Development of analysis tools for the MAGIC Telescopes and observation of the Segue 1 Satellite Galaxy with the MAGIC-I Telescope.

Coordinatore:  Prof. Attilio Stella  
Supervisore:  Prof. Mosè Mariotti  
Correlatore:  Dott. Villi Scalzotto  
Correlatore:  Dott. Michele Doro  

Dottorando: Saverio Lombardi  

Anno Accademico 2009-2010
A tutti coloro che mi vogliono bene
e mi hanno sempre sostenuto
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Il lavoro presentato in questa tesi è stato svolto nell’ambito dell’esperimento MAGIC durante i tre anni di Scuola di Dottorato all’Università degli Studi di Padova, Dipartimento di Fisica G. Galilei, in associazione con l’Istituto Nazionale di Fisica Nucleare, sezione di Padova, sotto la supervisione del Professor M. Mariotti, del Dott. Michele Doro e del Dott. Villi Scalzotto.

Il sistema stereoscopico di due telescopi MAGIC (Major Atmospheric Gamma-ray Imaging Cherenkov), situato nell’isola canaria di La Palma (Spagna), è basato sulla tecnica IAC (Imaging Atmospheric Cherenkov) per la rivelazione della luce Cherenkov emessa da cascate di particelle cariche che si originano nell’atmosfera terrestre. Tali cascate atmosferiche sono generate preminentemente dai raggi cosmici carichi, quali protoni e isotopi leggeri, che entrano isotropicamente nell’atmosfera terrestre, e da una percentuale di raggi gamma, la cui direzione di provenienza punta direttamente alla regione cosmica di emissione. Grazie a sofisticati algoritmi di riduzione dei dati è possibile estrarre efficacemente il segnale di natura gamma da quello di natura adronica per energie comprese tra ~50 GeV e ~10 TeV, permettendo studi in diversi settori scientifici quali l’Astrofisica galattica ed extragalattica delle alte energie, la Cosmologia e la Fisica delle Particelle Elementari.

Le attività scientifiche dell’esperimento MAGIC vengono portate avanti in sinergia e complementarità con gli esperimenti gamma posti su satellite, quali AGILE e Fermi-LAT, e quelli posti a terra, quali H.E.S.S., VERITAS e Milagro. I telescopi MAGIC, realizzati da una collaborazione internazionale di circa 150 fisici appartenenti ad una ventina di istituzioni di diversi paesi della comunità europea, rientrano nella seconda generazione di telescopi Cherenkov e presentano rispetto ai precedenti delle caratteristiche innovative legate per esempio alla grandezza della superficie riflettente (circa 250 metri quadri), all’abbassamento della soglia energetica al di sotto dei 100 GeV e alla velocità di puntamento nel caso dell’osservazione di fenomeni improvvisi, come ad esempio l’esplosione di raggi gamma (Gamma Ray Bursts). In particolare, la messa in opera del secondo telescopio MAGIC-II e la possibilità di condurre osservazioni stereoscopiche permettono un abbassamento della soglia energetica ed una sensibilità di circa un fattore 1.5-2 superiore a quella ottenuta con il singolo telescopio MAGIC-I. Attualmente, i telescopi MAGIC rappresentano il rivelatore posto a terra più sensibile al mondo per raggi gamma di energie comprese tra ~50 GeV e ~150 GeV.

Durante il lavoro di dottorato ho svolto la mia ricerca in due principali attività. La prima, di tipo tecnico, ha riguardato l’aggiornamento di una specifica parte del software dell’esperimento MAGIC necessario in vista della presa dati stereoscopica dei telescopi MAGIC (inizidata alla fine del 2009), ovvero il calcolo dell’area efficace collettrice. La seconda, di tipo più prettamente scientifico, è stata rivolta alla possibile rivelazione indiretta di Materia Oscura con il telescopio MAGIC-I. In particolare, ho contribuito all’analisi dati e alle relative pubblicazioni delle sorgenti osservate dal telescopio MAGIC-I candidate per la ricerca di possibili segnali gamma dovuti ad auto annichilazione di Materia Oscura, quali le galassie nane satelliti della Via Lattea Draco e Willman 1 e il cluster galattico Perseus. Durante i tre anni di dottorato ho passato complessivamente 4 mesi a La Palma, nel sito dei telescopi MAGIC, come operatore di presa dati e shift leader. Inoltre nel mese di Giugno del 2009 ho partecipato per 4 settimane alla fase di commissioning del secondo telescopio MAGIC-II.
Questo lavoro di tesi è suddiviso in otto capitoli.

Nel capitolo 1 verranno brevemente introdotti la fisica dei raggi cosmici e dei raggi gamma di natura astrofisica, gli attuali metodi sperimentali per la loro rivelazione, i meccanismi attivi nell’Universo per la produzione di raggi gamma di alte energie e le principali sorgenti cosmiche note di raggi gamma.

Il capitolo 2 sarà dedicato alla descrizione delle principali caratteristiche fisiche degli sciami atmosferici di natura adronica ed elettromagnetica, alla susseguente emissione di luce Cherenkov prodotta da essi e alla tecnica di rivelazione IAC, su cui i telescopi MAGIC si basano.

Nel capitolo 3 verranno descritte le principali componenti hardware dei due telescopi MAGIC e le innovazioni introdotte per il secondo telescopio.

Nel capitolo 4 si discuterà la catena di analisi standard del telescopio MAGIC-I per l’estrazione delle principali quantità fisiche di interesse, quali la significanza del segnale proveniente da una data sorgente e il suo flusso. Successivamente verranno brevemente descritte le principali novità introdotte per la corrente analisi delle sorgenti osservate stereoscopicamente. Le attuali performance del sistema stereoscopico (la cui fase di commissioning è stata portata a termine con successo durante il 2009) saranno inoltre presentate.

Nel capitolo 5 verrà introdotta la tematica riguardante la dipendenza alt-azimutale di una delle principali quantità che caratterizzano la rivelazione di luce Cherenkov da parte dei telescopi MAGIC ovvero l’area efficace collettrice. Infatti, se da un lato l’effetto zenitale su tale quantità è ben noto e correlato alla maggiore profondità atmosferica che gli sciami percorrono nel loro sviluppo per angoli zenitali di osservazione via via maggiori, una possibile dipendenza azimutale è associata agli effetti che il campo geomagnetico induce sullo sviluppo degli sciami stessi e alla particolare configurazione geometrica del sistema di telescopi. La direzione fissa tra di essi, infatti, rompe la simmetria circolare tipica delle osservazioni effettuate con un singolo telescopio. Alla luce della messa in funzione del secondo telescopio MAGIC-II e della osservazione stereoscopica delle sorgenti, l’introduzione della dipendenza azimutale dell’area efficace collettrice è stata dunque presa sistematicamente in considerazione e implementata efficacemente nel software di analisi dati dell’esperimento. Le principali modifiche apportate al software verranno illustrate assieme ai risultati di test effettuati su campioni di dati Monte Carlo e di dati reali.

Nel capitolo 6 verrà fatta una breve introduzione sulla Materia Oscura: saranno discusse le principali evidenze sperimentali, alcuni modelli che la descrivono e i principali candidati proposti in letteratura per spiegare la natura. L’attenzione sarà focalizzata sulla ricerca indiretta (tramite raggi gamma) di Materia Oscura che si basa sulla possibilità per i telescopi MAGIC di poter rivelare raggi gamma, prodotti da annichilazione o decadimento di particelle di Materia Oscura, provenienti da sorgenti caratterizzate da alte densità di tale tipo di materia quali, per esempio, galassie nane sferoidali satelliti della Via Lattea.

Il capitolo 7 sarà dedicato all’analisi dati dell’osservazione effettuata da parte del telescopio MAGIC-I della sorgente Segue 1, ritenuta essere una galassia nana sferoidale satellite della Via Lattea, la cui cinematica stellare sembra indicare un elevato rapporto massa-luminosità, rendendo tale oggetto celeste estremamente interessante dal punto di vista della ricerca indi-
retta di Materia Oscura. I dati di tale sorgente hanno richiesto particolare attenzione dovuta al fatto della presenza di una stella di magnitudo apparente 3.5 nel campo di vista della sorgente durante l’intera osservazione. Le tecniche adottate per trattare i problemi legati alla presenza della luce di tale stella saranno illustrati. L’analisi ha permesso di determinare, per energie maggiori di 100 GeV, limiti superiori sul flusso della sorgente assumendo diversi generici spettri di potenza. Un articolo sull’osservazione della sorgente Segue 1 condotta dal telescopio MAGIC-I, basato sui risultati di questa analisi, è in fase di preparazione.

Infine, nel capitolo 8, verranno riportate le conclusioni generali su questo lavoro di tesi.
Abstract

The work presented in this thesis has been carried out for the MAGIC experiment during three years of PhD student-ship at the University of Padova, Department of Physics G. Galilei, in association with the Padova section of the National Institute of Nuclear Physics (Istituto Nazionale di Fisica Nucleare, INFN), under the supervision of Professor M. Mariotti, PhD Michele Doro and PhD Villi Scalzotto.

The two MAGIC (Major Atmospheric Gamma-ray Imaging Cherenkov) telescopes, located in the Canary Island of La Palma (Spain), constitute a stereoscopic system based on the IAC (Imaging Atmospheric Cherenkov) technique and detect Cherenkov light emitted by atmospheric showers of charged particles that originate in the Earth’s atmosphere. These showers are predominantly generated by charged primary cosmic rays, such as protons and light isotopes, which impinge on the Earth’s atmosphere isotropically, and by a small percentage of gamma-rays, whose direction points back to the region of the cosmic emission. Thanks to sophisticated data reduction algorithms it is possible to efficiently extract the gamma-ray signal of the observed sources from the hadronic background in the energy range between \( \sim 50 \) GeV and \( \sim 10 \) TeV, allowing studies in different scientific fields such as galactic and extragalactic high energy Astrophysics, Cosmology and Particle Physics.

The scientific activities of the MAGIC experiment are carried out in synergy and complementarity with satellite experiments, like AGILE and Fermi-LAT, and ground-based experiments, like H.E.S.S., VERITAS, and Milagro. The MAGIC telescopes, operated by a collaboration of about 150 physicists from about 20 institutes, spread in several European countries, belong to the second generation of Cherenkov telescopes and have innovative features compared to previous IAC experiments related for example to the size of the reflecting surface area (about 250 square meters), to the lower energy threshold (below 100 GeV) and to a fast repositioning system in case of observations of transient phenomena such as the Gamma Ray Bursts. In particular, the start of the operations of the second telescope MAGIC-II and the possibility to perform stereoscopic observations allow a lower energy threshold and a better sensitivity (by a factor \( \sim 1.5-2 \)) compared to the single MAGIC-I telescope observations. Currently, the MAGIC telescopes are the more sensitive world-wide ground-based detector for gamma-rays in the energy range between \( \sim 50 \) GeV and \( \sim 150 \) GeV.

During the PhD student-ship, my research activity focused on two principal occupations. The first one, orientated on technical issues, concerned the upgrade of a specific part of the software of the MAGIC experiment required for the stereoscopic data taking (which started since fall 2009), i.e. the calculation of the Effective Collection Area. The second one, more focused on scientific topics, was addressed to the possible indirect detection of Dark Matter with the MAGIC-I telescope. In particular, I contributed to the data analyses and to the related publications of the observations of interesting sources for possible detection of gamma-ray signal from self-annihilation processes of Dark Matter, like the dwarf spheroidal galaxies (satellite of the Milky Way) Draco and Willman 1 and the galaxy cluster Perseus. During the three years of PhD student-ship, I spent, on the whole, 4 months at the site of the MAGIC telescopes, as data taking operator and shift leader. I also contributed in June 2009 for a period of 4 weeks to the commissioning phase of the second telescope MAGIC-II.
This thesis is divided into eight chapters.

Chapter 1 will be dedicated to a brief introduction on the physics of cosmic-rays and of astrophysical gamma-rays, on the current experimental methods for their detection, on the main mechanisms which take place in the Universe for the production of very high energy gamma-rays and on the main astronomical objects known as gamma-ray emitters.

In chapter 2, a description of the main physical characteristics of the hadronic and electromagnetic atmospheric showers, the subsequent Cherenkov light emission and the IAC technique will be given.

In chapter 3, the main hardware components of the MAGIC telescopes will be illustrated, together with the main innovations introduced for the second telescope.

In chapter 4, the standard analysis chain of the MAGIC-I telescope finalized to the extraction of the physical quantities of main interest, such as the significance and the gamma-ray flux of a given source, will be described. Subsequently, the new tools required for the stereo analysis of the MAGIC telescopes data will be introduced. The current performance of the stereoscopic system (whose commissioning phase was successfully accomplished during 2009) will be also shown.

Chapter 5 will be dedicated to the Alt-Azimuth dependence of one of the main quantities which characterize the detection of Cherenkov light by the MAGIC telescopes, i.e. the Effective Collection Area. Indeed, while the Zenith effect on this quantity is well known and related to the increased atmospheric depth the atmospheric showers must pass through for increasing Zenith angles of observation, a possible Azimuth dependence is associated to the geomagnetic effects induced to the development of the showers and to the particular geometric configuration of the two telescopes system. In fact, the fixed direction between them breaks the typical circular symmetry which characterizes the observations performed with a single telescope. In view of the start of the operations of the second MAGIC telescope and of the stereoscopic observations, the introduction of the Azimuth dependence of the Effective Collection Area has been therefore systematically taken into account and successfully implemented in the analysis software of the experiment. The main changes implemented in the software together with the results of tests performed on both Monte Carlo and real data will be shown.

In chapter 6, a brief introduction on the Dark Matter topic will be reported: the main experimental evidences and some of the models and candidates proposed in literature to describe the Dark Matter nature will be discussed. In particular, the attention will be focused on the indirect Dark Matter search (via gamma-rays) which is based on the possibility for MAGIC to detect gamma-rays as a result of annihilations or decays of Dark Matter particles. Gamma-ray signals are searched for in places where Dark Matter is concentrated, like, for example, the dwarf spheroidal galaxies satellite of the Milky Way.

Chapter 7 will be dedicated to the data analysis of the observation carried out by the MAGIC-I telescope of Segue 1, a source considered to be a dwarf spheroidal galaxy satellite of the Milky Way, whose stellar kinematics seems to indicate a high mass–to–light ratio, making this celestial object extremely interesting from the point of view of indirect Dark Matter searches. The data of this source required particular cares due to the fact that a 3.5 apparent magnitude star was present in the field of view of the source during the whole
survey. The adopted techniques used to face the problems related to the light of that star will be illustrated. The analysis allowed to determine, for energies above 100 GeV, upper limits on the flux emission derived from different assumed power law spectra. A paper on Segue 1 observation carried out by the MAGIC-I telescope, based on the results achieved by this analysis, is in preparation.

Finally, in chapter 8, the general conclusions of the work presented in this thesis will be given.
Acknowledgments

I would like to thank all the people who helped me during these years of PhD student-ship.

First of all, Professor Mosè Mariotti for all the opportunities he gave me and for all the experiences I could do thanks to him.

I also would like to thank all the senior Professors of the MAGIC Padova group: Luigi Peruzzo, Antonio Saggion and Donatella Pascoli.

Very special thanks to Villi, for his great comprehension and wisdom: he always helped and encouraged me during these years and I have learned several things from him, not only from the scientific point of view.

Other special thanks to Michele who explained me many things and helped me a lot with advices and suggestions.

I would like to thank Elisa, for her kindness, patience, availability and sincere friendship she always dedicated to me.

Many thanks to Francesco for his answers to my technical questions and to Marcos for several explanations he gave me.

Thanks also to all the other young people who belong or belonged to the MAGIC Padova group during these years of PhD student-ship: Fabio, Alessandro, Simona, and Cornelia.

I must say I spent very beautiful moments with all these people and I consider them as a second family.

I also would like to thank all the people of the MAGIC Collaboration (the “big family”) and the people who shared with me the data taking shifts I did at the MAGIC site.

I thank all my friends who always supported me, giving me fun and good cheer.

Last but not least, I would like to thank my family, my mother Silvia, my father Mariano and my sister Sonia for their love and support.

Thanks to all of you!
Introduction to Very High Energy $\gamma$-ray Astrophysics

Figure 1.1: Astroparticle physics with cosmic-rays.
Astroparticle physics is a relatively new branch of particle physics which is interested in elementary particles of astronomical origin and their relation to astrophysics and cosmology. The main searches carried out in this field are basically related to fundamental issues such as the contents and the history of the Universe, the nature of Dark Energy and Dark Matter as well as the nature of gravity, the stability of ordinary matter, the properties of neutrinos and the origin of cosmic-rays. Its rapid development has led to the design of new types of experiments which employ new detection methods to observe a wide range of cosmic particles including neutrinos, $\gamma$-rays and the other cosmic-rays up to the highest energies.

Cosmic-rays are energetic particles originating from outer space that constantly impinge on the Earth’s atmosphere. Their energy range is incredible wide (more than 13 orders of magnitude), between few MeV up to $10^{21}$ eV, much greater than the energies which can be investigated with the accelerator experiments. Even though the first evidences of these particles trace back to the beginning of nineteenth century, it is not yet completely clear where and how cosmic-rays are produced, what their composition is over the whole observed energy and how far they can propagate in space. Indeed, cosmic-rays are almost completely composed by charged particles which are deflected by cosmic magnetic fields and arrive almost isotropically to the Earth, losing the information on their place of production and acceleration.

Excluding neutrinos, only a tiny fraction of cosmic-rays is composed by neutral particles, the most suited ones for cosmic-ray studies being the $\gamma$-rays. In fact, $\gamma$-rays are always produced where high energy cosmic-rays are accelerated and travel undeflected from the regions of production to the Earth. They represent therefore a unique probe for understanding the mechanisms which rule the cosmic-ray physics and can cast light on fundamental issues of particle physics and cosmology.

In this chapter, after a brief introduction to the cosmic-ray physics, we will discuss the basic issues which characterize the $\gamma$-ray astronomy (i.e., the field which studies the $\gamma$-rays of cosmic origin) such as the experimental techniques, the mechanisms which can produce very high energy $\gamma$-rays, the main cosmic sources which are known to be $\gamma$-ray emitters and the connections with the fundamental physics.

### 1.1 Cosmic-rays and $\gamma$-rays

The term cosmic-rays (CRs), coined by the American physicist Millikan in the twenties of the last century, includes all the particles of extraterrestrial origin that constantly hit the Earth. The first experimental evidences of CRs were given by the Austrian physicist V. Hess in 1912 [1], by the discovery of a ionizing radiation permanently impinging on the Earth’s atmosphere which could not be explained by the Earth’s radioactivity. Hess indeed was observing the showers of charged particles induced by CRs in the atmosphere [2]. In the first decades after the discovery of extensive air showers many previously unknown particles could be identified ($e^+, \mu^\pm, \pi^\pm, K^\pm, \ldots$) as being part of the showers. The discovery of new particles gave rise to a completely new field in experimental and theoretical physics which nowadays is still a flourishing field of investigation.

The spectrum of the observed CRs is a remarkable power law extending in energy over 13 orders of magnitude, ranging from few MeV up to $\sim10^{21}$ eV, and with fluxes dropping from 1 particle/(cm$^2$ s) at $\sim100$ MeV to less than 0.01 particles/(km$^2$ century) for the highest observed energies. This spectral feature suggests that CRs have been emitted by objects that had not time to thermalize. For this reason, CRs are considered messengers of the most violent components of the Universe.

Many open questions are still debated about the origin, the composition and the propagation of CRs. Nevertheless, a general idea about the sources of galactic and extragalactic CRs has been established during last decades: the most important sources of galactic CRs are thought to be Supernova Remnants (shell type), Plerions, Pulsars, Microquasars, X-ray binary sys-
1.1. Cosmic–rays and $\gamma$–rays

tems, young star and young open star clusters, whereas, the main sources of extragalactic
CRs are Active Galactic Nuclei (in the beamed region, jet), Gamma Ray Burst, Starbust
Galaxies and Cluster of Galaxies.

Excluding neutrinos, which are weakly interacting particles, CRs are made for 99.9% of
charged particles, mainly protons ($\approx$89%), $\alpha$-particles ($\approx$10%), ionized nuclei of
heavier elements ($\approx$1%) and electrons/positrons ($\approx$1%), and only for a tiny fraction of
photons of energy greater than 1 MeV, called gamma–rays ($\gamma$–rays). A compilation of the
energy spectrum (multiplied by $E^3$) of CRs from 32 different experiments is shown in
Fig. 1.2. In the region of the energy spectrum which is unaffected by the Earth’s magnetic
field and the propagation of particles to the Earth through the solar wind (at energies
greater than $\approx$1 GeV), the differential flux of CRs follows a power law with spectral
index $\alpha$ = -2.7. The region around $10^{15.5}$ eV is called knee, because of the change in
the slope (the spectral index varies from -2.7 to -3), probably caused by a change in the
composition of the primary particles. Instead, the region around $10^{19}$ eV is referred to
as ankle, because of a second tilt in the slope: the spectral index for energies above the
ankle assumes a value of -2.6. The ankle is most probably caused by an
interplay between CRs of galactic and extragalactic origin. CRs with energies above
$10^{20}$ eV can interact with the Cosmic Microwave Background (CMB) radiation. This
leads to a cutoff in the CR–spectrum, also known as the GZK–cutoff $[3, 4]$. It is
generally believed that CRs below $10^{15}$ eV have a galactic origin and have been confined
inside our galaxy for at least $10^7$ yr. That is why these particles have been completely
isotropized before they arrive to the Earth. Particles above $10^{17}$ eV are believed to be
mostly of extragalactic origin, since the galactic magnetic field is not able to trap them
in our galaxy $[5]$. CRs up to about $10^{18}$ eV are isotropic, while a claim on anisotropy
at higher energies is still under debate $[6]$.

The main feature that distinguishes the charged component of galactic and extragalactic
CRs from the neutral component is its isotropic origin. In fact, the charged particles,
being deflected by the Lorentz force due to the galactic and extragalactic magnetic
fields, lose the information about the direction of origin and also about the temporal
structure of the emission. This is the main reason why, after one hundred years since their first experimental
evidence, CRs still keep many secrets and open questions. Many steps have been made to
understand their origin: we know for example the main objects involved in their acceleration, but we still do not know the precise emission mechanisms. The \(\gamma\)-rays and the neutrinos, conversely, being electrically neutral, do not suffer deflections along the travel from the place of their generation to the observer, thus they can provide valuable information on the astrophysical environments in which they were generated, on the characteristics of the medium they crossed through to reach the Earth and on the acceleration mechanisms at high energies which take place in the Universe.

Due to their extremely small cross section, the detection of neutrinos requires enormous experimental devices. To date, no sources of neutrinos were detected beyond the Sun and the supernova SN 1987A exploded \(1.7 \cdot 10^5\) ly far away from the Earth in the Magellanic Clouds (recorded simultaneously by the experiments Kamiokande and IMB [8]). Instead, many sources of \(\gamma\)-rays, both galactic and extragalactic, were detected, despite the very low fluxes which characterize the high energy emissions, compared to those at lower energies. For this reason \(\gamma\)-rays are currently the best window for exploring the high energy processes in action in the Cosmos.

The discipline that studies the \(\gamma\)-rays of cosmic origin is called \(\gamma\)-ray astronomy and it was born around the mid sixties of the last century with experiments carried out on the satellites, balloons, and, for very high energies, even on the ground. However, the first significant progress has been achieved in the last two decades, when the experiments on-board the satellites Granat, Compton Gamma Ray Observatory (CGRO) and BeppoSAX gave a completely new vision of the Universe at high energies, promoting the \(\gamma\)-ray astronomy to a new important branch of the astronomical and astrophysical searches related to the study of the most energetic and violent processes acting in the Universe. More recently, a new great progress in the knowledge of the \(\gamma\)-ray sky is being carried out by the Fermi Space Telescope, launched in 2008, which is currently discovering several new \(\gamma\)-ray sources and giving impressive new details about the previously detected \(\gamma\)-ray sources.

The energy spectrum studied by the \(\gamma\)-ray astronomy ranges from energies around some MeV to several tens of TeV, with fluxes rapidly decreasing with energy. Since this is a very large energy range, it can be divided into sub-domains according to the main physical interaction processes which take place in each particular energy region. \(\gamma\)-rays are detected through the production of secondary electrons which can occur through different processes. Therefore, each sub-domain of the \(\gamma\)-ray spectrum is characterized by its own detection technique. The energy bands can thus conventionally divided (even though the definitions are not standard and linked in part on the thresholds of the instruments) in [9]:

- **Medium Energy region (ME):** the involved energies are below 30 MeV. Compton scattering is the main interaction process used for the detection. Since cosmic \(\gamma\)-rays of this energy domain are completely absorbed in the atmosphere, satellite telescopes are the principal detectors in this energy band.

- **High Energy region (HE):** the energies are between 30 MeV and 100 GeV. Pair production is the main process on the basis of the detection of those photons. The produced leptonic pairs can be detected with satellites which are situated above the Earth’s atmosphere but also by atmospheric balloons.

- **Very High Energy region (VHE):** the energies range between 100 GeV and 100 TeV. These photons are enough energetic to produce, when entering the Earth’s atmosphere, electromagnetic showers which can be well detected by ground-based telescopes by the aid of secondary flashes of light produced during the showers. The Earth’s atmosphere acts therefore as a calorimeter and the detection of the primary \(\gamma\)-rays is indirect.
• Ultra High Energy (UHE): the involved energies are above 100 TeV. This is a rather unknown region. The experiments are placed on the ground and directly reveal the particles produced in the so called Extensive Air Showers (EAS). To date, a good discrimination between the γ or hadronic nature of the primary CR producing the shower is still difficult to achieve.

As in case of CRs, γ-rays have typical astrophysical spectra with rapidly falling fluxes. Therefore, single experiments cannot be sensitive in the entire γ-ray energy range and realistically only around 2-3 orders of magnitudes can be covered by one observational technique. In recent years, in order to effectively explore this vast energy window, which often presents phenomena which vary widely over time, different and sophisticated technologies have been developed. The MAGIC telescopes fit into this context as a ground-based detector designed for the detection of photons of energies between $\sim 50$ GeV to $\sim 10$ TeV, based on the Imaging Atmospheric Cherenkov (IAC) technique, which will be described in chapter 2.

1.2 Experimental approaches

In the last 40 years, several approaches have been used to detect CRs and γ-rays. The adopted technologies are many and very different one each other since the range of energies to be investigated is very wide.

Schematically, CRs can be detected

- directly by satellite and balloon experiments carrying compact particle detectors. Representative on-board satellite experiments are AMS 02 [10] and PAMELA [11], whereas important balloon-borne experiments are ATIC [12] and BESS [13].

- through the direct detection of the secondary particles generated in the CR induced air showers with large arrays of particle detectors on the ground. This technique works well for energies above $\sim 10^{14}$ eV. In case of lower energies of the primaries, shower particles do not reach the ground level. The energy threshold of an experiment can be thus lowered by building detectors at high altitudes. The most important examples of such technique are the KASCADE experiment [14] and AUGER [15] which is composed of 1600 water tanks distributed over 3000 km$^2$ in Argentina. AUGER also hosts 4 detectors for atmospheric fluorescence detection. To date, it is the biggest of all direct particle detection experiments.

CRs can be also detected indirectly on the ground by measuring

- the Cherenkov light emitted by the charged particles of the CR induced air showers (see chapter 2). Cherenkov light for air showers above 15 TeV can be detected directly by placing light detectors.

- the fluorescence light induced by the air showers. A small fraction of the energy of the shower particles is transferred to atmospheric nitrogen molecules and excites them which subsequently emit isotropically a distinct line spectrum in the near UV. Only showers with energy above $\sim 10^{15}$ eV produce a detectable amount of fluorescence light. An important example of experiments which employ such technique is the HiRes detector, located in the Utah desert, composed of two groups of telescopes at 12.6 km distance [16].

- acoustic waves or radio emissions of air showers using antennas. Recently, the feasibility of shower detection in radio for energies above $\sim 10^{17}$ eV was shown [17].
Concerning the detection of cosmic γ-rays, it can be broadly divided into two basic categories: the first one studies γ-rays in the HE domain, whereas the second one in the VHE domain. This distinction is made according to the type of experiment and the technology used to probe these energy ranges. The γ-rays subject of the study of the first category cannot get through the atmosphere without being completely absorbed and diffused, and thus cannot generate electromagnetic showers producing Cherenkov light suitable to be efficiently detected from the ground. Therefore, experiments at high altitude or on-board of satellites are needed to investigate the γ-ray HE domain. These experiments are basically limited by the small effective area (∼1 m²) which confines the investigation at the HE domain, due the rapidly falling down of the γ-ray fluxes. Conversely, as the energy of the primary γ-rays increases, the development of electromagnetic showers in atmosphere creates a quantity of Cherenkov light which turns out to be enough to be detected at the ground. Therefore ground-based telescopes are suitable for the investigation of the VHE γ-ray domain, since the effective collection area can reach values of the order of 10⁵ m².

It is worth mentioning that after the start of the operations of the Fermi Large Area Telescope (Fermi-LAT) and of the ground-based stereoscopic MAGIC telescopes system, the existing gap between ∼30 GeV and ∼100 GeV which was affecting the γ-ray astronomy studies has been finally overwhelmed. Indeed, simultaneous observations between the MAGIC telescopes and the Fermi-LAT detector will allow detailed studies of the high energy phenomena in the Universe in the wide energy range between few tens of MeV and ∼10 TeV.

In the next subsections, the techniques used for detecting the γ-rays of cosmic origin are briefly described.

1.2.1 Satellite-borne experiments

In order to avoid the opacity of the Earth’s atmosphere, the first instruments to detect γ-rays were installed on flying balloons. However their sensitivity was seriously limited by secondary γ-rays generated by the interaction of the charged CRs with the atmosphere above the balloon. It was not until the launch of the OSO-3 (Third Orbiting Solar Observatory) satellite [18], in 1967, that γ-ray astronomy with satellites started to obtain its first results. This satellite detected γ-ray emission of 100 MeV from the galactic plane. In the 1970s two other γ-ray satellites were launched: SAS-2 (Small Astronomy Satellite-2) in 1972 (which found evidence of γ-ray emission from single sources, the Vela pulsar and the Crab Nebula) and the ESRO’s COS-B in 1975 which identified more than 25 γ-ray sources. Since then many efforts have been carried out to improve the detection technique.

Modern satellite and balloon borne detector consists of a standard calorimeter adapted for the HE band with a segmented tracking system. The detection mechanism is pair production in the tracker and consequent γ-ray production into a calorimeter. The energetic range is basically defined by the width of the calorimeter and the telescope area (of the order of 0.5-1 m²), because γ-rays at high energies have both small fluxes and cannot be contained by a too short detector. Contrary to ground-based telescopes, these detectors are characterized by very large field of view but rather small angular resolution (depending on the energy, but of the order of few degrees)¹. Due to the rapidly falling flux of γ-rays, they are only marginally sensitive above few tens of GeV for point-like emissions, while the diffuse emission, that can be integrated for thousand of minutes, can be observed up to few hundreds of GeV. A famous representative experiment based on this technique was the EGRET experiment [19], on-board the CGRO satellite, which operated between the years 1991 and 2000. Currently the most

¹This can no longer be applied to the Fermi-LAT detector, for which the angular resolution can reach values even of the order of 0.15° (see Fig. 1.3).
important instruments are the ASI’s AGILE (Astro-rivelatore Gamma a Immagini LEggero), launched in 2007, and the NASA’s Fermi (formerly GLAST) $\gamma$-ray satellite, launched in 2008. Fermi satellite [20] carries two instruments, the Large Area Telescope (LAT) and the GLAST Burst Monitor (GBM). The energy range of LAT is between 20 MeV and 300 GeV. The Fermi–LAT detector, marking the footsteps of EGRET, is providing new insight on the HE $\gamma$-ray astronomy. So far many $\gamma$-ray sources have been detected and discovered by the Fermi–LAT detector [21,22] thanks to its much improved performance and sensitivity [23]. It represents a unique detector for triggering the observation of the ground-based telescopes and to give information up to the actual energy threshold of the Imaging Atmospheric Cherenkov Telescopes (IACTs), extending the spectrum down to the HE domain. In Fig. 1.3, a comparison between the performance of EGRET and Fermi–LAT detectors is reported.

### LAT Specifications and Performance Compared with EGRET

<table>
<thead>
<tr>
<th>Quantity</th>
<th>LAT (Minimum Spec.)</th>
<th>EGRET</th>
</tr>
</thead>
<tbody>
<tr>
<td>Energy Range</td>
<td>20 MeV - 300 GeV</td>
<td>20 MeV - 30 GeV</td>
</tr>
<tr>
<td>Peak Effective Area$^1$</td>
<td>&gt; 8000 cm$^2$</td>
<td>1500 cm$^2$</td>
</tr>
<tr>
<td>Field of View</td>
<td>&gt; 2 sr</td>
<td>0.5 sr</td>
</tr>
<tr>
<td>Angular Resolution$^2$</td>
<td>&lt; 3.5° (100 MeV)</td>
<td>5.8° (100 MeV)</td>
</tr>
<tr>
<td></td>
<td>&lt; 0.15° (&gt;10 GeV)</td>
<td></td>
</tr>
<tr>
<td>Energy Resolution$^3$</td>
<td>&lt; 10%</td>
<td>10%</td>
</tr>
<tr>
<td>Deadtime per Event</td>
<td>&lt; 100 μs</td>
<td>100 ms</td>
</tr>
<tr>
<td>Source Location Determination$^4$</td>
<td>&lt; 0.5'</td>
<td>15'</td>
</tr>
<tr>
<td>Point Source Sensitivity$^5$</td>
<td>&lt; 6 x 10$^{-9}$ cm$^{-2}$ s$^{-1}$</td>
<td>$\sim$ 10$^{-7}$ cm$^{-2}$ s$^{-1}$</td>
</tr>
</tbody>
</table>

$^1$ After background rejection  
$^2$ Single photon, 68% containment, on-axis  
$^3$ 1-$\sigma$, on-axis  
$^4$ 1-$\sigma$ radius, flux 10$^{-7}$ cm$^{-2}$ s$^{-1}$ (>100 MeV), high $|b|$  
$^5$ > 100 MeV, at high $|b|$, for exposure of one-year all sky survey, photon spectral index -2

**Figure 1.3:** Comparison of the performance between EGRET and Fermi–LAT. The improvements of Fermi–LAT detector are impressive. Image from http://fermi.gsfc.nasa.gov/.

### 1.2.2 Ground-based Atmospheric Cherenkov experiments

Ground-based atmospheric $\gamma$-ray astronomy is concerned with $\gamma$-rays of energies between around 100 GeV and several TeV, i.e. in the VHE domain. When the primary VHE $\gamma$-rays reach the top of the Earth’s atmosphere, they produce positron-electron pairs which then emit energetic $\gamma$-rays via Bremsstrahlung. The secondary $\gamma$-rays in turn emit positron-electron pairs giving rise to the so-called electromagnetic cascade where highly relativistic particles cause a flash ($\sim$3 ns) of UV-blue Cherenkov light which propagates in a cone with an opening angle of $\sim$1°. The resulting circle of projected light, at 2000 m a.s.l., has a radius of about 120 m. A IACT collects the Cherenkov light by a reflective surface which is then focused onto a multi-pixel camera which records the shape of the image produced by the shower which has an elliptical shape pointing to the nominal source position projected onto the camera. Since Cherenkov light is emitted also by charged particles produced in atmospheric showers
induced by charged isotropic CRs, an image reconstruction algorithm [24] is used in order to recover the energy and the direction of the primary particle and to determine whether it was more likely a hadron or a photon, allowing the rejection of more than 99% of the background. More details about the physics of the atmospheric showers and the IAC technique will be given in chapter 2.

The first IACT was the Whipple Telescope (Mount Hopkins, Arizona), which started the operations in 1982, with a multi-pixel camera of 37 photomultipliers and a spherical reflecting surface of 10 m of diameter. It discovered TeV emission from the Crab Nebula in 1989 [25] and the Active Galactic Nuclei Mkn 421 and Mkn 501, which, nowadays, are considered confirmed sources for the VHE γ-ray astronomy, often used to calibrate the performance of the telescopes. The current second generation of IACTs consists in arrays of telescopes. The most important telescopes in operation are H.E.S.S. (High Energy Stereoscopic System) [26], located in Namibia, VERITAS (Very Energetic Radiation Imaging Telescope Array System) [27], operating in Arizona (USA), CANGAROO III [28] (Collaboration of Australia and Nippon for GAamma Ray Observatory in the Outback) in Australia, and the MAGIC (Major Atmospheric Gamma-ray Imaging Cherenkov) telescopes [29], located in the Canary Island of La Palma (Spain), which will be described in chapter 3. Two projects for the next generation of IACT systems are already in their first stages of development: the CTA (Cherenkov Telescopes Arrays) Project [30], mainly promoted by the European scientific community, and the AGIS (Advanced Gamma Imaging System) Project [31], which is pursued by the American scientific community. The sensitivity of these new arrays of telescopes are foreseen to be at least 10 times better than the actual existing ones and they will be sensitive to γ-rays of energies between ~10 GeV and ~100 TeV.

An alternative technique to the IAC one is represented by the use of the large collection area of already existing solar power plants for reflecting the Cherenkov light produced by the atmospheric showers into photomultipliers placed in a tower at the center of the plants. The achievable energy threshold can be of the order of few tens of GeV but the sensitivity and the capability to discriminate between hadronic and electromagnetic showers are much lower than those reached by the IAC technique since the images of the showers are not reconstructed. This technique is used by experiments like STACEE [32], CELESTE [33] and GRAAL [34].

1.2.3 Water Cherenkov tanks experiments

The Water Cherenkov experiments take advantage of the Cherenkov light produced by atmospheric showers electrons and positrons in clear water. The water tanks must be located at high altitude to observe a large fraction of atmospheric products which are rapidly absorbed. The tanks are surveyed with several photomultipliers which collect the light signal. A water Cherenkov experiment is usually composed of several tanks distributed over a large area. The field of view is therefore very large, and cover almost the complete above sky, while the angular resolution is rather small. A representative of this technique is the Milagro experiment [35], situated in the Jemez Mountains near Los Alamos, New Mexico, which operated until the year 2008. It was primarily designed to detect γ-rays but also detected large numbers of CRs. Milagro measured the γ-rays at average energies of 15 TeV. There is a proposed follow up experiment called the High Altitude Water Cherenkov experiment (HAWC) to be located at the Sierra Negra Volcano, Mexico. HAWC should be 15 times more sensitive than Milagro [36].
1.3 Production mechanisms of VHE $\gamma$-rays

The thermal radiation emitted by stars is characterized by blackbody spectra that can be described by Planck’s radiation formula with temperatures spanning from $3000^\circ$ K to $50000^\circ$ K. This translates to photons in the range of visible to UV light. The hottest objects in the Universe emit radiation up to X-rays in the few KeV energy range. Higher $\gamma$-ray energies are reached in non-thermal processes. The emission of $\gamma$-rays is always associated to the presence of accelerated charged particles at the VHE regime. For this reason, the search for cosmic sources of $\gamma$-rays is strictly related to the identification of regions in the Cosmos where high energy hadrons and leptons (and possibly even exotic particles) are present. The knowledge of the production mechanisms is important in order to understand where and how these $\gamma$-rays can be observed.

There are several possible processes for the production of $\gamma$-ray radiation each of which takes place in specific energy ranges. Normally, several of these mechanisms interplay to shape the $\gamma$-rays outflow from a given source. In addition, the presence of gas, plasma and radiation fields in the vicinity of the emitters can strongly shape the observed emission.

The most relevant non-thermal $\gamma$-ray production processes which are commonly believed to take place are: synchrotron radiation, inverse Compton (IC) scattering, Bremsstrahlung radiation and neutral pion decay. The hadronic process of neutral pion decay and IC scattering are the most important sources of VHE $\gamma$-rays. In fact, it is a highly debated issue which of the two processes is the dominant one. In case of neutral pion decay, this would point to hadronic accelerators and thus explain at least part of the origin of CRs. Observation of VHE neutrinos from the same source would be the unambiguous proof of hadronic production. In case of IC scattering, leptonic accelerators would be favourite, leaving the origin of CRs an open issue.

In the following subsection the main mechanisms for VHE $\gamma$-ray production are briefly described. A more comprehensive review can be found e.g. in [37].

1.3.1 Synchrotron radiation

If particles are forced by magnetic field to follow a curved trajectory they emit synchrotron radiation. The radiated synchrotron power $P$ is [38]

$$ P = \frac{1}{6\pi\varepsilon_0} \frac{e^2a^2}{c^3} \gamma^4 $$

(1.1)

where $e$ is the charge of the particle and $a$ its centripetal acceleration. Since $P$ is strongly dependent on the mass $m$ of the particle ($\propto 1/m^4$), synchrotron radiation in most applications is only relevant for electrons/positrons. Indeed, synchrotron radiation induced by protons is considered an inefficient process. For protons and electrons with the same energy, the energy loss rate for protons appears $(m_p/m_e)^4 \simeq 10^{13}$ times lower than the energy loss rate for electrons.

When particles are relativistic, the emitted radiation is beamed in a cone centered on the particle and with an angular spread $\alpha \simeq mc^2/E$. The spectrum of synchrotron radiation of mono-energetic electrons is a continuum that peaks at

$$ E_\gamma = 1.5 \cdot 10^{-5} \left( \frac{E_0}{[\text{TeV}]} \right)^2 \cdot \left( \frac{B}{[\text{G}]} \right) \text{GeV} $$

(1.2)
where $E_e$ is the energy of the electron and $B$ is the magnetic field component perpendicular to the plane of the particle trajectory.

Synchrotron radiation of accelerated electrons is one of the most important process in the non-thermal Universe. In the context of VHE $\gamma$-rays, synchrotron radiation is the usual process for the generation of seed photons for Inverse Compton scattering. However, UHE CRs (electrons and/or protons) can emit synchrotron radiation directly in the VHE domain.

### 1.3.2 Inverse Compton scattering

In the Inverse Compton (IC) process a low energy photon is upscattered by a relativistic particle (especially electrons). In this process a considerable fraction of the electron energy (up to several TeV) can be transferred to the photon. One can distinguish two different regimes, the Thomson limit ($\gamma \epsilon \ll m_e c^2$) and the Klein-Nishina limit ($\gamma \epsilon \gg m_e c^2$), where $\epsilon$ is the photon energy before the scattering process and $\gamma$ is the Lorentz factor of the relativistic electron. The average energy of the photon after the scattering process is

$$\langle E_{\gamma} \rangle \simeq 4 \langle \epsilon \rangle \gamma^2 \quad \text{Thomson limit}$$

$$\langle E_{\gamma} \rangle \simeq \frac{1}{2} \langle E_e \rangle \quad \text{Klein-Nishina limit}$$

(1.3)

where $E_e$ is the energy of the electron. Typical target photon fields in IC scattering are the $2.7^\circ$ K CMB, synchrotron radiation and thermally generated photons from stars or clouds. The emitted spectrum depends on the spectrum and density of the target photons and on the velocity distribution of the involved electrons. This type of process becomes important in regions where there is a large density of photons and it is of great interest for astrophysics studies as it can produce very energetic photons and thus can explain the emission at high energies of some of the sources which will be described in section 1.5. In particular, the IC process is important in the production of high energy $\gamma$-rays in the jets of Active Galactic Nuclei. Indeed, the IC scattering of synchrotron photons by relativistic electrons is able to produce the VHE photons observed in TeV blazars.

### 1.3.3 Bremsstrahlung

This type of radiation occurs when charged particles are accelerated in an electric field. Since the acceleration is proportional to the particle mass, this radiation involves mostly the lighter charged leptons. When an electron passes close to an atomic nucleus it feels the strong nuclear charge, and thus it is subject to an acceleration with a consequent radiation of photons. When an electron passes through a medium, the overall effect of many of such processes is the conversion of the electron kinetic energy into radiation. The photons are emitted in the forward direction of the particle, within an angle $\theta \approx 1/\gamma$ (with $\gamma$ Lorentz factor) with a continuous spectrum which approximately flats up to an energy of the electron kinetic energy ($E_\gamma = (\gamma - 1)m_e c^2$) which follows a power law with the same spectral index as the one of the accelerated particle. Bremsstrahlung occurs when the energy of the charged CR $E = \gamma m_e c^2$ is above a critical energy $E_0$ which depends on the crossed material. Below this value, the most efficient energy loss mechanism is ionization. This mechanism is observed in astrophysical regions that contain ionized gas, such as gaseous nebulae, which emit in radiowaves and in the hot intra-cluster gas within cluster of galaxies, which emit in X-rays. If the electrons are accelerate to TeV energies or more, emission at $\gamma$-rays is also observed. The Bremsstrahlung
radiation plays an important role in the processes of atmospheric shower formation, as it will be described in chapter 2.

1.3.4 \( \pi^0 \) decay

Relativistic protons and nuclei interact with ambient gas through inelastic collisions and produce basically mesons, kaons and hyperons via strong interactions. One of the most common reactions is:

\[
pp \rightarrow pp\pi^+\pi^-\pi^0 \tag{1.4}
\]

The pseudoscalar \( \pi \)-mesons (pions) are the lightest mesons \( (m_{\pi^\pm} = 140 \text{ MeV}, m_{\pi^0} = 135 \text{ MeV}) \) and have the larger cross section for such processes. They are produced in equal amounts of positive, negative and neutral charged, thus one third of the produced pions are neutral. While the charged pions decay weakly (their lifetime is \( \tau = 2.6 \cdot 10^{-8} \text{ s} \) in the particle rest frame) basically in muons and neutrinos, the neutral ones decay electromagnetically in \( \gamma \)-rays with a short life time of \( 8.6 \cdot 10^{-17} \text{ s} \):

\[
\begin{align*}
\pi^0 & \rightarrow \gamma\gamma \quad \text{(B.R. 98.2\%)} \\
\pi^0 & \rightarrow e^+e^-\gamma \quad \text{(B.R. 1.2\%)}
\end{align*} \tag{1.5}
\]

The minimum kinetic energy of a proton to produce a \( \pi^0 \) is

\[
E_{\text{th}} = \frac{2m_{\pi^0}c^2(1 + m_{\pi^0})}{4m_p} \simeq 280 \text{ MeV} \tag{1.6}
\]

The energy of the photons emitted by a \( \pi^0 \) at rest is peaked at \( E_\gamma = m_{\pi^0}c^2/2 \simeq 67.5 \text{ MeV} \), whereas in the laboratory frame it depends on the emission angle (i.e., the angle between the photon direction with respect to the pion one) and on the initial energy of the meson. Its value is between

\[
\gamma_{\pi^0} \cdot (1 - \beta_{\pi^0}) \cdot m_{\pi^0}c^2/2 \leq E_\gamma \leq \gamma_{\pi^0} \cdot (1 + \beta_{\pi^0}) \cdot m_{\pi^0}c^2/2 \tag{1.7}
\]

The final energies of the emitted photons can thus reach the VHE regime.

\( \gamma \)-rays produced by pions are the so-called \( \gamma \)-rays of hadronic origin. They are distinguishable from those of electromagnetic origin because of their spectra, lack of correlation with X-rays, and presence of molecular clouds, and they occur together with a flux of neutrinos coming from the charged pion decays, whose spectrum is similar to that of \( \gamma \)-rays from \( \pi^0 \) decay.

1.4 \( \gamma \)-ray Absorption

\( \gamma \)-rays traveling cosmological distances from the source to the observer can suffer absorption losses by interaction with cosmic matter and with the low energy photon fields. If, on the one hand, the absorption of \( \gamma \)-rays emitted by distant sources by the interstellar and intergalactic matter is negligible (due to the low matter densities), on the other hand, the absorption due to the low energy photons belonging to the evolving extragalactic background light (EBL) can be relevant. Excluding the galactic plane, the Universe is almost isotropically filled with soft photons of the EBL. EBL is the second largest, in terms of contained energy, background after the CMB of 2.7° K. While the CMB conserves the structure of the Universe at the moment of decoupling of matter and radiation following the Big Bang (at redshift \( z \approx 1000 \)),
EBL is a measure of the entire radiant energy released by processes of structure formation that have occurred since the decoupling. Until now, several different models of the EBL have been proposed, which differ quite considerably. Nevertheless, it is commonly accepted that the energy density spectrum of the EBL is characterized by two pronounced peaks. The first peak at \( \sim 1\ \mu m \), known as the stellar component, is associated with light emitted by stars and it is redshifted through the history of the Universe. The second one at \( \sim 100\ \mu m \) comes from the re-processing of the starlight by dust and for this reason is called the dust component (see Fig. 1.4).

The process which VHE \( \gamma \)-rays undergo with EBL is the photon-photon electron-positron pair production

\[
\gamma_{\text{VHE}} + \gamma_{\text{EBL}} \rightarrow e^+ + e^-
\]

and it is an extremely relevant process for the observation of VHE \( \gamma \)-rays of far distant sources. The reaction can occur if the center-of-mass energy of the photon-photon system exceeds twice the squared rest energy of the electron. The cross section for this process peaks when [40]

\[
E_{\gamma_{\text{VHE}}}E_{\gamma_{\text{EBL}}}(1 - \cos \theta) \approx 2(m_e c^2)^2
\]

where \( \theta \) is the collision angle. For example, for VHE photons of energy about 100 GeV, the highest cross section occurs for the head-on collision with photons of few eV.

VHE \( \gamma \)-rays have high cross section for pair production with the EBL, therefore the Universe is not completely transparent to VHE \( \gamma \)-rays, at least beyond 100 GeV. This means that far distant sources cannot be observed, and therefore a \textit{Gamma Ray Horizon} is obscuring us what is beyond. This is particularly relevant for the observation of Active Galactic Nuclei. The effect of the EBL absorption is energy dependent, therefore original spectra get distorted and the measured spectrum must be deconvoluted.

The \( \gamma \)-ray observed flux \( F_{\text{obs}} \), for a given energy \( E \) and redshift \( z \), can be written as an exponential cutoff:

\[
F_{\text{obs}}(E, z) = F_{\text{int}}(E)e^{-\tau(E, z)}
\]
where $F_{\text{int}}$ and $\tau(E, z)$ are respectively the intrinsic flux of the source and the optical depth, as a function of the energy and the redshift $z$. For any given $\gamma$-ray energy, the *Gamma Ray Horizon* is the distance at which the Universe becomes optically thick to $\gamma$-rays, namely when the optical depth is 1. In Fig. 1.5, the behavior of the optical depth as a function of the energy at different redshifts, for a fixed EBL model, is shown: it is clear that for a given source the energy, at which the spectrum cutoff occurs, decreases rapidly with increasing redshifts. Once the opacity of the Universe to $\gamma$-rays is known, it is in principle possible to derive the unabsorbed intrinsic emission spectra of TeV sources and test the emission models. The EBL is difficult to measure directly due to strong foregrounds from our solar system and the Galaxy. Therefore, the observation of distant sources of VHE $\gamma$-rays using IACTs provides a unique indirect measurement of the EBL.

**1.5 Astrophysical Sources of VHE $\gamma$-rays**

In the following a brief summary of the major classes of known VHE $\gamma$-ray emitters is reported. These objects are naturally divided in galactic and extragalactic sources.

**1.5.1 Galactic Sources**

Galactic objects are of basic interest in $\gamma$-ray astronomy for several reasons. Due to their vicinity, they are often observed as extended objects in a vast range of wavelengths, and thus provide a unique opportunity to carry out studies on the morphology of the emission region. Moreover, the $\gamma$-ray emission does not suffer absorption of cosmological background light.
On the other hand, they could be absorbed locally in the vicinity of the emission region and thus give a clear indication on the distribution of matter and gas around the source.

- **Supernova Remnants and Plerions.** A Supernova Remnant (SNR) is the left-over of a supernova explosion of a massive star when it runs out of fuel necessary for the fusion reactions that counteract the gravitational pressure. When a star is massive enough (5-10 $M_\odot$, with nucleus with mass greater than $\sim1.5\, M_\odot$), its life ends in a cataclysmic explosion which blows out a huge amount of material expanding into the interstellar medium of a lower density building a shock wave. In the core a pulsar or a black hole forms, depending on the mass of the remaining object, while a gas nebula (the SNR), expands into the surroundings. About 99% of the initial gravitational energy is released into neutrinos, 1% into kinetic energy of the remnants particles and only 0.01% of the energy goes into radiation. The remnant consists of ejected material expanding from the explosion and the interstellar material it sweeps up and shocks along the way. Particles are accelerated in these shocks by means of collisionless interactions of the particles with magnetic clouds in the interstellar medium (Fermi-I acceleration mechanisms [41]) leading to steady VHE $\gamma$-ray emissions.

SNRs are believed to be the principal sites for CR acceleration below $\sim5$ PeV and their spectrum is thought to be composed by a superposition of a synchrotron spectrum, from radio to soft-$\gamma$ wavelengths, and an IC part, from soft-$\gamma$ to VHE-$\gamma$ wavelengths. The first part is due to the interaction of relativistic electrons with magnetic fields, while the IC peak comes from the scattering of relativistic electrons with target radiation fields, such as synchrotron, thermal infrared or CMB radiation. The mechanism of particle acceleration in the expanding SNR shocks is considered well understood. However, clear evidence for the production of high energy hadronic particles in supernova shells has proven to be remarkably hard to find. Only observation of VHE neutrinos coming from the same direction of the supernova shells would unambiguously prove the theory.

If a fast rotating magnetized neutron star (pulsar) remains in the system, it is referred to as a Plerion or Pulsar Wind Nebula (PWN). Plerions are amongst the most abundant sources of VHE $\gamma$-rays in our galaxy. The rotational energy of the neutron star is converted, by some not fully understood mechanisms, into a relativistic stream of particles (mainly electrons and positrons) generated near the pulsar which power up the VHE $\gamma$-ray emission. Indeed, compared to the case of SNRs, the upscattering process is very efficient in generating $\gamma$-rays, which explains the abundance of PWNs detected in HE $\gamma$-rays.

The best known and studied example of such systems is the Crab Nebula (see Fig. 1.6(b)). The Crab Nebula is the remnant of a supernova explosion that occurred in the year 1054. It is located at a distance of 2 kpc and it is one of the best studied non-thermal celestial objects in almost all wavelengths of the electromagnetic spectrum from $10^{-5}$ eV (radio) to nearly $10^{14}$ eV ($\gamma$-rays). The brightness and stability of its emission in the VHE $\gamma$-ray regime make this source one of the most important sources used as standard “calibration candle” for ground-based and satellite borne $\gamma$-ray experiments. The data of Crab Nebula recorded by the MAGIC telescopes will be presented during this thesis.

- **Pulsars.** Neutron stars with masses of the order of the solar mass and radii of some ten kilometers represent the densest state of stable matter known in the Universe, apart from the black holes. They are created in the aftermath of supernova explosions and are characterized by very short rotational periods (down to the order of milliseconds) and extremely intense magnetic fields (typically in the range $10^8$ to $10^{12}$ G) believed to be
the highest ones existing in the Universe. The accelerated particles near the pulsar are beamed forming outflowing jets. Since the rotation and the magnetic axes of the pulsar are misaligned, the observer only sees an emission from a pulsar when the beam crosses his line of sight. Pulsed VHE γ-ray emission is predicted to originate either from particle acceleration and subsequent synchrotron emission near the Polar Cap of the pulsar or in the Outer Gap regions of the magnetosphere by IC scattering. The emission can be also a mixture of both mechanisms. In Fig. 1.7 a sketch of the two models is drawn. Very recently, the MAGIC Collaboration could observe for the first time in the history of
IACT astronomy, pulsed \(\gamma\)-rays from the Crab pulsar above 25 GeV \([43]\) (see Fig. 1.8). This observation was possible thanks to a new electronic trigger which allowed to lower the energy threshold to 25 GeV for \(\gamma\)-ray pulsed studies. The observation reveals a moderately high cutoff energy in the phase-averaged spectrum. This indicates that the emission occurs far out in the magnetosphere, hence possibly excluding the Polar Cap scenario and favouring an Outer Gap scenario.

**Figure 1.8:** Pulsed emission of the Crab pulsar in different energy bands. The shaded areas show the signal regions for the main pulse (P1) and the inter pulse (P2). From top to bottom: (a) Evidence of an emission above 60 GeV for P2 measured by MAGIC-I; (b) Emission above 25 GeV measured by MAGIC-I; (c) Emission above 1 GeV measured by EGRET; (d) Emission above 100 MeV measured by EGRET; (e) Optical emission measured by MAGIC-I with an optical pixel.

- **Microquasars.** Microquasars are astronomical systems composed of a pair of objects: one very compact object, such as a black hole or a neutron star, and a relative normal companion star. Matter of the companion star is stripped off and forms an accretion disk around the compact object, much like the way a miniature Active Galactic Nucleus forms jets in which the high energy emission is thought to take place. An associated X-ray emission directly comes from the stream of collapsing particles. Phenomenologically, microquasars have many of the observational and morphological properties of Active Galactic Nuclei, namely strong emission over a wide energy range, rapid flux variability and the existence of jets of relativistic plasma along the rotational axis of the black hole \([44]\). However, there are also many differences between the two systems \([45]\). First of all microquasars are galactic sources and the compact objects are of stellar origin. However, the main difference consists on the presence of a companion star which interacts with the compact object by causing precession of the jets and hence a periodic variability of the non-thermal emission.

The typical microquasar spectrum is dominated by the emission from the jets. It is generally believed that relativistic charged particles interacting with magnetic fields produce synchrotron radiation (from radio to X-rays). Relativistic electrons can scatter the UV photons emitted by the companion star through the IC mechanism. The IC
emission peak can reach the GeV-TeV energies.

- **Galactic center.** The center of the Milky Way is crowded with many astrophysical sources where history and evolution have played an important role in shaping the $\gamma$-ray emission. Indeed, the Galactic Center (GC) has been found to emit steady VHE $\gamma$-rays signal up to 10 TeV as reported by Whipple [46], CANGAROO [47], H.E.S.S. [48] and MAGIC-I [49] Collaborations. A sketch of the $\gamma$-rays sky around the GC is shown in Fig. 1.9, where both point-like and diffuse emission are reported, as observed by H.E.S.S.. In particular, MAGIC-I and H.E.S.S. results agree on a quite hard $\gamma$-ray spectrum with index $\sim -2.2$ and a light curve which does not show any variability on time-scales from hours to years. The interpretation of the emission mechanism is difficult since the region is packed with different potential sources, and the angular resolution of the current instruments is not sufficient to disentangle the exact location of the VHE $\gamma$-ray emission. The most important GC $\gamma$-ray emitter candidates are the supermassive black hole SgrA*, the young SNR SgrA East and the PWN G359.95-0.04. The presence of dense molecular clouds can also act as a target for any shock wave, therefore acceleration can take place there. More exotic scenarios including Dark Matter particles annihilation have been also taken into account. However, only very massive neutralino of the order of 10 - 20 TeV could explain the results, for which the $\gamma$-ray yield is expected to be 2-3 orders of magnitude lower than the measured flux [50].

![Figure 1.9](image1.png)

(a) GC skymap.  
(b) GC skymap after point-like sources subtraction.

**Figure 1.9:** Galactic Center observed by H.E.S.S. [51].

- **Diffuse galactic emission.** The H.E.S.S. Collaboration reported a diffuse VHE $\gamma$-ray emission along the galactic plane after subtracting point sources [51]. Diffuse $\gamma$-ray emission is thought to come mainly from hadronic collision of CRs with dense molecular clouds and dust in the galactic plane creating, among other, neutral pions with a subsequent decay into VHE $\gamma$-rays. A special agreement between the observed VHE $\gamma$-ray excess and distribution of molecular clouds supports this scenario. IC scattering of electrons onto seed photons pervading the galaxy can also contribute to the VHE emission. An increased angular resolution could probably disentangle the diffuse emission by localizing eventual anisotropies that could be associated to singular emitters. Noteworthy, EGRET [52] and specially the Fermi-LAT [53] also observed a diffuse HE $\gamma$-ray excess in the galactic plane region. In Fig. 1.10, the Fermi-LAT released full-sky map after 1 years of data taking is shown.
1. INTRODUCTION TO VERY HIGH ENERGY $\gamma$–RAY ASTROPHYSICS

1.5.2 Extragalactic Sources

As already mentioned, the Universe is not completely transparent to $\gamma$–rays. Due to their interaction with the photons of the EBL, $\gamma$–rays disappear through pair production. For this reason, it is believed that we cannot observe $\gamma$–rays in the VHE domain from objects at high redshifts (let say greater than $z \approx 1$ for energies above 100 GeV). The current map of extragalactic emitters in the VHE domain above 100 GeV is shown in Fig.1.11. The class of extragalactic emitters is almost completely filled with Active Galactic Nuclei.

Extragalactic VHE $\gamma$–ray sources
($E_{\gamma} > 100$ GeV)

Figure 1.11: Extragalactic VHE $\gamma$–ray sources skymap above 100 GeV. The MAGIC Zenith angle ranges are superimposed. Figures from http://wwwmagic.mppmu.mpg.de/~rwagner/sources/ (R. M. Wagner).
• **Active Galactic Nuclei.** Active Galactic Nuclei (AGNs) are galaxies with an active core which is brighter and more energetic than the ones of usual galaxies. Currently, AGNs are believed to be the most powerful sources of non-thermal energies in the Universe.

In the unified scheme of AGNs [54], a supermassive black hole (SMBH) with around $10^6$ to $10^{10}$ solar masses is assumed to comprise the central region of AGNs, with an extension of often only 1 pc in diameter. The SMBH is surrounded by a disk of hot plasma being accreting by the SMBH which emits thermal radiation. Further out around the SMBH, a dusty torus can obscure the view of the central part. It is composed of relatively dense molecular clouds which are responsible for the broad emission lines detected in some spectra above all in the optical range. Accordingly the region is called Broad emission Line Region (BLR). Farther away from the SMBH, there is also a Narrow emission Line Region (NLR) filled with molecular clouds less dense than those in the BLR. Between these two regions, electron scattering becomes the basic physical mechanism. Perpendicular to the disc, two plasma jets are created carrying out part of the angular momentum of accreted material. The jets extend up to several kpc distance scale from the SMBH and seem to harbor conditions adequate for very efficient cosmic particle acceleration.

The currently believed classification scheme for the many observed subclasses of AGNs is based on the mass of the central black hole, its evolutionary status, its accretion speed and on the orientation of the galaxy and the emitting regions with respect to the line of sight (see Fig. 1.12). The AGN family includes a variety of objects with different observational features but equal underlying theory, such as Radio Galaxies, Quasars and Seyfert Galaxies. If by chance the observer is aligned with the jets, the emission is strongly boosted so that the chance of observation of a distant object is increased. An AGN observed from this point of view is called a blazar. This is the most common class of observed AGNs.

AGNs are of the most varying emitters in the sky: due to variation in the in-fall of materials, the flux can vary by up to two orders of magnitude in intensity. The observation of these flares are of utmost importance to study the source morphology and characteristics. The production process of VHE $\gamma$-rays is still under debate: both leptonic and hadronic models seem to be able to describe well the observational data.

![Figure 1.12: Classification of Active Galactic Nuclei.](image)

An AGN is constituted by an accreting torus composed of star and materials in-falling into the central supermassive black hole. In the perpendicular plane two co-aligned jets emerge and can extend up to several times the extension of the galaxy itself. When observing an AGN from the side of the jets, the obscuring torus shadows the black hole and the galaxy is either radio quiet or radio loud. When the jet is co-aligned with the observer the AGN is extremely VHE $\gamma$-ray emitting.
• **Gamma Ray Bursts.** Gamma Ray Bursts (GRBs) are the most energetic and violent short-term phenomena observed in the Universe. Their luminous flashes of photons last from fractions of seconds to minutes and they occur with a frequency of about 1-2 per day, as it was established by BATSE (Burst And Transient Source Experiment) [55] hosted in the CGRO satellite which observed about 2700 events (see Fig. 1.13). There are several models describing the phenomenon of GRB, but none of them can explain it in a complete and coherent way. The widely accepted current understanding of these highly transient bursts are asymmetric supernova or hypernova explosions. In the *Fireball* model [56], the explosion produces an ultrarelativistic outflow of an optically thick plasma shell, which emits the GRB as soon as it becomes optically thin. The short duration of GRBs (milliseconds to few hundreds seconds) hints at very compact progenitors. The mean distance of GRBs is about z = 2.8, which means that in most cases the VHE $\gamma$-rays are absorbed by the EBL and only HE $\gamma$-rays can reach the Earth. However, few GRBs occur at redshifts much closer than z = 1, which makes a detection possible. Typically $10^{51} - 10^{54}$ erg s$^{-1}$ are released within seconds up to tens of seconds. This *prompt emission* is followed by an *afterglow* of less energetic photons in all further domains of electromagnetic radiation. The delayed emission can last from days to years, according to the type of emission. X-ray afterglows were detected for the first time in 1996 by the Italian-Dutch satellite BeppoSAX [57]. Thanks to its accuracy in determining their position, they were detected also by ground-based telescopes, which allowed to establish the cosmological origin of such extreme phenomena. As shown in Fig. 1.13, GRBs are distributed uniformly in the sky and reveal no evidence for clustering.

A dedicate satellite for GRB studies, SWIFT [58], has been launched into space in September 2004 and it is able to send GRB alerts and their position within 10 seconds. Thanks to these alerts (and also to the alerts of the GLAST Burst Monitor (GBM) onboard the Fermi satellite) and to the fast repositioning system, the MAGIC telescopes have the possibility to observe GRBs in their prompt emissions. As of now, about 50 GRBs has been observed by MAGIC-I, however no significant VHE $\gamma$-ray emission was found.

![Figure 1.13: The 2704 GRBs detected by BATSE [55]. Note the isotropic distribution of this kind of phenomena.](image-url)
1.6 Connections to Fundamental Physics

The study of VHE cosmic γ-rays could probe fundamental issues and have a potential of great discoveries. In the next subsection the main topic which could be investigated by studies related to VHE γ-ray emissions are briefly presented.

1.6.1 Lorentz invariance

Lorentz Invariance (LI) was a fundamental finding in the history of physics and it has been repeatedly confirmed to an ever greater precision. So far, it is believed to be a symmetry of Nature. However, there are general Lorentz-violating extension of the standard model. Indeed, quantum gravity theories introduce quantum fluctuations on the Plank scale, characterized by energies of $E_{QG} = \sqrt{\hbar c^5/G_N} = \hbar c/L_{QG} \simeq 10^{28} \text{ eV}$, where $\hbar$ is Dirac’s constant, $c$ is the velocity of light, $G_N$ is Newton’s constant and $L_{QG} \simeq 1.6 \times 10^{-35} \text{ m}$ is the Planck’s length. These fluctuations result in a complex “foamy” structure of space-time at very small distances which gives rise to deformed dispersion relations which may result in observable energy dependent arrival time dispersions for high energy radiation, due to the fact that the speed of light could be no longer considered a constant of Nature. Any experimental evidence for such a violation of the LI would be a clear signature of new physics beyond General Relativity.

A natural place where to look for this effect is during short flares of far away AGNs. The large distance allows to increase the time discrepancy between photons and the flare allows to provide the time stamp of the produced photons. However, a possible detected delay of arrival times as function of the γ-ray energies could be intrinsically dependent on the source emission mechanisms. The MAGIC Collaboration reported a lower limit to the quantum gravity scale after a flare occurred in Mkn 501 (see [59] for more details).

1.6.2 Dark Matter

Nowadays there are compelling experimental evidences for a large non-baryonic component of the matter density of the Universe at all observed astrophysical scales, such as galaxies, galaxy clusters and cosmic background radiation [60]. The so-called Dark Matter (DM), believed to constitute about 80% of the matter of our Universe, makes its presence known through gravitational effects and it could be made of so far undetected relic particles from the Big Bang. One of the best theoretical candidates is the neutralino, the lightest supersymmetric weakly interacting massive particle. The neutralino, being a Majorana particle, can undergo a self annihilation process with subsequent (direct and indirect) emission of high energy γ-rays which could be potentially detected by the γ-ray instruments in operation, such as the Fermi-LAT detector and the IACTs. This topic will be further discussed in chapter 6. In chapter 7, the analysis of the data recorded by the MAGIC-I telescope of the Segue 1 Galaxy (a satellite galaxy of the Milky Way), which is supposed to be dynamically dominated by DM and characterized by a huge mass-to-light ratio of the order of $10^3 M_\odot/L_\odot$, will be reported.
Figure 2.1: Atmospheric windows for the observation of the Universe [37]. The continuous line indicates the height at which a detector can receive half of the total incoming radiation for each wavelength.
Electromagnetic waves, the main messengers from outer space, can be detected directly from the ground only in small energy bands in the optical and radio range, as shown in Fig. 2.1. Observations in nearly all other energy ranges require detectors to be placed outside the Earth’s atmosphere. In particular, VHE $\gamma$-rays cannot be detected directly by ground-based instruments, because they are absorbed in the atmosphere. In fact, they interact with nuclei of the upper layers of the atmosphere, initiating a shower of secondary particles called Extensive Air Showers.

Imaging Atmospheric Cherenkov Telescopes (IACTs) are ground-based detectors whose aim is to measure physical properties of primary VHE $\gamma$-rays by means of the Cherenkov light produced in the atmospheric showers induced by these $\gamma$-rays. In this chapter the physics of the atmospheric showers, the main characteristics of the subsequent production of the Cherenkov light, the technique used by the IACTs and the main sources of background for this kind of detectors are briefly summarized.

2.1 Atmospheric Showers

The first experimental evidences of the phenomenon of the atmospheric showers are dated 1938 and associated to the name of Pierre Victor Auger. The scientist, by performing experiments at high altitude carried out by the aid of Geiger counters, observed temporally coincident signals which could be attributed to bunches of particles belonging to the so-called Extensive Air Showers (EAS) [2].

Hadrons, charged light leptons and VHE photons when entering the atmosphere interact with atmospheric nuclei and lose their energy inducing the creation of secondary particles which in turn give rise to the production of an avalanche of particles which traverse the atmosphere with nearly the vacuum speed of light. The particles of the shower are strongly collimated along the direction of the incoming primary particle because of their relativistic energies. The physical processes involved in the development of the showers are somehow different depending on the nature and on the energy of the primary cosmic-ray (CR). In Fig. 2.2, a schematic view of air showers induced by a VHE $\gamma$-ray and a hadron is shown. The differences in the morphology and temporal feature of the shower developments can be used to discriminate the nature of the primaries.

2.1.1 Electromagnetic Atmospheric Showers

Electromagnetic (EM) showers are generated by VHE cosmic electrons, positrons and $\gamma$-rays. Two main processes contribute to the shower development induced by primary $\gamma$-rays with energy greater than $\sim 100$ MeV: the Bremsstrahlung, which leads to the radiation of new photons, and the pair production, in which a pair of $e^+ e^-$ is created\(^1\). The two processes are very similar: the radiation length of an electron or positron for the Bremsstrahlung is $X_{e0} = 37 \text{ g/cm}^2$ in air, and the mean free path of the photon before pair creation (interaction length) is $X_{0} = 7/9 X_{e0}$.

When a $\gamma$-ray impinges on the upper layers of the atmosphere, after interacting with the electric field of atmospheric nuclei, it produces a pair of $e^+ e^-$. Since the cross section of VHE $\gamma$-rays with air is weakly energy dependent, this first interaction occurs for different primary $\gamma$-ray energies typically at a height of about 20 km above the sea level. Each electron-positron in turn radiates new $\gamma$-rays via Bremsstrahlung which generate other pairs of $e^+ e^-$. The processes of Bremsstrahlung and pair production alternate, giving rise to the EM shower. During its development, the number of particles increases approximately in exponential way and the average energy per particle reduces accordingly. When the average

\(^1\)The cross section for muon-antimuon pair production is many orders of magnitude smaller due to the larger mass of the muon compared to the electron one.
energy per particle reaches a critical value $E_c$ (which is \( \approx 83 \) MeV in air), the energy losses per unit length by Bremsstrahlung and by ionization are equal. Almost contemporary, when the mean photon energy decreases below few MeV, the cross section for Compton scattering and photo-electric effect become dominant over pair production. At this stage of the shower development the number of secondary shower particles has reached its maximum. The height above the sea level where the number of particles in the shower is highest is defined as the shower maximum and occurs at \( \approx 13 \) to \( \approx 7 \) km above the sea level for energies of primary $\gamma$-rays between \( \approx 50 \) GeV and \( \approx 10 \) TeV. Below this height ionization losses dominate and the shower rapidly dies out.

In 1944 Heitler introduced a simple model which can describe in good approximation the characteristics of the development of an EM shower [37]. In this model, only the processes of pair production by energetic photons and of Bremsstrahlung emitted by the light charged leptons are considered. The model also makes two further assumptions, namely that at each vertex the energy is equally distributed between the produced particles and that the radiation length and the interaction length are equal (\( X_0^\gamma = X_0^e = X_0 \)) (see Fig. 2.3). According to this simple model, after $n$ radiation length $X_0$ the amount of particles is $N(n) = 2^n$, equally distributed between electrons, positrons and photons, whereas the mean energy per particle is $E(n) = E_0/2^n$, where $E_0$ is the primary $\gamma$-ray energy. The number of radiation length $n_{\text{max}}$ at which the maximum of the shower is reached (which occurs at the critical energy $E_c$) and the total number of produced particles at the depth $n_{\text{max}} \cdot X_0$ are thus

$$E_c = E_{n_{\text{max}}} = \frac{E_0}{2^{n_{\text{max}}}} \Rightarrow n_{\text{max}} = \frac{\ln(E_0/E_c)}{\ln 2} \quad (2.1)$$
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\[ N_{\text{max}} = 2^{n_{\text{max}}} \frac{E_0}{E_c} \] (2.2)

The number of particles with energy greater than \( E \) is

\[ N(> E) = \int_0^{n(E)} N \, dn = \int_0^{n(E)} 2^n \, dn = \frac{e^{n(E) \ln 2}}{\ln 2} = \frac{E_0/E}{\ln 2} \] (2.3)

where \( n(E) \) is the depth, in radiation length units, at which the mean energy of the particles is \( E \). From this last relation it is possible to infer that the differential energy spectrum of the particles decreases as

\[ \frac{dN}{dE} \propto 1/E^2 \] (2.4)

The results obtained by the Heitler model, namely that the depth at which the shower reaches its maximum is proportional to the logarithm of the primary \( \gamma \)-ray energy and that the number of particles in the maximum development of the shower is proportional to the energy of the primary \( \gamma \)-ray, are confirmed by more detailed studies based on Monte Carlo (MC) simulations.

More complicated models were developed by Rossi and Greisen in the 1940s [61]. It is possible to achieve analytical solutions of the shower equations if some approximations are made. In the so-called B approximation, Compton effect and photo-production mechanism are neglected. The number of electrons above the critical energy \( E_c \) in the shower can be thus

**Figure 2.3:** Heitler model. Each level corresponds to one radiation length \( X_0 \).
expressed as

\[ N_e(t, E_0) = \frac{0.31}{\sqrt{\ln(E_0/E_c)}} \cdot \exp[t(1 - 1.5 \ln s)] \] (2.5)

\[ s = \frac{3t}{t + 2 \ln(E_0/E_c)} \] (2.6)

\[ t = X/X_0 \approx X_{\text{air}}/X_0 \cos \theta_{\text{inc}} \cdot \exp\left(-\frac{h}{h_0}\right) \] (2.7)

with \( E_0 \) the energy of the primary photon, \( X_{\text{air}} = 1013 \text{ g cm}^{-2} \) the column height of air at ground level and \( h_0 \approx 8 \text{ km} \) the scale height of the atmospheric pressure. \( \theta_{\text{inc}} \) is the incident angle of the air shower. The slant depth \( t \) determines the thickness of the atmosphere in units of \( X_0 \). The parameter \( s \) is the so-called shower age which measured its temporal evolution. It goes from 0 at the first interaction, through 1 at the shower maximum (at \( t_{\text{max}} = \ln(E_0/E_c) \)), and reaches 2 at the point where the shower dies out. The longitudinal electron distribution given by equation 2.5 is shown if Fig. 2.4. The radiation length corresponding to the altitude above the sea level where the MAGIC telescopes are located (\( \sim 2230 \text{ m} \), corresponding to \( \sim 22 X_0 \)) is also drawn. Air showers induced by \( \gamma \)-rays of energies typically detectable by the MAGIC telescopes (\( \sim 30 \text{ GeV} \) to \( \sim 30 \text{ TeV} \)) develop their maxima well above the telescopes’ level.

The shower fluctuations (not including fluctuations of the height of the first interaction in atmosphere) can be expressed as

\[ \Delta N_e(s) \simeq \frac{9}{14} (s - 1 - 3 \ln s) \cdot N_e(s) \] (2.8)

Due to their statistical nature, the fluctuations from shower to shower can be rather large and thus can have important consequences e.g. for the energy estimation of the primaries.

An EM shower is strongly collimated along the incident direction because of the relativistic velocities of its particles, even though multiple Coulomb scattering and, in second order, the Earth’s magnetic field effect, broaden the shower transversely. The lateral distribution is explained by Molière theory which takes into account only multiple scattering and describes it as a gaussian for small deflection angles with a long tail. An analytic expression for the lateral distribution of the electrons, which is strictly valid for \( 1 \leq s \leq 1.4 \), can be parametrized by the Nishimura, Kamata, Greisen (NKG) formula [62]:

\[ \rho_e(r, t, E_0) = \frac{\Gamma(4.5 - s)}{2\pi \Gamma(s)(4.5 - 2s)} \cdot \frac{N_e(t, E_0)}{R_M^2} \cdot (r/R_M)^{s-2} \cdot (1 + r/R_M)^{s-4.5} \] (2.9)

where \( R_M = 21.2 \text{ MeV} \cdot (X_0/E_c) \approx 80 \text{ m} \) in air at the sea level is the Molière radius. About 99% of the shower energy is contained within \( 3.5R_M \), whereas only 10% are “lost” outside \( 1R_M \).

It is worth mentioning that air showers initiated by electron/positron primaries constitute an irriducible background for \( \gamma \)-ray shower detectors, as they develop in the same way the \( \gamma \)-ray showers do. However, due to the fact that the cosmic light leptons isotropically hit the atmosphere, this background can be neglected in most of the cases, since, on the contrary, the \( \gamma \)-ray showers point back to their astrophysical sources.
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Figure 2.4: Longitudinal development of EM showers. The curves correspond to different energies of the primary $\gamma$-rays. The values $\ln(E_0/E_C)$ are written above each curve and the values of the shower age $s$ is also reported. The atmospheric depth at the sea level is 28 radiation lengths, while at the MAGIC telescopes’ altitude it is nearly 22.

2.1.2 Hadronic Atmospheric Showers

Most of the primary CRs are composed by protons and nuclei of He. The strong interaction of these particles with the nuclei in atmosphere generates mainly pions ($\sim90\%$), kaons ($\sim10\%$) and light baryons (p, $\bar{p}$, n, $\bar{n}$). Since the energy of the primary can reach UHE, the production of heavier charmed mesons and baryons is also possible. These secondary particles, together with the primary CR, form the so-called core of the hadronic shower. The core develops with further interactions and decays, diluting gradually the energy between the produced particles. During the initial development of the shower, until the energy per nucleon is above the multiple production pion threshold ($\sim1$ GeV), the strong interactions are the dominant processes and hadronic particles are created, whereas in the final stage of the shower ionization and decays become the dominant processes and the shower starts to die out.

The fact that the radiation length $X^0_h$ for hadrons ($\approx 70$ g/cm$^2$ for protons) is higher than that for electrons $X^0_e = 37$ g/cm$^2$ has the consequence that the hadronic showers are more penetrating and extended than EM ones. Moreover, since the primary cosmic hadrons may lose in atmosphere up to half of their initial energy in the first shock, large fluctuations of the height of the maximum of the shower are present.

The air showers induced by hadrons are composed mainly by three components:

1. a hadronic core built up from high energy nucleons and mesons. Charged pions and kaons decay into muons and neutrinos through the channels:

\[
\pi^\pm \rightarrow \mu^\pm + \nu_\mu (\bar{\nu}_\mu) \quad (2.10) \\
K^\pm \rightarrow \pi^\pm + \pi^0 \quad (2.11) \\
K^\pm \rightarrow \mu^\pm + \nu_\mu (\bar{\nu}_\mu) \quad (2.12) \\
K_L^0 \rightarrow \pi^\pm + \mu^\pm + \nu_\mu (\bar{\nu}_\mu) \quad (2.13)
\]
2.2. Cherenkov Light

2. a muonic component which consists of muons produced by the meson decays. Muons basically lose energy via ionization, even if they can also decay in electrons:

\[ \mu^\pm \rightarrow e^\pm + \nu_e (\bar{\nu}_e) \]  

Only a small fraction of their energy is released into the electromagnetic component. Due to their large lifetime (\(2.2 \cdot 10^{-6} \text{ s}\)), muons can travel relevant portion of the atmosphere: they only rarely decay before reaching the ground due to their high initial energy. Isolated muons with large transversal momenta can travel far away from their parent shower.

3. an EM component mainly due to the neutral pion decay \(\pi^0 \rightarrow \gamma\gamma\). About one third of the produced \(\pi\)-mesons are neutral, so about one third of the energy in hadronic interactions is transferred to EM components. The \(\gamma\)-rays from pion decays in turn produce EM subshowers which constitute the main background for ground-based \(\gamma\)–ray detectors, as they are practically indistinguishable from showers initiated by primary \(\gamma\)-rays.

A hadronic shower is therefore characterized by a core which proceeds along the direction of the incidence primary CR and by a series of subshowers, both hadronic and electromagnetic, with high transverse momenta (produced by the decay of heavy hadrons generated with high transverse momentum) that make the lateral distribution of the hadronic showers much wider and irregular than the EM showers one (see Fig. 2.5). This is one of the main features which is used to discriminate \(\gamma\)-ray induced EAS from hadronic induced ones. However, subcas-cades of hadronic EAS, if only the EM subshower is detected, can mimic \(\gamma\)-ray induced EAS making them to an almost irreducible background. However, the most striking background suppression comes about by the fact that \(\gamma\)-ray showers point back to the source of their primary \(\gamma\)-rays, while the arrival direction of charged CR initiated air showers is isotropized.

The simplest hadronic shower model is the so-called Superposition Model described by Gaisser [64]. Assuming that a nucleus with atomic mass \(A\) and energy \(E_0\) is equivalent to \(A\) independent nucleons with energy \(E_0/A\), and that the height of first interaction is the same for all independent considered nucleons, the position of the shower maximum can be described by

\[ X_{\text{max}} \simeq \ln \left( \frac{E_0}{AE_c} \right) \cdot \xi_n \]  

where \(\xi_n\) is the nuclear interaction length in air. The equation shows that heavier nuclei are less penetrating than lighter ones. Moreover, the fluctuations of the position of \(X_{\text{max}}\) is smaller for heavier nuclei because each of them is equivalent to many lighter nuclei. In order to obtain a full determination of the hadronic shower characteristics in the atmosphere, MC simulations are necessary.

2.2 Cherenkov Light

In 1948, a decade after the discovery of the Cherenkov effect, Blackett advanced the hypothesis that a fraction of the night sky brightness (of the order of \(10^{-4}\)) was due to the Cherenkov light emitted in atmosphere by the secondary charged particles produced by CRs [65]. Indeed, air showers contain a large number of relativistic charged particles with a speed close to the speed of light in the vacuum which exceeds the local speed of light of the traversed medium and thus they can emit Cherenkov light.
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Figure 2.5: Development of showers induced by 1 TeV $\gamma$-ray (left) and 1 TeV proton (right). The different lateral distribution of the two showers is used to discriminate the nature of the primary. The pictures were created using CORSIKA [63].

IACTs detect this kind of light, mainly in the wavelength between 300 nm and 600 nm, by collecting it through large reflecting surfaces. The light is then focalized onto a multi-pixel camera made by photomultiplier tubes (PMTs) in order to reconstruct the shower development. The offline analysis of the recorded images gives then the possibility to discriminate the nature of the showers, allowing the determination of the amount of $\gamma$-rays coming from a certain observed source.

2.2.1 Cherenkov Effect

The Cherenkov effect was first discovered in 1934 by the Russian physicist during a research on the visible light induced in different liquids under the action of $\gamma$ radiation [66]. A complete theory of the effect was further derived by Tamm and Frank in 1937 [67].

Cherenkov light is produced whenever a charged particle travels through a transparent dielectric medium with refractive index $n$ with a velocity $\beta c$ exceeding the local phase velocity of light ($c/n$). The radiation is due to the in-phase reorientation of the electric dipoles induced by the charged particle in the medium. A polarization of the medium occurs along the particle trajectory because the particle is moving faster than the electromagnetic wave which induces the polarization itself. This radiation has an analogy in acoustics in the form of the shock waves produced by a projectile or an airplane traveling at an ultra-sonic speed (Mach waves).

The emission angle of the radiation, which defines the so-called Cherenkov light cone, can be
2.2. Cherenkov Light

inferred by simple geometrical arguments (see Fig. 2.6) and it turns out to be

$$\cos \theta = \frac{\Delta t \cdot c/n}{\Delta t \cdot \beta c} = 1/\beta n$$

(2.16)

The maximum Cherenkov angle is attained for $\beta = 1$, i.e. $\theta_{\text{max}} = \arccos(1/n)$. For a given medium, equation 2.16 determines a minimum particle velocity $\beta_{\text{min}} = n^{-1}$ which is required to produce Cherenkov light. The Cherenkov condition can be converted in a kinetic energy threshold for particles with rest mass $m_0$:

$$E > m_0 c^2 \cdot \frac{1}{\sqrt{1 - \beta_{\text{min}}^2}} = m_0 c^2 \cdot \frac{1}{\sqrt{1 - 1/n^2}}$$

(2.17)

The Cherenkov radiation, with spectrum mainly in the visible and near-UV, therefore propagates along a conical surface determined by the value of the angle $\theta$ with axis the trajectory of the particle (Cherenkov cone). Due to the dispersion (the refractive index depends on the wavelength of light) the cone of radiation actually has a certain thickness whose inner part is determined by the shorter wavelengths.

![Figure 2.6:](image)

Figure 2.6: Representation of the Cherenkov effect. (a) Symmetric polarization induced by a particle moving with velocity $v < c/n$. The orientation of the medium charges around the charged particle is symmetric so there are no effects at large distances and thus no emission of light takes place. (b) The net polarization due to a charge moving at $v > c/n$ creates a wave of radiation. (c) Geometrical description of the Cherenkov angle $\theta$ of emission.

2.2.2 Propagation of Cherenkov Light through the Atmosphere

The Cherenkov radiation produced in air has typical characteristics due to the refractive index of the medium close to 1. At the sea level, the refractive index of air is $n \approx 1.00029$. This implies that the maximum emission angle is $\theta_{\text{max}} \approx 1.3^\circ$ and that the energy threshold for Cherenkov light emission for electrons, muons and protons are 21.3 MeV, 4.4 GeV and 39.1 GeV respectively.

In order to compute the energy threshold for Cherenkov radiation at different atmospheric depths, it is necessary to take into account the exponential variation of air pressure as a function of the height $h$ [68]. The atmospheric density $\rho(h)$ can be written, in the isothermal
atmosphere approximation, as
\[ \rho(h) = \rho_0 \cdot e\left(-h/h_0\right) \] (2.18)
where \( h_0 \approx 8 \) km is the scale height of the atmospheric pressure and \( \rho_0 = 1.22 \cdot 10^{-3} \) g/cm\(^3\) is the density at the sea level. In first approximation the refractive index can be thus re-written as
\[ n(h) = 1 + \eta_0 \cdot e\left(-h/h_0\right) \] (2.19)
where \( \eta_0 = 2.9 \cdot 10^{-4} \).

Considering that the refractive index changes only about 2.5\% in the visible range at ground [69], the dependence of the refractive index on the wavelength \( \lambda \) can be neglected leading to the expression
\[ E_{th} = \frac{m_0 c^2}{\sqrt{1 - n^2}} \simeq \frac{m_0 c^2}{2 \eta_0 e\left(-h/h_0\right)} \] (2.20)
Since as height increases, \( \eta_h \) decreases, the energy threshold to emit Cherenkov light increases with height. For example at 10 km above the sea level \( E_{th} \) is about twice larger as the one at the sea level (~42 MeV for electrons). Since at that altitude \( E_{th} \) is smaller than the critical energy \( (E_c = 83 \) MeV), once an EM shower dies out, most of the electrons still emit Cherenkov light. For an EM shower, the number of electrons with energy greater than \( E_{th} \) as a function of the number \( N_e \) of electrons present in the shower maximum is given by
\[ N(E > E_{th}) \approx \frac{N_e}{1 + E_{th}/30} \] (2.21)
with energies expressed in MeV. Therefore, about half of the electrons at the shower maximum has energy greater than the minimum one necessary for emitting Cherenkov light.

The maximum angle \( \theta_{max} \) for particles with \( \beta \simeq 1 \) can be obtained from
\[ \cos \theta_{max} \simeq \frac{1}{1 + \eta_0 \cdot e\left(-h/h_0\right)} \simeq 1 - \eta_0 \cdot e\left(-h/h_0\right) \] (2.22)
and has an averaged over altitude value of about 1.2\%. By knowing \( \theta_{max} \), the distance from the emitted Cherenkov photons to the axis of the emitting charged particle at a given height \( h_{obs} \) can be derived:
\[ R_c = (h - h_{obs}) \cdot \tan \theta_{max} \] (2.23)

The Cherenkov radiation of an EAS consists of a cumulative Cherenkov light emitted by the charged shower particles. The light from the shower tail is emitted with larger angles, but its distance to the shower axis is smaller just because of its lower height (see Fig. 2.7(a)). In case of an EM shower, Cherenkov light illuminates typically a circle\(^2\), called Cherenkov light pool, with a radius of ~120 m if measured at about 2000 m above the sea level. A sketch of the Cherenkov light pool for 100 GeV \( \gamma \)-ray and for 400 GeV proton induced air showers is shown in Fig. 2.7(b). The energies of the two primaries were chosen in order to have a similar overall amount of Cherenkov light (see Fig. 2.8). The figure shows that for \( \gamma \)-ray induced showers the Cherenkov photon density is almost constant from the center to the rim at about 120 m. In this region the Cherenkov photon density is proportional to the energy of the primary. The region between the center and ~80 m is called plateau, whereas the rim with higher density is called hump. The higher density of the hump is due to the increasing opening angle \( \theta \) of the Cherenkov photons as the light emitting particles penetrate deeper.

\(^2\)A circle forms in case of vertical air shower. In case the EAS is inclined, the overlapping Cherenkov light illuminates an ellipse on the ground.
into the atmosphere. At distances larger than the hump the photon density decreases rapidly. Still, there is light emission even beyond 200 m impact distance, caused by the shower halo particles. The area hit by Cherenkov photons is very large, allowing the detection of photons with a large impact parameter. On the other hand, the photon density is very low (few photons per m$^2$ for a 100 GeV $\gamma$-ray induced shower), that is why ground-based telescopes need large light collection surfaces. Conversely, the lateral distribution of Cherenkov light density for hadron induced air showers shows a completely different behavior related to the intrinsic differences in the shower development with respect to the EM shower case: it is peaked at small distances to the shower axis and decreases constantly. The photon density inside the Cherenkov light pool as a function of the energy of the primary particle is given for different particle types in Fig. 2.8. For $\gamma$-rays, an almost constant fraction of the primary energy is converted into Cherenkov photons. Therefore a measurement of the Cherenkov light intensity is a good measure of the primary $\gamma$-ray energy. This means that the atmosphere behaves almost like an ideal calorimeter for $\gamma$-ray induced showers. On the contrary, the same cannot be said for hadron induced showers. Moreover, as the energy decreases the hadronic showers produce increasingly less Cherenkov photons.

### 2.2.3 Intensity and Spectrum of the Cherenkov Radiation

The number of emitted Cherenkov photons per unit track length for a particle with one elementary charge and per wavelength of the photon is [72]:

$$\frac{d^2N}{dx d\lambda} = \frac{2\pi\alpha}{\lambda^2} \cdot \left(1 - \frac{1}{\beta^2 n^2(\lambda)}\right)$$  \hspace{1cm} (2.24)

$$\frac{d^2N}{dE d\lambda} \approx 370 \cdot \sin^2\theta(E) \ [\text{eV}^{-1} \text{cm}^{-1}]$$  \hspace{1cm} (2.25)
where $\alpha$ is the fine-structure constant. In an EM air shower about 500 Cherenkov photons are created per GeV of primary $\gamma$–ray energy between 300 nm and 600 nm for $\beta \simeq 1$. The spectral intensity has a strong $\lambda^2$ dependency of the photon wavelength $\lambda$. As shown in Fig. 2.9, most of the photons are emitted in the UV range. However, the emitted photons suffer transmission losses in the atmosphere mainly caused by Rayleigh scattering of air molecules with diameter of the order of few nm. This process has a cross section with a $\lambda^4$ dependence and affects the UV/blue part of the spectrum. It represents the dominant contribution to light transmission losses during good atmospheric conditions. Another more complex process which affects the light transmission is the Mie scattering of aerosols particles (also water droplets, dust) which shows no strong dependence on the photon wavelength. The Mie scattering is difficult to model because the presence of aerosol particles is subject to continuous variations on a hour scale. Light with $\lambda < 280$ nm is strongly attenuated due to UV absorption by ozone molecules. Infrared absorption occurs for wavelengths above 800 nm caused by H$_2$O and CO$_2$ molecules, and does not play an important role for Cherenkov telescopes. In Fig 2.9 all these effects are evident from the comparison of the Cherenkov spectrum around the shower maximum with that measured at 2200 m above the sea level for different $\gamma$–ray primary energies. The measured Cherenkov photon spectrum mainly ranges from $\sim$300 nm and $\sim$600 nm with a peak at about 320 nm. It is worth mentioning that the Cherenkov spectrum is Zenith angle dependent with a shift of the peak position towards larger wavelengths for an increasing Zenith angle of the atmospheric showers.

The typical duration of the Cherenkov flash produced by a $\gamma$–ray initiated air shower is of the order of 3 ns for impact parameters below $\sim$120 m. On the contrary, for a hadron induced shower the time spread is slightly wider ($\sim$10 ns). These extremely short flashes are very difficult to measure due to a high background of the night sky. The light of the night sky (LONS) consists of light of bright and faint stars, diffuse light of the galactic plane, zodiacal light, ionospheric airglow, polar light, Cherenkov light, moonlight and artificial light (see [73] for more details). The LONS is not isotropically distributed and is higher towards e.g. Galactic plane than for a typical extragalactic field of view. Moreover it increases with increasing Zenith distance. The typical night sky photon flux between 300 and 600 nm has

Figure 2.8: Cherenkov photon yields for different particle species within 125 m of the shower core for vertically incident showers. In case of $\gamma$–ray induced air shower the atmosphere behaves almost like an ideal calorimeter. Plot taken from [71].
been measured to be \((1.75 \pm 0.4) \times 10^{12}\) photons m\(^{-2}\) sr\(^{-1}\) s\(^{-1}\) at the MAGIC telescopes' site \cite{74}. This translates to an average single photoelectron rate due to the night sky background in one 0.1° diameter pixel of the MAGIC-I camera of \(~10^8\) MHz. The given number are valid for extragalactic dark areas of the sky with no bright star present, and for clear moon-less nights. In fact, the presence of the moon in the sky can deteriorate the Cherenkov light detection. Since the moon contribution is uniform in the MAGIC telescopes’ cameras, the pedestal fluctuations are larger, but this does not make impossible the measurements. However some precautions must be followed: the measurements are taken only with an angular distance to the moon above 50° up to three or four days before and after full moon.

One of the key element of the Imaging Atmospheric Cherenkov technique, which will be presented in next section, is therefore the discrimination between light originating from EAS and LONS, which requires an integration time of the device of the order of the typical duration of the EAS temporal development. This makes, for example, an usage of CCD cameras impossible because they need milliseconds for readout. Instead, the usage of PMTs to measure Cherenkov light of atmospheric showers has been proven to be efficient, due to their very fast (few ns) response time.

2.3 The Imaging Atmospheric Cherenkov Technique

The discrimination between atmospheric showers induced by primary hadrons and VHE \(\gamma\)-rays is one of the most important task for the ground-based experiments in the \(\gamma\)-ray astronomy. In fact, the ratio between the number of showers generated by hadrons and that induced by \(\gamma\)-rays is always typically greater than \(10^4\), even during the observation of the most bright \(\gamma\)-ray sources, such as the Crab Nebula. Among the different developed techniques, the Imaging Atmospheric Cherenkov (IAC) technique is so far the most efficient and reliable one, in the energy range between few tens of GeV and several TeV. Currently, four
big Collaborations operate world-wide four IACTs which are shown in Fig 2.10. The IAC technique consists in collecting a sample of the Cherenkov light produced by an atmospheric shower through large reflecting surfaces placed inside the Cherenkov light pool, and then in focusing it in a multi-pixel camera equipped by PMTs which record the intensity and the temporal development of the Cherenkov flash. Short exposure times (~10 ns) and a trigger concept that exploits spatial and temporal coincidences between different pixels allow for the discrimination between EAS events and statistical fluctuations of the steady LONS flux. A sketch of the IACT detection of atmospheric showers is shown in Fig. 2.11. The camera is placed in the focal plane of the reflector, where parallel light rays are focused to the same point (see Fig.2.12(a)). Photons which reach the telescope under an angle $\beta$ with respect to the telescope axis are focused, under parallax approximation$^3$, at a distance $\rho$ from the focal point (see Fig. 2.12(b)). Fixing a polar coordinate system in the focal plane, the focused position $(\rho, \phi)$ can be calculated from the incoming direction of the photon $(\beta, \Phi)$ by

$$\rho = \text{sen}(\beta) \cdot f \simeq \beta \cdot f$$

$$\phi = \Phi$$

where $f$ is the parabolic mirror surface focal length of the reflecting surface (17 m in case of the MAGIC telescopes). Therefore, the distance from the focal point (the telescope camera center) goes linearly with the incoming angle, i.e. there is a univocal relation between illuminated pixels in the camera and (almost) parallel beams of Cherenkov photons with a certain incoming angle $\beta$. The bigger the reflectors, the more Cherenkov photons can be collected from a single shower and, hence, the better image of the EAS can be recorded. In order to achieve high sensitivity, high Quantum Efficiency (QE) fast photo detectors and fine pixelized camera are required. The subsequent analysis of the recorded images allows to discriminate, on a statistical basis, the major part of the hadronic background from the $\gamma$-ray events.

$^3$The “small angle approximation” is justified by the limited angular acceptance of the IACTs’ cameras (typically $<5^\circ$).
2.3. The Imaging Atmospheric Cherenkov Technique

Figure 2.11: Schematics of the IACT detection of atmospheric showers. When the primary particle interacts in the top layers of the atmosphere, an air shower of particles is generated, characterized by a head (dark blue) and a tail (light blue). From the shower, the Cherenkov photons (blue lines) propagate to the ground at increasing angle with increasing shower development. The photons are reflected onto the focal plane at a distance from the center of the camera which reflects the shower impact parameter, i.e. the distance from the telescope axis. The camera is pixelized and the image can be reconstructed. Courtesy of M. Lopez [75].

Thus, the image formed on the camera have approximately an elliptical shape whose edges along the major axis represent the head and the tail of the shower, while the inner pixels correspond to its core. The shape, the orientation and light content of the image can be used to infer physical pieces of information about the particle producing the air shower, such as the energy, the incident direction and the particle nature (hadron, muons or $\gamma$-ray). In Fig. 2.13, some examples of different EAS events recorded by the MAGIC-I telescope are shown.

The images recorded by the PMTs of the camera can be parametrized by a set of parameters, the so-called image parameters. The basic ones were originally introduced by A. M. Hillas in 1985 [24] and are related to the distribution of the photons in the different pixels which constitute the image. The most used parameters are derived basically from the moments
Figure 2.12: Focusing of photons in a parabolic mirror. The camera lies in the focal plane \( f \) of the parabolic mirror surface. (a) Photons coming from a perpendicular direction with respect to the telescope axis are focused on the center of the camera. (b) Photons which reach the telescope under an angle \( \beta \) with respect to the telescope axis are focused at a distance \( \rho \approx \beta \cdot f \) from the camera center.

up to the third order of the photons distribution in the pixels of the image\(^4\). A detailed mathematical definition of the image parameters can be found in [76]. In the following the main parameters are briefly described. For a graphical interpretation of some of them see Fig. 2.14.

- **SIZE**: overall number of photoelectrons in the shower image. In first approximation and for fixed values of the impact parameter and Zenith angle, the number of detected photoelectrons is proportional to the primary particle energy. For the MAGIC-I telescope, the conversion factor between photons and photoelectron is about 6 (i.e. 1 photoelectron each \( \sim 6 \) photons).

- **CoG**: the so-called center of gravity of the image. The CoG consists of a pair \((X,Y)\), which is the position in the camera of the weighted mean signal along the X and Y axis respectively. The \((X,Y)\) are the first moments of the charge distribution in the image and are called **MEANX** and **MEANY** respectively.

- **WIDTH**: half width of the minor axis of the shower ellipse. This parameter is correlated to the transversal development of the shower. As the transversal development is larger in hadronic cascades with respect to \( \gamma \)-ray induced one (see section 2.1.2), this parameter is important for the discrimination of the nature of the primary.

- **LENGTH**: half length of the major axis of the shower ellipse. This parameter is correlated with the longitudinal development of the shower. Note that LENGTH is generally larger for hadron induced showers than for \( \gamma \)-ray induced showers.

- **CONC\(n\)**: fraction of photoelectrons contained in the \( n \) brightest pixels. They provide information about the shower core since they describe the compactness of the shower

\(^4\)The zero-order momentum gives the image amount of light, the 1st-order momenta give the image gravity center, the 2nd-order momenta give the shape and direction of the ellipse and the 3rd-order momenta give the image asymmetry.
2.3. The Imaging Atmospheric Cherenkov Technique

Figure 2.13: Air showers as imaged with an IACT (the examples are taken from real MAGIC-I telescope data). (a) \(\gamma\)-ray candidate event. (b) Hadron candidate event. (c) Hadron candidate event characterized by multiple islands. (d) Muon ring caused by an isolated muon.

- **LEAKAGE**: fraction of signal distributed in the outermost camera ring with respect to the total image size. This parameter estimates the fraction of signal loss because of a too large impact parameter, thus it allows to reject images which cannot be reconstructed correctly.

- **M3LONG**: third longitudinal momentum of the image along the major axis of the ellipse. It gives a measure of the asymmetry in the signal distribution along this axis. It is useful to determine the head-tail discrimination of the shower, since it is expected that the head of the image has a higher charge concentration than the tail. This parameter is positive when the shower head is closer to the camera center than the tail, otherwise maximum region. For \(\gamma\)-ray induced air showers this region ought to be very compact, therefore CONC can be helpful for the \(\gamma\)-hadron separation. We will refer to CONC without any index for the case \(n = 2\).
• **NUMBER OF ISLANDS**: number of distinct islands in the shower image. This parameter can be correlated to the fragmentation of the air showers. Hadron induced showers are generally characterized by a larger number of distinct islands as compared to γ-ray induced showers.

Some Hillas parameters are calculated with respect to a defined position in the camera which is normally related to the expected source position. In the ON-OFF data taking mode the source position usually coincides with the camera center, whereas for WOBBLE data taking mode the source position is displaced from the camera center. The main features of these two data taking modes will be described in section 4.2.1. The most common source position dependent parameters (i.e. calculated with respect to a certain reference point in the camera) are:

- **ALPHA**: angle between the major axis of the ellipse and the direction from the image CoG to the reference point. This parameter has the highest γ-hadron separation power, since the images from the γ-rays point toward the position of the source in the camera, thus they are characterized by a small ALPHA angle. On the other hand, hadron showers are distributed isotropically in the sky implying a rather flat ALPHA distribution\(^5\).

- **DIST**: distance of the image CoG to the reference point in the camera. It provides information about the distance of the shower maximum from the telescope axis, correlated to the impact parameter.

Some parameters can be defined also taking into account the arrival time of the Cherenkov signal in the camera. In case of the MAGIC telescopes, the main used time parameters are [77]:

- **TIME GRADIENT**: it measures the magnitude of the time profile of the event. To compute it, first the major axis of the image is found, then the pixels coordinates are projected along the axis so that the problem is reduced to one dimension. Finally, using only the space coordinate along the major axis, a graph of the arrival time is built and fitted with a linear function \(y = mx + q\). The linear coefficient \(m\) represents the TIME GRADIENT image parameter.

- **TIME RMS**: it measures the arrival time spread of the Cherenkov photons in the pixels belonging to the image. It is defined by

\[
\text{TIME RMS} = \sqrt{\sum_{i=1}^{k} (t_i - \bar{t})^2}
\]  \hspace{1cm} (2.28)

where \(k\) is the number of pixels which form the image, \(t_i\) is the arrival time of the \(i\)-th pixel and \(\bar{t}\) is the mean arrival time of the pixels of the image. Due to the time structure of the events, this parameter is also correlated with the TIME GRADIENT. The RMS of the arrival times would be less correlated if computed respect to a fit along the mean axis of the image, but for sake of simplicity the normal RMS is chosen.

\(^5\)Due to the trigger region geometry, the ALPHA distribution for background is not perfectly flat.
2.3. The Imaging Atmospheric Cherenkov Technique

Once the images are parametrized, they can be tagged as hadron-like or \( \gamma \)-like, depending on the values of their image parameters. This procedure, called \( \gamma \)/hadron separation, will be explained in section 4.3.5.

Observations at high Zenith angles imply a larger path from the point of the first impact of the primary particle with the top layers of the atmosphere to the detector. The air shower can be subject to larger fluctuations and light losses compared to observations with the telescope pointing vertically (i.e., with Zenith angle \( = 0^\circ \)). Thus the Zenith angles of the observations have to be taken into account when analyzing the image parameters of a recorded shower. This topic will be further discussed in chapter 5.

2.3.1 Main sources of background for the IACTs

Once sensitive measurements of the Cherenkov light produced in the EAS induced by the VHE \( \gamma \)-rays of a given source are taken, most of the underlying physical pieces of information, such as the origin of the source, its extension and its spectral and temporal properties, can be extracted from the data. However, strong backgrounds reduce the sensitivity of the measurements. The most important sources of background for the IACTs are:

- **LONS**
  The LONS is always contaminating the Cherenkov light of EAS and is one of the limiting factor at trigger level. As mentioned before, the LONS must be rejected by the aid of special trigger system data acquisition. This contamination can be indeed reduced to an acceptable level by requiring a multi-fold next neighbour coincidence of several pixels.
with a very short coincidence time of the order of 5 ns. Another source of noise can be due to bright stars in the field of view of the telescope which can illuminate one or few pixels, thus producing a higher pedestal RMS in the corresponding pixels which affects the pixel signal to noise ratio.

- **Hadrons**
  Hadrons induced showers are the main background of the IACTs’ measurements. Indeed, more than 99% of the recorded images has a hadronic origin, even during the observation of the brightest $\gamma$-ray sources. The energy distribution of the hadron background follows a power law with spectral index $\Gamma = -2.7$. This background can be highly suppressed by an image analysis based on the image parameters, because the hadron induced showers form images quite irregular compared to those induced by $\gamma$-rays. The so-called $\gamma$/hadron separation efficiency (see section 4.3.5) can be larger than 99%. However, the hadronic background cannot be completely removed.

- **Muons**
  Close-by single muons are easily recognized by the ring or arc image they produce in the camera. However, Cherenkov light from single high altitude muons with high impact parameter (>50 m) can mimic $\gamma$-ray images of low energies and are often difficult to reject. Nevertheless, muons contamination can be strongly suppressed with the use of stereoscopic observation and partly by image analysis.

- **Cosmic Electrons and Positrons**
  At energies below about a few 100 GeV cosmic light leptons become a non-negligible background, since they produce air showers with exactly the same features of the $\gamma$-ray induced one. It is an irreducible but isotropic background. The only way to estimate it is via MC simulation.

- **Diffuse $\gamma$-rays**
  Diffuse VHE $\gamma$-rays coming from unresolved galactic and extragalactic sources has been detected by H.E.S.S. [51], EGRET [52] and Fermi–LAT [53] [78]. The diffuse extragalactic $\gamma$-ray background is believed to be smaller by several orders of magnitude than the galactic one. Diffuse $\gamma$-rays represent another source of an irreducible background when searching for $\gamma$-ray emitting point sources. Fortunately, this background is very weak even in the galactic plane where its level seems to be highest.

- **Electronic Noise**
  All the devices and the readout chain have an intrinsic electronic noise that could not be eliminated. The photon sensors in the focal plane instrumentation produce the so-called dark noise which is the output current when the device is not illuminated. Dark noise has continuous and pulsed components which can vary in time. Other sources of noise can be due to the Vertical Cavity Surface Emitting Lasers (VCSELs) (a device used by the MAGIC telescopes) which convert the currents into a light signal, the receiver circuits and the FADCs. The continuous component is eliminated by an AC coupling of the signal but the pulsed one can be confused with the signal. These background are treated with the use of special data runs, taken before each observation (see section 4.2.2).

### 2.3.2 Energy Threshold for the IACTs

The energy threshold for an IACT is strictly connected to the possibility to trigger small signals over the Night Sky Background (NSB). Indeed the Cherenkov light signals are acquired
by the PMTs over the NSB signals. A useful, although simplified, expression for the energy threshold of an IACT can be written as

$$E_{th} \propto \sqrt{\frac{\Phi_B \Omega \tau}{\epsilon A}}$$

(2.29)

where $\Phi_B$ is the night sky background light flux, $\Omega$ is the solid angle seen by a single PMT, $\tau$ is the integration time of the front-end electronics, $\epsilon$ is the QE of the PMT, and $A$ is the reflecting surface area. According to this equation, it is clear that the reflecting surface area, as well as the QE of the PMTs and the fast integration time of the signals, turn out to be key elements for reducing the energy threshold of an IACT.
2. THE IMAGING ATMOSPHERIC CHERENKOV TECHNIQUE
The MAGIC Telescopes

Figure 3.1: Picture of the MAGIC telescopes during data taking at the Astronomic Observatory of Roque de Los Muchachos, located in the Canary Island of La Palma (~2230 m above the sea level). MAGIC-I telescope is on the left, whereas the MAGIC-II telescope is on the right. The smaller dismissed IACT on the right is the CT-3 telescope which was operating during the HEGRA experiment.
The Atmospheric Imaging Cherenkov MAGIC telescopes for ground-based $\gamma$-ray astronomy consist of an array of two telescopes operating in stereoscopic mode since fall 2009 at the Canary Island of La Palma. MAGIC-I telescope has been operating since late 2003, whereas the second telescope, dubbed MAGIC-II, has been successfully completed and commissioned during 2009. Both telescopes have a 17 m diameter tessellated parabolic shaped reflector and are currently the biggest world-wide (see Fig. 3.1). Already for MAGIC-I telescope, the large reflective surface permitted to lower the energy threshold to a value far beyond that of the past generation of telescopes (as Whipple and HEGRA), covering the existing energy gap between 100 GeV and $\sim$300 GeV.

The two telescopes’ system is designed to achieve an improved sensitivity in stereoscopic/coincidence operation mode and simultaneously lower the energy threshold, compared to the single telescope observation mode. The stereoscopic observation mode brings a significant improvement of the shower reconstruction and of the background rejection and consequently better angular and energy resolutions, a lower energy threshold (up to $\sim$30%) and a $\sim$1.5-2 times higher sensitivity.

In the first part of this chapter the main hardware features of the MAGIC-I telescope will be presented. Then, the main new hardware components introduced in the construction of the MAGIC-II telescope will be briefly summarized.

3.1 Introduction

The MAGIC (Major Atmospheric Gamma-ray Imaging Cherenkov) telescopes are an array of two IACTs located in the Astronomic Observatory of Roque de Los Muchachos, at the Canary Island of La Palma (N 28°45'43", W 17°53'24", ~2230 m a.s.l.). The site is well known for its very good atmospheric conditions for most of the year and it is considered to be one of the best observatory sites in the Northern hemisphere: the site benefits from about 1000 hours of excellent conditions for astronomical observations per year. It was previously used for cosmic-ray and $\gamma$-ray experiments, such as the HEGRA CT system (five IACTs, each one with 5 m diameter reflective surface [79]).

The MAGIC telescopes have been designed and are operated thanks to a collaboration of about 150 physicists and roughly 100 technicians from about 20 institutes spread in several European countries. The first telescope has been fully operational since fall 2003 and since then, roughly 40 VHE $\gamma$-ray sources have been observed. The design report of the first telescope was published in 1998 [70] and the financial support was allocated between late 2000 and early 2001. In late 2001 the mechanical structure of the telescope was installed and in 2002-2003 it was equipped and put into operation. From fall 2003 until the early fall 2004 the telescope was in the commissioning phase. After autumn 2004 MAGIC-I started to regularly observe $\gamma$-ray source candidates and to collect data. The construction of the second telescope, 85 m away from the first one, started in summer 2005 and all the technical elements were installed by the end of 2008. Its commissioning phase has been accomplished during 2009.

The HEGRA experiment was the first system of IACTs which successfully used the stereoscopic observation of the air showers [79] (see Fig. 3.2). The principle of a stereoscopic observation is to place several IACTs in the light pool of an air shower and to operate them as a system. That means e.g. that a trigger will only be created when at least two telescopes of the system have a certain signal over a given threshold. In each telescope, inside the light pool, a shower image is recorded. The combination of the recorded images improves the reconstruction of the impact point of the shower and of the height at which the shower reached its maximum and thus the incoming direction of the air shower. These pieces of information significantly improve the energy reconstruction of the shower and the hadron background rejection.

Nowadays, MAGIC-I and MAGIC-II operate in stereoscopic mode, although the single tele-
3.1. Introduction

Figure 3.2: Sketch of the measuring principle of stereoscopic observations of the 5 IACTs HEGRA system. The orientation of the shower axis in the atmosphere can be unambiguously reconstructed through the intersection point of the main image axes of the images recorded by the single telescopes.

scope observation mode is still possible and used for particular purposes. All aspects of the physics program addressed by the MAGIC Collaboration, ranging from fundamental physics to galactic and extragalactic astrophysics, will greatly benefit from the increased sensitivity of the stereoscopic observation of the γ-ray sources. In particular, the expected lower energy threshold of the MAGIC telescopes can have a strong impact on pulsar studies and will extend the accessible redshift range, which is limited by the energy dependent absorption of VHE γ-rays by extragalactic background light (see section 1.4). Moreover, a special synergy is also foreseen with Fermi–LAT, which, being a survey instrument, has a large field of view (FoV) and a very low background counting rate at energies above few GeV: once the position of the sources discovered by the Fermi–LAT detector is available, the MAGIC telescopes (in case those sources are characterized by hard spectra and sufficiently high fluxes) can deliver spectra above ~50 GeV with higher accuracy and better time resolution for variability studies, thanks to their large collection area. The MAGIC telescopes and the Fermi–LAT detector can thus be considered as complementary instruments in the overlap region, from ~30 GeV to 300 GeV, with the possibility to produce highly accurate spectra spanning an energy range of several orders of magnitude (from tens of MeV to several TeV, or wherever a possible source cutoff is present).

The basic aims of the MAGIC-I telescope were the achievement of the lowest energy threshold among all operating IACTs (down to 30 GeV) and the possibility to re-position the telescope 5-10 times faster than other similar experiments. The former goal was pursued following the idea that some important astrophysical processes were happening in the energy gap not accessible to any of the known techniques of the previous generation instruments and thus it was intended to bridge the observational energy gap between the upper limit of satellite
borne detector experiments, which was situated at around 10 GeV before Fermi–LAT started to be operating (2008), and the lower limit of ~300 GeV of the previous generation of ground based detectors. The latter intent was instead required to quickly react on Gamma Ray Burst (GRB) alerts and point the telescope as fast as possible to the sky position of the burst. In order to achieve these characteristics, several novel technologies were used for the design of the MAGIC-I telescope:

- A 17 m tessellated parabolic shaped reflector composed of 964 light weight, square aluminum mirrors with a total area of 234 m². The parabolic shape conserves the time structure of the Cherenkov pulses and the large diameter of the reflective surface permits to lower the energy threshold.

- A telescope frame made of light carbon fiber tubes. The low inertia momentum allows a very fast repositioning of the telescope (on average <40 s), making MAGIC-I the fastest moving Cherenkov telescope in the world and the most suitable ground-based instrument for the observation of the GRBs in their prompt emission.

- An Active Mirror Control system monitored by software which allows the online reorientation of the panels hosting the mirrors. In this way, the focusing properties of the detector can be efficiently monitored and the telescope can always operate in the best optical conditions.

- A 3.5° FoV camera equipped with 577 pixel photomultiplier tubes (PMTs) with an enhanced Quantum Efficiency (QE).

- An analog optical transport of the signal from the camera to the readout system.

- A 300 MHz Flash Analog to Digital Converters (FADCs) data digitization system which has been upgraded in 2007 with ultra-fast FADCs capable to digitize the signals at 2 GSamples/s.

- A fully programmable two-level trigger system which is able to reject the major part of the signals due to the Night Sky Background (NSB).

The basic structure of MAGIC-II telescope is almost identical to the first telescope (for this reason it is often called as “clone”), although many subsystems have been substantially improved thanks to the experience made inside the MAGIC Collaboration during the construction and the operation of MAGIC-I and to novel available technologies. The major hardware modifications introduced for MAGIC-II will be presented after the description of the main hardware components of MAGIC-I. An upgrade of the MAGIC-I telescope, which will basically involve the camera and the readout chain, is foreseen for summer 2011.

### 3.2 The MAGIC-I Telescope

All the subsystems of the MAGIC-I telescope are controlled by dedicated control software programs and coordinated by a Central Control (CC) system, which also collects reports from the subsystems. Besides the regular raw data files which store the pieces of information of the recorded air showers, other files are created and stored in ASCII files. In the following the main features of the MAGIC-I telescope hardware are described.
The MAGIC-I Telescope

3.2. The MAGIC-I Telescope

3.2.1 The telescope frame and the drive system

The reflecting surface is mounted on a robust, light weight structure made by carbon fiber reinforced plastic (CFRP) tubes which weighs only ~5 tons. The tubes are joined with aluminum knots as shown in Fig 3.4(a). The structure allows for fast repositioning: the telescope can perform a complete turn-around in about only a minute. The CFRP tubes construction is about three times stiffer and has less than a third of the weight of an equivalent steel construction.

The camera is at a distance of around 17 m from the reflector and it is placed inside a metallic box. The total weight is around half a ton. The camera is sustained by a single aluminum tubular arc, which is supported by narrow steel cables connected to the main structure designed to minimize shadow effects on the reflecting surface (see Fig. 3.3). The frame of the telescope, including mirrors and the camera support, weighs less than 20 tons, while the whole structure plus the undercarriage amounts to about 65 tons.

MAGIC is an Alt-Azimuth mount telescope: the whole structure can move in the Azimuth and in the altitude direction and it is mounted on a circular rail of 19 m . This means that in order to track a source in the sky, the telescope has to be moved around two axes (Azimuth and elevation) with variable speed along each of them. One of the design goals of the drive system was to allow the continuous observation of a given source without reaching one of the end positions in Azimuth. Therefore, the allowed range of movements in Azimuth spans from to (where denotes Geographic North), whereas the range of movements in elevation is between and . The accessible range in both
directions is limited by software and by two mechanical end-switches. While tracking a source, tracking loop algorithms transform celestial coordinates of the target (RA, DEC) into telescope coordinates.

Two servo-motors control the motion in Azimuth and one servo-motor in the Zenith motion with a peak power consumption of 11 kW each (see Fig. 3.4(b) and 3.4(c)). The angular position of the telescope is controlled by three absolute 14-bit shaft encoders. One of them is in Azimuth axis and two in the declination to control twists of the mirror dish. With this configuration it is possible to measure the telescope position with an accuracy of about 0.02°.

In order to improve the pointing accuracy, the optical axis of the telescope is calibrated by taking pictures of stars at different Zenith and Azimuth pointing angles using a CCD camera (the so-called T-Point camera) placed in the middle of the reflective surface. Thanks to these pictures, accurate bending models of the telescope can be produced, allowing for a tracking accuracy of the order of 0.01°. The position of the telescope is further constantly monitored using a sensitive 4.6°×4.6° FoV CCD camera mounted on the center of the mirror dish and aligned with the pointing axis of the telescope (starguider camera). Six markers, created by six LEDs, indicate the reference system of the PMT camera while the stars are recognized thanks to a dedicated software which compares the images of the starguider camera with star catalogs. The information recorded by the starguider camera are used to correct offline any misalignment of the telescope. Moreover, the comparison of the number of identified stars in the FoV with the expected one from the star catalog can give information about the weather conditions (in particular about the atmospheric transmission) during the observation.

More details about the drive system, the pointing models and the starguider system of the telescope are given in [71,80,81].

3.2.2 The reflector dish and the mirrors

The reflector follows a parabolic profile with \( f \) (focal length) \( \simeq D \) (diameter) = 17 m, which was chosen to minimize the time spread of the Cherenkov light flashes on the camera plane, thanks to the isochronous properties of parabolic dishes. The preservation of the time structure of the Cherenkov flashes is indeed important to minimize the trigger gate, reducing the number of the triggered random coincidence due to the NSB and thus to increase the signal to noise ratio of Cherenkov flashes with respect to the NSB. Another advantage of preserving the intrinsic time structure of the Cherenkov pulse is to get a possible improvement in the separation power between hadronic and electromagnetic showers as well as isolated muons with large impact parameters which can look like \( \gamma \)-ray showers. A drawback of a parabolic dish is mainly due to the coma aberration, which makes the images to look extended (blurred) if looking off-axis. The effect has a radial symmetry to the camera center resulting in an image stretching, which is in first order directly proportional to the angle under which the shower is seen by the telescope. In case of the MAGIC reflector the coma aberration effect amounts to \( \sim 7\% \), i.e an image point which should have a distance \( d \) from the camera center has an effective distance of \( 1.07 \cdot d \).

The reflector of MAGIC-I telescope is tessellated and comprises 956 mirrors with a total area of 234 m². Each mirror is a square of 0.495 m side length and has a spherical profile whose radius of curvature (ranging from 36.6 m to 34.1 m) is optimized for each position in the reflector to best approximate the overall parabolic shape. MAGIC I mirrors are generally grouped onto panels of 4 elements each. At the borders of the reflecting surface the panels are composed by 3 mirrors.

The mirrors are made of 5 mm thick AlMgSi alloy plates glued on aluminum honeycomb inside a thin aluminum-box [82] (see Fig. 3.5). The reflectivity of the mirrors ranges between
3.2. The MAGIC-I Telescope

Figure 3.4: Figure (a) shows the joints between the CFRP tubes and the aluminum knots. In figure (b) one can see one Azimuth servo-motor, mounted on the circular rail. In figure (c), the Zenith servo-motor drives the elevation and is mounted on a aluminum arc. Figure (d) shows the open box of the SBIG camera, which is used for the active mirror control (AMC) repositioning system.

~80% and ~90% over a broad range of wavelengths (~250 nm to ~750 nm), whereas the average reflectivity is about 85%. The surface of the mirrors was coated with a thin 100 nm layer of quartz (with some admixture of carbon) for protection against corrosion and acid rain. The mirrors are equipped with internal heating wires for de-icing and the removal of dew. The total power consumption for heating the entire reflector is 40 KW.

During the observations the mirror elements are focused to a distance of 10 km, the typical height of the shower maximum of a 100 GeV γ-ray induced shower at vertical incidence. In this configuration 80% of the light of a point like source is focused into a circle with a radius of ~17 mm (roughly the size of a camera pixel). More details about the properties of the reflecting surface of the MAGIC-I telescope can be found in [83].

3.2.3 The Active Mirror Control

Degradations of the optical performance due to deformations of the telescope frame and of the mirror panels, caused by the effect of the gravity when tracking a source, are compensated by an Active Mirror Control (AMC) [84]. The AMC is continuously monitoring the declination angle of the telescope and accordingly adjusting the individual mirror elements. The system consists of two actuators per mirror panel and a red laser pointer installed in the center of each
3. THE MAGIC TELESCOPES

Figure 3.5: Sketch of the structure of the mirrors of the MAGIC-I telescope: reflecting surface, internal heating wires, aluminum honeycomb and aluminum-box.

panel. A control software can switch on the lasers which project the red spots on the camera cover which are viewed by a CCD camera (see Fig. 3.4(d)). This CCD camera determines the laser spots with respect to some reference markers whose positions are defined in respect to several LEDs mounted on the PMT camera cover. According to the deviation to the nominal positions of the red spots, the software activates the actuators in order to properly realign the mirror panels. This operation is done every night before starting observations and take around 5-10 minutes. The alignment can be also done with stars: the CCD camera can identify reflected star light of each panel whose image is used to adjust the panel such that the spot is focused in the camera center. The corresponding actuator positions are stored in lookup-tables (LUTs) as a function of the Zenith angle pointing position, so that absolute actuator positions are stored. The LUTs are very useful since allow for a fast readjusting of the panels in the correct (Zenith dependent) positions whenever it is necessary, for instance during the repositioning or even tracking of the telescope. Typically, it is automatically requested by a control software to adjust the mirror segments using LUTs when the telescope orientation changes by more than 5° in Zenith.

More details about the AMC system are given in [85].

3.2.4 The PMT Camera

The camera represents a key element to improve the γ-ray sensitivity and the γ-hadron separation of any IACT. The MAGIC-I camera is designed to record the Cherenkov light from the atmospheric showers with the help of PMTs, amplify the signal, re-convert it again into an analog light signal and transmit it to the counting house (where the electronic and control rooms of the telescopes are placed) via optical fibers.

The MAGIC-I camera has a diameter of \(\sim1.2\) m, a FoV of 3.5° and it is equipped with 577\(^1\) hemispherical PMTs that convert the Cherenkov photons into a measurable electrical

\(^1\)This number includes the Central Pixel that is not used to record shower images but is reserved to perform optical measurements, particularly important for pulsar studies.
3.2. The MAGIC-I Telescope

The camera is divided in an inner (where the trigger system is operating) and in an outer part. The inner part is equipped with 397 1" PMTs from Electron Tubes Ltd of type 9116A each with a pixel diameter of 0.1° (inner pixels). The PMTs are combined into hexagonal arrangement for a total inner FoV of roughly 2° (see Fig. 3.6(a)). The inner pixels are surrounded by 180 1.5" outer PMTs of type ET9117A each with a pixel diameter of 0.2° (outer pixels). While a fine pixelization (0.1°) in the camera center is mandatory for observing low energy small showers, the outer part of the camera generally records parts of large energy showers or shower tails in which larger statistical fluctuations in the shower development are expected. Moreover, the outer part of the camera is generally affected by the coma aberrations of the reflector. These conditions justified therefore the use of larger diameter (0.2°) pixels in the outer part of the camera. In addition, this solution permitted to reduce the costs and the overall camera weight.

The PMTs are equipped with dedicated hexagonal light collectors (glued in front of the PMTs) which guides the light inside the PMT, maximizing the double-crossing probability of photon trajectory at large incidence angles. Thanks to their hexagonal arrangement, the PMTs can uniformly cover the camera plane with minimum blind areas. For protection purposes pixels and light guides are protected by a 2 mm thick UV transmitting plexiglass window.

The hemispherical bialkali photocathode PMTs were enhanced in QE with a diffuse lacquer doped with a wavelength shifter (P-Terphenyl) to enhance the sensitivity for the UV photon component of the Cherenkov light [86]. The obtained final QE of the PMTs has a typical peak of 30% at ~400 nm and it is of the order of 20% for a wide useful wave length range (~300 nm to ~550 nm).

Each PMT has 6 dynodes: three of the 6 dynodes are stabilized to a fixed voltage, by Zener diodes (dynode number 1) or by active loads (dynodes number 5 and 6), while the remaining three can be regulated online. The High Voltages (HV) regulation for each PMT fully covers the 0-2000 V range. The PMT current and HV are read out for each pixel, multiplexed in groups of 96 and digitized by a 12-bit ADC. In order to compensate with the aging and the possible degradation of PMTs, a readjustment of HV is done a few times per year (flat-fielding procedure). The dynodes operate at a gain of roughly 3·10^4 (inner PMTs) and 2·10^4 (outer PMTs), instead of the standard gains of 10^6-10^7, to prevent fast aging and damage from high currents by light during moonshine observations [87]. The PMTs typically have a signal risetime of ~800 ps, a full width at half maximum (FWHM) of 1.5 ns and a tail extending to ~4 ns. A sketch of the camera layout and the two PMT types are shown in Fig. 3.6.

The camera has two metallic lids which are open only during data taking and protect the camera from daylight. In addition, there are water and air cooling and heating systems, which prevent the camera from getting too hot or reaching the dew point. The whole camera housing can be moved forward and backward with respect to the mirror dish in order to obtain different focuses, ranging from ~1 km to infinity.

A detailed description of the MAGIC-I camera and its performance can be found in [88,89].

3.2.5 The Readout chain

The readout chain has the task to record the events that were triggered by air showers. In Fig. 3.7 a sketch of the readout chain is shown.

Once the Cherenkov photons are reflected into the camera, they produce photoelectrons in the PMTs. At the base of the PMTs the signal is AC coupled to a fast transimpedance preamplifier with a gain of about 6. The fast preamplified PMT signals are then converted back into light by means of fast current driver amplifiers coupled to vertical cavity surface emitting laser diodes (VCSELs). The signal is then transported via 160 m long optical fibers
into the counting house. The conversion of the electrical signal into an optical one has several advantages. First of all, the dispersion and the attenuation of the signal during the transmission over 160 m are minimized. Then, additional benefits are the reduced weight and diameter of the signal carrying cables as well as high immunity against electromagnetic interference. Details of the characteristics of the optical transmission system can be found in [88].

In the counting house the optical signals are received and converted back into electronic ones by fast GaAs PiN photodiodes in the receiver boards. After a further amplification the signal is split in two. One part of the signal is routed to the trigger branch and the other part is routed to the FADC system, which digitizes the signal and writes it into a ring-buffers. When the trigger sends a signal, the data stream is copied into a first-in-first-out (FiFo) buffer. From there it is read out and stored to disk. Data are then sent over day by Gbit connections.
3.2. The MAGIC-I Telescope

Figure 3.7: Sketch of the MAGIC-I readout chain. PMT signals are reconverted into optical pulses, sent to the counting house via optical fibers, and fed into the trigger and DAQ system. Figure taken from [71].

to some nodes in Europe.

The Trigger System

The trigger of the MAGIC-I telescope is a two-level system following a pipeline philosophy, similar to those adopted in high energy physics experiments, and makes the decision if an event takes place so that the camera should be read out. The trigger system consists in a programmable logic multi level decisional system which has the task to discriminate the signals arriving from Cherenkov flashes from those related to the NSB, moonlight and bright stars in the FoV of the camera, which increase the probability of accidental coincidences. For this purpose, an efficient trigger strategy has been organized according to timing and topology considerations which allows to get an energy threshold of the order of 50 GeV at the trigger level. A detail description of the MAGIC-I Trigger system can be found in [90,91].

The are three trigger decisional levels, the first one being just a discriminator threshold system for the individual channels:

- **Level 0 Trigger (L0T).** Its task is to check if the signal coming from a camera pixel is greater than a fixed discriminator threshold (DT). The DTs are adjustable for each pixel via software in order to take into account different individual pixel FoV light content and different levels of the NSB. An almost flat trigger rate over the camera is therefore achieved by adjusting the discriminator level of the individual pixels. Thanks to dedicated channel rate counters the Individual Pixel Rate (IPR) information is available and can thus be monitored online preventing high trigger rates e.g. caused by bright stars in the individual pixels. In this way, accidental trigger due to stars in the FoV of the camera can be evaded. The IPR control is always active during the observation and thus the DT values for any given pixel vary with time. If the signal of the pixel exceeds the corresponding threshold, the pixel is tagged as “fired” and a very fast digital signal
is generated with a standard duration of 6 ns (the typical duration of a Cherenkov flash) which is then processed by the Level 1 Trigger. The L0T is restricted to only 325 inner pixels (see Level 1 Trigger section), which correspond approximately to 1° radius from the camera center. Most of the γ-ray induced shower images have in fact its center of gravity within 1° radius from the nominal source position in the camera, whereas most of the images contained in the outermost part of the camera are not likely to be γ-ray induced and thus they will be not triggered. However this configuration has some disadvantages for the extended sources and for sources observed off-axis. As it will be shown, for the MAGIC-II telescope the trigger region has been increased. In Fig 3.8 the sketch of the L0T is shown.

![Schematic view of the Level 0 Trigger system of MAGIC-I.](image)

**Figure 3.8:** Schematic view of the Level 0 Trigger system of MAGIC-I.

- **Level 1 Trigger (L1T).** This level requires fast temporal and spatial coincidences of the digital signals of the “fired” pixels. All signals from the L0T are grouped into 19 hexagonal overlapping cells (macrocells) of 37 pixels (36 active and one blind) each, which cover 325 inner pixels of the camera (see Fig. 3.9(a)). The L1T gives a trigger signal if a certain amount of next neighbour (NN) pixels fulfill a temporal and spatial compact configuration which can be set by software. When the selected multiplicity condition is satisfied by any of the trigger macrocells, the signals in a 20 ns window around the original trigger time are stretched to 50 ns width and transmitted to the Level 2 Trigger for further selection. The decision of the L1T is taken in less than 80 ns with a total time jitter of the order of 1 ns. Several option are available: 2NN, 3NN, 4NN and 5NN. During normal operation, the 4NN condition is used leading to a typical trigger rate of about 200-250 Hz (in moon-less nights), depending on the Zenith angle. In Fig. 3.9(b) a schematic view of the L1T system is shown.

- **Level 2 Trigger (L2T).** The L2T looks for more sophisticated topologies in the overall trigger region. The information of each trigger macrocell is distributed over three different 12-bit lookup-tables (LUTs). Inside each LUT any possible pixel combinations can be checked and then merged into a fourth LUT for a global analysis of the image over the macrocell. Up to now the L2T has been used just in flag-mode and not implemented in the usual data taking.
3.2. The MAGIC-I Telescope

(a) Macrocells. (b) L1T system.

Figure 3.9: (a) Sketch of the trigger topology in the MAGIC-I camera. (b) Schematic view of the Level 1 Trigger system of MAGIC-I.

The FADC system

After passing the trigger conditions, the events are digitized using FADCs, and stored to disk. The time and the trigger information for each event are also recorded by dedicated digital modules which are read out together with the FADC modules.

- The old 300 MSamples/s FADCs
  The original DAQ system of the MAGIC-I telescope was based on custom-made dual gain 300 MSamples/s 8-bit FADC system [92]. This rather low sampling rate required additional pulses stretching to ensure proper sampling of the signals.
  In order to increase the limited dynamic range, the part of the signal going from the receiver to the FADCs was split into two (see Fig. 3.7). One half of the signal was amplified by another factor of ten, called high-gain, the other half of the signal, called low-gain, was delayed by ~55 ns. If the high-gain signal exceeded a certain amplitude (i.e. it was likely saturating the dynamical range), a GaAs switch was activated and also the delayed low-gain signal was digitized. In this way the usable dynamic range of the FADCs was extended to a factor about 1000.
  Just before the digitalization, the short electronic pulses (~2.5 ns FWHM) needed to be stretched to a FWHM larger than 6 ns before they were passed to the FADCs, just to make possible the sampling at 3-4 different positions (which is the minimum required sampling to get all the information contained in a pulse).
  The FADCs continuously wrote the digitized amplitude information into a ring buffer. In case of a trigger the digitization stopped and the corresponding part of the ring buffer was written onto a disk. The dead time introduced by the readout was 25 μs, corresponding to a dead time of 1.6%, thus negligible compared to an average trigger rate of 200-250 Hz.

- The current operating 2 GSamples/s MUX FADCs
  In order to enhance the suppression of the NSB contribution and to increase the detector sensitivity thanks to a better determination of the timing structure of the signals (which
helps to increase the discrimination power between hadronic events and γ-ray ones) an upgrade of the digitizing system by a 2 GSamples/s FADCs was performed in 2007. To reduce the number of these new FADCs, for reasons of costs and power consumption, a new fiber-optic multiplexing (MUX) technique was developed and tested. The MUX feeds 16 signals into one FADC consecutively (see Fig. 3.10), which is possible because of the typical relatively low trigger frequency (below 1 kHz). These signals are delayed using optical fibers of different length. Thanks to good time resolution of about 200 ps and the higher dynamical range, the stretching of the pulses and the splitting into high-gain and low-gain are no longer needed. A sustainable data acquisition rate up to 100 MBytes/s (corresponding to a trigger rate of 1 kHz) has been achieved. A complete description of the MUX FADC system can be found in [94].

The performance of the MUX FADC system showed an increase of the best sensitivity of the MAGIC-I telescope by up to ~40% [95], which also includes a new analysis chain using image parameters related to the timing information of the events. In this thesis, all the considered data were taken after the installation of the MUX FADCs and thus benefit of the enhanced performance due to the better arrival time determination of the recorded air showers.

**Figure 3.10:** Concept scheme of the Multiplexed readout system.

3.2.6 The Calibration system

Due to gain fluctuations, changing behavior of several components in the readout chain with changing extrinsic influences (e.g. the temperature) the PMT signals have to be calibrated. In order to obtain calibration constants for converting FADC charge into physical quantity of photoelectrons and FADC timing into an absolute signal timing, an optical calibration system providing independent calibration methods was installed [96] (see Fig. 3.11). The calibration box is located in the center of the reflecting surface dish and contains ultra-fast (~3 ns FWHM) emitting LEDs at different wavelengths (370 nm, 460 nm and 520 nm) which allow for an uniform illumination of the PMT camera with light pulses of different intensities (which can be varied from 4 to 700 phe per inner pixel) and different frequencies. Different

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2The MUX technique allowed a cost reduction of about 85%.
distinct methods can be used to calibrate the whole readout chain: the F-factor method [97],
the positive intrinsic negative (PIN) diode method and the so-called blinded pixels method.
Special calibration runs are taken before each observation. Thanks to a dedicated trigger

![Figure 3.11: Schematic view of the MAGIC-I calibration system setup.](image)

signal, the calibration system is working also during normal data acquisition: interleaved
calibration events are recorded as to trace gain variations of the PMTs and the VCSELs. A
precise description of the calibration system and the different methods which can be used are
given in [89,98,99].

### 3.3 The MAGIC-II Telescope

Following the results of a dedicated Monte Carlo study, showing moderate dependence of
the sensitivity on the distance of the two telescopes [100,101], the second telescope has been
installed at a distance of 85 m from the first one. In order to minimize the time, the costs
and the resources required for design and production, the MAGIC-II telescope is structurally
a clone of the first telescope: the light weight CFRP tubes frame, the drive system and
the AMC have been developed with marginal improvements following the design used for
MAGIC-I. Nevertheless, many hardware improvements have been implemented in the second
telescope whenever they allowed cost reduction, improved reliability or increased physics
potential of the new telescope with reasonable efforts. In particular, the main innovations
deal with the reflecting surface, the PMT camera and the electronics readout. Moreover, a new
Central Control (CC) program has been developed to communicate with all the subsystems
of both telescopes and to coordinate their functionality, thus easing the stereo data taking
procedure [102].

In the following the basic hardware innovations of MAGIC-II are briefly described.

### 3.3.1 The reflector dish and the mirrors

The MAGIC-II reflector follows the same parabolic profile of the MAGIC-I dish \( f \) (focal
length) \( \simeq D \) (diameter) = 17 m) and is composed of two types of spherical mirrors with
area of 1 m\(^2\): 143 full-aluminum mirrors and 104 glass-aluminum mirrors. The full-aluminum
The MAGIC-II telescope is a clone of the MAGIC-I telescope. The main innovations introduced with respect to the first telescope concern the reflector surface, the PMT camera, and the readout electronic chain.

The mirrors consist of a sandwich of two 3 mm thick aluminum plates and a 65 mm thick aluminum honeycomb layer in the center, protected with a quartz coating, resulting in a ~85% reflectivity. The remaining 104 mirrors are produced as a 26 mm thick sandwich of 2 mm glass plates inter-spaced by a honeycomb layer using a cold slumping technique [103]. The frontal glass surface is coated with a reflecting aluminum layer and a protecting quartz coating. These mirrors are placed in the outermost part of the reflector.

The use of large mirrors and the removal of the intermediate panel as used for small mirrors of MAGIC-I carried a general improvement in terms of a decrease of the number of items to be produced and installed, and in general a decrease in costs and manpower.

As in MAGIC-I, the 247 1 m^2 mirror panels can be adjusted by the AMC system, according to the pointing position of the telescope. The full-aluminum mirrors have a hole drilled in the middle to support the AMC laser needed to track the panel position. In the case of glass-aluminum mirrors, the lasers are mounted at the corner of the mirror.

The overall optical performance of the reflector surface of MAGIC-II gives a total Point Spread Function (PSF) of the same order of the PSF of MAGIC-I reflector (~10 to ~13 mm).

Details about the mirrors of the MAGIC-II telescope, the performed quality tests, the performance, and the mounting on the reflector structure are reported in [83].

### 3.3.2 The PMT Camera

A major effort has been made to improve the MAGIC-II camera photon detection system. The design of the camera is improved with respect to that of MAGIC-I, while maintaining a FoV of 3.5° and a total weight of about 0.6 tons. The camera has a round shape and has been uniformly covered with 1039 identical 0.1° FoV pixels equipped with hemispherical superbialkali Hamamatsu R1408 6 stage PMTs, of a 25.4 mm diameter (the same size as the small ones in the MAGIC-I camera), with a QE of 32% at the peak wavelength (~330 nm) and...
3.3. The MAGIC-II Telescope

(a) Single pixel configuration.

(b) Cluster configuration.

Figure 3.13: (a) Single pixel configuration of the MAGIC-II camera. (b) Cluster configuration of the MAGIC-II camera.

without lacquer coating. The smaller pixelization leads to an enhancement of the sensitivity compared to MAGIC-I (see Fig 3.13(a)). The 6 dynodes of the PMTs are operated at a rather low gain of typically $3 \times 10^4$ in order to allow observations under moderate moonlight. The camera is sealed by a 3 mm thick, UV transparent plexiglas window.

Each pixel of the camera consists of a hexagonal Winston cone type light guide glued on top of the PMT (in order to minimize the dead areas between adjacent pixels) followed by a Cockcroft-Walton type HV generator, an amplifier and a VCSEL. In addition, a monitoring circuitry is included which provides readout of relevant operating parameters such as average PMT current, PMT operating voltage, and pixel temperature. The PMT signals are amplified by 700 MHz band-width AC coupled preamplifiers placed in one of the circuit boards. The amplified PMT signals are converted back to optical signals by VCSELs, which transmit the analog signals by multi-mode optical fibers of 160 m length to the readout system in the counting house.

Pixels in the camera are arranged in cluster units. In total 169 clusters are installed: 127 of them are fully assembled with PMTs, while 42 clusters in the circumference of the camera are only partially equipped with PMTs in order to approximate the overall round configuration. Each cluster comprises 7 pixels grouped in a hexagonal configuration (see Fig. 3.13(b)). This modular design allows easier control and maintenance of the camera and easier upgrade with improved photo detectors which is foreseen for the near future. A new photodetector type, the Hybrid Photo Detector (HPD), will substitute part of the PMTs in the camera [104]. HPDs consist of a vacuum tube with a GaAsP photocathode set at 6-8 kV and an avalanche diode acting as an electron bombarded anode with internal gain. The HPDs have a much higher QE than PMTs (~50%), better single photon resolution and are almost free from after-pulses. This will significantly increase the sensitivity for low energy showers.

The steering and monitoring of the parameters of each of the 7 pixels in each cluster is done by a Slow Control Cluster Processor (SCCP). The SCCP controls the readout and setting of all parameters of the pixels of the cluster: the HV setting, the HV reading, the anode current reading, the VCSEL bias current setting and VCSEL temperature reading. It is in turn controlled by a PC in the counting house over a custom made RS485 and VME optical link. In addition, there are three SCCPs installed in the camera to control the power of the low voltage (LV) power supplies, for steering the lids and the starguider target of the camera.
and to read the humidity and the temperature in different points of the camera. They also control the intensity of the AMC and starguider LEDs. The camera control software [105] is programmed in Labview and can be remotely steered by the central control system.

A new feature of the MAGIC-II camera is a pulse injector which can be used to inject charge pulses (adjustable between 0 and 1.6 V amplitude) into the input of the amplifier, without needing to have HV present. This device has been particularly useful during the commissioning phase in order to perform linearity tests of the readout chain during daylight, without wasting precious observation night time and reducing the systematics related to the fact that no signal passing the PMTs is needed.

Since the MAGIC-II camera have more pixels than the MAGIC-I one, also the cooling system has been much improved. Two cooling plates have been installed which serve also as a support structure for the clusters (clusters are inserted into holes between the two cooling plates) in order to avoid space problems in the camera housing. The cooling system keeps the camera at a stable temperature of 27°C during operation. The total power consumption of the camera is \( \sim 1 \text{ kW} \).

The trigger area of the camera has been increased: the trigger region includes 91 cluster modules covering a FoV of 2.5°. This is an advantage for the WOBBLE mode observations (see section 4.2.1), for a scan of a sky region with unknown source position as well as for extended sources.

More details about the hardware components, the structure and the performance of the MAGIC-II camera can be found in [106] (and reference therein).

### 3.3.3 The Readout chain

The very fast analog signals (\( \sim 2.5 \text{ ns FWHM} \)) from the PMT MAGIC-II camera are transmitted via optical cables to the counting house, where they are routed to new high bandwidth and fully programmable receiver boards which convert back the signals from optical to electrical ones. The signal is then split in two branches. One branch is further amplified and transmitted to the digitizers while the other branch goes to a discriminator with a software adjustable threshold. As in MAGIC-I, if an analog pulse is above a certain threshold, a digital signal is generated and shaped with a software controllable width. The signal is then sent to the trigger system with a software adjustable time delay. Scalers measure the trigger rates of the individual pixels: this allows to use the online IPR control, as in case of MAGIC-I.

The new digitization and acquisition system is based upon a low-power consumption analog sampler called Domino Ring Sampler (DRS) [107, 108]. The fast Cherenkov pulses are sampled and temporarily stored in a multi capacitor bank (1024 cell in DRS version 2, currently used) that is organized as a ring buffer, in which the single capacitors are sequentially enabled by a shift register driven by an internally generated 2 GHz clock locked to a common synchronization signal. When a trigger signal arrives, the sampling is stopped and the signals currently stored in the capacitors are frozen. Those signals are then read out and digitized at 40 MHz rate by a 12-bit resolution external ADC. From there, the digital information is transmitted via optical links to a single computer which is the final point of the data taking chain.

Signals are sampled at 2 GSamples/s, which allows a precise measurement of the signal arrival times in all pixels. As mentioned in section 3.2.5, a fast sampling of the pulses allows one to minimize the integration time and thus to reduce the influence of the background from the Light of the Sky (LONS) and, in addition, to exploit the arrival time information in order to improve the hadron background suppression. The Domino samplers are integrated in newly designed mezzanines which equip a set of 14 multi-purpose PULSer And Recorder
3.3. The MAGIC-II Telescope

(PULSAR) boards. Finally, the data are sent through a S-LINK optical interface to a single computer. The entire DAQ hardware is controlled through a VME interface and steered by a slow control software program called MIR (Magic Integrated Readout). The data acquisition software program proceeds finally to the event building and data storage.

Among the characteristics of the DRS, the most important ones are the ability to host up to 10 channels in one chip, its high analog bandwidth (150 MHz), excellent time resolution and very low power consumption. Nevertheless, the current DRS-2 chips have some significant disadvantages. Namely, they are sensitive to temperature changes and highly non-linear devices. Each capacitor of each channel has a different response to a given input signal. Because of this, a calibration that corrects these non-linearities of the DRS-2 has to be done regularly during data taking (at least 2 times per night). An upgrade to the newest version of the DRS chip (DRS version 4), which have several advantages with respect to DRS-2, are foreseen in the very near future.

For a complete description of the new readout chain of the MAGIC-II telescope see [109–111].

3.3.4 The Trigger system

The trigger system of MAGIC-II is a slightly modified version of the trigger system of MAGIC-I. It is based on the same compact next neighbour logic as the MAGIC-I trigger system. One significant change is the distribution of the macrocells which has been necessary to increase the trigger region of the MAGIC-II camera up to 2.5° diameter FoV (see Fig.3.13), covered by 599 pixels. The enhanced effective trigger area permits a higher sensitivity over MAGIC-I, and thus increases the potential to study extended sources and to perform sky scans.

When the two telescopes are operated in stereo mode a coincidence trigger (so-called Level 3 Trigger) between the two telescopes rejects events which only triggered one telescope. In order to minimize the coincidence gate in the Level 3 Trigger, the triggers produced by the individual telescopes will be delayed in a time which depends on the geometry of the telescopes. This will reduce the overall trigger rate to a rate which is manageable by the data acquisition system.

3.3.5 The Calibration system

In order to calibrate the PMT camera of the MAGIC-II telescope, instead of LEDs as light emitters, a new system consisting of a subnanosecond laser has been developed. The advantages of the subnanosecond laser system is that the width of the light pulses is below than the typical width of the PMT pulses. The stability of the laser system is also higher than the one of the LED system.

The calibration box of MAGIC-II has been installed in the center of the reflecting surface and is based on a frequency tripled passively Q-Switched Nd-YAG laser, operating at the third harmonic at 355 nm and producing pulses with duration of ~700 ps FWHM. For providing a large dynamic range, the output beam of the laser goes through two filter wheels with different attenuation filters. After the laser beam is attenuated it goes through an integrating sphere (Ulbricht sphere) to diffuse the light exiting the box, thus providing a homogeneous illumination of the camera. Due to 16 possible combinations of the attenuation filters, it is possible to adjust the pulse energy in a wide range of intensities within 100 steps from single to 1000 photoelectrons (per camera pixel). The wide range of possible intensities of the light coming from the calibration box enables the precise monitoring of the linearity of the complete signal chain. It is foreseen to equip MAGIC-I with a similar calibration system in the next future.

More details about the calibration system of MAGIC-II can be found in [112].
3.3.6 Upgrade of MAGIC-I

The smaller trigger area and somewhat lower light conversion efficiency of MAGIC-I telescope can limit the overall performance of the MAGIC stereo system. Therefore it has been recently decided to upgrade the MAGIC-I PMT camera as to be a clone of the MAGIC-II PMT camera. However its inner section (about 400 pixels) may be readily equipped with HPDs. The camera frame and electronics are already under construction. The readout will also be upgraded to a digitizing system similar to that of MAGIC-II. The finalization of the MAGIC-I upgrade is foreseen for summer 2011.
The Analysis Chain of MAGIC data

In this chapter, the main features and performance of the analysis chain of the MAGIC telescopes data (referring to spring 2010) are described. Since the start of the new operations and data taking of the stereoscopic system, an extension of the official software package for the analysis of MAGIC data has been developed and it is now in the final stage of implementation and testing. The description of the analysis chain of MAGIC-I mono data will be taken into account first since it can be used as a valid guideline for the introduction of the new tools required for the analysis of the stereoscopic data. Then, the main new features of the stereoscopic analysis will be introduced.

After a short general introduction to the software framework and to the MAGIC-I analysis chain, in section 4.2 the observation modes currently used for the data taking and the different types of data will be presented. In section 4.3 the main analysis steps needed to cover all MAGIC-I data processing, starting from the uncompressed raw data and extending to the calculation of energy spectra and light curves for the detected $\gamma$-ray sources (or to the calculation of the flux upper limits) will be illustrated. Finally in sections 4.4, the current analysis chain of the MAGIC stereoscopic data will be illustrated. In this section the current performance of the stereoscopic system will be presented.

4.1 The MARS software and the MAGIC-I analysis chain overview

The official software package for the reduction of MAGIC data is called MARS (MAGIC Analysis and Reconstruction Software). It is a collection of ROOT-based [113] programs (classes and executables) written in C++ which includes all the necessary algorithms to transform the raw data recorded by the telescopes into information about the physics parameters of the observed targets.

MARS has been developed during the last decade within the MAGIC Collaboration [114], and is currently the official analysis package of MAGIC data. The data analysis chain implemented in MARS is divided into several steps, each of which is performed by an independent program which takes as input the output of one or more of the previous stages. The initial input to MARS are the raw data recorded by the telescopes, consisting of binary files containing the full information available per pixel (digitized signal amplitude vs. time) for every triggered event, plus ASCII files containing regular reports from the different telescope subsystems.
(like the telescope drive system, the trigger system or the weather station). Throughout the analysis chain, the data are organized in ROOT trees containing a set of parameter containers for every entry. Typically, the core of a MARS program is an event loop which executes an ordered list of tasks (the task list) for every event in the input file. Besides the task list, every loop is associated to a parameter list. It contains pointers to all the parameter containers holding the input data needed by the tasks, and to the containers where the tasks store the results of their calculations.

Since no natural or artificial calibrators of VHE $\gamma$-rays are available for ground based $\gamma$-ray detectors, the analysis chain banks on Monte Carlo simulated $\gamma$-ray events (MC-$\gamma$) to resemble $\gamma$-rays excesses in MAGIC data. MC-$\gamma$ events are indeed essential for the $\gamma$/hadron separation (that is for the suppression of the hadronic background in the data) and for the energy estimation of the events and the calculation of the Effective Collection Area. Several studies have been performed in order to check the correspondence between $\gamma$-ray simulations and real $\gamma$-like events [42, 115, 116], proving the good behavior of the simulations. In order to estimate the background, MAGIC data with no $\gamma$-ray source in the analyzed region are used. In case the given source is observed in the so-called WOBBLE mode (see section 4.2.1), background data are extracted from the same field of view (from a region safely far away from the expected signal region). Whereas, in case the source is observed in ON mode (see section 4.2.1), extra OFF data are necessary to estimate the background.

The main basic steps of the MAGIC-I analysis can be summarized as follows:

- **Data preparation**: the binary raw data, coming out directly from the DAQ system, are properly organized into the custom MARS file format. The MARS program in charge of this task is called MERPP (MERging and Pre-Processing program).
- **Calibration**: the information coming from the FADC are properly calibrated for each pixel of the camera and converted into number of photoelectrons, including timing information of the signal. The MARS program used for accomplishing such a task is called CALLISTO (CALibrate LIght Signals and Time Offsets).
- **Image cleaning and image parametrization**: the aim of this analysis step is to remove, for each triggered event, those pixels containing light not likely related to a Cherenkov event and subsequently to perform the calculation of the parameters associated to the image of each event (see section 2.3). The program which takes care of this task is called STAR.
- **Data quality check**: the quality of the data can be checked by the aid of the pieces of information provided by the telescope subsystems and from the image parameter distributions. During this step, standard quality cuts related to some image parameters (and their combinations) are generally applied. In addition a good indicator for the quality of the data is the nominal event rate. Lower rates (than those expected at a specific Zenith angle) generally indicate bad atmospheric conditions, while rate fluctuations that exceed the statistical level of fluctuations can be most likely related to technical problems during the data taking.
- **$\gamma$/hadron separation and energy estimation**: once for each triggered event the set of image parameters is available, an algorithm is performed in order to create matrices for estimating the nature of the event (hadron-like or $\gamma$-like) and to reconstruct its energy. For the former task, a subsample of MC-$\gamma$ events is used vs. a subsample of background hadronic events extracted from the real data. For the latter task, only a subsample of MC-$\gamma$ events is used. The program used for performing these calculations is
called OSTERIA (Optimize STandard Energy Reconstruction and Image Analysis). The determined \( \gamma \)/hadron matrices and the energy estimation matrices are then applied to the signal data, to the background data and to a test (statistically independent) sample of MC-\( \gamma \) events. A new parameter called Hadronness, related to the probability for each event to have a hadronic nature, is thus calculated, as well as the Estimated Energy parameter. The program which allows these calculations is called MELIBEA (MErge and Link Image parameters Before Energy Analysis).

- **Calculation of the flux and of the light curve**: several calculations are done during this step of the analysis. The effective time of the observation is estimated from the data sample. Through the application of \( \gamma \)/hadron separation cuts to the data (signal and background) and to the test sample of MC-\( \gamma \), as function of bins of Estimated Energy, the number of excesses events are extracted. The Effective Collection Area is determined from the corresponding MC-\( \gamma \) sample. The spectrum and the light curve are then computed as function of estimated energy and time. In case the observation does not present significant \( \gamma \)-ray excesses, the calculation of the flux upper limits is computed. All these calculations are performed by the MARS program called FLUXLC.

- **Unfolding of the spectrum**: the energy spectrum is unfolded taking into account the energy resolution of the telescope and other analysis effects. The energy spectrum in bins of true energy is thus obtained. The program used for this task is called UNFOLDING.

- **Sky-maps**: there are two independent programs in MARS which allow the production of sky-maps for a given detected source which indicate the arrival directions of the incoming \( \gamma \)-rays and allow to determine the excess position or to study extended sources. The name of these programs are CASPAR and ZINC.

For each step of the analysis, the different programs can be steered by inputcards which allow the analyzers to set the most important parameters used by the algorithms.

4.2 Data taking modes and classification of the MAGIC-I data

4.2.1 Data taking modes

The MAGIC telescopes observe in the so-called tracking mode, i.e. the telescopes follow a given position in the sky while recording air shower images. There are three different data taking modes: the ON mode, the OFF mode and the so-called false-source tracking mode \cite{117}, usually named WOBBLE mode. These data taking modes are used for the single telescope observations as well as for the stereoscopic observations.

- **ON mode.** The potential \( \gamma \)-ray source is tracked in such a way that its nominal position is always located in the center of the camera. The trigger acceptance for the \( \gamma \)-ray showers is the maximum achievable one for constructive reasons.

- **OFF mode.** In case ON data are taken for a given source, additional dedicated data must be taken in order to obtain reliable background control data. Therefore, a close-by sky region without any known \( \gamma \)-ray sources is selected and tracked. The OFF data should have very similar observation conditions as the corresponding ON data, such as the star field, the Night Sky Background (NSB) brightness, the Zenith and the Azimuth distributions, the moon phase (in case of moon observation) and the overall hardware setup. If the above conditions are fulfilled, the image parameter distributions of the OFF
sample (before the analysis cuts or after the analysis cuts outside the expected signal region) should match the ON data image parameter distributions. Any strong deviation indicates that most likely the two data sample were taken in different observation and/or hardware conditions.

- **WOBBLE mode.** In this observation mode, two directions on opposite sides and 0.4° far away from the potential γ-ray source are tracked alternately for 20 min each. The two position are called W1 and W2 positions. The main consequence is that the nominal source position in the camera is always 0.4° far from the center of the camera. This procedure provides a simultaneous measurement of the signal and the background. Therefore, there is no need a priori of additional time which must be dedicated to OFF observations. Indeed, the background can be determined by calculating the shower image parameters with respect to the so-called anti-source position, which is 0.8° symmetrically far (with respect to the camera center) from the source position. Additional background regions (for example at ±90° from the source position) can also be defined (see Fig.4.1).

The fact that two opposite directions in the sky (W1 and W2) are tracked each for 20 min is done in order to avoid systematic effects due to efficiency differences in the camera. Indeed, the event acceptance is not just a smooth function of the camera geometry and the trigger region: the inhomogeneities are mostly due to the presence of dead pixels in the camera, different rates of the individual trigger macrocells and a possible insufficient flatfielding of the camera. Therefore, by swapping each 20 min the nominal source and anti-source positions (such as to be again 0.4° from the camera center but 180° rotated with respect to the previous positions), the source and the anti-source positions spend almost the same amount of time in the same camera regions (if a reasonable amount of data is taken and a roughly equal amount of observation time is spent in both W1 and W2 WOBBLE positions). An additional reduction of possible systematic errors is given by the fact that the nominal source and anti-source positions continuously rotate around the so-called WOBBLE circle, due to the Alt-Azimuth mount of the telescope.

Currently, the major part of the MAGIC data are taken in WOBBLE mode. The most important advantage of this observation mode, compared to the ON-OFF one, is the simultaneous measurement of the background. This permits to increase the duty cycle of the observations and, at the same time, to reduce the systematic effects due to different weather conditions and NSB levels which, conversely, could strong affect the ON-OFF observations. In addition, the frequent wobbling between the two WOBBLE positions (W1 and W2) allows the averaging over systematic effects in the camera acceptance. Anyway, the WOBBLE mode has as well some disadvantages, mainly related to the reduction of the γ-ray trigger acceptance (of the order of 15-20%, depending on the energy [118]) due to the offset of the γ-ray source from the center of the camera. Moreover, in case the γ-ray signal extraction is needed for a short data sample (e.g. a data sample with only few WOBBLE pointings), the background determination can suffer high uncertainties due to the quite high camera inhomogeneities for lower energies (of the order of 20% after the analysis cuts) not sufficiently averaged by the wobbling procedure.

\[1\text{This effect is highly reduced in case of MAGIC-II because of the larger trigger area of the PMT camera (see section 3.3.2).}\]
4.2. Data taking modes and classification of the MAGIC-I data

Figure 4.1: Sketch of the possible observation modes. In the ON-OFF observations the $\gamma$-ray source (ON data) and the sky position for the determination of the background (OFF data) are tracked such that their nominal positions are always at the center of the camera (green spot). On the contrary, during WOBBLE observations, the nominal source and the anti-source positions are displaced $0.4^\circ$ far from the camera center and observed simultaneously (red and blue spots). As a consequence of the telescope mount, during the data taking these positions rotate along the so-called WOBBLE circle (black $0.4^\circ$ radius circle). Every 20 min the WOBBLE pointing position on the sky is changed, usually placing the nominal source and anti-source positions on the opposite side of the camera center respectively. These conditions allow to reduce the systematic errors introduced by the intrinsic camera inhomogeneities. More control background region can be defined along the WOBBLE circle, for example at $\pm 90^\circ$ from the source position (yellow and purple spots), increasing the statistics of background data. The number of allowed background regions are basically limited by the extension of the $\gamma$-ray source and by the fact that additional OFF regions different from the anti-source are not totally equivalent to the anti-source position, i.e. they have less WOBBLE symmetries compared to those the nominal source and anti-source positions have. Therefore for the lowest energy range ($E < 200$ GeV), where the camera inhomogeneities are more pronounced, only the anti-source position is used as OFF control region. Above $\sim 200$ GeV the use of 3 OFF background regions generally starts to be reliable.

4.2.2 Real data

During the data taking, all the digitized information related to the observation conditions and to the trigger events are stored in files called runs. Unique run numbers are assigned by the DAQ system. Three basic kinds of run are recorded:

- **Pedestal runs**: the background light which is not related to the Cherenkov flashes and which is mainly caused by the NSB and the intrinsic noise of the readout chain must be taken into account and subtracted from the Cherenkov light content. To achieve this, special pedestal runs with the trigger activated randomly (which minimizes the probability to trigger air shower events) are taken before the start of each observation. A pedestal run contains typically 1000 events. These events contain the signal baseline and its fluctuations. In this way the mean pedestal signal of each pixel of the camera can be estimated.
4. THE ANALYSIS CHAIN OF MAGIC DATA

- **Calibration runs**: in order to be able to calculate the conversion factor from FADC counts into number of photoelectrons and the arrival time offsets, special runs are taken containing events triggered by the standard calibration light pulses from 10 UV LEDs generated by the calibration system (see section 3.2.6). These runs contain usually 4096 events.

- **Data runs**: these files contain the information, coming from each pixel, about the events that triggered the telescope while tracking a given sky position. Thanks to a dedicated trigger signal, these runs contain also interleaved calibration events at a fixed rate (typically 25-50 Hz), which are used to monitor and update the conversion factors and the arrival time offsets, initially calculated from the calibration run, in order to correct possible fluctuations and drifts. Also the pedestal levels are properly extracted directly from the data runs. The data runs contain 10775 events (53521 events for the old 300 MSamples/s FADC system) corresponding to 1 GB of disk space. When this disk size is reached, the subsequent data run is started automatically with no dead time. According to the type of data taking, the data runs are classified in three main typologies: ON data, OFF data or WOBBLE data. In case of WOBBLE data, each runs is divided into a number of subruns as to cover 20 minutes of observation for each WOBBLE position. Each subruns has the typical dimension of the runs taken in ON and OFF mode.

All runs are grouped into sequences, each of which contains at least one pedestal run and one calibration run which are used to estimate the pedestals and the pixel responses used for calibrating the data runs belonged to the same sequence. The duration of a sequence can vary according to the data taking mode. During the data taking, also ASCII files containing regular reports from the different telescope subsystems are recorded. The aim of these files is to give direct information about the correct behavior of the main subsystems during the data taking and they represent useful tools for the subsequent quality data monitoring and selection. These pieces of information are incorporate to the regular data runs in the first steps of the analysis, and organized in ROOT trees containing a set of parameter containers for every entry.

4.2.3 Monte Carlo data

As already mentioned, MC-$\gamma$ events are essential for the reduction of MAGIC data since they allow the calculation of several physical quantities which cannot be measured directly, as the expected morphology of $\gamma$-ray images as a function of the image parameters. This is used to provide tools for the $\gamma$-hadron separation and energy reconstruction of the real events, and the effective $\gamma$-ray Effective Collection Area for a given observation.

Despite the fact that the shower events recorded by an IACT have mostly a hadronic nature, and thus a proper simulation of those events should be desirable, the simulation of hadronic showers is not used for the extraction of the physical quantities of a given observation. One of the reason can be addressed to the difficulty in generating enough amounts of hadronic MC events. Indeed, the Cherenkov hadronic events, being generated by a much more complicated physical interactions in atmosphere compared to the $\gamma$-rays induced showers, and being isotropic, are widely more difficult to simulate. For this reason, the hadronic background, for a given observation, is estimated by the aid of proper real data depending on the observation mode.

The MC-$\gamma$ simulation for the MAGIC-I telescope is divided in three main steps and the final available files have the same formats and typologies of the real data:
4.2. Data taking modes and classification of the MAGIC-I data

- The simulation of the $\gamma$-ray shower development in air are performed by a dedicated software called CORSIKA [63] (version 6.019) using the US standard atmosphere. The input parameters of the software are specifically set for the MAGIC telescope site. Binary files store all the relevant parameters of the Cherenkov photons arriving on the ground around the telescope location of the simulated showers. Typically, each file contains the information of $10^5$ $\gamma$-ray simulated events.

- A program called REFLECTOR [119] processes the binary output of CORSIKA, containing information on the Cherenkov photons direction and their position on the ground. This program accounts also for the Cherenkov light absorption and scattering in the atmosphere (using the US standard atmosphere to compute the Rayleigh scattering as well as the Mie scattering losses using the Elterman model [120] for the distribution of aerosols and ozone). During this step, the reflection of the survived Cherenkov photons on the mirror dish is also simulated and their location and arrival time on the camera plane are obtained.

- A program called CAMERA [121] processes the output of the REFLECTOR program simulating the MAGIC-I PMT camera as well as the complete readout chain including the PMT signal response, the VCSELs, the amplifiers, the optical cables, the trigger system and the FADC readout. Realistic pulse shapes, noise levels and gain fluctuations obtained from the real MAGIC-I data have been implemented in the simulation.

Depending on the telescope performance and the observational conditions, many parameters can be tuned during the MC simulation chain in order to get a MC-$\gamma$ sample better reproducing the real data. Once the telescope conditions are set, the most important quantities to be considered for the simulation are:

- The energy range. Typically it is set from 10 GeV to 30 TeV. The energy distribution is a pure power law with a spectral index of -2.6 (for the so-called standard MC-$\gamma$ production) and -1 (for the so-called high energy MC-$\gamma$ production).

- The Zenith angle range. This is one of the most important parameters since the telescope performance is strictly related to it (see chapter 5).

- The Azimuth range. In MAGIC-I simulated data this parameter is fixed to two specific values equal to $0^\circ$ and $90^\circ$ (where $0^\circ$ indicates the Geomagnetic North). As it will be shown in chapter 5, for MAGIC stereo data all the Azimuth values have been considered in the simulation.

- The maximum impact parameters of the showers. It is commonly set to be 300 m for the $\gamma$-ray events with Zenith angles below $<30^\circ$ and 450 m for higher Zenith angles.

- The simulated optical Point Spread Function (PSF, sigma of the gaussian). In MC-$\gamma$ there are two contributions to the optical PSF: a fixed one with sigma equal to 7 mm at the REFLECTOR simulation level, and an additional one that can be set in the CAMERA simulation program to tune the MC-$\gamma$ to the data. These two contributions add in quadrature. Typically, set of MC-$\gamma$ samples with PSF spanning a range of values between 10.5 mm and 15 mm are generated\(^2\).

\(^2\)See http://magic.pic.es/priv/data/psf_plot.html for an exhaustive explanation of the two terms quadratically summed used for the computation of the final effective MC-$\gamma$ PSF.
The observation mode of the simulation. Depending on the data taking mode, the nominal source position in the camera frame can be set to the center (ON-OFF mode data) or 0.4° far from the center (WOBBLE mode data).

All the analysis chain steps performed on the real data are as well applied to the simulated data. Thus, after processing the full simulation chain from CORSIKA to CAMERA, also particular pedestal and calibration runs are simulated for the MC-\(\gamma\) data and used for estimating the proper pedestals and pixel responses.

### 4.3 MAGIC-I standard analysis

In this section the main steps of the MAGIC-I standard analysis introduced in section 4.1 are described.

#### 4.3.1 Data preparation

As first step of the analysis, the MARS program MERPP takes care to translate the binary raw data of recorded atmospheric showers, coming out directly from the DAQ system and containing the full information available per pixel (digitized signal amplitude vs. time) for every triggered event, into the custom MARS file format. During this operation the program merges the raw data with all the information stored in the report files coming from the major telescope subsystems (like the camera control system, the tracking system, the starfield monitor, the trigger system, the central control system and the weather station). All these pieces of information are then used in the subsequent analysis steps for the data selection.

#### 4.3.2 Calibration

The aim of the calibration process of the raw data is to extract for each triggered events of the raw data and for each channel the signals from the FADC slices and to convert the extracted information back to number of photoelectrons. In this process also the arrival times of the signals are determined. The calibration of MAGIC-I data is performed by the MARS program CALLISTO and it is done basically in two steps: in the first step the signal of each pixel is extracted from the data, then a conversion factor is calculated and applied to the extracted signal to obtain the absolute number of photoelectrons.

A complete description of the MAGIC calibration system and the possible calibration methods can be found in [89,98,122].

#### Signal extraction

The recorded FADC slices comprise not only the Cherenkov photons but also fluctuation of the NSB and noise induced by the readout chain. After the subtraction of these pedestal offsets, the signals coming from the FADCs are reconstructed in a fixed window smaller than the full digitized information. This is done because a small size of the window allows better noise suppression, since in this way less background fluctuations enter the reconstruction algorithm. The signal and the arrival time extraction can be done using several different methods [122]. Since the upgrade of the MAGIC-I data acquisition to 2 Gsamples/s in 2007 (see section 3.2.5), the so-called spline method method has become the standard one. It deals with the integration around the peak of a cubic spline built from the raw digitized pulse (pedestal subtracted). Besides the integrated signal, the arrival time of the pulse is computed,
as the position of the rising edge of the spline at 50% of the peak value. Therefore, the signal extractor returns two parameters:

- The signal amplitude in FADC counts that is converted into photoelectrons in the subsequent step.
- The arrival time of the signal. This time information can be used to identify pixels that very likely do not contain a signal belonging to the triggered shower.

**Conversion of Signal Amplitude into Photoelectrons**

In order to convert for each channel the extracted signal amplitude in FADC counts into a number of photoelectrons (assuming a linear relation between both quantities), calibration factors are needed. These calibration factors $C_i$ are in general different for each channel $i$. Several methods were developed to determine the conversion factors from the dedicated calibration runs and from the interleaved calibration events simultaneously recorded in the data runs (see 4.2.2). The current used method is the so-called excess–noise factor method ($F$-factor method) [97,99,123]. This method is based on the deviation from the poissonian statistics in the pixel signals. For each channel, the number of photoelectrons are assumed to be Poisson distributed, i.e. the distribution of the photoelectrons has a mean of $N$ photoelectrons and a root mean square (RMS) of $\sqrt{N}$. The measured amplitude in FADC counts after pedestal subtraction has a mean value $Q$ and a RMS $\sigma_Q$. However, the RMS $\sigma_Q$ is wider than a pure Poissonian one. The relation between the relative width of the distribution of the photoelectrons ($\sqrt{N}/N$) and the relative width of the distribution of the extracted signals ($\sigma_Q/Q$) can be thus written as

$$ F \cdot \frac{1}{\sqrt{N}} = \frac{\sigma_Q}{Q} \quad (4.1) $$

where the $F$-factor accounts for the additional broadening of the measured distribution beyond what is expected by Poisson statistics. This broadening is mainly originated in the statistical fluctuations in the amplification of the electrons in the PMT dynode system$^3$. The $F$-factor has to be quantified in laboratory for each individual PMT by evaluating single photoelectron spectra. In case of MAGIC-I PMTs, an average $F$-factor of 1.15 is used for all PMTs. Since $Q$ and $\sigma_Q$ are determined by the calibration events and the $F$-factor is known, the conversion factor $C$ can be thus derived

$$ C = \frac{N}{Q} = F^2 \frac{Q}{\sigma_Q^2} \quad (4.2) $$

The estimated (and constantly updated) conversion factors $C_i$ of each channel are then applied to the data events to obtain the number of the photoelectrons $N_i$ from the measured number of FADC counts $Q_i$

$$ N_i = C_i \cdot Q_i \quad (4.3) $$

It is worth noting that this method provide the relation between FADC counts and the number of photoelectrons released in the PMT, but not the photon to photoelectron conversion efficiency and losses due to the plexiglass entry window of the PMT camera and the light collectors. In order to get these absolute calibration information the light input to the PMT camera has to be known. Therefore the PIN diode calibration method or the blinded pixels

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$^3$In principle, small fluctuations of the signal introduced by the amplifiers, optical transfer, FADC and instabilities of the calibration light pulser should be taken into account as well, but they turn out to be negligibly small compared to the intrinsic $F$-factor of a PMT.
one can be used (see section 3.2.6). It has been estimated that, considering the averaged QE of the PMTs plus plexiglass window and light guides, the global efficiency is about 18%.

### Bad pixel treatment

During the calibration process, the pixels showing malfunctioning hardware behaviors (related to the PMTs themselves or the corresponding receivers or the FADCs) are found out and tagged as bad pixels. Typically, the amount of unsuitable pixels is of the order of 5% (10-15 pixels). The signal of bad pixels is replaced by the linear interpolation of the signals of the properly working neighbour pixels (in case they are more than three).

### 4.3.3 Image cleaning and image parametrization

When a shower event triggers the telescope, the signals coming from all pixels are stored by the DAQ system, although most of them contain light which is most likely related only to fluctuations of the NSB rather than to the Cherenkov light of the shower. Pixels that presumably do not contain a shower signal or contain signals from small tracks far outside the shower must be therefore rejected in order to allow a robust and stable determination of the image parameters (see section 2.3) associated to the event. Indeed, any disturbing pixel far from the pixels which are involved in the shower image can severely mislead the image reconstruction. On the other hand, any unnecessary tightening of the exclusion limits will lead to signal loss at low γ-ray energies. For this reasons, clever algorithms are needed, which have to be tuned such as to provide a low energy threshold and, at the same time, ensure that the vast majority of the images are unaffected by noise.

The MARS program which perform the image cleaning algorithm and the subsequent calculation of the image parameter is called STAR. In principle, three main input quantities can be employed in setting the image cleaning algorithm. These quantities (given by the output of the calibration process) are the number of photoelectrons of each pixel (already pedestal subtracted), the relative arrival time of the measured signal in FADC slices and the RMS of the NSB fluctuations per pixel.

Before the update of the MAGIC-I 300 MSamples/s FADC system to the 2 GSamples/s MUX FADC system (see section 3.2.5), several studies were made to find out the best image cleaning procedure for the MAGIC-I data considering the three above mentioned quantities [89, 124–126]. As a compromise between a low energy threshold of the analysis and the robustness of the method, the so-called Absolute 10-5 cleaning was made standard by the MAGIC Collaboration. In this method, only the charge (in photoelectrons) contained in the pixels is used as criteria for keeping or rejecting a given pixel (hence the name Absolute). The Absolute cleaning compares the number of reconstructed photoelectrons (phes) per pixel with two charge reference level values \(Q_{\text{core}} = 10\) phes and \(Q_{\text{boundary}} = 5\) phes. The algorithm proceeds in three steps:

1. All pixels with a number of photoelectrons above \(Q_{\text{core}}\) are selected.

2. A selected pixel is marked as core pixel, if it has at least one direct neighbour also selected in the previous step. Note that an event can have several isolated groups of core pixels.

3. Pixels with direct neighbours having survived the previous steps and which have a charge greater than \(Q_{\text{boundary}}\) are also included and marked as boundary pixels.

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4 time slice = 3.3 ns for the old 300 MSamples/s FADC system whereas 1 time slice = 0.5 ns for the new MUX FADCs.
those pixels which are not core or boundary pixels are excluded (i.e., their charge are set to zero, see Fig. 4.2).

![Image](image.png)

(a) Event before image cleaning.  
(b) Event after the image cleaning.

**Figure 4.2:** Shower image before and after the application of the image cleaning algorithm. The ellipse superimposed on the camera corresponds to the parametrization of the event.

This algorithm is simple and robust if initialized with the standard levels $Q_{\text{core}} = 10$ phes and $Q_{\text{boundary}} = 5$ phes. Indeed it is not strongly affected by different levels of the NSB since the two levels are much higher than typical values of the pedestal RMS. The Absolute cleaning does not use the arrival time information, thus it does not allow to reach the lowest possible energy threshold. Moreover, in this cleaning different levels of noise between individual pixels (e.g. due to stars in the field of view of the camera) are not taken into account. For these reasons also other cleaning algorithms based on the arrival time and/or the pedestal RMS information were studied. However, only after the installation of the new 2 GSamples/s MUX FADC system (in 2007) the time information has been fully exploited in the image cleaning algorithm for the MAGIC-I data. Indeed, the better determination of the temporal evolution of the individual pixel signals of the recorded images allows to significantly improve the performance of the image cleaning, thanks to the addition of proper time constraints to the selection of the core and boundary pixels. These further time constraints allow to relax the standard cleaning levels $Q_{\text{core}} = 10$ phes and $Q_{\text{boundary}} = 5$ phes, lowering in this way the energy threshold. For these reasons, the MAGIC Collaboration has decided to make the so-called Time-constrained 6-3 cleaning as the current standard image cleaning for the MAGIC-I data. The new algorithm, which has the two charge reference level values $Q_{\text{core}} = 6$ phes and $Q_{\text{boundary}} = 3$ phes, can be summarized as follows:

- After selecting the core pixels in the same way as in the Absolute cleaning (see above), it rejects those core pixels, whose arrival time is not within a fixed time constraint of 4.5 ns with respect to the mean arrival time of all core pixels.

- In the selection of the boundary pixels, it also requires that the time difference between the boundary pixel candidate and its neighbour core pixels is less than 1.5 ns.

The two time constraints (4.5 ns and 1.5 ns) have been optimized on MC-\(\gamma\) basis [77]. In Fig. 4.3 a comparison between the old Absolute 10-5 and the new Time-constrained 6-3 image
cleaning is reported. A detailed comparison between the different image cleaning techniques and their performance can be found in [95]. Once the images have been cleaned, the image parameters are calculated. The most impor-

Figure 4.3: Illustration of the improvement in the shower parameter reconstruction using the Time-constrained 6-3 cleaning compared to the non-time Absolute 10-5 image cleaning and to the non-time Absolute 6-3 image cleaning. The image refers to a MC-\(\gamma\) events with energy 71 GeV and impact parameter of 111 m. The simulated \(\gamma\)-ray source is located in the center of the camera (yellow star). First row: display of raw recorded data (left) and arrival times information (right). Second row: comparison of standard cleaning with 10-5 phes minimum charge levels (left) and 6-3 minimum charge levels (right). The Absolute 10-5 non-time cleaning loses information whereas lowering the charge levels to 6-3 (without time constraints) can add noise to the shower image which could lead to wrong image parameter reconstruction. Bottom: image obtained with the Time-constrained 6-3 cleaning (with 4.5 ns and 1.5 ns time constrains). The reconstruction of the image is much improved. Image taken from [95].
tant ones have been already introduced in section 2.3. The shower images are well characterized by these parameters and all the other information are no longer used in the subsequent analysis steps. Indeed, the image parameters are the key tool for the whole data analysis: it is just thanks to them that it is possible to separate the $\gamma$ events from the dominating hadron background, to estimate their incoming direction and their energy.

Just after the image parameter calculation some cuts (the so-called quality cuts) are applied to the data. The aim of these cuts is to remove cleaned events which are most likely not produced by air showers or which can be related to showers whose parametrization is probably affected by large uncertainties (as in case of large values of the LEAKAGE parameter or NUMBER OF ISLANDS). In particular, the so-called spark events (originating from occasional discharges between the photocathode and the light guide of the PMTs, which can trigger the readout of the camera) and the so-called car-flash events (due to rare artificial illumination of the camera during the data taking) can be easily identified and removed.

4.3.4 Data selection and quality cuts

Before starting the $\gamma$-hadron separation, a check on the quality of the runs is useful to reject runs not suitable for further analysis. In this step, the information coming from the major subsystem of the telescope are used to found out possible bad behaviors during the data taking. The most important information come from the nominal trigger rate, the drive system, the weather station (which provide e.g. the humidity value during observations) and the pyrometer which allows to measure the sky quality inside the field of view of MAGIC-I\(^5\) and provides a percentage quantity called cloudiness. The less the value, the better the atmospheric conditions. Typically values of cloudiness smaller than 40% indicate good sky quality. Moreover, the image parameter distributions are usually checked run per run in order to reject runs with anomalous distribution behaviors.

To get information day by day about the hardware conditions and the status of the readout software, the online run book is a useful source for manifold information. The electronic run book is maintained by the shift crew during the observations. Another useful tool is represented by the so-called automatic dailycheck. It consists in a program automatically executed every morning after the data taking to extract information from the subsystems through the central control and to produce plots which can be easily checked by the analyzer in order to find out possible problems in the data. For more detail about the automatic dailycheck see [127–129].

4.3.5 $\gamma$/Hadron separation

After the image cleaning and the data selection, the events are highly dominated by background events not induced by $\gamma$-rays. These background events are mostly originated by cosmic hadrons (mainly protons, but also nuclei of helium and of heavier elements) and in a smaller part by isolated muons which are produced in hadron induced showers through $\pi^\pm$ and $K^\pm$ decays.

Before additional analysis cuts, even during the observation of bright $\gamma$-ray sources like the Crab Nebula, the ratio of recorded background events over the $\gamma$-ray ones is of the order of $10^4$. Therefore, to obtain a sensitive measurement of a VHE $\gamma$-ray source, an efficient discrimination between the $\gamma$ events from the background is one of the most important task to be performed during the analysis reduction of the data. Since the major contribution to the

\(^5\)see http://www.magic.iac.es/subsystems/pyrometer/
background is due to the hadronic showers, this discrimination procedure is called γ-hadron separation.

The basic idea to get an efficient suppression of the background is to take into account the morphological and temporal differences which distinguish the γ-ray events from the background events. Indeed, images produced by hadronic showers are typically longer, wider and have more irregularities and larger arrival times than those produced by γ-rays. The physical reason for this difference is the larger longitudinal and lateral particle distribution in the hadron induced air shower developments (see section 2.1). Thus, the image parameters whose distributions show significant dependences on the γ-ray or hadronic nature of the showers can be used to estimate the nature of the events in the real data.

At the beginning, IACTs applied fixed cuts, called static cuts, to the image parameters for the background rejection. Since the shape of the shower images depends on other variables, as the energy of the showers, the impact parameter and the Zenith angle of the observation, the so-called dynamic cuts were successively developed for IACTs’ data analysis. These cuts take into account the different parameters correlations, and turned out to provide a big achievement in the background rejection [130]. Alternative γ-hadron separation technique have been also used, such as the so-called SuperCuts optimization [131] or neural networks [132].

The standard γ-hadron separation algorithm used for the MAGIC-I data is performed by a multidimensional classification technique based on the Random Forest (RF) method [133–135] and is performed by the MARS program OSTERIA. The basic difference between the RF method and the methods using static or dynamical cuts is that the cuts are not parametrized. The RF does not apply independent cuts on the discriminator image parameters but employs a “forest” of decisonal trees to classify each event, using all the parameters simultaneously, whereby correlations among them are taken into account automatically during a training section. The RF method needs training samples to find a set of classification trees in the space of the image parameters. The training sample representing γ-ray showers is a sample of MC-γ events⁶, whereas the training sample for the hadronic events is constructed by selecting randomly events from the real experimental data. The input parameters used to train the method can be selected by the analyzers. In each step of the tree growth process one image parameter is chosen at random and a cut is found in this parameter such as to minimize the so-called Gini index [136]:

\[ Q_{Gini} = 2 \left( \frac{N_{left}^{signal} \cdot N_{left}^{background}}{N_{tot}^{left}} + \frac{N_{right}^{signal} \cdot N_{right}^{background}}{N_{tot}^{right}} \right) \]  

(4.4)

where \( N_{left} \) (\( N_{right} \)) are the events of the subsamples that are below (above) the cut and which constitute new subspaces. Subsequently in each subspace an image parameter is chosen at random and the Gini index minimized. This procedure is repeated until, in each subspace at the end of the tree (terminal node), the Gini index is zero or the number of remaining events is less than a certain threshold (usually 1-10 events) or belonging to only one category. In order to avoid over-fitting, several (usually 100) random trees are generated in this way each from different random samples of training events. The final sample of trees forms the Random Forest. The combination of random forests constitutes an ensemble of uncorrelated trees, which are combined to form a more generalized predictor. In order to classify an event, each of the trees is followed until its terminal node. At the end of the process, it is counted

⁶In order to minimize the systematic errors in the RF method, the γ-ray events of the train sample used for the algorithm should resemble real γ-ray events as closely at possible. Comprehensive studies [42,115] have shown that there is a reasonable agreement between the observed γ-ray events and the simulated ones.
in how many trees the event ends up at a node that consists of a hadron and a $l = 1$ value is assigned to the events. In case the event ends up at a node that consists of a $\gamma$ a $l = 0$ value is assigned to the events. As there is a whole forest, each random tree $i$ assigns a value $l_i$ to the event, and the parameter $\text{Hadronness}$ can be calculated as

$$\text{Hadronness} = \frac{\sum_{i=0}^{N_{\text{trees}}} l_i}{N_{\text{trees}}}$$  \hspace{1cm} (4.5)

The RF algorithm produces $\gamma$-hadron separation matrices which are then applied to the real data (ON, OFF or WOBBLE) as well as to a sample of MC-$\gamma$ statistically independent from the one used to train the RF. In this way, the new image parameter $\text{Hadronness}$ is added to the events. The $\text{Hadronness}$ is a real number ranging from 0 to 1 and yielding values distributed close to 1 for hadron-like and close to 0 for $\gamma$-like event (see Fig. 4.4(a)). A cut in $\text{Hadronness}$ is then equivalent to a cut in the multidimensional parameter space of the image parameters included in the RF training. The MARS program which take care of the application of the RF matrices is called MELIBEA.

The set of input image parameters for the RF training depends on whether the source dependent or source independent approach is used to extract the signal. Typical parameters used in the training of the RF methods are SIZE, ZENITH, WITH, LENGTH, DIST, CONC, M3LONG and (after the update of the MAGIC-I 300 MSamples/s FADC system to the 2 GSamples/s MUX FADC system) the temporal parameters TIME GRADIENT and TIME RMS, which have demonstrated to provide a significant improvement in the $\gamma$-hadron separation performance [95]. In case a source independent approach is used to extract the signal, all source dependent parameters (e.g. the DIST parameter) are not used as input parameters for the RF method. This is done in order not to bias the training towards a certain sky position. In any case, the parameter with the highest discrimination power, ALPHA, is usually not included in the list of parameters for the RF training, as it is used to extract the $\gamma$-ray signal (see section 4.3.7). A detailed study on the possibility to develop new analysis methods with the parameter ALPHA used as input parameter for the RF training can be found in [125].

In Fig. 4.4 the so-called Gini plot, which shows the relative contribution to the $\gamma$-hadron separation of the parameters used for the RF training is reported together with the distribution of the $\text{Hadronness}$ for normalized samples of $\gamma$-ray events and hadron events (statistically independent from those used to train the RF method).

It is worth noting that before the training, the SIZE distributions of the $\gamma$ and hadron training data set are made equal (re-sizing) such that the RF training does not train on possible differences in the SIZE distribution but only exploits correlations of SIZE with other parameter. Since the distribution of the Zenith parameter (for the same range of Zenith values) is typically different for MC-$\gamma$ events and hadron events, the same equalization between the $\gamma$ and hadron training data set is done also with respect to the Zenith (re-zenithing). Anyway, individual $\gamma$-hadron separation processes are performed in case of observations extending in a wide range of Zenith angle above $\sim 30^\circ$.

The separation power between $\gamma$-ray and hadron induced events is clearly a function of the energy of the event. For very low energetic events the distinction between $\gamma$-ray induced showers and hadronic showers becomes more and more difficult. The reasons are related to the poor determination of the conventional image parameters and the resembling of their distributions at low energies. This leads to a higher value for the energy threshold of the analysis (after proper $\text{Hadronness}$ cuts) compared to the trigger threshold. For low Zenith angle observation ($\lesssim 30^\circ$), the typical analysis threshold is around 100 GeV.

Since WOBBLE data provides a simultaneous measurement of the signal and the background,
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Figure 4.4: (a) Hadronness distributions for normalized subsamples of MC-\(\gamma\) and hadron events. (b) Typical Gini plot which shows the relative importance of the different parameters used to train the RF algorithm. The higher is the value for a given parameter, the higher is the separation power of that parameter.

The source dependent distributions calculated with respect to the source position and with respect to the anti-source position (and possibly with respect other background control regions defined in the camera, as explained in Fig 4.1) are highly correlated because they contain the same events. Therefore, if source dependent parameters are used in the training of the RF method, the algorithm is performed not only with respect to the source position (ON sample) but also with respect to each defined background positions (OFF samples). This is an essential feature of the WOBBLE data which guarantee (after the application of the same (in value) Hadronness cuts) to treat the ON and OFF distributions as independent samples, as in case of the ON-OFF observation mode. For more detail about this issue see [137,138].

4.3.6 Energy estimation

At first order, the recorded light content of \(\gamma\)-ray induced air showers (with impact parameter within \(\sim 120\) m, for vertical incidence) is proportional to the initial energy. However, additional dependencies exist and must be properly taken into account in order to allow for a more precise estimation of shape of the spectra of the observed sources.

The first parameter which has to be taken into account is the impact parameter of the shower, which can be characterized by the DIST parameter, together with the shape parameters WIDTH and LENGTH. Another important parameter is the Zenith angle of observation: showers observed at large Zenith angles travel a longer distance, as they have to pass through a larger layer of atmosphere. At the ground, they arrive therefore spread over a larger area and showers with larger impact parameters can be detected. On the other hand, the photon density decreases by reducing the number of showers detectable at low energies. The global effect is thus a shift of the energy threshold to higher energies. This topic will be further discussed in chapter 5.

Due to the large number of required parameters which have to be considered, a slightly modified version of the RF method is used to estimate the energy of the events\(^7\). The parameters used for the energy estimation are usually the same as for the \(\gamma\)-hadron separation. Only

\(^7\) The MARS program which performs these calculations is OSTERIA, as in case of the RF calculation for the \(\gamma\)-hadron separation.
one training data set of MC-\(\gamma\) showers with known simulated energy (true energy) is used to train the method. The data set is divided in many bins of logarithmic energy. For each of those bins, a forest is trained that separates \(\gamma\)-rays with energies that fall within the bin from those with energies outside of the bin. Subsequently, the combination of image parameters determines the probability of an event to belong to a given energy bin and the one with the highest probability is defined as the Estimated Energy of the event. As in case of \(\gamma\)-hadron separation, the MARS program MELIBEA takes care to apply the calculated estimated energy parametrization to the real data (ON, OFF or WOBBLE) as well as to a sample of MC-\(\gamma\) statistically independent from the one used to train the RF. From this latter sample it is possible to infer the energy resolution and its bias of the energy reconstruction method and the actual migration from the true energy of the MC-\(\gamma\) events \(^8\) into the Estimated Energy (the so-called Migration Matrix). In Fig. 4.5, the typical performance of the energy reconstruction method (in case of source dependent analysis and for Zenith angles of observation \(\lesssim 30^\circ\)), after all analysis cuts, are shown. The energy resolution has a clear dependence on the Estimated Energy: above 100 GeV resolution between 25% and 15% can be achieved, whereas for lower energies the reconstruction gets considerably worse due to the largest fluctuations of the recorded showers. Note that the values of the energy resolution reported in Fig. 4.5(b) have been obtained by using source dependent parameter in the RF method. In case no source dependent parameter are used, the achievable energy resolution is typically 30-40% worse in the whole energy range. The finite energy resolution and bias must

\[E_{\text{true}} \quad 10^2 \quad 10^3 \quad 10^4 \quad 10^5 \]

\[E_{\text{est}} \quad \text{(a) Energy Migration Matrix.} \]

\[\frac{E_{\text{true}}}{E_{\text{est}}} \quad \frac{E_{\text{true}}}{E_{\text{est}}} \quad \text{(b) Energy resolution and bias.} \]

**Figure 4.5:** (a) Example of Migration Matrix between reconstructed and true energy. The matrix is normalized to 1 separately for each bin in the reconstructed energy. The red dashed line indicates the perfect correlation at which the true energy of the simulated \(\gamma\)-ray events is equal to the reconstructed energy. (b) Quality of the energy reconstruction defined by the resolution (red points) and the bias (blue) points. Note that below \(\sim 100\) GeV the energy resolution and its bias get considerably worse.

be taken into account and correct in order to calculate accurately the final energy spectra of the observed sources. This is done in the so-called unfolding procedure (see section 4.3.11). As in case of the \(\gamma\)-hadron separation, for WOBBLE data, if source dependent parameters are used as input parameters for the energy reconstruction process, the algorithm is performed with respect to the source position as well as with respect to the chosen background regions.

\(^8\)The true energy is given only for the MC-\(\gamma\) events, since the information of the energy value used for the simulation is known.
4.3.7 $\gamma$-ray signal extraction and optimization of the cuts

The extraction of the $\gamma$-ray signal from a given source is done after the application of a proper *Hadronness* cut. Since the *Hadronness* parameter is energy dependent, the signal is usually extracted in bin of *Estimated Energy*. However, even after a *Hadronness* cut, there are always background induced images (background events) surviving which are mixed with the $\gamma$-ray induced images (signal events). This is the reason why it is mandatory to have a set of proper OFF data in order to evaluate the irreducible background of the observation.

As explained in section 4.2.1, in case of ON-OFF observation mode the control background is extracted from dedicated OFF data, whereas in the WOBBLE mode, the background can be evaluated from different region of the camera from the same data set. Shower images that are produced by $\gamma$-rays from the source under observation are pointing with their major axis to the source position in the camera. Therefore, the possible $\gamma$-ray signal in the data sample is investigated through the directional information of the parameter $|\text{ALPHA}|$ (which ranges from $0^\circ$ to $90^\circ$). The $\gamma$-ray events should peak at lower values of $|\text{ALPHA}|$, whereas the background events are instead characterized by an almost flat distribution, since they have an isotropic incoming direction. This approach to extract the signal from the data is called ALPHA approach.

The number of excess events can thus be calculated (after an appropriate *Hadronness* cut) by subtracting the normalized number of OFF events, surviving the same *Hadronness* cut applied to the ON data, from the number of ON events in the signal region of the $|\text{ALPHA}|$ distribution. This signal region is dependent on the *Estimated Energy* and it is usually lower than $15^\circ$.

The distribution of the $|\text{ALPHA}|$ parameter for the ON and OFF samples is called *Alphaplot*. The normalization between the ON and OFF in the *Alphaplot* is necessary only in case of ON-OFF observation mode (since the exposure time of the two sample are in general different) and is performed in the so-called *tail region* (also called background normalization region), i.e. a region where no signal is expected and where the trend of the ON and OFF distribution should match. The *tail region* is typically chosen to be between $20^\circ$ and $80^\circ$.

The number of excess events in the signal region is then defined as

$$N_{\text{exc}} = N_{\text{ON}} - \eta \cdot N_{\text{OFF}}$$  \hspace{2cm} (4.6)

where $N_{\text{ON}}$ are the entries in the signal region of the ON sample, $N_{\text{OFF}}$ are the entries in the signal region of the OFF sample and $\eta$ is the normalization factor, i.e. the ratio between the entries of the ON sample and the OFF sample in the *tail region*.

The found number of $\gamma$-ray excess may be caused by fluctuations of the irreducible background. The hypothesis, that all excess events are in fact background fluctuations can be tested by different methods which reject this hypothesis with a certain significance level. For the IACT experiments, the widely accepted and used method for calculate the significance of an observation is given by the Li and Ma formula number 17 [139]:

$$\sigma_{\text{LiMa}} = \sqrt{2 \cdot \sqrt{N_{\text{ON}} \ln \left( \frac{(1 + \eta)N_{\text{ON}}}{\eta(N_{\text{ON}} + N_{\text{OFF}})} \right) + N_{\text{OFF}} \ln \left( \frac{(1 + \eta)N_{\text{OFF}}}{N_{\text{ON}} + N_{\text{OFF}}} \right)}}$$  \hspace{2cm} (4.7)

In $\gamma$-ray astronomy, it is commonly accepted that a source is detected if its significance is greater than 5. This means that the probability that the excess would be compatible with background fluctuation is lower than $\sim 10^{-40}\%$.

In Fig 4.6, an example of *Alphaplot* is reported. It refers to a selected sample of roughly

\footnote{It should be mentioned that the $|\text{ALPHA}|$ distribution widens with the extension of the $\gamma$-ray source.}
5 hours of Crab Nebula data taken in WOBBLE mode by the MAGIC-I telescope between November 2008 and March 2009. This sample will be used during the analysis of the Segue 1 source reported in chapter 7. The Zenith angle of the observation is between $\sim 13^\circ$ and $\sim 33^\circ$. The Alphaplots is calculated for estimated energies above 200 GeV. Standard quality cuts were applied and the Hadronness cuts were optimized in logarithmic energy bins. A fixed $|\text{ALPHA}|$ cut of $8^\circ$ (corresponding to a cut efficiency of the order of 70%) was applied for extracting the excesses and the significance.

There are basically two main strategy for the optimization of the Hadronness cuts and the size of the signal region (as a function of the Estimated Energy). Either the optimization is based on quantities which can be calculated from the MC-$\gamma$, or, alternatively, one can optimized the cuts on an independent real data sample (observed in the same period and under the same experimental conditions) containing strong $\gamma$-ray signal (like the Crab Nebula).

In the former case, the so-called quality factor $Q$ ($Q$-factor) can be calculated and optimized. The $Q$-factor is a measure of the effectiveness of a cut and give a quantitative measure of the $\gamma$/hadron separation power [130]:

$$ Q = \frac{\epsilon_\gamma}{\sqrt{\epsilon_{bg}}} $$

where $\epsilon_\gamma$ is the is the fraction of MC-$\gamma$ events that are retained after a certain cut ($\gamma$ efficiency) and $\epsilon_{bg}$ is the hadron efficiency. To compute the $\epsilon_{bg}$ the background events are extracted from the real data sample. The higher $Q$ the better is the $\gamma$/hadron separation power. Therefore, the $Q$-factor describes the enhancement of a given signal due to a certain cut. Usually, for the background rejection a cut in the $|\text{ALPHA}|$ parameter is applied in combination with a cut in the Hadronness parameter in order to maximize the corresponding $Q$-factor, with a further condition of a a minimum efficiency for the MC-$\gamma$ events.

Figure 4.6: Alphaplot above 200 GeV for a Crab Nebula data sample observed in WOBBLE mode. The black histogram is the $|\text{ALPHA}|$ distribution (after Hadronness cuts optimized in logarithmic energy bins) calculated with respect to the nominal source positions, whereas the red histogram is the $|\text{ALPHA}|$ distribution calculated with respect to the nominal anti-source (after the same Hadronness cut values, but computed with respect to the anti-source position). No normalization between the two distribution was applied.
In the latter case the cuts of the analysis are optimized in bin of estimated energy on a well established source (test source) ad then apply to the source under investigation. Usually, the cuts are selected in order to get the highest significance of the signal for the test source. Also in this case, a minimum efficiency of the cuts for the MC-γ events is required. The minimum MC-γ events efficiency required for the applied cuts determines two basic kinds of cuts: tight cuts and loose cuts. Tight cuts are obtained by requiring MC-γ efficiencies of the order of 40-50%, whereas the MC-γ efficiencies are of the order of 70-80%. Tight cuts are generally used when searching for new VHE γ-ray emitters, whereas loose cuts are applied to determine the energy spectrum and the light curve of already established γ-ray sources. The reason is that once the γ-ray signal is established, the total number of excesses events and the energy threshold become more important than the significance. Thus, looser cuts in Hadronness are applied in order to increase the γ-ray acceptance.

4.3.8 Disp method and sky maps

![Disp Method Diagram](image)

**Figure 4.7:** Left: A schematic view of two showers with the same estimated source position, but with different Disp values: images closer to the source position are more roundish. Right: The Disp parameter provides two possible solutions, the correct one is estimated by using a head-tail discriminator as the M3LONG parameter.

If the location of the γ-ray source is not precisely known or the source has an extended morphology one has to estimate the incoming direction of each detected γ-ray. In telescope arrays, multiple images of the same shower are recorded from different positions (see Fig. 3.2) and the incoming direction can be derived by the point of intersection of the major axes of the images (see also section 4.4). In order to reconstruct the impact point of γ-ray shower if the shower is imaged by only one telescope, the so-called Disp method is used [117, 140]. The basic idea behind this method is to use the information of the shape of the shower image to reconstruct the incidence direction of each detected shower. In fact, the feature that images of showers with a given direction become more roundish the closer the impact point of the shower is to the telescope (see the left sketch of Fig. 4.7) can be used to estimate the angular distance between the image center of gravity (CoG) and the unknown source position (i.e. the quantity Disp) which is assumed to lie on the prolongation of the major axis of the image. This assumption is plausible since
the major axis is a projection of the shower axis onto the camera. The \(Disp\) can be estimated for each event by using the so-called elongation of the image, defined as the ratio of the shape parameters \(\text{WIDTH} \) and \(\text{LENGTH}\). The first parametrization for the \(Disp\) was originally proposed by the Whipple Collaboration \cite{117}

\[
Disp = \xi \cdot \left(1 - \frac{\text{WIDTH}}{\text{LENGTH}}\right) \tag{4.9}
\]

where \(\xi\) is a factor dependent on the image parameter \(\text{SIZE}\).

A more general parametrization has been adopted in case of the MAGIC-I data \cite{141}

\[
Disp = A(\text{SIZE}) + B(\text{SIZE}) \cdot \frac{\text{WIDTH}}{\text{LENGTH} + \rho(\text{SIZE}) \cdot \text{LEAKAGE}} \tag{4.10}
\]

where \(A\), \(B\) and \(\rho\) are second order polynomial function of \(\log(\text{SIZE})\) and are optimized from MC-\(\gamma\) simulations or real data from a strong and well-known point-like source. The \(\text{LEAKAGE}\) parameter is used in the parametrization in order to take into account the images which are truncated at the edge of the camera. This allows a better reconstruction of the arrival direction of the more energetic showers.

The \(Disp\) algorithm has an important degeneracy: there are two possible solutions for a source location along the major shower axis (see the right sketch of Fig. 4.7). Therefore the use of a head-tail discriminator (as the M3LONG parameter) is essential in order to judge which of these solutions is most probably the correct one, placing the source position closer to the head of the image. Recently, it has been implemented in MARS the possibility to calculate the \(Disp\) parameter by using the \(\text{RF}\) method \cite{142} in a similar way as it is done for the energy reconstruction, which has demonstrated to provide a significant improvement in terms of angular resolution and sensitivity of the order of \(\sim 30\%\) (for \(\text{SIZE}>400\) phes) compared to the determination of the \(Disp\) parameter with the parametrization in equation 4.10.

The \(Disp\) method allows one to reconstruct arrival directions of the events on the camera plane and thus provides a reconstructed source position along the major axis for each event. In case of \(\gamma\)-ray events, these reconstructed source positions should be close to the real position of the \(\gamma\)-ray source projected on the camera, whereas for hadrons the reconstructed sources are isotropically distributed. Therefore, the \(Disp\) method can also be used for the detection of a signal using the so-called \(\theta^2\)-plot. The angle \(\theta\) is defined as the angular distance between the position of the \(\gamma\)-ray source projected onto the camera and the reconstructed source position, as shown in the left sketch of Fig. 4.7\textsuperscript{10}\textsuperscript{10}. For convenience, it is used the squared quantity \(\theta^2\) because the distribution of \(\theta^2\) values is, for geometrical reasons, nearly flat for background events, whereas it is peaked at \(\theta^2 = 0\) [deg\(^2\)] for \(\gamma\)-ray events. Hence, the \(\theta^2\) approach provides an alternative to the ALPHA approach in order to estimate the number of excess events for a given observation. In Fig. 4.8 an example of \(\theta^2\)-plot is reported. It refers to the same Crab Nebula data used for the \(\text{Alphaplot}\) of Fig. 4.6.

\textsuperscript{10}The sketch shows the situation for the ON-OFF observation mode. In case of WOBBLE data the nominal source position is displaced from the center of the camera and rotate along the WOBBLE circle, nevertheless the basic concepts of the \(Disp\) method are still the same as in case of ON-OFF mode.

\(Disp\) method is also used for the calculation of the sky maps of a given source: sky maps are computed by reconstructing the arrival direction of each events in the camera, transforming these into sky coordinates, filling the sky coordinates of all \(\gamma\)-like events into histograms and subtracting the expected background. The transformation (de-rotation) from local camera coordinates to celestial ones is possible because the current pointing position and the respective event time are known \cite{143}. Fig. 4.9 shows the sky map above 200 GeV obtained for the
4. THE ANALYSIS CHAIN OF MAGIC DATA

Figure 4.8: \(\theta^2\)-plot above 200 GeV for a Crab Nebula data sample observed in WOBBLE mode. The black histogram is the \(\theta^2\) distribution (after source independent Hadronness cuts optimized in logarithmic energy bins) calculated with respect to the nominal source positions, whereas the red histogram is the \(\theta^2\) distribution (after the same Hadronness cuts) calculated with respect to the nominal anti-source, which always lies on in the opposite position of the camera. No normalization between the two distribution was applied. It is worth noting that the performance of the source independent analysis is generally slightly worse with respect the source dependent analysis.

The same Crab Nebula sample so far considered and calculated with the MARS program CASPAR\textsuperscript{11}. The content of each bin has been weighted (smoothed) with a bidimensional gaussian with a \(\sigma\) taken as large as the telescope resolution at 200 GeV (~0.08\degree). For the sky maps it is important that no source dependent parameters are used in the \(\gamma\)-hadron separation algorithm in order not to bias certain camera areas.

### 4.3.9 Spectrum and light curve calculation

The differential \(\gamma\)-ray energy flux is defined as

\[
\frac{dF}{dE}(E) = \phi(E) = \frac{dN_\gamma}{dEdA_{eff}dt_{eff}}
\]  

(4.11)

where \(dN_\gamma\) is the number of excess events in bins of estimated energy \(dE\), \(A_{eff}\) is the Effective Collection Area and \(t_{eff}\) is the effective observation time. The number of excess events is extracted from the Alphaplots (or \(\theta^2\)-plots) calculated in different bins of estimated energy. The effective observation time is evaluated directly from the event rate, which is obtained from the distribution of the time intervals between successive events (and is the same for all energy bins). The Effective Collection Area of the telescope characterizes the detection efficiency for \(\gamma\)-ray induces air showers, i.e represents the area in which \(\gamma\)-ray air showers can be observed by the telescope folded by the detection efficiency \(\epsilon_\gamma\) after all cuts. All the characteristics and performance of the detector are included in the calculation of the Effective Collection Area. Due to the absence of a calibrated \(\gamma\)-ray source the Effective Collection Area can only

\textsuperscript{11}http://magic.pic.es/priv/wiki/index.php/MAGIC\_software:Caspar
4.3. MAGIC-I standard analysis

Figure 4.9: Sky map of a Crab Nebula data sample observed in WOBBLE mode for energies above 200 GeV. The bins of the original map are smoothed out with a bidimensional gaussian with $\sigma$ taken as large as the telescope resolution. The color code represent the significance of the signal. Note that the signal is well peaked at the nominal sky coordinates of the Crab Nebula. The PSF is also reported in the lower left corner of the plot.

be determined from MC-$\gamma$ events (statistically independent from those used for the training of the RF for the $\gamma$-hadron separation and the energy reconstruction). For this reason the Effective Collection Area is determine in bins of true energy.

The final Effective Collection Area depends on several quantities related to the hardware components of the detector (trigger efficiency, reflector surface, ...) and to the observation conditions (such as the Zenith angle pointing position) and the applied analysis cuts. The most evident dependence is the energy one: the Effective Collection Area has a steep rise at lower energies, related to the trigger efficiency and to the $\gamma$-hadron separation power (which improve with rising energy), and then reaches a plateau region above $\sim$ 1 TeV of the order of $10^5$ m$^2$ (for lower Zenith angle observations), as shown in Fig. 4.10(a). Another dependence of the Effective Collection Area is represented by the Azimuth angle of the observation, due to the Geomagnetic field. In case of stereoscopic observations, another pointing dependent effect must be considered and it is related to the configuration of the array. This topic will be discussed in chapter 5. In Fig. 4.10(b), the spectrum of the MAGIC-I Crab Nebula sample, so far considered, is shown.

The errors on the spectral points include statistical errors on the number of excess events and the uncertainty of the Effective Collection Area. Note that the spectrum is calculated in bins of estimated energy, which might be differ from the true energy, especially close to the energy analysis threshold (as can be noted from the Migration Matrix reported in Fig. 4.5(a)). The calculation of the energy spectrum in bins of true energy is performed in a subsequent analysis step called unfolding (see section 4.3.11).

The presence of a signal in a sample of data from a variable source requires the study of possible variations of the signal in time intervals. When the $\gamma$-flux is determined in bins of
4. THE ANALYSIS CHAIN OF MAGIC DATA

Figure 4.10: (a) Typical MAGIC-I Effective Collection Area before and after analysis cuts. (b) Spectrum of a Crab Nebula sample before the unfolding. The red dashed line represents the parametrization published in [42].

time, the so-called light curve is created. The light curve is defined as

$$\frac{dF}{dE}(E, t) = \frac{dN_{\gamma}}{dEdA_{\text{eff}}(t)dt_{\text{eff}}(t)}$$

(4.12)

Usually the statistics of the signal does not allow to make light curves in many energy bins for several time intervals. Therefore, light curve are practically obtained in bins of time calculating the number of the excess events in each time bin above a certain Estimated Energy.

The time binning can be of order of days, hours or minutes according to the strength and the variability of the source. The Effective Collection Area is calculated separately in each time bin since it depends on the pointing position of the observation and thus, implicitly, on time.

The MARS program which allows the determination of the spectrum of a given source and its light curve is called FLUXLC.

4.3.10 Upper limit calculation

In case no significant signal is found for a given observation, an upper limit on the emission flux of the source can be calculated. In the following, $E$ represents the Estimated Energy, whereas $E_{\text{true}}$ the true energy (which is given only for the MC-$\gamma$, from which the Effective Collection Area is derived).

The number of excess events in a particular estimated energy bin $\Delta E$, after the analysis cuts, is extracted from the corresponding Alphaplot (or $\theta^2$-plot) and is given by

$$N_{\Delta E} = t_{\text{eff}} \cdot \int_{0}^{\infty} \phi(E_{\text{true}})A_{\text{eff}}(E_{\text{true}} | \Delta E)dE_{\text{true}}$$

(4.13)

where $\phi(E_{\text{true}})$ is the supposed differential flux of the source $[\text{ph cm}^{-2} \text{s}^{-1} \text{TeV}^{-1}]$, $A_{\text{eff}}(E_{\text{true}} | \Delta E)$ is the Effective Collection Area after all cuts, including the estimated energy dependent cut in $\Delta E$, and $t_{\text{eff}}$ is the total effective time of the observation. It is important to note that the integral is performed in true energy $dE_{\text{true}}$ and that the differential flux as well as the Effective Collection Area depends on the true energy, while the cuts for the selection of
the excess events $N_{\Delta E}$ and for the Effective Collection Area $A_{eff}(E_{\text{true}} \mid \Delta E)$ depend on the estimated energy.

The method used in MAGIC for deriving the upper limits on $N_{\Delta E}$ (i.e. $N_{\Delta E}^{UL}$) is based on the prescriptions of Rolke et al. [144, 145], which computes confidence intervals for the number of excess events assuming a Poisson distribution of ON events, in the presence of background approximated as gaussian. The method includes nuisance components in the used statistic, like for example the uncertainty on the background and on the efficiency. It is usually assumed a 30% of systematic uncertainty and a confidence interval width of 95%.

If it is assumed that the $\gamma$-ray spectrum is a general function of the energy, the differential flux can be written as

$$dF/dE_{\text{true}} = \phi(E_{\text{true}}) = K \cdot S(E_{\text{true}})$$ (4.14)

where K is a normalization factor [ph cm$^{-2}$ s$^{-1}$ TeV$^{-1}$]. From equation 4.13, and once the number of excess events $N_{\Delta E}^{UL}$ is computed, this constant can be re-expressed as:

$$K_{UL}^{N} = \frac{N_{\Delta E}^{UL}}{t_{eff}} \frac{1}{\int_{0}^{\infty} A_{eff}(E_{\text{true}} \mid \Delta E)S(E_{\text{true}})dE_{\text{true}}}$$ (4.15)

The integral flux upper limit above a certain estimated energy $E_{0}$ is therefore calculated as

$$\Phi_{UL}^{E}(E > E_{0}) = K_{UL}^{N} \int_{E_{0}}^{\infty} S(E)dE \quad [\text{ph cm}^{-2} \text{s}^{-1}]$$ (4.16)

where $S(E)$ is expressed in estimated energy and is derived by smearing $S(E_{\text{true}})$ with the energy resolution (by the aid of the energy Migration Matrix) and $\Delta E$ of equation 4.15 is to be considered as the range of estimated energies greater than $E_{0}$.

It is possible to derive differential upper limits by assuming that locally, between the estimated energies $E_{1}$ and $E_{2}$, $S(E_{\text{true}})$ follows a power law ($E_{\text{true}}/E_{0}^{\text{true}})^{\Gamma}$, where $\Gamma$ is the spectral index and $E_{0}^{\text{true}}$ is the pivot energy (e.g. 100 GeV). The mean $\gamma$-ray energy $E^{*}_{\text{true}}$ between $E_{1}$ and $E_{2}$ is given by

$$E^{*}_{\text{true}} = \frac{\int_{0}^{\infty} E_{\text{true}} S(E_{\text{true}})A_{eff}(E_{\text{true}} \mid E_{1}, E_{2})dE_{\text{true}}}{\int_{0}^{\infty} S(E_{\text{true}})A_{eff}(E_{\text{true}} \mid E_{1}, E_{2})dE_{\text{true}}}$$ (4.17)

where $A_{eff}(E_{\text{true}} \mid E_{1}, E_{2})$ represents the Effective Collection Area after the estimated energy dependent cuts between $E_{1}$ and $E_{2}$. The differential upper limit can thus be expressed (in units of [ph cm$^{-2}$ s$^{-1}$ TeV$^{-1}$]) as

$$\phi_{UL}^{E}(E_{1} < E < E_{2}) = \frac{N_{E_{1}, E_{2}}^{UL}}{t_{eff}} \frac{1}{\int_{0}^{\infty} A_{eff}(E_{\text{true}} \mid E_{1}, E_{2})(E_{\text{true}}/E^{*}_{\text{true}})^{\Gamma}dE_{\text{true}}}$$ (4.18)

The calculation of the flux upper limits is performed by the MARS program FLUXLC and is done, in practice, in bins of estimated energy (set by the analyzer) where the analysis cuts have been optimized. Therefore, in each estimated energy bin the Alphaplot (or $\theta^{2}$-plot) is computed and from it the number of the ON and OFF entries in the corresponding signal region is derived. Then, these quantities are summed in order to get the final ON and OFF events in the energy range where to compute the (differential or integral) upper limit. From the overall quantities the number of excess events and its upper limit are derived in the energy range under investigation and the upper limit on the flux is calculated.
4.3.11 Unfolding the spectrum

Since the detector has a finite energy resolution (which depends on the energy itself, i.e. it gets worse for the low energy range) and since the true energy is not directly measured, the measurements of the energy of the $\gamma$-like event extracted from the data of a given source are systematically distorted, particularly at the energies close to the threshold. The distortions due to biases and finite resolution can be written in the form:

$$Y(y) = \int M(y, x) S(x) dx$$

or

$$Y_i = \sum_j M_{ij} S_j$$

or

$$Y = M \times S$$ (4.19)

where $y$ is the estimated energy, $x$ is the true energy, $M$ describes the detector response (i.e. the Migration Matrix), $Y$ is the measured distribution and $S$ is the true undistorted distribution. The aim is to determine $S$, given $Y$ and $M$.

There are various approaches to solve this problem. The simplest solutions (called deconvolution) is to invert the matrix $M$ and it is technically correct. However, this is often useless due to large correlations between adjacent bins, which imply large fluctuations of their contents. Another option is to use an unfolding with regularization:

$$\chi^2 = \frac{w}{2} \times \chi_0^2 + \text{Reg}(S)$$ (4.20)

where $w$ is the regularization parameter, steering the strength of the regularization, $\chi_0^2$ expresses the degree of agreement between the prediction $M \times S$ and the measurement $Y$, and $\text{Reg}(S)$ is a measure of the smoothness of $S$. A solution for $S$ is obtained by minimizing expression 4.20 for a fixed regularization parameter $w$. In the MAGIC software, a variety of methods is available, such as the Tikhonov [146], the Bertero [147] and the Schmelling [148] methods, which differ in the way the regularization is implemented. In chapter 7 the unfolding of the spectrum of the MAGIC-I Crab Nebula data sample used to optimized the analysis of Segue 1 data will be shown.

The various unfolding methods are used for each observation, and the consistency of the results is checked. Only when results of different unfolding methods agree, the result of the unfolding is considered trustworthy. More detail about the unfolding procedures used in the MAGIC analysis can be found in [149,150].

4.3.12 Sensitivity

The integral flux sensitivity of an IACT is the minimum flux of $\gamma$-ray events (per unit time and area) which, in a given observation time, results in a statistically significant excess above the isotropic background of cosmic-ray initiated showers above a certain energy $E_0$. The flux is integral in the sense that it refers to all the $\gamma$-rays above a given energy $E_0$. When comparing different instruments, it is most often assumed that the source is point-like, and that its energy spectrum is a pure power law of differential index -2.6 (which is the spectral index for Crab around 1 TeV).

A common sensitivity unit for different IACTs is the flux which will be measured with significance of 5$\sigma$ in 50 hours of observation. In the estimation of the sensitivity the significance is not generally computed from equation 4.7 but, instead, by requiring that observed excess events are five times the RMS of the background ($N_{\text{exc}} = 5 \times \sqrt{N_{\text{bg}}}$). This definition does not take into account fluctuations of the uncertainty of the mean number of the background and therefore gives an optimistic estimation of the sensitivity. The sensitivity can be derived by using MC-$\gamma$ events [100] or directly from an observation of
a standard candle, like the Crab Nebula. When the sensitivity is computed from a real Crab Nebula observation, the analysis cuts are optimized above the energy \( E_0 \) in order to get the best possible significance, i.e. the highest achievable \( \frac{N_{\text{exc}}}{\sqrt{N_{\text{bg}}}} \), with the further constraint to get at least 10 \( \gamma \)-ray events in 50 hours of observation. Generally, the cuts which guarantee the best significance are very tight and thus not suitable for the spectrum calculations. If the time of the observation is \( T_{\text{obs}} \), it follows that the integral flux sensitivity in terms of the Crab nebula flux is:

\[
\Phi^{\min}_{E > E_0} = 5 \times \frac{\sqrt{N_{\text{bg}}}}{N_{\text{exc}}} \times \sqrt{\frac{T_{\text{obs}}}{50 \text{ [h]}}} \times \Phi^{\text{Crab}}_{E > E_0} \tag{4.21}
\]

In order to express the integral sensitivity in percentage units of the Crab Nebula flux (C.U.), equation 4.21 can be rewritten as

\[
\text{Sensitivity (50h - 5\sigma) [C.U.]} = 100 \times 5 \times \frac{\sqrt{N_{\text{bg}}}}{N_{\text{exc}}} \times \sqrt{\frac{T_{\text{obs}}}{50 \text{ [h]}}} \tag{4.22}
\]

The best integral sensitivity is not reached for \( E_0 \) close to the absolute energy threshold (\( \sim 100 \text{ GeV} \)) of the telescope, since the background discrimination degrades rather fast for decreasing energy, and therefore, on reducing \( E_0 \), the amount of background after all cuts grows much faster than the signal. For higher \( E_0 \) there is a competing effect, which is the fast reduction of the event statistics (both of signal and background), and therefore the best discovery performance of the telescope is expected to be found at intermediate energies. Currently, the best integral sensitivity of the MAGIC-I telescope is about 1.6% of the Crab Nebula flux for \( E_0 > 280 \text{ GeV} \) (and \( \text{SIZE > 400 phes} \)) and for source dependent analysis cuts [137].

In case the excess events and the background events are extracted in bins of estimated energy, a differential sensitivity can be derived.

### 4.3.13 Systematic effects

Apart from statistical errors, the determination of the \( \gamma \)-ray spectral features of the observed sources are affected by large systematic uncertainties which are often difficult to evaluate. One of the main problems is related to the fact that a calibration of the instrument is not possible due to the lack of a natural or artificial calibrators of VHE \( \gamma \)-rays. The ground-based \( \gamma \)-ray experiments therefore have to rely on MC simulations which are particular susceptible to systematic uncertainties. Other important contributions to the systematic errors are related to uncertainties of the hardware components of the detectors.

The most important systematic uncertainties for MAGIC-I are:

- **Atmospheric model:**
  The MC data are simulated using the US standard atmosphere to compute the Rayleigh scattering as well as the Mie scattering losses using the Elterman model [120] for the distribution of aerosols and ozone. This model represent an ideal atmosphere which does not take into account atmospherical changes as humidity, temperature, high clouds, haze and calima (a fine dust blown from the Sahara Desert). These effects may lead to an underestimation of the energy of the showers. Moreover, the model is not perfectly reproducing the atmosphere at the latitude of La Palma, the effect being estimated to be of the order of 10%. In addition, the average atmospheric values between Summer and Winter change the atmospheric transmission by about 15%.
• **Hardware components of the detector:**
  The main systematic errors due to hardware components are related to:
  
  - the effective reflectivity of the mirrors and the Point Spread Function (PSF). The reflectivity is assumed to be 85% and is monitored by reflectivity measurements [85]. It can be affected by dust which can deposit on the mirrors and also by the aging. The quality of the optical PSF mainly depends on the accuracy of the Active Mirror Control. The uncertainties are estimated to be of the order of 5% for both contributions and affect the energy estimation of the γ-ray showers.
  
  - photon losses on the camera entrance. The protective plexiglass in front of the PMT camera and the Winston cone can yield less transmission due to dust pollution. The effect is estimated to be 3-5%.
  
  - conversion of photons to measurable photoelectrons. The effect introduces a systematic error of 5-10% and is dominated to the uncertainty of the light collection efficiency of the first dynode of the PMTs. Further contribution of about 2% comes from the equalization of the PMT gains. These errors affect the energy estimation.
  
  - trigger inefficiencies, which are most pronounced for primary γ-rays below 150 GeV (effect of the order of 20%) while the effect is much smaller at higher energies (5%). The trigger inefficiency affect both the flux and the energy estimation.
  
  - non-linearity in the PMT gains, in the amplifiers, in the optical signal transmission and in the FADC system lead to an overall uncertainty of about 10%.

• **Analysis chain:**
  The error related to the calculation of the effective observation time of a given observation is of the order of 2% and is caused by dead times of the system which are not considered in the analysis. This error affects the flux and is independent on energy. Different methods for the signal extraction, the image cleaning, the γ-hadron separation and the unfolding lead to a systematic error of 3% [126]. The calibration chain using the F-factor method introduces an overall systematic error of 8% [89]. This error affects the energy estimation of the events.

• **MC simulation:**
  Uncertainties introduced by the overall simulation chain, from the production of the showers to the simulation of the different components of the detectors, lead to an overall systematic error estimated to be 10%. This error affects the determination of the Effective Collection Area and thus the absolute energy scale and the flux level.

The overall systematic errors have been estimated by considering separately the different contributions and then by adding these individual errors in quadrature. The overall systematic uncertainty in the energy scale is estimated to be about 30%, assuming a spectral index of $\Gamma = -2.6$. For steeper slopes the uncertainties become higher. The systematic uncertainty on the spectral index is conservatively estimated to be $\Delta \Gamma_{\text{sys}} = \pm 0.2$. More detailed discussions of the systematic errors affecting the MAGIC-I telescope can be found in [42,126].

### 4.4 MAGIC stereo analysis

Since December 2009, the default observation mode for the MAGIC telescopes is the so-called full stereo: the Level 1 Trigger of MAGIC-I and MAGIC-II are set with a 3NN logic (see section 3.2.5) and the events are recorded only when Level 3 Trigger identifies the coincidence.
The typical rate of this configuration is about 200 Hz. Concerning the data taking mode, the two telescopes can observe the sources in ON-OFF mode or in WOBBLE mode (see section 4.2.1).

As already mentioned, the stereoscopic observation mode allows a more precise reconstruction of the shower parameters as well as a more efficient suppression of the hadronic showers and other background events (as single muon events). However, the global trigger logic can be also set in order to record the events independently for each telescope. Therefore, the two MAGIC telescopes can as well operate separately. The independent operations are used for particular purposes and are useful to monitor the individual performance of the two telescopes.

When the data of the MAGIC-II telescope are analyzed independently, the same analysis steps described in section 4.3 are applied, with specific modifications implemented in MARS which take into account the different hardware components of the two telescopes, such as the cameras and the trigger and readout systems. The main modifications have been introduced for the MERPP and CALLISTO programs. In particular special calibration runs for the MAGIC-II readout system, called domino calibration runs [110], are used during the calibration of the MAGIC-II data in order to account for the high non-linearity of the current DRS-2 chips used in the FADC system (see section 3.3.3). For the MAGIC-II data calibration, the signal is not extracted with the spline method (as in MAGIC-I) but using a sliding window algorithm [89] with a pedestal estimation from the beginning of the pulse. This is basically due to the different FADC system of the second telescope. The same image cleaning algorithm (Time-constrained cleaning) is used for the showers recorded by the second telescope: up to now, the charge levels of the cleaning are set to $Q_{\text{core}} = 10$ phes and $Q_{\text{boundary}} = 5$ phes \[12\], whereas the time constrain values are the same as those used for MAGIC-I, namely 4.5 ns for the core pixels and 1.5 ns for the boundary pixels. In Fig. 4.11, an example of a same shower recorded by both telescopes is shown.

The MC-\(\gamma\) simulation chain for the stereo production basically follows the same steps described in section 4.2.3. The program REFLECTOR and CAMERA have been updated in order to simulate the telescopes’ configuration as well as to the new hardware components of the MAGIC-II telescope. At the end of the simulation two set of MC-\(\gamma\) data are available, one for each telescope, and are analyzed as the real data. It is worth noting that in the standard

\[12\] These two charge level values have been chosen after dedicated MC studies.
MC-\(\gamma\) production for the stereoscopic system, the showers are simulated with all the Azimuth values (between 0° and 360°), whereas in the standard MAGIC-I MC-\(\gamma\) production only two values of Azimuth were simulated (at 0° and 90°). The extension of the simulated showers to all the Azimuth range was done in order to better take into account the Geomagnetic effect on the shower developments, which mainly affects the low energy showers, and also to take into account the fixed configuration geometry of the stereoscopic system, which breaks the azimuthal symmetry the single telescope observation has (see chapter 5). For more detail about the MC production for the stereoscopic system see [151].

Up to the image parametrization calculation, the analysis chain runs over the data of each telescope separately. At this point, two sets of files, one per telescope, which contain two different views of the same cleaned showers are available\(^\text{13}\). The stereo analysis then proceeds with the aid of a new MARS program called SUPERSTAR whose aims are to identify the matching pairs of images belonging to the same event from the two data streams, based on the Level 3 Trigger numbering, to update the names of the image parameters of each telescope according to the telescope number (e.g. SIZE1, SIZE2, WIDTH1, WIDTH2, etc.) and to calculate further image parameters which refer to the stereoscopic view of the showers (see Fig. 4.12), such as the primary incoming direction (see below), the ground impact point with respect to the two telescopes (IMPACT1 and IMPACT2) and the height of the shower maximum (MAXHEIGHT). The output files of the SUPERSTAR program contains the information of each single telescope parameters and the new calculated stereo parameters.

\begin{figure}[h]
\centering
\includegraphics[width=\textwidth]{figure412.png}
\caption{Sketch of the stereoscopic technique with two telescopes. The single image parametrizations are geometrically combined to give the incoming direction of the shower and core distance to each telescope.}
\end{figure}

\(^{13}\text{In an IACT array, the image parameters of multiple images of the same shower are generally combined in order to reduce the number of parameters. However, the MAGIC Collaboration preferred to keep the parameters separated as they are only two and the two cameras have different designs.}\)
4.4. MAGIC stereo analysis

4.4.1 Arrival Direction estimation

In the single telescope mode, the shower direction is estimated through the Disp method (see section 4.3.8). With two telescopes, the primary direction (and hence the location of the reconstructed source on the cameras) and the ground impact point can be reconstructed by stereoscopy with the major axis of the 2 images: the direction is determined from the intersection of two planes, each of them defined by one of the recorded images, plus the position and orientation of the corresponding telescope. In case of an array of two telescopes, like MAGIC, there is only one solution for the geometry of the shower axis, and its accuracy depends on the relative positions of the telescopes and the shower: the more parallel the two images on the camera planes are, the larger the uncertainties in the reconstructed parameters. This method provides good results for angle $\Delta \delta$ greater than $\sim 30^\circ$ (see upper part of Fig. 4.13(a)) and it is independent on the MC simulations. In order to improve the shower parameter reconstruction, the single telescope Disp method, calculated with the RF algorithm using MC-$\gamma$ samples (one for each telescope) [142] (see lower part of Fig. 4.13(a)), and the stereo reconstruction based on the intersection of the major axes (stereo reconstruction) can be combined. The sketch of the combined reconstruction method (called stereo-Disp method) is shown in Fig. 4.13(b). The values of the different quantities reported in the sketch have been optimized on MC-$\gamma$ simulations and real Crab Nebula stereo data. For angles $\Delta \delta$ smaller than $15^\circ$ the stereo reconstruction is not used and the two single telescope Disp determinations are considered in order to estimate the direction of the showers or to reject the events. For angles $\Delta \delta$ wider than $15^\circ$, the stereo reconstruction is considered together with the two single telescope Disp determinations and a set of decisional constraints is applied in order to estimate the direction of the events or to reject them.

The arrival direction is related to the angular resolution. Its definition, here, is given by the radius of the circle centered on the simulated source containing 68% of the reconstructed events. In Fig.4.14 it is reported a comparison between the angular resolutions achieved on MC-$\gamma$ studies for the single telescope (with the Disp RF method) and for the stereo reconstruction and the stereo-Disp reconstruction. As shown in figure, the stereo-Disp reconstruction method provide a $\sim 30\%$ improvement of the angular resolution.

Once the shower axis is determined, an estimate of the height of the shower maximum is defined as the angle at which the image center of gravity is viewed from each telescope.

4.4.2 Energy estimation

As of now, the reconstruction of the energy of the stereo events is not performed by using the RF method as in case of MAGIC-I data. Instead, for each telescope the energy is reconstructed using lookup-tables based on the image size, impact parameter, the height of the shower maximum and the Zenith angle. Combination of these two energy estimations provides the final reconstructed energy of the event as well as a parameter describing the compatibility between these two energies. The lookup-tables are computed from a train sample of MC-$\gamma$ and then are applied to the real data, as well as to an independent MC-$\gamma$ sample, by the MARS program SUPERSTAR. Although the lookup-table method is simpler than the RF method, it provides a better energy resolution with respect to the typical one achieved for the MAGIC-I using a source independent estimation with the RF method. This is due to the 3D-parameters of the shower (i.e. the impact parameter and the height of the shower maximum) which are well reconstructed by stereoscopy. In Fig. 4.15 the energy resolution obtained from the lookup-table method applied to stereo MC-$\gamma$ events is compared to the energy resolution obtained from the RF method applied to MAGIC-I MC-$\gamma$ events. Note that the two curves are obtained without any assumption on the source position, i.e. a source independent analysis.
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(a) Direction reconstruction of the showers recorded by the two telescopes.

(b) Sketch of the combined method using stereo reconstruction and single telescopes reconstructions using the Disp method.

Figure 4.13: Sketches of the current method used to estimate the direction of the showers recorded by both telescopes. The method combines the information of the stereo reconstruction (major axes crossing point) and of the single telescopes reconstructions, obtained by the Disp method (with the RF algorithm), through a series of decisional constraints.

was applied in both cases. The improvement is particularly evident for the lower energies. Spectral analysis performance of extended or poorly localized sources will strongly improve
4.4. MAGIC stereo analysis

**Figure 4.14:** Angular resolution as a function of the energy for the MAGIC telescopes with the stereo-Disp method (red points) and the axis-crossing method (black squares). The single MAGIC-I telescope angular resolution obtained with the Disp method (with the RF algorithm) is also shown (blue triangle). The improvement of the angular resolution with the stereo analysis is achieved in all the energy range. In particular, the stereo-Disp method provides a $\sim$30\% improvement, especially at the lowest energies. Courtesy of P. Colin.

with the stereo observations. The energy reconstruction for the stereo analysis using the

**Figure 4.15:** Energy resolution as a function of the estimated energy for the stereo analysis using lookup-tables (black points) and for the stand alone MAGIC-I using the RF method with source independent parameters (grey squares). Image taken from [152].

MARS program OSTERIA with the RF method is currently under development and should improve the energy resolution even more.
4.4.3 $\gamma$-hadron separation

The $\gamma$-hadron separation in the stereo analysis is based on the RF method, as in case of MAGIC-I data. Up to now, the algorithm is performed by an independent macro but the update of the OSTERIA program in order to embed the calculation for stereo data is ongoing. Samples of SUPERSTAR real data (containing hadrons) and MC-$\gamma$ data are used as input for the training of the RF algorithm. The current default input parameters are SIZE1, SIZE2, WIDTH1, WIDTH2, LENGTH1, LENGTH2, IMPACT1, IMPACT2, TIME GRADIENT1, TIME GRADIENT2 and MAXHEIGHT, which are all source independent parameters. In Fig. 4.16 the typical Gini plot for the stereo $\gamma$-hadron separation algorithm is shown. The most discriminant parameters are the stereo parameter MAXHEIGHT, and the single telescope parameters WIDTH1 and WIDTH2. The MELIBEA program then applies the $\gamma$-hadron separation matrices to the real stereo SUPERSTAR data as well as to an independent stereo SUPERSTAR MC-$\gamma$ sample in order to add the Hadronness parameter to the data.

![Gini plot](image)

**Figure 4.16:** Typical Gini plot for the $\gamma$-hadron separation of the stereo analysis. The most discriminant parameters are the stereo parameter MAXHEIGHT and the single telescope parameter WIDTH1 and WIDTH2.

4.4.4 Spectrum, light curve calculation and unfolding

The FLUXLC and UNFOLDING programs have been updated in order to allow the determination of the spectrum and the light curve of the sources observed in stereoscopic mode. In particular the programs have been changed in order to recognize the stereo parameters and the new names of the single telescope parameter, which now are tagged with the number of the corresponding telescope. The basic calculations are the same as those described in section 4.3.9. An important difference is that for stereo analysis the $|\text{ALPHA}|$ approach is no longer valid and only the $\theta^2$ approach is used to infer the excesses. Another important difference is how the Effective Collection Area is calculated (see chapter 5).

In Fig. 4.17 the unfolded spectrum (with the Bertero method [147]) of the Crab Nebula observed in stereo mode between 50 GeV and 8 TeV is shown. The observation was carried out
in November 2009 (13-14-15/11/2009) for a total observation time of roughly 190 min. The data are in good agreement, within the systematic errors, with the results achieved with the MAGIC-I telescope [42]. The parametrization assuming a curved power low function is (in units $[\text{TeV}^{-1} \text{cm}^{-2} \text{s}^{-1}]$)

$$dF/dE = (5.2 \pm 0.2_{\text{stat}}) \times 10^{-10} \times (E/300\text{GeV})^{-2.32 \pm 0.04_{\text{stat}}} + (-0.27 \pm 0.09_{\text{stat}}) \log(E_{300\text{GeV}})$$ (4.23)

**Figure 4.17:** Unfolded spectrum of the Crab Nebula observed in stereo mode in November 2009 (13-14-15/11/2009). The effective observation time is ~190 min. The red, green and yellow dashed line are the parametrizations published respectively by the MAGIC Collaboration [42], by the HEGRA Collaboration [153] and by the H.E.S.S. Collaboration [154]. The blue solid line represents the curved power law fit of the data, which is in good agreement with that published in [42]. Image taken from the official MAGIC released plots web page http://wwwmagic.mpp.mpg.de/documents/released.

### 4.4.5 Current integral sensitivity of the MAGIC telescopes

The improvement for stereo observations of the $\gamma$/hadron separation, of the angular resolution and of the energy reconstruction compared to the observation carried out with the single telescope improves significantly the sensitivity and spectral analysis performance. In Fig. 4.18 the preliminary integral sensitivity curve (see section 4.3.12) achieved from real Crab Nebula data taken in WOBBLE mode during the end of 2009 is shown. The improvement with respect
to the integral sensitivity achieved with MAGIC-I is a factor \( \simeq 1.5-2 \) between \( \sim 200 \) GeV and \( \sim 2 \) TeV, whereas for low energies the sensitivity enhancement is even larger, due to a more precise reconstruction of the shower parameters as well as a better suppression of the hadronic showers and other background events (like single muon events). Up to know, the MAGIC telescopes are the most sensitive world-wide ground-based instrument for \( \gamma \)-ray astronomy between \( \sim 50 \) GeV and \( \sim 150 \) GeV. Above few TeV, the MAGIC performance suffers the relatively small trigger region of the MAGIC-I camera (which is foreseen to be upgraded in summer 2011).

A great effort is ongoing inside the MAGIC Collaboration in order to further improve the sensitivity of the MAGIC telescopes by means of a better understanding of the new telescope hardware components and, at the same time, of the development of new high-performance analysis tools. With the upgrade of the PMT camera and the readout system of MAGIC-I, the sensitivity of the MAGIC telescopes will further improve. A paper dedicated to the performance and the evaluation of the systematic errors of the new stereoscopic system is in preparation.

![Figure 4.18: Preliminary integral sensitivity of the MAGIC telescopes compared with the MC expectations and the MAGIC-I achieved sensitivity. The best sensitivity of the stereoscopic system is about 1% of the Crab Nebula flux. Courtesy of E. Carmona.](image-url)
Azimuth implementation in the Effective Collection Area calculation

One of the main estimated quantities needed for the calculation of a spectrum (or its flux upper limit) is the Effective Collection Area for primary $\gamma$-rays. For a single Imaging Atmospheric Cherenkov Telescope (IACT), besides the features and performance of the detector, the value of the Effective Collection Area mostly depends on the energy of the primary $\gamma$-rays and on the Zenith angle of the observation. A possible Azimuth angle dependence due to the Geomagnetic field (GF) should also be considered, since the GF affects the development of the Extensive Atmospheric Showers (EAS) generated by primary $\gamma$-rays and thus of the recorded images, as already shown in previous studies for MAGIC-I [155–157]. In case of observations with an array of two IACTs, like the MAGIC telescopes, a further Alt-Azimuth effect is expected, in particular for high ranges of Zenith angle. In fact, unlike the single telescope observations, the fixed position of the two MAGIC telescopes breaks for the stereoscopic observations the symmetry under Azimuth rotations.

In this chapter, the main reasons for taking care of the Azimuth dependence of the Effective Collection Area, in light of the MAGIC Phase-II, and the main changes implemented in the MAGIC analysis software (MARS) in order to account for this new dependence are described. Results of some studies on stereo Monte Carlo $\gamma$-rays (MC-$\gamma$) and of some tests performed on a sample of Crab Nebula MAGIC-I data are also reported.

5.1 Introduction

A basic quantity needed for the computation of the $\gamma$-ray flux (or its upper limit) for a given source observed by IACTs, once the number of the excesses and the effective observation time are extracted from the data, is the Effective Collection Area of the telescopes for the $\gamma$-ray detection related to that observation. This quantity represents the estimation of the effective area within which the telescopes can detect primary $\gamma$-rays coming from a certain source during its observation. We define it as:

$$A_{eff} = \frac{N_\gamma}{\Phi T}$$ (5.1)
where \( N \) is the number of detected \( \gamma \)-ray events, \( \Phi \) [ph cm\(^{-2}\) s\(^{-1}\)] is the \( \gamma \)-ray flux (in a certain energy range) from a given source observed for a period of time \( T \). The Effective Collection Area is basically related to the size of the Cherenkov light pool (see section 2.2.2) on the ground and to the efficiency of the detector, reaching a maximum value, for the MAGIC telescopes’ observations, of the order of \( 10^5 \) m\(^2\) at low Zenith angles.

Since no real pure \( \gamma \)-ray beams are available for calibrating the ground-based \( \gamma \)-ray detectors, the Effective Collection Area must be derived by means of dedicated MC-\( \gamma \) simulations: for given energy \( E' \), Zenith angle (ZA) \( \theta' \) and Azimuth angle (AZ) \( \phi' \), the Effective Collection Area can be written as

\[
A_{\text{eff}}(E', \theta', \phi') = \int_{0}^{IP_{\text{max}}} \varepsilon(E', \theta', \phi', r) \, rdr
\]  

where the variable \( r \) represents the impact parameter (IP) used for the simulation of the primary \( \gamma \)-rays and

\[
\varepsilon(E', \theta', \phi', r') = \varepsilon_{\text{trigger}}(E', \theta', \phi', r') \cdot \varepsilon_{\text{IC}}(E', \theta', \phi', r') \cdot \varepsilon_{\text{cuts}}(E', \theta', \phi', r')
\]

\[
= \frac{N_{\text{MC surviving}}(E', \theta', \phi', r')}{N_{\text{MC simulated}}(E', \theta', \phi', r')}
\]  

represents the efficiency for the MC-\( \gamma \) events simulated with the given values \( E', \theta', \phi' \) and \( r' \), passing the trigger conditions and surviving the image cleaning (IC) and the further applied analysis cuts (see the sketch in Fig. 5.1). The global \( \gamma \)-ray detector efficiency depends on the features and on the performance of the telescopes such as the reflector collection surfaces, the efficiency of the PMTs in collecting the Cherenkov light, the extension and the design of the trigger region of the cameras, the data transmission, the readout system and the overall analysis chain. Finally, in case of stereo observations, it depends also on the configuration of the array. The \( \gamma \)-ray detector efficiency increases quickly with the energy of the primary \( \gamma \)-rays until it starts to reach a plateau for the higher energies.

\[\text{Figure 5.1: Sketch showing the main quantities involved in the definition of the differential efficiency for } \gamma \text{-rays characterized by the set of values } E', \theta', \phi' \text{ and } r', \text{ as written in equation 5.3.}\]

In order to perform the simulation for the Effective Collection Area for different incoming direction of the primary \( \gamma \)-rays the so-called geometric area must be considered. The geometric
area is defined as the area, on the level of the detector, perpendicular to the shower axis, which is illuminated by the Cherenkov light coming from an electromagnetic atmospheric shower. It depends on the physics of the shower development in the atmosphere, on the multiple scattering of the charged particles emitting the Cherenkov light and on the incoming direction of the $\gamma$-ray shower (in particular the Zenith angle). A good expression for the geometric area (where the second part of the expression is valid in the flat Earth approximation, i.e. up to $ZA \approx 70^\circ$) can be written as

$$A_{\text{geom}} = \pi \cdot [D(h(\theta), \theta) \cdot \tan(\alpha)]^2 \approx \pi \cdot [(h(\theta) - h_{\text{obs}}) \cdot \tan(\alpha)]^2 \cdot \cos^{-2}(\theta)$$  (5.4)

where $\theta$ is the ZA of the observation, $h(\theta)$ is the height of the maximum of the shower, $h_{\text{obs}}$ is the height above the sea level (a.s.l.) of the detector, $D(h, \theta)$ is the distance between the detector and the maximum of the shower and $\alpha$ is the typical Cherenkov emission angle ($\sim 1^\circ$) (see the sketch shown in Fig. 5.2). Typical values of the geometric area for low ZA observations are of the order of $10^5$ m$^2$ at the MAGIC site ($\sim 2230$ m a.s.l.), corresponding to $D(h, \theta) \approx 10$ km.

Indeed, as shown in Fig. 5.3, for vertical showers induced by primary $\gamma$-rays, the Cherenkov light distributes roughly uniformly at the ground in an area with radius of the order of 120 m (at the MAGIC site), which defines the so-called plateau region, with a fast decreasing light density after the plateau border in the so-called halo region, leading to an overall value of the geometric area of the order mentioned above. All the primary $\gamma$-rays passing through this area, emit Cherenkov light which is potentially detectable by a single telescope. In the figure is also shown that even if the Cherenkov light density is dependent on the energy of the primary $\gamma$-ray ($\propto E$), the radius which defines the plateau region (which is related to the value of the geometric area), can be considered in first approximation as independent on the energy and, thus, the geometric area dependences expressed in equation 5.4 can be considered valid, in first approximation, for a large range of energies.

Another important quantity affecting the Effective Collection Area is represented by the energy threshold ($E_{\text{th}}$) of the IACTs for $\gamma$-ray showers. The $E_{\text{th}}$ is usually defined as the
peak value of the γ-ray energy distribution of the triggered events, although there are different ways to define it. For instance, it can be also defined as the energy peak of the events after the image cleaning (and the application of some quality cuts). The $E_{th}$ (which is spectrum dependent) has a strong dependence with respect to the ZA of the observation, since it is correlated to the atmospheric depth the Cherenkov light must pass through to reach the detector. For a given primary γ-ray energy and for observations under increasing ZA, the light concentration decreases, due to the enhancement of the geometric area ($\cos^2(\theta)$) and of the atmospheric depth (and thus of the light absorption factor due to the Rayleigh and Mie scatterings) and therefore less light is collected by the reflector surfaces. All these factors lead to the rise of the $E_{th}$ of the detector for observation with increasing ZA. Empirically it is found that the $E_{th}$ scales with the ZA as

$$E_{th}(\theta) \approx E_{th}(\theta = 0^\circ) \cdot \cos^{-2.7}(\theta)$$ (5.5)

The main source of uncertainties for the Effective Collection Area values are related to all possible mismatches between the MC-γ simulations and the real shower developments in atmosphere. The most problematic factor is the intrinsic transparency of the atmosphere which depends on the location on the Earth and also on the season. Moreover, the clouds in the troposphere and some weather phenomena, like the calima (a fine dust blown from the Sahara Desert), can also modulate the atmospheric transparency within a time range of the order of the day, or even less (see section 4.3.13).

Once the telescopes’ location, configuration and hardware performance are fixed in the MC-γ simulation, the Effective Collection Area for a given analysis chain can be studied as a function of few main parameters, namely the energy of the primary γ-rays and their incoming direction in the sky, i.e. the ZA and AZ of the observation. In case of MAGIC-I, even if previous dedicated studies (see [155–157]) have shown that the Geomagnetic field (GF), whose strength module (proportional to the Lorentz force) has a clear Alt-Azimuth dependence, can significantly affect the orientation and the shape of the recorded MC-γ images and their
trigger efficiency, the AZ dependence of the Effective Collection Area has not explicitly been taken into account. In order to deal somehow with it, the MC-$\gamma$ events used for the analysis of the sources observed by MAGIC-I have been always simulated with two different AZ values: 0° and 90° (in the CORSIKA program [63] reference frame, see section 5.1.1). In this way, one can expect that an average fair account for the AZ dependence can be achieved. Obviously, this is a first order solution which indeed seems to work properly above the $E_{th}$ of the telescope, particularly at the lower ZA ($< 30^\circ$).

On the other hand, in light of MAGIC Phase-II, we need to start considering the AZ dependence of the Effective Collection Area since, in addition to the GF effect, the stereoscopic system introduces a new Alt-Azimuth dependence due to the fixed direction between the two telescopes which breaks the rotational symmetry which characterizes the single telescope configuration: the intersection of the fields of view of the two telescopes changes for each pointing direction, for example it achieves its maximum along the direction which joins the two telescopes. Therefore, for a fixed range of ZA, different AZ pointing positions can lead to different Effective Collection Area values, besides the possible contribution due to the GF effect. This issue should be particularly important for the higher values of ZA. Moreover, this effect combines in a nontrivial way with the GF one, since the fixed direction between the telescopes is not along the direction of the GF (see Fig. 5.4). In light of these considerations, the account for the overall Alt-Azimuth dependence of the Effective Collection Area for the stereoscopic system becomes important and it has been systematically taken into account by upgrading the parts of the MAGIC software involved in the calculation of the Effective Collection Area.

![Figure 5.4: Birds-eye view of the MAGIC telescopes’ site. The fixed direction between the two telescopes forms an angle of about 38° with respect to the Geographic North direction and an angle of about 45° with respect to the Geomagnetic North direction, since the magnetic field lines at La Palma are tilted by ~7° westwards with respect to the meridian [158].](image)
5.1.1 The reference frame systems

Before describing the effects which can produce non-negligible AZ dependences of the Effective Collection Area, for given ZAs and energy ranges, it is worth introducing the possible reference frames which can be used.

There are basically two coordinate systems which can be considered. The first one is the CORSIKA program [63] reference frame and it is related to the MC-γ simulation (see Fig. 5.5(a)):

- the x-axis is aligned with the Geomagnetic North pole
- the y-axis points to the West

The AZ is counted counterclockwise from the positive x-axis and refers to the incoming direction of the primary γ-rays. The AZ values range in the interval [0, 2\pi] radians. However, we will always refer to these values in degrees.

The second coordinate system, instead, is the MAGIC telescopes’ drive reference frame and it is related to the real data taking:

- the x-axis is aligned with the Geographic North (which is pointing \(\sim 7°\) eastwards with respect to the Geomagnetic North)
- the y-axis points to the East

The AZ in the drive reference frame is counted clockwise from the positive x-axis and ranges in the interval \([-90°, 318°]\). The range is extended below AZ = 0° for technical reasons (see section 3.2.1). Indeed, the sources which culminate in the sky (at MAGIC site) in a position located in the shortest imaginary line projected in the sky joining the Polar star and the Zenith have AZ = 0° at the culmination. Since the sky rotates counterclockwise with respect to the celestial North pole, the AZ values corresponding to increasing ZAs, are positive. Therefore just after the culmination, the AZ values should be decreasing, starting from AZ = 360°. In order to avoid this AZ value discontinuity, the telescopes do not stop to track the source and thus, after the culmination, the AZ values are negative. On the contrary, for sources which culminate at AZ = 180° (i.e. in a position located in the shortest imaginary line joining the celestial South pole and the Zenith at the MAGIC site), such kind of AZ discontinuity does not occur and the AZ values are always positive (before and after the culmination).

In Fig. 5.5 the CORSIKA program reference frame and the conversion function between it and the drive reference frame are shown. For clarity, the AZ values are displayed within 0° and 360°: in case of AZ values between \(-90°\) and 0° in the drive reference frame, the situation is the same as for the \([270°, 360°]\) case. For this reason we decide to keep the range for the AZ values of the real data, during the new version of the Effective Collection Area calculation, to \([0°, 360°]\), which means that in case of negative AZ values, a factor equal to 360° is added\(^1\).

Throughout this chapter, both the coordinates systems introduced above will be used: when we will deal with stereo MC-γ studies we will use the CORSIKA program reference frame, whereas when we will introduce the main changes implemented in MARS and the performed tests on MAGIC-I data samples we will use the drive reference frame, since this last frame is used in MARS for the Effective Collection Area calculation.

\(^1\)This does not represent a problem for the mispointing correction, through the starguider information (see section 3.2.1), since it is applied before the Effective Collection Area calculation steps.
5.1. Introduction

(a) CORSIKA reference frame.

(b) Conversion function between CORSIKA and telescopes’ drive reference frames.

Figure 5.5: (a) Definition of the CORSIKA program coordinate system. \( \alpha \) denotes the angle between the direction of the EAS and the direction of the GF. The AZ refers to the momentum of the incoming \( \gamma \)-ray and is counted counterclockwise from the positive \( x \)-axis towards West. (b) Conversion function between the CORSIKA program reference frame and the MAGIC telescopes’ drive one. The \( x \)-axis of the drive reference frame is aligned with the Geographic North (which is pointing \( \approx 7^\circ \) eastwards with respect to the Geomagnetic North) and the \( y \)-axis points to the East. The AZ in the drive reference frame is counted clockwise from the positive \( x \)-axis.

5.1.2 The Geomagnetic field effect

During the development of the showers induced by primary \( \gamma \)-rays, the separation between positrons and electrons due to the component of the GF normal to the shower axis can affect the shape, the orientation and the concentration of the overall Cherenkov light flash event at the ground, and, hence, the capability to trigger the event (particularly at the lowest energies) and also to discriminate it against the hadronic-induced ones. Detailed studies on these issues can be found in [156,157] and particularly in [155].

The GF at the MAGIC telescopes’ site is described in the MC simulations according to the epoch 2005 International Geomagnetic Reference Field (IGRF) model [158]. The GF value is calculated at 10 km a.s.l. where the mirrors of the MAGIC telescopes are focused, i.e. likely the location of the shower maximum for 100 GeV \( \gamma \)-ray induced air shower at small ZAs.

Figure 5.6(a) shows the absolute value of the GF component \( |\vec{B}_\perp| \) normal to the direction of the EAS versus AZ and ZA (in the CORSIKA program reference frame) for the MAGIC telescopes’ site (N 28\(^\circ\)45\(^\prime\)43\(\prime\)s, W 17\(^\circ\)53\(\prime\)24\(\prime\)s, \( \approx \)2230 m a.s.l.). In the MAGIC site, at 10 km a.s.l., the GF has a strength module \( |\vec{B}| \approx 38.4 \, \mu T \), being the Geomagnetic vector field \((B_x, B_y, B_z) \approx (29.9,0.0,24.1) \, [\mu T] \) [158]. As shown, the module of the Lorentz force (which is proportional to the component of the GF normal to the shower axis) has a symmetric behavior with respect to the plane perpendicular to the Geomagnetic North-South direction. Since the magnetic field lines at La Palma are tilted by \( \approx 7^\circ \) westwards with respect to the meridian, the trajectories of the sources in the sky are asymmetric with respect to 180\(^\circ\) Azimuth angle (see the path covered in the sky by the Crab Nebula source, observed at the Roque de los Muchachos site, superimposed in the plot shown in Fig. 5.6(a)). At La Palma, the minimum influence of the GF is expected to occur for EAS developing in direction of the Geomagnetic North at ZA = (90\(^\circ\) - I) \( \approx \) 51\(^\circ\) and AZ = 0\(^\circ\), where the angle \( \alpha \) between the shower axis and the GF becomes the smallest (see Fig. 5.5(a)), i.e. for EAS developing along the GF lines. \( I \) denotes the angle under which the GF lines dip into the Earth’s surface, which is around 39\(^\circ\) for La Palma [158]. The maximum influence is therefore expected for EAS developing per-
pendicularly to the direction of the GF lines, i.e. for \(ZA \approx 39^\circ\) and \(AZ = 180^\circ\) (see Fig.5.6(a)).

### 5.1.3 The stereo configuration effect

In the stereoscopic system observations, the Effective Collection Area must be computed considering the MC-\(\gamma\) events triggered by both the telescopes. The distance between the two detectors is therefore an important quantity to take into account because it determines the overlap of the fields of views (FoVs) of the two telescopes. The best relative distance between the MAGIC telescopes for low ZA was determined by dedicated MC studies to be around 85 m [100,101].

For observations very close to the Zenith, the distance between the axes of the two telescope FoVs coincides with the physical distance between the two telescopes, i.e. 85 m, and it is AZ independent. For higher ZAs of observation, the relative distance starts to show a dependence on the AZ, simply due to a geometric effect. In fact, the relative distance between the axes of the two telescopes’ FoVs can be written (in the CORSIKA program reference frame) as

\[
d(\theta, \phi, \eta) = d(0^\circ, \phi, \eta) \cdot \left[ (\cos^2 \theta \cos \eta - \sin^2 \theta \sin \phi (\cos \phi \sin \eta - \sin \phi \cos \eta))^2 + (\cos^2 \theta \sin \eta + \sin^2 \theta \cos \phi (\cos \phi \sin \eta - \sin \phi \cos \eta))^2 + (- \cos \theta \sin \theta (\sin \phi \sin \eta + \cos \phi \cos \eta))^2 \right]^{\frac{1}{2}} 
\]  

(5.6)

where the angle \(\theta\) and \(\phi\) are respectively the ZA and the AZ, while \(\eta\) is the angle between the \(x\)-axis and the direction joining MAGIC-I to MAGIC-II, which has a value of about 135\(^\circ\).

\(d(0^\circ, \phi, \eta)\) is the distance between the two telescopes for Zenith observations and hence it is equal to 85 m.

Figure 5.6(b) shows the value of the relative distance between the axes of the two telescope FoVs as a function of the Alt-Azimuth pointing position (according to equation 5.6): the stereo configuration effect has a periodicity of 180\(^\circ\) in AZ (the maximum distance of 85 m being achieved for all ZA for AZ\(\approx 45^\circ\) and AZ\(\approx 225^\circ\)) and it reaches a difference between the maximum and minimum distances of the order of 30\% for ZA\(\approx 40^\circ\).

It is worth mentioning that, since the fixed direction between the telescopes is not along the direction where the GF vector lies (the angle between the two direction being around 45\(^\circ\)), the stereo configuration effect and the GF one combine each other in a nontrivial way.

### 5.2 Azimuth dependence of the stereo \(\gamma\)-ray Effective Collection Area

In order to investigate the dependence of the \(\gamma\)-ray Effective Collection Area with respect to the AZ due to the GF and the stereo configuration effects, we used a sample of MAGIC stereo MC-\(\gamma\) available when this study was done (autumn 2009). Details of the simulation steps and of the characteristics of this sample can be found in [151]. Here we report the main characteristics of the used sample:

- The energy ranges from 10 GeV to 30 TeV. The energy distribution is a pure power law with a spectral index of -1.6. This spectral index was set a unit higher than the real Crab Nebula observed spectrum to get more statistics at higher energies.
- The ZA ranges between 0\(^\circ\) to 45\(^\circ\).
- The AZ covers uniformly the whole circle, from 0\(^\circ\) to 360\(^\circ\).
5.2. Azimuth dependence of the stereo $\gamma$–ray Effective Collection Area

Figure 5.6: (a) The absolute value (in [$\mu$T]) of the component of the GF normal to the direction of the EAS versus AZ and ZA for the Roque de los Muchachos observatory on La Palma. (b) The relative distance (in [m]) between the axes of the FoVs of the two MAGIC telescopes as a function of the AZ and ZA. Since the fixed direction which joins MAGIC-I to MAGIC-II forms an angle of about $135^\circ$ with respect to the $x$-axis (i.e. with respect to the Geomagnetic North) the maximum relative distance for ZA $> 0^\circ$ is achieved for AZ values around $45^\circ$ and $225^\circ$. The two plots are displayed in CORSIKA program reference frame.

- The maximum IP was fixed to 300 m for the $\gamma$-ray events with ZA $< 30^\circ$ and 450 m for $\gamma$–ray events with $30^\circ < $ZA$< 45^\circ$. The IP has a value equal to 0 m for EAS with propagation axis passing through the point at the middle of the line which joins the two telescopes.

- The simulated optical point spread function (sigma of the gaussian) is for both telescopes’ surfaces set to $\sqrt{(7 \, \text{mm})^2 + (8 \, \text{mm})^2} = 10.6 \, \text{mm}$ (see section 4.2.3).

- The observation mode is ON-OFF: the nominal source position is always placed in the center of the of the MAGIC-I and MAGIC-II cameras, that is the telescopes’ optical axes have always been set parallel to the direction of the primary $\gamma$-rays.

- The overall number of simulated stereo MC-$\gamma$ events of the sample is about $6 \times 10^6$.

The MC-$\gamma$ generated showers were analyzed using the MARS analysis chain explained in chapter 4. At the image cleaning level a soft SIZE cut of 30 phes was applied for both telescope MC-$\gamma$ separately. After the image parameters were calculated for each event surviving the image cleaning (performed separately for each telescope), the program SUPERSTAR recognized the events triggered by both telescopes. For those events, the two different images recorded by the two telescopes were then considered to obtain the stereoscopic image parameters of the shower (see section 4.4). A set of standard quality cuts were applied at this stage: SIZE1(2)$> 60$ phes, LEAKAGE1(2)$< 0.2$ and NUMBER OF ISLANDS1(2)$\leq 2$, which are typical quality cuts used for the MAGIC analysis. No further cuts were applied, in particular no $\gamma$/hadron separation cuts were taken into account. Thus, the calculated values for the Effective Collection Area reported below must be considered independent on the Hadronness efficiency. Moreover, since no energy reconstructed algorithm was performed, the energy of the events refers to the true simulated energy.

Since the Effective Collection Area and the $E_{th}$ strongly depend on the ZA of the observation,
we decide to define the ZA bins used in this analysis in the following way: given that the bins in ZA used for the MAGIC-I MC-\(\gamma\) simulations (zbins) are defined equally spaced in \(\cos(\theta)\) with a step of 0.01, we redefined new larger ZA bins corresponding each one to 6 zbins (see the first two columns of Table 5.1). This was done in order to keep the \(\cos(\theta)\) binning dependence and also to get a good balance between the amount of statistics for each ZA bin and the width of the bins. From now, we will refer to them as Zbins. In Fig. 5.7 the energy thresholds and the Effective Collection Area for the different Zbins are shown: the ZA dependence of those quantities is evident.

---

**Figure 5.7:** (a) Energy thresholds for the different Zbins defined in table 5.1. For each Zbin, the \(E_{th}\) is defined as the peak of the stereo MC-\(\gamma\) distribution (with a given power law of spectral index -1.6) after the quality cuts. As the ZA increases, the \(E_{th}\) increases, roughly as \(\cos^{-2.7}(\theta)\), as explained in the text. (b) Effective Collection Area for the different Zbins defined in table 5.1. The \(y\) axis is not plotted in logarithmic scale in order to emphasize the different behavior of the Effective Collection Area values as a function of the Zbins. For higher ZAs the Effective Collection Area starts to rise at higher energies and then it reaches higher values, due to the increasing of the \(E_{th}\) (see expression 5.5) and of the \(\gamma\)-ray detection efficiency. Note that the decreasing of the Effective Collection Area at the higher values of energy is due to the applied LEAKAGE quality cuts.
5.2. Azimuth dependence of the stereo $\gamma$-ray Effective Collection Area

In order to have, for each Zbin, a reasonable energy binning where to estimate the AZ dependence of Effective Collection Area, we defined four basic energy bins (Ebins) according to the following choice:

- The first bin ranges from about $2/3$ of the value of the $E_{th}$ to the $E_{th}$ value (very low energy bin).
- The second bin ranges from the $E_{th}$ to the double of its value (low energy bin).
- The third bin ranges from the double value of the $E_{th}$ to 6 times its value (medium energy bin).
- The fourth bin ranges from 6 times the value of the $E_{th}$ to the maximum energy value of the simulation, i.e. 30 TeV (high energy bin).

In table 5.1 the definition of the Zbins used for the analysis and of the corresponding Ebins (according to the different energy thresholds) are shown, whereas in table 5.2 the amount of statistics for the simulated events and for the events surviving the selection (i.e. the trigger, the image cleaning, the coincident trigger in both telescopes and the quality cuts discussed above) are reported for all the defined Zbins and Ebins.

In Fig. 5.8 the Effective Collection Area values for the different Zbins as a function of the corresponding Ebins and 12 AZ bins are plotted. In order to show in a better way the AZ dependence of the Effective Collection Area values, in Fig. 5.9 the values with respect to the 12 AZ bins of each Zbin - Ebin combination are displayed (in CORSIKA program reference frame).

As expected, the AZ dependence of the Effective Collection Area is not negligible particularly at the energies close to the threshold (for a given Zbin) and for the higher Zbins. Moreover, the GF effect, particularly for the lowest Zbins, where the stereoscopic effect is not strong, appears to be the main reason for the AZ dependence of the Effective Collection Area, and it is mainly characteristics by a GF North-South dependence (i.e. a symmetry with respect to $AZ = 180^\circ$). Indeed, the AZ value where, for a given ZA, the GF component normal to the shower is minimum (i.e for $\gamma$-rays coming from the Geomagnetic South), clearly show a higher value of the Effective Collection Area. This is particularly evident for the lower energies, whereas for the higher ones the AZ dependence due to the GF effect seems to vanish. It is also worth mentioning that, in case of evident Geomagnetic effect, the differences between consecutive AZ bins increase toward AZ values approaching the Magnetic North (i.e. decreasing values toward $AZ = 0^\circ$ and increasing values toward $AZ = 360^\circ$), whereas around $AZ = 180^\circ$ a sort of plateau is present. This means that the higher density of Cherenkov light for the events coming from the Geomagnetic South (which are less affected by the GF), compared to the light density for those coming from the Geomagnetic North, has a clear effect on the Effective Collection Area values. The difference between the light collected at North and South directions should decrease with increasing IPs [157], but this behavior was not investigated since our results are integrated in the IP.

Concerning the stereoscopic effect, it is not possible to clearly distinguish it in the plots shown in Fig. 5.9, also because, as said above, it combines with the GF effect in a nontrivial way. Possibly, the best evidence of this effect is related to the differences which appear in the highest Ebin for the different Zbins: at those energies, the GF is no longer strongly evident and, as the ZA increases, the AZ distributions become less flat. Also the Effective Collection Area values shown in some plots seem to show a modulation of the GF effect which could be related to the stereo configuration effect. Anyway, the stereoscopic effect should become stronger and stronger, as the ZA increases and it could become comparable or even stronger.
Table 5.1: Definition of the Zbins and of the Ebins for each Zbin used throughout the stereo MC-$\gamma$ analysis. The Ebins are defined according to the different values of the $E_{th}$ estimated from the energy distribution of each Zbin (see Fig. 5.7(a)).

than the GF effect for $Z_A > 45^\circ$ also for the lowest energies bins. For checking this behavior, simulations at $Z_A$ higher than 45$^\circ$ are needed.

In table 5.3, the maximum AZ difference values (in percentage) for each Zbin and Ebin combination, estimated by the plots reported in Fig. 5.9, are shown$^2$. Considering the width of each Ebin and that the power index of the used stereo MC-$\gamma$ simulation is -1.6, we would expect even a larger Azimuth dependence in case of a typical source spectrum (for instance -2.6 for the Crab Nebula case). Anyway, for a given Zbin, this should not invalidate the general behavior found as a function of energy (see discussion above).

The results obtained in this section justify the implementation of the Azimuth information in the analysis chain.

$^2$Note that, of course, the reported results are AZ binning dependent.
<table>
<thead>
<tr>
<th>Zbin</th>
<th>simulated events</th>
<th>surviving events</th>
<th>Ebin1</th>
<th>Ebin2</th>
<th>Ebin3</th>
<th>Ebin4</th>
</tr>
</thead>
<tbody>
<tr>
<td>Zbin1</td>
<td>1498942</td>
<td>110596</td>
<td>111615</td>
<td>128209</td>
<td>129171</td>
<td>125295</td>
</tr>
<tr>
<td></td>
<td>1536178</td>
<td>125062</td>
<td>94894</td>
<td>124754</td>
<td>117467</td>
<td>112787</td>
</tr>
<tr>
<td></td>
<td>1156971</td>
<td>68104</td>
<td>72293</td>
<td>85048</td>
<td>79593</td>
<td>76259</td>
</tr>
<tr>
<td></td>
<td>954586</td>
<td>41795</td>
<td>52569</td>
<td>62062</td>
<td>58203</td>
<td>54343</td>
</tr>
<tr>
<td></td>
<td>853023</td>
<td>41043</td>
<td>41912</td>
<td>50137</td>
<td>46664</td>
<td>43440</td>
</tr>
</tbody>
</table>

Table 5.2: Number of simulated and of surviving stereo MC-γ events (after trigger, image cleaning, coincident trigger and quality cuts) as a function of the Zbins and Ebins defined in table 5.1.

<table>
<thead>
<tr>
<th>Zbin</th>
<th>simulated events</th>
<th>surviving events</th>
<th>Ebins [GeV]</th>
<th>difference between maximum and minimum</th>
</tr>
</thead>
<tbody>
<tr>
<td>Zbin1 (0° - 19.09°)</td>
<td>very low (60 - 90)</td>
<td>24%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>low (90 - 180)</td>
<td>10%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>medium (180 - 540)</td>
<td>6%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>high (540 - 30000)</td>
<td>3%</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Zbin2 (19.09° - 27.75°)</td>
<td>very low (75 - 110)</td>
<td>29%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>low (110 - 220)</td>
<td>18%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>medium (220 - 660)</td>
<td>8%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>high (660 - 30000)</td>
<td>3%</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Zbin3 (27.75° - 34.41°)</td>
<td>very low (85 - 130)</td>
<td>41%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>low (130 - 260)</td>
<td>21%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>medium (260 - 780)</td>
<td>13%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>high (780 - 30000)</td>
<td>5%</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Zbin4 (34.41° - 40.09°)</td>
<td>very low (105 - 160)</td>
<td>47%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>low (160 - 320)</td>
<td>21%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>medium (320 - 960)</td>
<td>11%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>high (960 - 30000)</td>
<td>9%</td>
<td></td>
<td></td>
</tr>
<tr>
<td>Zbin5 (40.09° - 45.17°)</td>
<td>very low (125 - 190)</td>
<td>45%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>low (190 - 380)</td>
<td>24%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>medium (380 - 1140)</td>
<td>11%</td>
<td></td>
<td></td>
</tr>
<tr>
<td></td>
<td>high (1140 - 30000)</td>
<td>8%</td>
<td></td>
<td></td>
</tr>
</tbody>
</table>

Table 5.3: Maximum AZ difference value (in percentage) for each Zbin and Ebin combination, estimated from Fig. 5.9. The used number of AZ bins for each Zbin-Ebin combination is 12, which corresponds to 30° AZ steps.
Figure 5.8: Effective Collection Area computed for the different Zbins with the energy binning given by the corresponding Ebins values, using 12 AZ bins for each Zbin - Ebins combination.
Figure 5.9: Azimuthal distributions (using 12 AZ bins) for the Effective Collection Area computed for the different Ebins (columns) for each Zbin (rows). See table 5.1 for their definitions.
5.3 Main changes implemented in MARS

As explained in chapter 4, the MAGIC analysis software (MARS) is a collection of ROOT-based [113] programs written in C++. The package is divided in class and executables (programs), each of which has specific tasks for the computation of all the quantities needed to reduce the MAGIC data.

The implementation in MARS of the Azimuth dependence of the Effective Collection Area required a deep modification and updating of some basic classes and executables. In the following, we will enter into technical code details. The given results shown in the figures are nevertheless of general interest.

We started the modifications from the change of the two main classes used for the computation of the Effective Collection Area, namely \textit{MHMcCollectionArea} and \textit{MMcCollectionAreaCalc}. The main intent was to introduce the AZ parameter for the calculation of the main quantities involved in the Effective Collection Area determination. After that, we consequently updated all the other classes and executables which needed the introduction of the new AZ dependent quantities. In table 5.4 the list of the main modified MARS elements is reported\(^3\). We will explain in the next subsections the most relevant modifications.

<table>
<thead>
<tr>
<th>MARS location</th>
<th>name</th>
</tr>
</thead>
<tbody>
<tr>
<td>Mars/mhistmc/</td>
<td>\textit{MHMcCollectionArea}</td>
</tr>
<tr>
<td>Mars/mmontecarlo/</td>
<td>\textit{MMcCollectionAreaCalc}</td>
</tr>
<tr>
<td>Mars/mh\textit{eft}/</td>
<td>\textit{MHEffectiveOnTime}</td>
</tr>
<tr>
<td>Mars/mh\textit{eft}/</td>
<td>\textit{MHThetaTime}</td>
</tr>
<tr>
<td>Mars/mhistmc/</td>
<td>\textit{MHMcEnergyMigration}</td>
</tr>
<tr>
<td>Mars/</td>
<td>\textit{FLUXLC}</td>
</tr>
</tbody>
</table>

\(^3\)Table 5.4: List of the main MARS classes and executables modified in order to introduce the AZ dependence of the Effective Collection Area in the analysis chain.

5.3.1 Changes implemented for the calculation of the Effective Collection Area

The two main classes involved in the computation of the Effective Collection Area are called \textit{MHMcCollectionArea} and \textit{MMcCollectionAreaCalc}.

For a clear explanation of what follows in this subsection, we introduce here the two different kinds of binning used throughout the analysis chain: the \textit{fine binning} and the \textit{coarse binning}. The \textit{fine binning}, made up by a large number of bins (with respect to the \textit{coarse binning}), is used to store into some histograms the energy, the ZA and the IP (as well as the AZ, in the new MARS version) values of the events belonging to the MC-γ sample and the real data.

The \textit{coarse binning}, made up by a lower number of bins, refers, instead, only to the energy and the ZA, since no coarse bin are defined for the IP (in case of no AZ dependence implemented in the software) or for the AZ (in the new MARS version with AZ dependence). This last feature will be explained later. Generally speaking, the main histograms needed to evaluate the Effective Collection Area are computed in fine bins. Then, in order to infer the flux of a given source, the Effective Collection Area computed in fine bins is properly re-weighted with some additional pieces of information which are the supposed source spectrum (typically a power law of index \(Γ\)) and the effective observation time of the real observation as a function of the zenith angle.\(^3\)

\(^3\)Some minor changes in other classes not shown in table 5.4 were also done.
of the telescope pointing position. The final result is the Effective Collection Area expressed as a function of energy and $Z_A$, in coarse bins.

In MARS the different fine bins for the involved quantities are defined as follows:

- **Energy ($E$)**: 125 logarithmic bins between 2 GeV and 200 TeV. We will use the index $i$ when referring to this binning.

- **$Z_A$ ($\theta$)**: 60 bins, equally spaced in $\cos(\theta)$ (with step of 0.01). The $\theta$ parameter therefore ranges from 0° to 66.73°. We will use the index $j$ when referring to this binning.

- **IP ($ip$)**: 52 bins between 0 m and 520 m (the maximum IP of the MC-$\gamma$ simulations never exceeds the maximum edge of the IP fine bins). We will use the index $k$ when referring to this binning.

- **AZ ($\phi$)**: the *fine binning* in Azimuth has been introduced by the work described here. The $\phi$ angle ranges from 0° to 360° and it is linearly divided in a number of bins which can be set by the analyzer\(^4\). By this definition, for example, in case only one AZ bin is defined, the results should be the same than those achieved in case of no AZ dependence. Ideally, a good number of AZ fine bin can be 12, corresponding to 30° per bin, but this choice is strictly dependent on the number of events of the MC-$\gamma$ sample. We will use the same index $k$ as the IP when referring to the AZ binning.

All the fine bins (apart from the number of AZ bins) are fixed in the analysis, whereas the *coarse binning* must always be set by the analyzers. As said above, the coarse bins refer only to the energy (let use the index $I$) and to the $Z_A$ (index $J$). The proper setting for the coarse bins is mainly dependent on the $Z_A$ of the given observation and the strength of the source under investigation. Normally, in case of observations performed below $\sim 30°$, only one $Z_A$ coarse bin is set. Concerning the energy, 30 logarithmic coarse bins defined between 5 GeV and 50 TeV are typically used.

An important choice we made at the beginning of the implementation of the AZ dependence in the Effective Collection Area calculation was to avoid the introduction of coarse bins for the AZ parameter (i.e. to have always only 1 AZ coarse bin in the final results). This means that the AZ parameter is properly taken into account during the calculation of the Effective Collection Area in fine bins, but, at the end of the calculation, the final Effective Collection Area is only dependent on ($E$, $\theta$) coarse bins, as before the implementation of the AZ dependence in the software. This choice did not deprive the results to properly account for the AZ dependence and, in addition, it allowed an easier implementation in MARS of the AZ parameter, avoiding deeper changes in some of the classes we had to modified. Indeed, most of the classes which need the Effective Collection Area values for further calculations require those values in coarse bins, so, technically speaking, nothing changes for these classes after the introduction of the AZ dependence (whatever the number of AZ fine bin is).

In the following, we describe briefly how the $\text{MHMcCollectionArea}$ and $\text{MMcCollectionAreaCalc}$ classes work without the AZ dependence. The calculation involves four main steps which are performed by $\text{MHMcCollectionArea}$. $\text{MMcCollectionAreaCalc}$ is used to define the parameters and to fill the basic histograms. Each step of the calculation is characterized by a histogram:

- **$fHistAll (A)$**: this histogram is filled with all the events generated in the MC-$\gamma$ simulation taken into account for the analysis of a given source (which are the MC-$\gamma$ events

\(^4\)The AZ values in use refer to the MAGIC telescopes’ drive reference frame and not to the CORSIKA program one. This has been chosen because the AZ values of the real data events refer to the drive coordinate system. All the MC-$\gamma$ AZ values are hence properly reconverted in this reference frame (see Fig. 5.5(b)).
statistically independent from those used for training the $\gamma$-hadron separation and the energy reconstruction). It is a 3D fine bin histogram in $(E, ip, \theta)$ with entries for each fine bin combination $(i, k, j)$ given by

$$ A_{ikj} = \sum_{Q} F(Q) \quad (5.7) $$

where $Q$ represents the index for the simulated events of the considered MC-$\gamma$ sample and $F(Q) = 1$ if $Q \in (i, k, j)$ and zero otherwise.

- **fHistSel** ($S$): this histogram is filled by all the events surviving the trigger selection, the image cleaning, and all analysis cuts (optimized in each estimated energy coarse bin). It is a 3D fine bin histogram in $(E, ip, \theta)$, as $fHistAll$, with entries for each fine bin combination $(i, k, j)$ given by

$$ S_{ikj} = \sum_{q} f(q) \quad (5.8) $$

where $q$ represents the index for the surviving events among the simulated ones and $f(q) = 1$ if $q \in (i, k, j)$ and zero otherwise.

- **fHistCol** ($C$): this histogram represents the Effective Collection Area in fine bins. In order to get it, the first step is to find out for each index $j$ (i.e. for each ZA fine bin), the maximum IP (as needed for subsequent calculations). Thus, the 2D projection of $fHistAll$ with respect to $E$ is used: for each $j$ the IP histogram is taken into account and the bin corresponding to a huge decrease of the entries is found out. From this bin the maximum $ip$ value of the simulation is extracted and stored for its use in the next calculations (let call it $M_j$). In case more than one homogeneous sample of MC-$\gamma$ is taken into account, i.e. with different simulated maximum IPs, this procedure also avoid possible problems: all the events within the $j$-th ZA fine bin with IP greater than $M_j$ (if any) are no longer taken into account in the next calculations. After the determination (for each $j$) of $M_j$, $fHistCol$ is calculated by dividing the 2D projection (along the $ip$) of the 3D $fHistSel$ histogram by the 2D projection (again, along the $ip$) of the 3D $fHistAll$ histogram. In this way, we obtain the efficiency of the surviving events in fine bins of energy and ZA. Finally, to obtain the Effective Collection Area we multiply the result, for each couple $(i, j)$, by a number representing the suitable area defined in the simulation (which is independent on the energy index $i$):

$$ Area_j = \pi \cdot (M_j)^2 \quad (5.9) $$

$fHistCol$ is a 2D histogram in $(E, \theta)$, since the IP variable is no longer needed. Thus, resuming, the values of this histogram, for each fine bin combination $(i, j)$ is given by

$$ C_{ij} = Area_j \cdot \frac{\sum_k S_{ikj}}{\sum_k A_{ikj}} \quad (5.10) $$

The errors are computed taking into account a binomial statistics, since the events in $fHistSel$ are a subsample of those in $fHistAll$.

- **fHistColCoarse** (CC): this histogram is the Effective Collection Area in the coarse bins selected by the analyzer. It represents the quantity which is eventually used for the evaluation of the flux (or its upper limit, in case of no significant detection) of a given source. It is important to note that this histogram is not a simple re-binning of the
5.3. Main changes implemented in MARS

The $fHistCol$ histogram. Indeed, in order to calculate it, two further quantities (in fine bins) are taken into account: the assumed source spectrum ($F(E)$, provided by the analyzer) and the effective observation time as a function of the ZA (in fine bins) related to the observation of the given source ($T_j$) (calculated by a MARS class called $MHEffectiveOnTime$). These two quantities are used to calculate the proper weight $O_{ij} = F_i \cdot T_j$ for each $(i, j)$ (where $F_i$ represents the integral of $F(E)$ in the i-th energy fine bin) to be used to compute the final Effective Collection Area values in $(E, \theta)$ coarse bins. Thus for each $(I, J)$

$$CC_{IJ} = \frac{\sum_{i \in I} \sum_{j \in J} C_{ij} \cdot (F_i \cdot T_j)}{\sum_{i \in I} \sum_{j \in J} (F_i \cdot T_j)}$$

(5.11)

In Fig. 5.10 the main quantities used to get the final Effective Collection Area in coarse bins as a function of the true MC-$\gamma$ energy are displayed. Since they refer to a Crab Nebula subsample (see section 5.4) performed in the second Zbin ($19.09 < \theta < 27.75$), only one ZA coarse bin was set, whereas the estimated energy coarse bins were set to 30, in a logarithmic scale between 5 GeV and 50 TeV. The assumed spectrum was set to be a power law with spectral index equal to -2.6.

Concerning the introduction of the AZ dependence, the first main problem we had was to understand how to include that new degree of freedom avoiding to introduce quantities (histograms) with more than 3 dimensions, since $fHistAll$ and $fHistSel$ were already 3D histograms, and the increasing of one dimension, in principle, might have increased the computing time and the CPU memory requirements (apart the fact that in the ROOT environment 4D histograms are not defined). The best solution was to replace the IP variable with the AZ one, since the IP was basically used only to find out the proper values of the maximum IP in each ZA fine bin. In the following, the index $k$ will be thus related to the AZ fine bin, instead of the IP ones. The AZ information of each event is extracted from the container $MMcEvt.fTelescopePhi$ and properly converted in the telescopes’ drive system coordinates.

In order to avoid the problems connected to the determination of the maximum IP, a weight

$$W_r = \frac{\text{AreaMax}}{\pi \cdot (M_r)^2}$$

(5.12)

is computed, for each simulated event taken into account, where

- $r$ refers to the MC-$\gamma$ run index
- $\text{AreaMax}$ is a unique and arbitrary number set to $\pi \cdot 10^6 \text{ m}^2$ (corresponding to an arbitrary maximum IP equal to $10^3 \text{ m}$, thus, greater than all maximum IPs of the MC-$\gamma$ simulations)
- $M_r$ is the highest value of IP found out among all the events belonging to a given MC-$\gamma$ run with index $r$ (for this reason we need now also the run index $r$). This value is properly stored in each MC-$\gamma$ run ($McRunHeader.fImpactMax$) and it is unique, therefore, $W_r$ is equal for all the simulated events inside a particular MC-$\gamma$ run, independently on the other quantities, i.e. independent on the indexes $(i, k, j)$.

The calculation of $W_r$ for all the considered MC-$\gamma$ events is performed by the new version of $MMcCollectionAreaCalc$ class. After this, the following histograms are calculated by the new version of $MHMcCollectionArea$ in order to get the final Effective Collection Area in coarse bins, with the AZ dependence properly taken into account:
Figure 5.10: Main quantities involved in the computation of the Effective Collection Area in coarse bins in case of no AZ dependence. (a) \( f_{HistAll} \) projection along the \( ip \) axis. (b) \( f_{HistSel} \) projection along the \( ip \) axis. (c) \( f_{HistCol} \) (d) Effective observation time of the considered Crab Nebula observation (between 19.09 < \( \theta \) < 27.75) as a function of the ZA fine bins. (e) \( f_{HistColCoarse} \) as a function of the true MC-\( \gamma \) energy, with one ZA coarse bin, 30 logarithmic estimated energy coarse bins between 5 GeV and 50 TeV and a spectral index of -2.6 for the assumed spectrum.

- the new \( f_{HistAll} \): it is now a 3D histogram in \((E, \phi, \theta)\). Each considered event belonging to the \( r \)-th MC-\( \gamma \) run is rescaled with its corresponding value of \( W_r \). In this way, all the events can be thought as they were simulated with a maximum IP equal to \( 10^3 \) m, avoiding the problems which can arise in case of different simulated maximum IP for a
given ZA index \( j \). The entries for each fine bin combination \((i, k, j)\) are hence given by

\[
A_{ikj} = \sum_r \sum_{S_r} G(S_r)
\]  

(5.13)

where \( S_r \) represents the run dependent index for the simulated events and and \( G(S_r) = W_r \) if \( S_r \in (i, k, j) \) and zero otherwise.

- the new \( fHistSel \): it is now a 3D histogram in \((E, \phi, \theta)\). It is filled, as in the old case, with all the surviving events, but now according to the values of the energy, AZ (instead of IP) and ZA of each event (see equation 5.8).

- the new \( fHistCol \): this histogram is now a 3D histogram in \((E, \phi, \theta)\) fine bins (i.e., \( C_{ikj} \)). \( fHistCol \) is calculated for each fine bin combination \((i, k, j)\) dividing \( fHistSel \) by \( fHistAll \), and multiplying the result for a number representing the appropriate rescaled quantity due to the introduction of the IP weights for the calculation of \( fHistAll \). The errors are properly evaluated considering the binomial statistic and the introduction of the IP weights.

- \( fHistColCoarse \): it is a 2D histogram in \((E, \theta)\) coarse bin (as in the old case), since, as said before, no AZ coarse bin has been introduced. Again, the assumed source spectrum and the effective time of the given observation are taken into account. However, now, the effective observation time is a function of the ZA fine bins and also of the AZ fine bins (so we have \( T_{kj} \)). Therefore, the weights calculated to get the final Effective Collection Area in coarse bins are also AZ dependent: \( O_{ikj} = F_i \cdot T_{kj} \). The final Effective Collection Area values in \((E, \theta)\) coarse bins for each \((I, J)\) are thus

\[
CC_{IJ} = \frac{\sum_{i\in I} \sum_{k \in K} \sum_{j\in J} C_{ikj} \cdot (F_i \cdot T_{kj})}{\sum_{i\in I} \sum_{k \in K} \sum_{j\in J} (F_i \cdot T_{kj})}
\]  

(5.14)

In Fig. 5.11 the main quantities used to get the final Effective Collection Area in coarse bins (as a function of the true MC-\( \gamma \) energy) with AZ dependence are displayed. They refer to the same Crab Nebula sample shown before (for the previous standard case of no considered AZ dependence). In this case, 6 AZ fine bins are set. Note that the final \( fHistColCoarse \) displayed in Fig. 5.11(h) has values very close to those calculated in case of no AZ dependence (see 5.10(e)), the differences being due to the AZ weights introduced by the Azimuth dependence of the effective time of the observation. In case only 1 AZ fine bin is set, the final values and errors of the Effective Collection Area in coarse bins have been checked to be nearly identical, as expected.
Figure 5.11: Main quantities involved in the computation of the Effective Collection Area in coarse bins in case of AZ dependence. The number of AZ fine bins is set to 6. (a) $fHistAll$ projection along the $\theta$ axis. (b) $fHistAll$ projection along the $\phi$ axis. (c) $fHistSel$ projection along the $\theta$ axis. (d) $fHistSel$ projection along the $\phi$ axis. (e) $fHistCol$ projection along the $\theta$ axis. (f) $fHistCol$ projection along the $\phi$ axis. (g) Effective observation time of the considered Crab Nebula observation (between 19.09 < $\theta$ < 27.75) as a function of the ZA and AZ fine bins. (h) $fHistColCoarse$ as a function of the true MC-$\gamma$ energy, with one ZA coarse bin, 30 logarithmic estimated energy coarse bins between 5 GeV and 50 TeV and a spectral index of -2.6 for the assumed spectrum.
5.3.2 Changes implemented for the calculation of the effective observation time of the real data

If the number of events in a time range follows the poissonian statistics, the distribution of the time difference between consecutive events ($\Delta t$) is an exponential (with index equal to the poisson parameter)

$$P(\Delta t) = \lambda \cdot e^{-\lambda \cdot \Delta t}$$  \hspace{1cm} (5.15)

Hence, performing an exponential fit of the $\Delta t$ values, it is possible to estimate the poisson parameter, corresponding to the mean rate of the events (which would be actually observed in absence of deadtime). From that value and from the total number of observed events used for deriving it, the effective time of the observation can be calculated.

In MARS, the class in charge to perform this calculation is called $MHEffectiveOnTime$. Besides the information of the effective observation time as a function of the Coordinated Universal Time (UTC), the class gives also information about the effective observation time in fine bins and coarse bins of ZA. With the introduction of the AZ dependence for the Effective Collection Area, we thus had to update that class, relatively to the fine bin calculation. This was needed to provide the class $MHMcCollectionArea$ with the proper AZ and ZA weights, since these values are exactly the effective observation times, as a function of the AZ and ZA fine bins (i.e. $T_{kj}$).

Before the changes we implemented, the effective observation time was calculated only as a function of the ZA of the observation, giving, for each ZA fine bins, the exponential fit $\lambda_j$ of the corresponding $\Delta t$ distribution (see figure 5.10(d)). The $\Delta t$ binning is defined by 500 bins between 0 and 0.2 s, while the exponential fit is evaluated in the range between 0.003 and 0.095 s. The overall effective time for a given observation was thus (using, as before, the index $j$ to indicate the ZA fine bins)

$$T_{eff} = \sum_j \frac{\sum_q h(q)}{\lambda_j}$$  \hspace{1cm} (5.16)

where $q$ is the index for the surviving events, $\lambda_j$ is the value of the mean rate for the $j$-th ZA fine bin calculated from the exponential fit and $h(q) = 1$ if $q \in j$ and zero otherwise.

After the introduction of the AZ dependence, the effective observation time is dependent also on the $k$-th AZ fine bin, therefore

$$T_{eff} = \sum_{k,j} \frac{\sum_q l(q)}{\lambda_{kj}}$$  \hspace{1cm} (5.17)

where $q$ is the index for the surviving events, $\lambda_{kj}$ is the value of the mean rate for the $k$-th AZ fine bin and $j$-th ZA fine bin calculated from the exponential fit and $l(q) = 1$ if $q \in (k, j)$ and zero otherwise (see Fig. 5.11(g)).

All the other quantities computed by the class $MHEffectiveOnTime$, such as the probabilities and the $\chi^2$ of the fits, have been properly extended in order to include the new AZ dependence. A new plot, displaying the mean rate as a function of the AZ, was also added in the output, beside the rate as a function of the ZA. An example is shown in Fig. 5.12.

Another change was introduce in another MARS class called $MHThetaTime$. Before this work, this class was used by the program FLUXLC (see section 4.3.9) to extract a histogram of the real events as a function of the ZA and of the UTC (for the calculation of the light curve of the observed source). After the AZ implementation, that class was updated to $MHPhiThetaTime$, which has now the task to compute a 3D histogram in ($\phi$, $\theta$, $T$), where
Figure 5.12: (a) Rate of a Crab Nebula sample between 10° and 30° as a function of the ZA. (b) Rate of the same Crab Nebula observation as a function of the AZ. Note that the observation was performed for different AZ corresponding to the same ZA. This sample was selected from the overall sample of the Crab Nebula data explained in subsection 5.4.

the bins for ZA and AZ are fine bins (see Fig. 5.15).

5.3.3 Changes implemented for the calculation of the energy Migration Matrix

As already explained in section 4.3.11, since the detector has a finite energy resolution and since the true energy cannot be directly measured, the measurements of the energy of the γ-like events extracted from the data of a given source are systematically distorted, particularly at the energies close to the threshold. The spectrum calculated in bins of estimated energy (which is the output of the FLUXLC program) must be thus corrected by computing the energy spectrum in bins of true energy. This task is accomplished by the subsequent unfolding procedure which employs the fact that the MC-γ events have both the true and the estimated energy. The distortions due to biases and finite resolution can be written in the form:

$$Y(y) = \int M(y, x)S(x)dx \quad \text{or} \quad Y_L = \sum_I M_{LI} S_I$$

(5.18)

where $y$ is the estimated energy ($E_r$), $x$ is the true energy ($E$), the index $L$ refers to the estimated energy coarse bins, the index $I$ refers to the true energy coarse bins, $M$ describes the detector response (i.e. the energy Migration Matrix), $Y$ is the measured distribution and $S$ is the true undistorted distribution. The aim is to determine $S$, given $Y$ and $M$.

There are various approaches to solve this problem, the best one being the so called unfolding with regularization. In what follows, we will not explain in details the unfolding procedure but rather report how the Migration Matrix is calculated in MARS and describe the main modifications implemented in order to introduce the AZ dependence (detailed explanations of the unfolding procedure and methods can be find in [149, 150]). In MARS, the Migration Matrix, which is used as input by the unfolding program, is defined as a $L \times I \times J$ matrix. $L$ is the number of the estimated energy coarse bin, $I$ is the number of the true energy coarse bin (which is generally a factor 1.4 smaller than $L$, as shown in the last row of table 5.5) and $J$ is the number of the ZA coarse bins. The Migration Matrix is calculated by the class `MHMmcEnergyMigration`.

In order to get it, several operations in fine bins must be applied, in a way quite similar
5.3. Main changes implemented in MARS

to what is done for the calculation of the Effective Collection Area. Concerning the case of no AZ dependence, the fine bin calculation has mainly the aim to use the basic weights for the calculation of the proper values of the Migration Matrix in coarse bins. These weights are indeed carrying also the information of the assumed spectrum of the source and of the observation time spent in each bin of ZA covered by the source. For each couple of indexes \((i, j)\), where the index \(i\) refers to the true energy fine bins and the index \(j\) to the ZA fine bins, we define the weights as

\[
W_{ij} = C_{ij} \cdot F_i \cdot T_j \tag{5.19}
\]

where \(C_{ij}\) is the Effective Collection Area in fine bins, \(F_i\) is the integral in the \(i\)-th fine bin of the assumed spectrum \(F(E)\) chosen by the analyzer and \(T_j\) is the effective time of the given source observation for the \(j\)-th fine bin. For each couple of indexes \((I, J)\), where the index \(I\) refers to the true energy coarse bins and the index \(J\) to the ZA coarse bins, the covariance matrix between the weights and their normalized values are also calculated, taking into account all the \(i \in I\) and \(j \in J\): the number of calculated covariance matrixes are \(I \times J\) and each one has a dimension equal to \((i \in I) \times (j \in J)\). These matrixes are used to compute the proper errors for the final normalized energy Migration Matrix in \((E_r, E, \theta)\) in coarse bins, performing a loop over the estimated energy coarse bins.

The algorithm explained above needs to use 3D histograms in estimated energy, true energy and ZA (all these parameters must be considered), so by adding the AZ we would need a further dimension. For technical reasons, this was a problem, since in in ROOT it is not possible to define a histograms with a dimension higher than 3. Therefore, the implementation of the AZ dependence required to pass from the usage of a histogram, to the definition of a N-tuple (a sort of matrix of numbers) in \((E_r, E, \phi, \theta)\). These N-tuples are used to fill an array of 3D histograms in \((E_r, E, \theta)\). Each element of the array is then a 3D histogram, containing the parameters of the events within a certain AZ range, as defined by the corresponding \(k\)-th AZ fine bin. Therefore the dimension of the array is given by the number of the AZ fine bins set by the users. In this way it is possible to use simply a loop over the AZ fine bins and to calculate the proper weights with also the AZ dependence of the observation:

\[
W_{ikj} = C_{ikj} \cdot F_i \cdot T_{kj} \tag{5.20}
\]

In the end of the calculation, the Migration Matrix is expressed as a function of \((E_r, E, \theta)\) in coarse bins, as it was in the case without taking the AZ into consideration \(^5\). Note that the N-tuple is stored, since this object is required by the subsequent unfolding procedure. In Fig. 5.20 three projections along the ZA coarse bins of three different Migration Matrices are shown. All are related to the Crab Nebula sample already mentioned in the previous subsections (see also section 5.4). In particular since only 1 ZA coarse bin is used, the three projections can be thought as the final Migration Matrices in coarse bin. The first plot represents the case of no AZ dependence algorithm, whereas the other two are calculated with the new AZ dependence algorithm, respectively with 1 and 6 AZ fine bins. In case only 1 AZ fine bin is set, the final Migration Matrix is equal to the case of no AZ dependence (a part from differences of the order of \(10^{-5}\) related to the different algorithm for the calculation of the Effective Collection Area in fine bins with AZ dependence). The means of the three distributions agrees very well and the energy dispersion seems to be slightly smaller when the 6 AZ fine bins are used. However, this feature can be probably related to the fact that when more than 1 AZ fine bin are used, the available MC-\(\gamma\) statistics is reduced. This is due to the

\(^5\)This situation is similar to the implementation of the AZ dependence in the Effective Collection Area calculation in coarse bins.
fact that all the MC-\(\gamma\) events with AZ values outside the AZ bins covered by the real Crab Nebula data are no longer taken into account since their AZ weights are equal to zero (and thus also the overall weights). Actually, the errors of the Migration Matrix values (which are not drawn) for the case of no AZ dependence (and 1 AZ fine bin case), are smaller than those of the 6 AZ fine bin case. Further studies on the energy resolution as a function of the AZ parameter are foreseen.

5.3.4 Changes in the executable for the calculation of the flux

The main changes performed in the executable FLUXLC have been essentially related to the proper call of the classes modified for taking into account the AZ dependence (see table 5.4) and to the drawing of the new displays showing the AZ dependent quantities. In the following section, which concerns the test of the new implementations on a MAGIC-I Crab Nebula sample, the new displays will be shown. A new flag in the inputcard of the FLUXLC program (\textit{FluxLC.finebinsphi}) has also been added in order to let the analyzer to chose the number of AZ fine bins.

5.4 Tests on MAGIC-I Crab Nebula data

In order to check all the changes done in MARS for the implementation of the AZ parameter, we performed an analysis of a sample of Crab Nebula source observed in WOBBLE mode by MAGIC-I\(^6\). First of all, we needed to get a WOBBLE MC-\(\gamma\) sample for MAGIC-I simulated with all the AZ values \([0^\circ, 360^\circ]\) (since, as already mentioned, the standard MC-\(\gamma\) production for MAGIC-I accounts only for two AZ values). This sample (provided us by our colleague J. Sitarek) has the following main features:

- The energy ranges from 30 GeV to 30 TeV. The energy distribution is a pure power law with a spectral index of -1.6.
- The ZA ranges from 10° to 45°.
- The AZ covers the whole circle, from 0° to 360°.
- The maximum IP was fixed to 500 m.
- The simulated optical point spread function (PSF) was set to 10.6 mm \(^7\).
- The observation mode is WOBBLE, i.e. the nominal source position is always placed 0.4° far from the center of the camera.
- The overall number of simulated \(\gamma\)–ray events is \(2.1 \times 10^6\). The statistics is actually quite small, nevertheless this MC-\(\gamma\) events turned out to be sufficient in order to accomplished the tests.
- The adopted cleaning is the Time-constrained 6-3 (see section 4.3.3)

\(^6\)At the time this test was carried out no Crab Nebula data observed in stereoscopic mode were available. Nevertheless, the current analysis chain for the stereo data fairly account for the Effective Collection Area calculation with the AZ dependence.

\(^7\)Actually this value for the MC-\(\gamma\) PSF underestimates the real values of the telescope PSF of the real Crab Nebula sample here considered (which belong to the range from 12.5 mm to 14.5mm). Nevertheless, since here we want only to demonstrate that the new code is working properly, it should be not considered as a compromising problem.
The Crab Nebula sample was selected from all the WOBBLE data taken between October 2008 and March 2009. The data were analyzed up to the image cleaning as explained in chapter 4. Then, the following criteria were applied to the Crab Nebula subruns in order to get a good data sample, according also to the characteristics of the MC-\(\gamma\) files available for the analysis:

- Rate within the 20\% of the nominal ZA dependent rate value given by
  \[ \text{Rate} \ [\text{Hz}] = 250 \times \cos(\theta)^{0.5} \]  
  (5.21)

- ZA ranges from 10\° to 45\°.

- Mean values of cloudiness lower than 40\%.

- Only the AZ branch with values greater than 180\° (in drive reference frame) was taken into account (note that for a given source observation and for a given ZA value there are two possible AZ values, see Fig. 5.6(a)). This choice was taken because the surviving runs with AZ below 180\° were too few and belonging only to ZA smaller than 20\° (see Fig. 5.13).

![Figure 5.13: ZA vs AZ distribution of the events surviving to the selection criteria applied to get the final selected Crab Nebula sample. The AZ branch below 180° was not taken into account in the following analysis.](image)

11 hours of good Crab Nebula data were selected. In Fig. 5.14 the ZA and AZ distributions of the sample are reported. Before starting the analysis, we decided to split the Crab Nebula sample in different ZA bins: we used the same Zbins criteria applied in the previous sections\(^8\), therefore 5 subsamples were created. The \(\gamma\)-hadron separation and the energy estimation were performed in each subsample according to the standard analysis explained in section 4.3. For each subsample, the proper ZA cuts were obviously applied. The MC-\(\gamma\) data were equally divided in two subsamples in order to get the training and testing MC-\(\gamma\) samples.

The main checks we did with this Crab Nebula data samples were the comparison between the results achieved using the executable FLUXLC and the subsequent UNFOLDING (for

\(^8\)Only Zbin1 has a different ZA range, since here the minimum ZA is equal to 10\°.\)
deriving the spectra in true energy) with the MARS version without the AZ parameter implementation and the results we could get with the new version of MARS with the AZ parameter implementation\(^9\). Imposing, in the new MARS version, the number of AZ fine bins to be 1, the results should be equal to those found with the standard MARS version. The setting used for the FLUXLC inputcards are the standard ones normally used for the WOBBLE data. In table 5.5 the main settings are reported. Concerning the unfolding process we tested successfully the different methods (Tikhonov, Bertero and Schmelling, see section 4.3.11). For each subsamples, we also performed an analysis with a number of AZ fine bin equal to 6. In this way also the nontrivial case with AZ fine bins greater than 1 was tested and found out to work properly.

It is worth mentioning that, during the Hadronness and Estimated Energy calculations, the AZ values of the real data and MC-\(\gamma\) ones were not taken into account, since the possibility to use the AZ parameter as a discriminator parameter in the MARS program OSTERIA is not yet implemented. Indeed, the implementation in OSTERIA of the AZ as a possible discriminator parameter would be an interesting issue to investigate, and it is foreseen for the next feature. For all the 5 Zbin subsamples, a complete analysis (using both source dependent (ALPHA approach) and source independent (\(\theta^2\) approach) analyses) was thus performed. In Figures from 5.16 to 5.21 the main output plots are reported for the case of no implemented AZ dependence and for the case of implemented AZ dependence (for both 1 and 6 AZ fine bins). All the plots refer only to the Crab Nebula Zbin2 subsample (2.9 hours), since the other subsamples showed the same generally results.

The MARS version with the AZ implementation gave the right expected results: in case of only 1 AZ fine bin, the achieved results are practically identical (for both the values and the corresponding errors) to those achieved by the version with no AZ dependence, the small differences being due only to the different algorithm for the computation of the Effective Collection Area in the two software versions (see section 5.3.1).

In the nontrivial case of 6 AZ fine bin, where the AZ is actually used to re-weight the Effective Collection Area calculation through the AZ values of the real data, the results are slightly different but perfectly compatible within the errors with the standard ones.

These general results were obtained for all the 5 Zbin subsamples. Even if not shown in the

\(^9\)This comparative analysis is not expected to show improvements of the performances of the analysis, but only to test the consistence of the changes involved the Azimuth parameter and to show some new implemented output plots.
5.4. Tests on MAGIC-I Crab Nebula data

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Table 5.5: Standard settings applied to the FLUXLC input cards used for the analysis. The ZA coarse bin was always set to be 1. The ZA ranges were selected according to the corresponding Zbin sample. The last row shows the factor applied for the ratio between the number of used estimated energy coarse bin and the number of true energy coarse bin. In case of the analysis performed with the MARS version with the AZ implementation, there is a further flag allowing to set the number of AZ fine bin (FluxLC.finebinsphi).

In the figures, we have also verified that the full source independent analysis was working properly. We report in the following the resulting power law fits between 200 GeV and 2 TeV of the Crab Nebula spectrum before and after the unfolding procedure, for the case of the Zbin2 subsample (i.e., the subsample considered for the Figures from 5.16 to 5.21) which demonstrate the right behavior of the implementation of the Azimuth dependence in the calculation of the Effective Collection Area:

- Resulting fits found with the MARS version with no AZ dependence:
  Not-unfolded spectrum fit:
  \[
  \frac{dF}{dE} = (2.7 \pm 0.2_{\text{stat}}) \times 10^{-11} \times (E/\text{TeV})^{-2.47\pm0.12_{\text{stat}}}\text{[TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}] \\
  \chi^2/\text{NDF} = 5/6
  \]  
  (5.22)

  Unfolded (Tikhonov method) spectrum fit:
  \[
  \frac{dF}{dE} = (2.4 \pm 0.3_{\text{stat}}) \times 10^{-11} \times (E/\text{TeV})^{-2.54\pm0.14_{\text{stat}}}\text{[TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}] \\
  \chi^2/\text{NDF} = 0.05/2
  \]  
  (5.23)

- Resulting fits found with the MARS version with AZ dependence and 1 AZ fine bin:
\[ \frac{dF}{dE} = (2.7 \pm 0.3_{\text{stat}}) \times 10^{-11} \times (E/\text{TeV})^{-2.47^{+0.11}_{-0.11}\text{stat}} [\text{TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}] \]

\[ \chi^2/\text{NDF} = 5/6 \quad (5.24) \]

Unfolded (Tikhonov method) spectrum fit:

\[ \frac{dF}{dE} = (2.4 \pm 0.3_{\text{stat}}) \times 10^{-11} \times (E/\text{TeV})^{-2.54_{-0.14}^{+0.14}\text{stat}} [\text{TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}] \]

\[ \chi^2/\text{NDF} = 0.05/2 \quad (5.25) \]

• Resulting fits found with the MARS version with AZ dependence and 6 AZ fine bins:

\[ \frac{dF}{dE} = (2.7 \pm 0.2_{\text{stat}}) \times 10^{-11} \times (E/\text{TeV})^{-2.46_{-0.13}^{+0.13}_{\text{stat}}} [\text{TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}] \]

\[ \chi^2/\text{NDF} = 5/6 \quad (5.26) \]

Unfolded (Tikhonov method) spectrum fit:

\[ \frac{dF}{dE} = (2.1 \pm 0.2_{\text{stat}}) \times 10^{-11} \times (E/\text{TeV})^{-2.58_{-0.11}^{+0.11}_{\text{stat}}} [\text{TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}] \]

\[ \chi^2/\text{NDF} = 1.00/2 \quad (5.27) \]
5.4. Tests on MAGIC-I Crab Nebula data

Figure 5.15: (a) ZA vs UTC plot computed by the class *MThetaTime* before the AZ dependence implementation. (b) ZA vs AZ vs UTC plot computed by the class *MPhiThetaTime* after the AZ dependence implementation (and 1 AZ fine bin). (c) ZA vs AZ vs UTC plot computed by the class *MPhiThetaTime* after the AZ dependence implementation (and 6 AZ fine bins).
Figure 5.16: *MHEffectiveOnTime* class plots for the Zbin2 Crab Nebula subsample obtained before the AZ dependence implementation.

Figure 5.17: *MHEffectiveOnTime* class plots for the Zbin2 Crab Nebula subsample obtained after the AZ dependence implementation and with 1 AZ fine bin.
Figure 5.18: *MHEffectiveOnTime* class plots for the Zbin2 Crab Nebula subsample obtained after the AZ dependence implementation and with 6 AZ fine bins. Note (from figures 5.16, 5.17 and 5.18) that the two effective observation time values, calculated as a function of the UTC (left plot at the bottom) and as a function of the pointing position (right plot at the bottom in Fig. 5.16, and center plot at the bottom, for figures 5.17 and 5.18), are the same for all the three different analyses (standard, 1 AZ fine bin and 6 AZ fine bin). (The former value is $10374.0 \pm 22.3$ s, the latter $10327.0 \pm 22.1$ s.)
Azimuth Implementation in the Effective Collection Area Calculation

Figure 5.19: (a) Effective Collection Area as a function of the true energy in coarse bins computed by the class \textit{MHMC\textunderscore Collection\textunderscore Area} before the AZ dependence implementation. (b) Effective Collection Area as a function of the true energy in coarse bins computed by the class \textit{MHMC\textunderscore Collection\textunderscore Area} after the AZ dependence implementation (and 1 AZ fine bin). (c) Effective Collection Area as a function of the true energy in coarse bins computed by the class \textit{MHMC\textunderscore Collection\textunderscore Area} after the AZ dependence implementation (and 6 AZ fine bins). The first two plots are almost identical, whereas the Effective Collection Area calculated with 6 AZ fine bins is compatible with the other plots but has higher errors since the Azimuth weights are actually used.
5.4. Tests on MAGIC-I Crab Nebula data

Figure 5.20: (a) Energy Migration Matrix computed by the class \textit{MHMcEnergyMigration} before the AZ dependence implementation. (b) Energy Migration Matrix computed by the class \textit{MHMcEnergyMigration} after the AZ dependence implementation (and 1 AZ fine bin). (c) Energy Migration Matrix computed by the class \textit{MHMcEnergyMigration} after the AZ dependence implementation (and 6 AZ fine bins).
Figure 5.21: (a) Crab Nebula spectrum (not unfolded) calculated by the MARS version before the AZ dependence implementation. (b) Crab Nebula spectrum (not unfolded) calculated by the MARS version after the AZ dependence implementation (and 1 AZ fine bin). (c) Crab Nebula spectrum (not unfolded) calculated by the MARS version after the AZ dependence implementation (and 6 AZ fine bins). The data refers to the Zbin2 Crab Nebula sample.
5.5 Discussion

As reported in [155, 157], the GF affects the shape and the orientation of the shower images in a complex way, giving different effects for each different energy and coming direction of the \( \gamma \)-ray showers. The MAGIC telescopes’ sensitivity makes these effects not negligible such to require the work made and reported in this chapter.

Nowadays, in light of the MAGIC Phase II, a further reason for the need of the AZ as a new degree of freedom of the analysis has been the incoming change from the single telescope observations to the new stereoscopic ones. Indeed, this new stereo data taking procedure combines the effect of the telescopes’ array fixed configuration with the GF effect in a non-trivial way.

The intent of our work has been to implement in the software the possibility to consider the Azimuth information of the observations, and to give a former look to the resulting effects, at the level of the Effective Collection Area and of the subsequent spectrum estimation. Just giving the software the information of the relative Azimuth of the observation, it is so possible to take into account at the Effective Collection Area level the right AZ to consider in order to have a better matching between the real data and the MC-\( \gamma \) simulation. Moreover, this new introduction allows to estimate the Effective Collection Area area in case of stereoscopic observations, where, at the level of the stereoscopic trigger (Level 3 Trigger), we have a selection of the events depending on the Azimuth.

From this work we have obtained, as shown in Fig. 5.9, the Effective Collection Area of the stereoscopic system as a function of the Azimuth angle for different Zenith bins and energy bins, by using the first version of stereo MC-\( \gamma \) simulation, available at the time of this work. We could see that the Effective Collection Area depends in a nontrivial way on the Azimuth, as a function also of the energy range and the Zenith position. Nevertheless, giving a look to the behavior of these collection areas, the symmetry around 180° in Azimuth would emphasize that the main contribution of this dependence is given by the GF (at least for ZAs below 45° and for the applied quality cuts). Indeed, the fixed configuration of stereoscopic system (that is the position of the two telescopes), would have shown clearly some kind of asymmetry with respect to the AZ in those plots. This result does not exclude the possibility that the configuration effect contribution could be stronger at higher Zenith angle of observation. This study will be made as soon as the MC-\( \gamma \) simulation at these Zenith values (i.e. higher than 45°) will be produced.

In any case, thanks to our implementation of the Azimuth in the code, it is now possible to directly look at the global resulting effect, obtained by the combination of the two contributions (i.e., the GF effect plus the telescopes’ configuration effect), which would be otherwise quite difficult to evaluate.

Concerning the flux estimation, we made several tests to check the proper functioning of the analysis chain. Since at the moment of the work reported in this chapter no stereoscopic data were available, we used a sample of MAGIC-I data from the Crab Nebula, and we studied the changing in the FLUXLC outputs, under different configurations of the AZ binning. In particular, by changing the number of AZ bins, we obtained results in agreement with what expected. For example, the case with 1 single AZ fine bin reproduced almost exactly the analysis with the old version of MARS with no Azimuth information. Also by increasing the number of AZ bins, we have obtained results in agreement with those expected, even if in this case the error bars (for example those of the Effective Collection Area in coarse bins) are larger, due to the reduction of simulated events given by the azimuthal weighting. This suggests that a larger amount of MC-\( \gamma \) data is needed, in order to efficiently use the Azimuth information in our analysis.
A possible solution might be the production of a dedicated MC-γ simulation for each observed source, where the MC-γ sample follows the same ZAs and AZs of the observation itself, in order to increase the statistics of only the events we are interested in. The new code presented in this chapter, would then give the proper weights to that simulation, according to the effective time of the observation in each value of ZA and AZ for the global data taking. On the other hand we have to consider that this idea in any case would require a dedicated production for each source, a issue which is not so trivial to carry out. After minor changes in some of the MARS classes involved in the source independent analysis, we also concluded that the software is working properly, both at the level of the spectra obtained by the usage of the $\theta^2$ approach, and at the level of the skymap. All these tests endorse that the implementation of the AZ dependence of the Effective Collection Area is properly working and is ready to be used for the stereoscopic observations. As shown in this chapter, the Azimuth parameter was never used in the $\gamma$-hadron separation and energy estimation algorithms. Further studies are needed to understand if the implementation of the AZ dependence in the $\gamma$/hadron separation and energy reconstruction procedures can give some improvements in terms of sensitivity and smaller systematic errors. If on one side, the study of the effects obtained by introducing the Azimuth as input parameter for the RF algorithm for the estimated energy calculation is in principle pretty simple, on the other side, in case of the $\gamma$/hadron separation, the correct use of the Azimuth information needs to properly weight the Azimuth distribution of our simulation, in order to reproduce the Azimuth distribution of the real data observation, otherwise we could introduce a strong bias in the analysis (re-azimuthing). This problem is well known and has been considered already for the case of ZENITH (the so-called re-zenithing) and SIZE (the so-called re-sizing) parameters, as explained in section 4.3.5.
Nowadays there are compelling experimental evidences for a large non-baryonic component of the matter density of the Universe at all observed astrophysical scales, such as galaxies, galaxy clusters and cosmic background radiation. This matter, which accounts for about 23% of the energy budget of the Universe, makes its presence known through gravitational effects and could be made of so far undetected relic particles from the Big Bang.

Although plenty of experimental and theoretical efforts have taken place so far and despite recent exciting and controversial results which can be interpreted as possible Dark Matter (DM) detection [159–163], the nature of DM has not yet been clarified, keeping open one of the most important fundamental questions in modern physics.

In this chapter, the DM paradigm will be briefly introduced. The major experimental evidences will be presented together with the principal particle candidates among the huge plethora proposed in literature, with particular emphasis to the supersymmetric neutralino. A brief review of the experimental techniques for DM detection will be introduced, in particular for indirect DM searches in γ-rays, which can be suitably carried out by ground-based experiments, such as the IACTs, and satellite experiments, as Fermi–LAT. The places in the Universe where large amounts of DM is expected will be also shortly discussed.

A recent and comprehensive review on DM physics, with a vast bibliography, describing the most compelling evidences coming from different research areas obtained with different experimental techniques and introducing the most important particle candidates and experimental results, can be found in Bertone et al. [60].

6.1 Introduction

Despite earlier evidences of an amount of invisible matter based on the study of the mass estimated by star luminosity of the galaxy Andromeda (known also as M31, at 725 kpc from the Milky Way) [164] and of the dynamical density of matter present in the solar neighbourhood [165], the modern meaning of DM was firstly adopted by Zwicky in the 1930s in his work on the dynamics of the Coma galaxy cluster [166]. By studying the motion of galaxies in the cluster and using the virial theorem (assuming that the galactic motion was virialized) he determined the mass distribution in the cluster and reported that the matter associated with luminous objects was not accounting for the total mass as calculated from the gravitational balance through the virial theorem. Zwicky called it “missing mass”. In his calculations, the
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mass-to-light ratio (M/L) of the system should have been of the order of $400 \, M_\odot/L_\odot$, where $M_\odot$ and $L_\odot$ are respectively the mass of the Sun and its luminosity.

These early studies did not awake the attention of the astronomical community for decades until, in the late 1970s, strong evidences that spiral galaxies contain a substantial component of un-associated form of mass [167, 168] were accepted by the scientific community. Since then, additional astrophysical evidences of the presence of some mass excess with respect to the visible part have accumulated with independent experimental techniques at all scales, from cosmological distances and all the way down to galactic scales.

The most accepted theoretical foundation of a DM dominated Universe was provided by Blumenthal et al. in 1984 [169]. Studying the formation processes of galaxies, clusters and filamentary superclusters, the hypothesis of non-baryonic, cold DM (CDM) was established against the so-called hot DM (HDM) scenario. CDM is indeed needed to explain how galaxies form from the very smooth (one part in $10^5$) baryonic matter distribution deduced from the isotropic microwave background radiation at high redshift ($z \approx 10^3$).

The most recent years have been devoted to a detailed elaboration of the concept of the CDM dominated Universe, i.e. to determine how much DM exists in the Universe, how it is distributed in galaxies and clusters, and to attempt to constrain its nature, both from the experimental and theoretical sides. There is an very large amount of literature on experimental techniques, candidates and proposed solutions. In addition, the discovery (using Type I supernova data [170] as well as precise mapping of the Cosmic Microwave Background radiation [171, 172]) that a new unknown kind of energy (the Dark Energy) is contributing to the overall energy amount of the Universe for $\sim 70\%$ and is responsible for the observed accelerated expansion of the Universe, has led to the widely accepted $\Lambda$CDM cosmological scenario, whose main theoretical basis are explained e.g. in [173].

So far, despite huge efforts have taken place both from the theoretical and experimental point of view, the nature of DM is still unknown and represents one of the most important fundamental questions in modern physics.

6.2 Main Dark Matter Evidences

In the next subsection the most important experimental evidences of DM are briefly summarized.

6.2.1 Evidences from the dynamics of spiral galaxies and galaxy clusters

One of the most classical and direct evidence for DM comes from the observation of the rotation curves of spiral galaxies, i.e. the circular velocity of star and gas as a function of the distance to their galactic center $r$. The rotation curves of the galaxies can be obtained by combining observations of neutral Hydrogen 21cm-line with optical surface photometry. By using Kepler’s law, the total mass $M(r) = 4\pi \int \rho(r)r^2 dr$ enclosed within a radius $r$ from the center and the rotation speed at this radius $v(r)$ are related by

$$v(r) \propto \sqrt{\frac{M(r)G_N}{r}}$$

where $G_N$ is Newton’s constant. If only the visible mass is present in galaxies, then, in the outer region, outside the radius within which this luminous contribution is distributed, the velocity should fall as $\sqrt{1/r}$, since moving to larger distances no more mass is included.
However, this expectation is not supported by observations [174]. In particular, rotation curves show an unexpected flat behavior beyond the edge of visible stars (see Fig. 6.1), indicating that a nearly spherical “halo” of DM, in which the galactic disk is embedded, extends farther than the visible component, with $M(r) \propto r$ and $\rho(r) \propto 1/r^2$.

After 30 years, the study of rotational curve is still a very active field. A large amount of data is already available for a large number of objects. A compilation of more than 1000 rotational curves can be found for example in [175].

Strong indications of DM halos come also from the dynamic studies of clusters of galaxies, which are the largest and most massive gravitationally bound systems in the Universe, with radii of the order of the Mpc and total masses around $10^{14} - 10^{15} M_\odot$ [177]. When studying the dynamics of the clusters, the mass inferred by photometry is hundreds times smaller than the virial mass used to fit the observed velocities of the galaxies. The experimental data suggest that in these large systems most of the mass resides in a relatively smooth DM halo with substructures and filaments [178].

### 6.2.2 Gravitational Lensing

In General Relativity, the light-rays travel into geodesic and can be deflected when they propagate through gravitational field which create local curvature of the space-time. Starting from the 1980s, a flourishing field has grown around the studies of the images of background astrophysical objects distorted by foreground energy-mass distributions (such as galaxies and galaxy clusters). These foreground mass distributions (deflectors) act on the light in the same way of a lens, and both displacements and deflections can be used to infer properties on the mass of the deflectors.

Gravitational lensing is nowadays a powerful tool for several astrophysical and cosmological studies [179]. The entity of the distortions of the images leads to different regimes of gravitational lensing. Strong gravitational lensing occurs when multiple images of the same
background object are created and evident arcs are present (see Fig. 6.2). Conversely, when weakly coherent distortion of single images of a large number of background faint galaxies (the so-called Arclets) occurs the gravitational lensing is in its weak regime, for which a significant signal can be obtained only statistically from the analysis of a large number of images. The gravitational lensing permits to perform mass distribution measurements, to constrain the mass concentrations and also to estimate cosmological parameters. In particular, both strong and weak lensing phenomena have been observed for clusters of galaxies. The distortion of the images of background objects due to the gravitational mass of a cluster can be used to infer the shape of the potential well and thus the overall mass of the clusters. The experimental results show that the mass of the intercluster gas (which account for \(~85\%\) of the baryonic mass of the cluster and can be deduced by X-ray measurements [180]) and the mass of the galaxies and stars belonging to the cluster (\(~15\%\) of the baryonic mass) can account only for \(~10\%\) of the total mass derived from gravitational lensing. Thus, about \(~90\%\) of the mass of the clusters are composed by DM [181].

![Gravitational Lens in Abell 2218](image)

**Figure 6.2:** Strong gravitational lensing as observed by the Hubble Space Telescope in Abell 1689. The studies of the distorted images and the arcs indicates the presence of DM.

### 6.2.3 Evidences from CMB and Large Scale Structure

The amount of DM present on cosmological scales has recently been accurately measured by the WMAP mission [182] through the analysis of the Cosmic Microwave Background (CMB) anisotropies.

In the early Universe, before the recombination of electrons and nuclei into neutral atoms (a process called recombination), photons were tightly coupled to the particles. With the cooling down of the Universe during the expansion the photons decoupled from the matter and started their free travel through the Universe. Due to the expansion of the Universe these photons have propagated till us becoming cooler and cooler. Today, this relic radiation (i.e. the CMB) follows a perfect blackbody spectrum peaked at \(T \approx 2.73^\circ\) K, isotropically distributed over the sky within fluctuations of the order of \(\delta T / T \approx 10^{-5}\) (see Fig. 6.3). The properties of these photons carry a record of the conditions at the time of decoupling. Since the initial mass fluctuations are conserved during inflation [183] they are still observable in the tiny fluctuations of the CMB.
6.2. Main Dark Matter Evidences

Figure 6.3: The Cosmic Microwave Background full-sky measured by the WMAP mission [182]. Fluctuations are of the order of $10^{-5}$, the radiation is a perfect black-body emission peaked at $2.73 ^\circ$ K. Image taken from http://map.gsfc.nasa.gov/media/080997/index.html.

Today, the analysis of CMB anisotropies enables accurate testing of cosmological models and puts stringent constraints on cosmological parameters. The observed temperature anisotropies in the sky can be expanded as

$$\frac{\delta T}{T}(\theta, \phi) = \sum_{l=2}^{\infty} \sum_{m=-l}^{+l} a_{lm} Y_{lm}(\theta, \phi)$$

(6.2)

where $Y_{lm}(\theta, \phi)$ are spherical harmonics. The variance $C_l$ of $a_{lm}$ is given by

$$C_l \equiv \langle |a_{lm}|^2 \rangle = \frac{1}{2l+1} \sum_{m=-l}^{+l} |a_{lm}|^2$$

(6.3)

The acoustic vibration modes of the coupled photon-matter fluid of the early Universe appear as peaks in the angular power spectrum shown in Fig 6.4. The first three acoustic peaks are well determined, allowing an overall fit in the multiparameter space of cosmological parameters [172]. In particular, the location of the first peak probes the spatial geometry, while the relative heights of the peaks probes the baryon density. The achieved results are consistent with the $\Lambda$CDM cosmological model and an overall fit in the multiparameter space has provided accurate determinations on many of the involved parameters.

The contribution of the $i$-th component of the Universe is expressed in terms of the ratio $\Omega_i$ between the relative density $\rho_i$ and the critical density $\rho_c = 3H^2/8\pi G_N$ (where $H$ is the Hubble’s constant) which today has a value of $\sim 1.38 \times 10^{-7} \text{ M}_\odot \text{ pc}^{-3}$ [173]. The precise measured position of the location of the first peak yields a total density $\Omega_{\text{total}} = 1.02 \pm 0.02$ (i.e. compatible with a flat geometry), whereas the best-fit parameters for the abundance of baryons and matter are

$$\Omega_b h^2 = 0.023 \pm 0.001 \quad \Omega_m h^2 = 0.14 \pm 0.02$$

(6.4)

The derived baryon density is consistent with that coming from Big Bang nucleosynthesis [72]. These results imply the need for both DM and Dark Energy, with no evidence for the dynamics of Dark Energy (i.e, consistent with a pure cosmological constant).
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Therefore, the CMB anisotropies results show that today the energy budget of the Universe is composed for 72% of an exotic form of energy, i.e. the Dark Energy, for 23% of an unknown non-baryonic DM component and for only 4.6% of visible and invisible baryonic matter (see Fig.6.5). The non-baryonic DM accounts thus for 83% of the total matter in the Universe.

Besides the cosmological measurement provided by CMB studies, the matter distribution in the Universe can be used to constrain the cosmological parameters. Indeed, the growth of structures is thought to be governed by the DM. The entire history of the Cosmos in all its complexity, as governed by a handful of cosmological parameters, imprints its influence into the large scale structure, from which we can derive their values.

Two huge galaxy surveys have been taking place: the Sloan Digital Sky Survey (SDSS) [184] (which is still operating) and the 2-degree Field Galaxy Survey (2dFGRS) [185] (whose activity ended in 2002). Both large astronomical surveys measured the angular position, the distance and the luminosity of galaxies and galaxy clusters, yielding a 3-dimensional map of the visible Universe which can be used to extract the 3-dimensional power spectrum of matter fluctuations. The power spectrum reflects the distribution of matter at different epochs, leading to the extraction of several pieces of information on the cosmological parameters, e.g.

Figure 6.4: Angular power spectrum of anisotropies in the CMB radiation. Data refers to the 5th data release from WMAP [172]. Image taken from http://map.gsfc.nasa.gov/news/index.html.

Figure 6.5: The relative contribution of matter and radiation fields at the recombination epoch (left) and today (right).
6.2. Main Dark Matter Evidences

the ratio of the baryon to matter density and also the matter density itself. The information coming from the CMB analysis and the power spectrum of matter fluctuations are combined to narrow down the uncertainty on the determination of the cosmological parameters. An estimate of the cosmological parameters combining the SDSS (using over 200,000 galaxies) and WMAP measurements can be found in [186].

6.2.4 The Bullet Cluster

In the galaxy clusters the luminous galaxies, the interstellar medium and the DM halo, are characterized by a mass distribution that peaks in the same point at the center of the cluster, since the clusters generally present a well-relaxed configuration. In case two clusters are colliding, this is no longer true: in that case the galaxy population and the DM halo would react in a similar way as a collisionless fluid, almost not influenced by the collision, whereas the interstellar medium (hot plasma that emits X-rays) experiences ram pressure and its distribution is strongly influenced.

An extremely exciting result of recent years, sometimes considered the first direct evidence of DM, is the observation of spatial segregation between the interstellar X-ray emitting plasma and the weak lensing reconstruction total mass distribution in two compenetrating galaxy clusters in the merging system 1E0657-558 at a redshift $z = 0.296$ (the so-called Bullet Cluster). The mass surface density of the system has been inferred by weak gravitational lensing analysis [187]1. As shown in Fig. 6.6, the luminous component of the clusters almost follows the overall mass distribution of the colliding clusters, whereas the hot plasma mass distributions (which account for most of the baryonic content of the systems) are separated (with a significance of 8 and 12σ for the two merging clusters) from the source of the gravitational potential detected through gravitational lensing. This provides a strong indication in favour of the DM scenario, since, if no DM is present, the mass distribution peak should be almost coincident with the hot X-ray emitting plasma which contributes for $\sim$85% of the total baryonic matter in the clusters. The Bullet Cluster data are also one of the main argument used to rule out theories of modified gravity, although some groups claimed the possibility to explain the Bullet Cluster experimental data with particular alternative theories of gravitation without the need of a non-baryonic DM component [189,190].

6.2.5 Recent controversial evidences

Annual modulation of DAMA data

The DAMA experiment at Gran Sasso laboratories published in 2008 evidences of regular modulations signals. The results have been interpreted as the modulation of the diffuse DM halo of our galaxy (which local density, at our galactocentric distance $R = 8.5$ kpc, is estimated to be in the range $\rho_0 \simeq 0.2 - 0.5$ GeV cm$^{-3}$) due to combined motion of the Sun around the galactic center and the annual Earth motion around the Sun [159]. However, despite the sensitivity of this result is very high, the DM interpretation is in contradiction with the negative detection of other similar underground experiment. Future experiments, in particular with cryogenic noble gas detector, will soon clarify this claim.

1The results were subsequently confirmed by an independent analysis based on both strong and weak gravitational lensing [188].
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Figure 6.6: (a) Color image in the visible range from the Magellan telescope of the 1E0657-558 cluster. (b) A 500 ks Chandra image of the same cluster. The green contours in the pictures correspond to weak lensing maps marking the gravitational potential. The white contours show the errors on the positions on the peaks of the mass distribution deduced by the weak lensing analysis and correspond to 68.3%, 95.5% and 99.8% confidence levels. The blue crosses in the left picture locate the mass peak of the measured baryonic plasma clouds which are shown in bright colors in the right picture. The white line on the right bottom indicates 200 kpc size scale. Images taken from [187].

Figure 6.7: (left) The scheme of the DM signal modulation in the annual motion of the Earth around the Sun. (right) The modulation in the range 2–6 keV observed in DAMA results.

Positron excess in PAMELA data

The PAMELA satellite experiment reported in 2008 an excess on the positron abundance in the cosmic radiation in the energy range 10 – 100 GeV [160] as shown in Fig.6.8(a) (note that up to ~10 GeV the results of PAMELA are consistent with the earlier experiments taking into consideration the solar modulations) compared to cosmic-rays (CRs) diffusion model predictions, basically obtained with the GALPROP code [191]. This results has been considered as the first indirect evidence of DM particle annihilation, since the data deviate significantly from predictions of secondary production models based on the interactions between CR nuclei and interstellar matter. However, these excess of positron abundance could be explained as positron production from nearby pulsars.

Electron–Positron excess in ATIC, Fermi–LAT and H.E.S.S. data

The Advance Thin Ionization Calorimeter (ATIC) Collaboration, reported an excess of electrons–positrons in the energy range 300-800 GeV than that expected from the diffuse electron-
6.3. Dark Matter Candidates

As shown, the evidence for non-baryonic DM is compelling at all observed astrophysical scales. Since last decades, a wide range of DM candidates has been proposed in the literature (see [195–197] for recent reviews). Some of them have an \textit{ad hoc} nature, i.e. they were created specifically to solve the DM problem. Others were instead developed independently to solve theoretical problems of the Standard Model of Particle Physics (SM), but they provided, in a natural way, good candidates for DM. Different candidates can have very different characteristics, with masses that go from the sub-eV scale to values larger than the TeV scale.

Experimental data and phenomenological considerations are available and can constrain to some extent the properties of a good DM candidate. Indeed, we know that the non-baryonic particles related to DM should be long-lived enough for being present in the Universe since their decoupling at the early Big Bang epoch (period in which we assume they were created). They should be weakly interacting to ordinary matter and neutral (both for electrical charge...
and color) and they should be within the limits posed by astrophysics experiments results. A recent review of those experimental constraints are given in a “ten-point test” described in [198], where the authors propose a list of 10 characteristics that a good DM candidate should have.

So far, the most studied DM candidates are the so-called Weakly Interacting Massive Particles or WIMPs. A WIMP is a generic, stable particle with a mass of the order of the GeV-TeV and a very weak cross section with SM particles ($\sigma v \approx 10^{-24} - 10^{-28} \text{ cm}^3 \text{ s}^{-1}$) that has been in thermal equilibrium with the primordial photons in the early Universe. The WIMP paradigm is particularly appealing for the so-called “WIMP miracle”: when particles of a certain type drops out of equilibrium, they have a fixed comoving density (excluding the secondary interactions) and their energy is redshifted and cooled down. By knowing the particle density at the equilibrium drop (also called the particles freeze–out), and scaling the volume of the Universe, one can estimate the relic density particle now and their energy/temperature. When making an order-of-magnitude estimate for a massive particle $\chi$ of the order of GeV-TeV, the resulting energy density now is [199]:

$$\Omega_\chi h^2 \approx \frac{3 \times 10^{-27} \text{ cm}^3 \text{ s}^{-1}}{<\sigma v>}$$

(6.5)

where the cross section times the velocity is averaged over the temperature. Therefore in order to obtain the DM relic density predicted today, the particle should have a cross section of the order of the weak interactions, making this scenario experimentally accessible to the present and future generation of experiments that are looking for DM. Moreover for WIMPs, at the freeze–out, the temperature is typically $\sim 5\%$ of the particle mass and the particle is therefore nonrelativistic, fulfilling the condition to be cold DM.

A complete list of DM candidates is out of the target of this chapter. In this section we will briefly mention the most important candidates reported in literature, with particular emphasis (after a short introduction to the SuperSymmetry) on the probably most studied candidate: the neutralino, a particle which appears to be a suitable candidate in the supersymmetric extension of the Standard Model of Particle Physics (SUSY) [200]. A complete collection of DM candidates can be found in [60] (and references therein). Some of the basic candidates are schematically represented in the plot by Roszkowski [201] in Fig. 6.9. As shown, the range of masses and cross sections are extremely wide.

- **Neutrinos**

  Neutrinos have been considered as excellent candidates, until very recently, since we know they exist, they pervade the entire Universe in large quantity, they interact very little with matter and, from neutrino oscillation experiments, they have a mass. In particular, from tritium $\beta$-decay experiments at Troitsk and Mainz we know they have a mass smaller than $\sim 2 \text{ eV}$ [202]. However, this mass cannot explain the total amount of the DM present in the Universe: indeed the relic density is predicted to be

$$\Omega_\nu h^2 = \sum_{i=1}^{3} \frac{m_i}{93 \text{ eV}}$$

(6.6)

where $m_i$ is the mass of the $i$-th neutrino. Therefore, neutrinos are simply not abundant enough to be the dominant component of DM since the upper limit on their mass implies an upper bound on the total neutrino relic density of $\Omega_\nu h^2 < 0.07$. Taking into account the CMB anisotropies measurements the neutrino density is further constrained to $\Omega_\nu h^2 < 0.0067$. The only possible solution is that there is a fourth family of neutri-
6.3. Dark Matter Candidates

Figure 6.9: A schematic representation of some well-motivated DM candidate particles. \( \sigma_{\text{int}} \) represents a typical order of magnitude of interaction strength with ordinary matter. The box marked “WIMP” stands for several possible candidates, e.g., from Kaluza–Klein scenarios. Figure from [201].

• Axions
The axions were theoretically motivated to solve the problem of CP violation in particle physics. These are chargeless spin 0 particles associated with the spontaneous breaking of the Peccei-Quinn symmetry. Laboratory searches, stellar cooling and dynamics of supernova 1987A, constrain the axions to be very light \( (\leq 0.01 \text{ eV}) \). They are expected to be weakly interacting with ordinary particles, which implies that they were not in thermal equilibrium in the early Universe. So, the calculation of their relic density is very uncertain and depends on some assumptions on their production mechanism. Anyway, it is possible to find an acceptable range were axions satisfy all present-day constrains and represent a possible viable DM candidate [203, 204].

• Extra-dimension, Kaluza Klein State
Many of the current researches toward a Grand Unification Theory (one to describe all known interactions, including gravity) postulate the existence of some hidden additional space dimensions, accessible only at very small length or very high energy scales. In this framework, the Unified Extra Dimensions (UED) theories predict the existence of a tower of Kaluza-Klein (KK) states for each SM particles which can provide suitable DM candidates. The most favoured candidate as a stable lightest Kaluza-Klein particle...
(LKP) is the first KK excitation of the hypercharge gauge boson, usually referred to as $B^{(1)}$. The $B^{(1)}$ can account for the observed quantity of DM if its mass lies in the range of 400 to 1200 GeV [205].

- **MeV DM**
The INTEGRAL satellite has observed an excess of 511 keV positron annihilation emission toward the galactic center, which is difficult to reconcile with astrophysical sources [206]. It has been suggested that this emission might be due to $e^+e^-$ pairs that are the annihilation products of DM particles in the MeV range [207]. However, the particle physics motivation for this scenario is weaker than for other candidates.

- **Super-heavy DM**
Rejecting the hypothesis that DM particles were in thermal equilibrium with other species before decoupling allows one to think of superheavy (mass greater than $10^{10}$ GeV) candidates dubbed WIMPzillas. These massive particles are extensively believed to be created by gravitational effects at the end of inflation epoch [208]. They present an interesting phenomenology, including a possible solution to the problem of CRs observed above the GZK cutoff, since they could decay into standard model particles with extremely high energies (for example, directly into UHE or EHE CRs).

As already mentioned, there are several more DM candidates which can be found in literature. The reader is referred to dedicated reviews as for example [195].

In the next subsection we briefly introduce the SuperSymmetry and its most relevant DM candidates, in particular the neutralino.

### 6.3.1 Short introduction to Supersymmetry

The SM of fundamental interactions consist on an explicit description of the strong, weak and electromagnetic interactions that governs the dynamics of the elementary particles. This theory is based on the Gauge principle, in which all forces are driven by Gauge fields of the local symmetry group $SU(3)_\text{color} \otimes SU(2)_\text{gauge} \otimes U(1)_\text{hypercharge}$. In the SM, the field contents can be described by 3 sectors: the bosonic sector (spin 1), the fermionic sector (spin 1/2), and the Higgs boson sector (spin 0). The last is imposed in the theory *ad hoc* and it is necessary to allow the spontaneous electroweak symmetry breaking and allows the elementary particles to get a physical mass.

Although the SM of particle physics is a theory that received excellent confirmations from decades of experimental data, it has some theoretical and experimental (e.g the oscillation of neutrinos and the anomalous magnetic moment of the muon) problems and thus it is well known that it should be considered as an effective low–energy theory.

The supersymmetric theory (SUSY) [200] was developed to explain few of the theoretical problems of the SM, in particular:

- The hierarchy problem: all interactions are thought to be unified at the typical unification scale of the grand unified theory (GUT) $\Lambda_{\text{GUT}} \sim 10^{14}$-$10^{16}$ GeV. GUT is based on the idea that at high energies, all symmetries have the same gauge coupling strength, which is consistent with the speculation that at the EW scale all of them are different manifestations of a single overarching gauge symmetry. The hierarchy problem arises in the radiative corrections to the mass of the Higgs boson. All particles get radiative corrections to their mass, but while fermion masses increase only logarithmically, scalar
masses increase quadratically with energy giving corrections at 1-loop of:

\[ \delta m_{\text{scalar}}^2 \sim \left( \frac{\alpha}{2\pi} \right) \Lambda^2 \]  
(6.7)

The radiative corrections to the Higgs mass (which is expected to be of the order of the electroweak scale \( \sim 100 \text{ GeV} \)) will destroy the stability of the electroweak scale if \( \Lambda \) is near the grand unified theory (GUT) scale.

- The couplings of the three gauge interactions of the SM do not unify at some high energy scale, so that the SM cannot easily be included in GUT (see Fig. 6.10).

![Figure 6.10: The measurements of the gauge coupling strengths at LEP do not (left) evolve to a unified value if there is no supersymmetry but do at \( \sim 2 \cdot 10^{16} \text{ GeV} \) (right) if supersymmetry is included. Image taken from [60].](image)

The above mentioned problems can be solved if the SM gauge group is enlarged to a new symmetry group named supersymmetry where bosons and fermions are coupled in common multiplets. Every known particle of the SM is then provided a superpartner sharing the same quantum numbers, except the spin which differs for \( 1/2 \). As superparticles (sparticles) are expected to have mass at the TeV scale they can only be tested at powerful particle accelerators, like the Large Hadron Collider (LHC). Moreover, the SM cannot provide any suitable DM candidate, whereas its SUSY extension can provide, in a natural way, good candidates for DM.

There are many ways in which the SUSY can be carried out. Normally, only the minimal extension is considered (Minimal SUSY Model) [209]. The minimal supersymmetric extension of the SM (MSSM) is minimal in the sense that it contains the smallest possible field content necessary to give rise to all the fields of the SM. The MSSM requires a doubling of the SM degrees of freedom including two complex Higgs doublets giving mass to the down-type and up-type fermions, respectively. In Fig. 6.11 the SM particles and their superpartners in the MSSM are shown. Out of the particle content of the MSSM there are six classes of physical particles that superpartners of the SM fall into: squarks, sleptons, sneutrinos, gluinos, charginos and neutralinos.

In the MSSM has been introduced \textit{ad hoc} a new discrete symmetry, the so-called \( R \)-parity [210]. It is defined as \( R = (-1)^{3B+2S+L} \), where \( B \) is the baryon number, \( S \) is the spin and \( L \) is the lepton number, and distinguishes SM particles \( (R = 1) \) and their SUSY partners \( (R = -1) \). In \( R \)-parity conserving models the sparticles can only be produced/annihilated in pairs, so that the lightest supersymmetric particle (LSP) is stable and can represent
6. THE DARK MATTER PARADIGM

<table>
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<th>Standard Model particles and fields</th>
<th>Interaction eigenstates</th>
<th>Supersymmetric partners</th>
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<tbody>
<tr>
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</tr>
<tr>
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<td>( \tilde{q}_L, \tilde{q}_R )</td>
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<td>( l = e, \mu, \tau )</td>
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<td>( \tilde{l}_L, \tilde{l}_R )</td>
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<td>( H^0_2 )</td>
<td>Higgs boson</td>
<td>( H^0_2 )</td>
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Figure 6.11: Standard Model particles and their superpartners in the MSSM. Adapted from [60].

an excellent DM candidate. Another consequence of a conserved \( R \)-parity is the prevention of the proton decay (which was actually the basic motivation for the introduction of the \( R \)-parity).

Unfortunately, the number of free parameters (masses and mixing angles) in a general MSSM extension is large and can exceed 100, which makes any phenomenological study based on a blind approach to the MSSM parameter space very problematic. The usual way of studying SUSY effects is therefore to assume specific frameworks, where further well-motivated assumptions are introduced.

**mSUGRA**

Since no SUSY particle with the same mass as its SM partner has been seen by accelerator experiments, SUSY has to be broken. It is typically assumed, that the breaking takes place at a high energy scale (the GUT scale). There are several models with different messenger particles (gravitons, gauge bosons, etc.) mediating the SUSY breaking effects from the GUT down to the EW scale. If the breaking information are mediated to the electroweak scale by gravitational interactions, the model is called minimal supergravity (mSUGRA) [211]. The key point of these models is the unification at the Grand Unification scale \( \Lambda_{GUT} \) of the bosonic masses to the common scalar mass \( m_0 \), of the fermionic masses to the common gaugino mass \( m_{1/2} \) and of the trilinear scalar couplings to \( A_0 \) in addition to gauge coupling unification. As a consequence, the number of free parameters in mSUGRA is dramatically reduced and the whole mSUGRA scenario can be described by only five additional parameters to the SM ones defined at the GUT scale. These parameters are: the universal scalar (sfermion and Higgs boson) mass \( m_0 \), the unification of the gaugino masses \( m_{1/2} \), the ratio of the Higgs vacuum expectation values \( \tan \beta = \langle H^0_2 \rangle / \langle H^0_1 \rangle \), the universal trilinear coupling \( A_0 \) and the sign of the Higgsino mass parameter \( \mu \). All parameters of the MSSM can be derived by renormalization group equations (RGE) from the values of these five input parameters (which define a specific mSUGRA model) at the GUT scale, as illustrated in Fig. 6.12. This means that by solving the RGE, these five input, defined at the GUT scale, can be translated into masses and couplings at the electroweak scale.

Despite the reduced number of parameters, the parameter space of mSUGRA is still quite wide, and provides a rich phenomenology. Important experimental constraints can help in
6.3. Dark Matter Candidates

Figure 6.12: Unification of the sparticle masses at the GUT scale. The common gaugino mass $m_{1/2}$ and the common scalar mass $m_0$ are input parameters of mSUGRA. Image taken from [212].

excluding parts of the parameter space, such as the relic DM density derived from WMAP data, the absence of new particle at LEP below $\sim 100$ GeV, the agreement of $b \to s\gamma$ decays with predictions of the SM and the measurements of the anomalous magnetic momentum of the muon.

pMSSM

There exists another popular SUSY model, the so-called phenomenological SUSY extension, or pMSSM. The main difference compared to the mSUGRA model is that it is not based on particular theoretical assumptions but on few phenomenological considerations that reduce the number of free parameters at 10 - 20. These free parameters are described at the electroweak scale, and not renormalized from the unification scale, as in case of mSUGRA.

In particular, the most common free parameters are $\tan\beta$, $\mu$, $m_0$, $M_A$, $M_2$, $A_b$ and $A_t$, where the first three terms were already presented for mSUGRA, while $M_A$ is the mass of the pseudo-scalar Higgs boson, $M_2$ of the second gaugino and $A_b$, $A_t$ are trilinear coupling appearing in SUSY breaking terms.

6.3.2 Supersymmetric Dark Matter Candidates

- Sneutrinos
  The supersymmetric partner of the standard neutrino would represent an interesting DM candidate if its mass lies between 0.5 and 2.3 TeV. Such a particle, however, has a quite large cross section for scattering off nucleons, and hence should have been already observed by direct detection experiments [213].

- Gravitinos
  Gravitinos are the superpartners of the (so far undetected) graviton in supersymmetric models. In some supersymmetric scenarios, like the Gauge Mediated Supersymmetry Breaking (GMSB) models, gravitinos can be the LSP and be stable. Despite being
theoretically well-motivated, they would interact only gravitationally, making them very difficult to observe [214].

- **Axinos**
  In many ways, axinos, the superpartner of the axion, and gravitinos share similar phenomenological properties, in particular very low interaction with matter.

- **Neutralinos**
  Neutralinos are by far the most studied DM candidate particle. In the MSSM there are four neutralinos $\tilde{\chi}_i^0$, results from the physical superpositions of the fermionic partners of the neutral electroweak gauge bosons, called Bino ($\tilde{B}$) and Wino ($\tilde{W}^\pm$), and of the fermionic partners of the neutral Higgs bosons, called Higgsinos ($\tilde{H}_1^0$ and $\tilde{H}_2^0$) (see Fig. 6.11). The lightest neutralino of the four neutralinos is $\tilde{\chi}_1^0$ and is simply dubbed $\chi$. All SUSY particles can therefore decay into $\chi$ (if it is the LSP) while $\chi$ itself is stable if the $R$-parity is a conserved symmetry. The neutralino can be expressed as

$$\chi \equiv \tilde{\chi}_1^0 = N_{11} \tilde{B} + N_{12} \tilde{W}^+ + N_{13} \tilde{H}_1^0 + N_{14} \tilde{H}_2^0 \quad (6.8)$$

The coefficients $N_{ij}$ are obtained by diagonalizing the neutralino mass matrix and are mainly function of the Bino and Wino masses and of the parameter $\mu$. It is commonly defined that the lightest neutralino is mostly gaugino-like if $P = |N_{11}|^2 + |N_{12}|^2 > 0.9$, Higgsino-like if $P < 0.1$, and mixed otherwise. The relative composition of the neutralino (i.e. the factors $N_{ij}$) strongly affects the cross sections of the different annihilation processes. If the neutralino is the lightest SUSY particle and $R$-parity is conserved then it must be stable and it can represent an excellent cold DM candidate with a relic density compatible with the WMAP bounds, a mass at the GeV-TeV scale and a typical cross section of the order of the weak interactions.

### 6.4 Dark Matter Experimental Searches

Searches for non-baryonic, cold DM exist in a large variety of techniques. One can classify them in three basic categories: direct production in accelerator experiments, direct detection through their scattering on target nuclei (measuring nuclear recoils) and indirect detection of self-annihilation products of these particles in high density DM regions. It is commonly accepted that a multidisciplinary approach of all the three mentioned types of DM search must be carried out in order to efficiently constraint the nature of DM [215].

#### 6.4.1 Direct production

The Large Hadron Collider (LHC), which started its operations in 2010 at CERN, will allow to explore possible extensions of the Standard Model of Particle Physics, reaching a center-of-mass energy of about 14 TeV. Despite the fact that the discovery of new particles would be of paramount importance also for Astrophysics and Cosmology, it will not be easy to extract, from accelerator experiments alone, enough information to unambiguously identify DM particles [215], since the constraints that can be placed on a DM candidate from Collider experiments are strongly model dependent.

SUSY particles are expected to be produced in pairs at the LHC, with cross sections around
6.4. Dark Matter Experimental Searches

The particle pairs will subsequently decay to quark or gluon jets plus a WIMP that can escape the detector without interactions. Therefore, the LHC experiments could observe events with many hadronic jets and a missing energy event which should be signature of SUSY.

6.4.2 Direct detection

Direct detection experiments are based on the assumption that the Earth moves within the DM halo of our galaxy (the estimated density of DM in the solar system is around 0.3 GeV/cm$^3$) and that DM particles have a non-zero cross section for scattering off ordinary matter. DM can thus be observed indirectly through the outcomes of nuclear recoils when crossing dense targets. A nuclear recoil determines measurable signals from ionization, scintillation and phonon release. The predicted energy depositions are very small, ranging from a few keV to a few tens of keV with typical cross section of the order of $10^{-43}$ cm$^2$ [217]. Such small energy scales and cross sections oblige direct detection experiments to be performed underground, in order to be shielded from CRs which would be an overwhelming background. Moreover, a great effort is done against backgrounds, trying to reduce natural radioactivity sources in the surrounding environment and in the materials employed for the experiment construction. Also the electronic of the detectors must be as much as possible noise-free.

More than 20 direct DM detection experiments are either now operating or are currently in development. In these many experiments, numerous techniques have been developed to measure the nuclear recoil produced by DM scattering. Some of these methods include the observation of scintillation (used by DAMA, ZEPLIN-I, NAIAD, LIBRA), photons (used by CREST and CUORICINO) and ionization (used by HDMS, GENIUS, IGEX, MAJORANA and DRIFT). Some experiments use multiple techniques, such as CDMS and EDELWEISS which use both ionization and photon techniques, CRESST-II and ROSEBUD which use both scintillation and photon techniques and XENON, ZEPLIN-II, ZEPLIN-III and ZEPLIN-MAX, which use both scintillation and ionization techniques. The use of such a large array of techniques and technologies is important not only to accelerate the progress of the field, but also to vary the systematic errors from experiment to experiment, allowing for a critical assessment of a positive signal.

The basic output of such experiments is an upper limits in the rate of recoils per kg of detector, which is interpreted within a given theoretical model in terms of limits in the WIMP–nucleon interaction cross section.

There is a wide literature on direct detection and the field is currently very active and developing. Two recent reviews can be found in [218,219].

6.4.3 Indirect detection

Indirect detection of DM is the technique of observing the products of self-annihilation processes or decays of DM particles, such as energetic leptons, hadrons, neutrinos, synchrotron radiation and $\gamma$–rays, emerging in the follow up hadronization and fragmentation of the involved processes. The flux of such products is directly proportional to the annihilation rate, which further is dependent on the DM density. Since these products usually mix with the CRs background, it is often very complicate to disentangle the origin$^2$. Indirect DM searches are performed by satellite or balloon–borne experiments, and ground–based telescopes.

Restricting our discussion to the indirect DM detection in $\gamma$–rays, the IACTs (see chapter 2)

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$^2$As the $\gamma$–rays are not deflected by magnetic fields, they point to the DM dominated regions and should act as tracers of the DM density.
and the Fermi–LAT detector are well suited for DM searches since the WIMPs have masses expected at the order of the GeV-TeV scale, i.e. the energies where these experiments are sensitive.

Considering the case of the neutralino as lightest stable particle in mSUGRA framework, many different processes are involved in the neutralino self-annihilation. The cross section strongly depends on the relative composition of the lightest neutralino defined by the factor $N_{ij}$ of the equation 6.8. The most relevant neutralino interaction for the purposes of indirect DM searches in $\gamma$-rays is the self annihilation in fermion-antifermion pair-production, primarily heavy fermions (top, bottom, charm quarks and tau leptons), gauge bosons pair ($W^+W^-$ and $Z^0Z^0$) and final states containing Higgs bosons. The subsequent hadronization results in a $\gamma$-ray power law spectrum with an exponential cutoff at the neutralino mass (expected to be between 50 GeV and several TeV). In figure 6.13, the predicted $\gamma$-ray spectra for different specific annihilation channels is shown.

A direct annihilation in $\gamma$-rays, such as $\chi\chi \rightarrow \gamma\gamma$ or $\chi\chi \rightarrow Z^0\gamma$, (which result in $\gamma$-rays of energies $E_\gamma = m_\chi$ and $E_\gamma = m_\chi - M_{Z^0}^2/4m_\chi$ respectively) or $\chi\chi \rightarrow H\gamma$ provides line emission (i.e., a smoking gun signature). The $\gamma$-ray flux originating from the line emissions is 2 to 3 orders of magnitude lower than the continuum power law spectrum mentioned above since no tree level process is possible and the loops, visible in Fig. 6.14, introduce two more vertices into the calculation of the corresponding cross section. The basic modes of annihilation of the neutralino are treated in a more extensive way in [60].

Recently Bringmann et al. [225] showed that in some regions of the mSUGRA parameter space, a hitherto neglected contribution to $\gamma$-ray emission comes directly from charged sparticles mediating the annihilation into $\chi\chi \rightarrow l^+l^-\gamma$. They defined this intermediate state radiation as Internal Bremsstrahlung (IB). The IB mechanism permits to restore the helicity

![Image](image_url)

**Figure 6.13:** Energy spectra of photons per annihilation for different channels. The solid [220] and dotted [221] lines both correspond to the $b\bar{b}$ annihilation channel for the supersymmetric neutralino, the differences are due to different quark fragmentation and different DM particle masses: $m_\chi = 1$ TeV and $m_\chi = 100$ GeV, respectively. The short-dashed line corresponds to the spectra for annihilation through the WW and ZZ channels for the neutralino [222]. The long-dashed line shows the spectrum for annihilation of Kaluza-Klein DM [223]. Image taken from [224].
6.5 Expected γ-ray flux from Dark Matter self-annihilation

The γ-ray flux originating from DM particle annihilations can be factorized into a contribution called the astrophysical factor \( J(\Psi) \) related to the morphology of the emission region and
a contribution called the particle physics factor $\Phi^{PP}$ depending on the candidate particle characteristics:

$$\Phi(E > E_0) = J(\Psi) \cdot \Phi^{PP}(E > E_0)$$

(6.9)

where $E_0$ is the energy threshold of the detector and $\Psi$ is the angle under which the observation is performed.

The astrophysical factor can be written as

$$J(\Psi_0) = \frac{1}{4\pi} \int_V d\Omega \int_{\text{l.o.s.}} d\lambda \rho^2 \ast B_{\theta_0}(\theta)$$

(6.10)

where $\Psi_0$ denotes the direction of the target. The first integral is performed over the spatial extension of the source, the second one over the line-of-sight variable $\lambda$. The square of the DM density $\rho$ is convoluted with a gaussian function $B_{\theta_0}(\theta)$ in order to consider the telescope angular resolution ($\sim 0.1^\circ$), where $\theta = \Psi - \Psi_0$ is the angular distance with respect to the center of the object. It is worth mentioning that the integration of equation 6.10 involves foreground (MW halo) and extragalactic background whose contributions can be substantial [226]. The astrophysical factor depends on the DM morphology and the distance of the emission region (as well as the PSF of the telescope), but for a given DM profile it does not depend on the particular DM candidate.

The particle physics factor can be expressed as a product of two terms. The first one depends only on the DM candidate mass and cross section, whereas the second term depends on the annihilation $\gamma-$ray spectrum and must be integrated above the energy threshold $E_0$ of the telescope

$$\Phi^{PP}(E > E_0) = \frac{\langle \sigma v_{\chi\chi} \rangle}{2m_\chi^2} \int_{E_0}^{m_\chi} S(E) dE$$

(6.11)

where $\langle \sigma v_{\chi\chi} \rangle$ is the total averaged cross section times the relative velocity of the particles, $m_\chi$ is the DM particle mass, and the factor 2 takes into account that the neutralino annihilates with itself. The $\gamma-$ray annihilation spectrum is composed of different contributions: $S(E) = \sum_i dN_i^\gamma/dE$ where $dN_i^\gamma/dE$ is the differential flux of the $i$-th annihilation mode (such as monochromatic $\gamma$ lines, secondary photons produced in the hadronization and further decay of the primary annihilation products (mainly through the decay of neutral pions) and the Internal Bremsstrahlung (see section 6.4.3)).

The indirect search for DM is nowadays affected by large uncertainties in the flux prediction which put serious hindrances to the estimation of the observability. On the one hand, the astrophysical factor uncertainties can raise up to typically several order of magnitude, depending on the target and the profile model and on other uncertainties due to the presence of substructures (expected to be present in any DM halo [227]) and possibly to adiabatic compression of the DM in the innermost regions of the halos [228]. On the other hand, the allowed parameter space for the mass and the annihilation cross section of the DM particle spans many orders of magnitude giving rise to flux estimations which can differ up to six orders of magnitude (or even more).

### 6.6 Dark Matter density profiles

The formation of large scale structures is a highly complex process. An essential tool to study the evolution of such structures are N-body simulations (see [229, 230] for recent reviews). Nowadays, numerical simulations play a key role in the interpretation of observational data since high resolution simulations are possible thanks to the increasingly computational power.
Since the amount of DM exceeds the normal matter by far, N-body simulations with the entire matter in form of DM are used as a good approximation for the distribution of matter. Moreover, gravity is the dominant force at large scales and normal matter is expected to follow DM at these scales. The inclusion of baryonic matter in the simulations is usually very complicated, because of the additional electroweak and strong forces between the SM particles.

In general, only the CDM simulations reproduce the observables at a large-scale, while at local scales there are some failures. For example, they predict more expected substructure orbiting within galactic halos [231].

One of the important outcomes of N-body simulations is the suggestion of a universal DM density profile with the same shape for all masses, epochs and input power spectra. A widely used parameterization for such a DM halo is

$$\rho_{DM}(r) = \frac{\rho_s}{(r/r_s)^\gamma[1 + (r/r_s)^\alpha]^\frac{(\beta-\gamma)}{\alpha}}$$

(6.12)

where \(r\) is the distance to the center of the halo, \(\rho_s\) is linked to the DM density at the center and \(r_s\) is called the scale radius of the profile and represents the distance where the slope is equal to -2. Various groups have obtained different results for the spectral shape in the innermost regions of galaxies and galaxy clusters. These regions are the hardest regions to be simulated because of the short dynamical time-scales and strong gravitational forces, normally dominated by the resolution of the simulation, i.e. the computational power. The parameters for common used profiles are listed in Fig. 6.16.

Other popular DM density profile are the Burkert profile [236,237]

$$\rho_{Burkert}(r) = \frac{\rho_s}{(1 + r/r_s)(1 + (r/r_s)^2)}$$

(6.13)

and the Einasto profile

$$\rho_{Einasto}(r) = \rho_s e^{-2n[(r/r_s)^{1/n} - 1]}$$

(6.14)

which has been shown to be a good fit to CDM halos with the Einasto index, \(n\), ranging from \(\sim 3 - 7\) [238].

Although it is definitely clear that the slope of the density profile should decrease as one moves from the center of a galaxy to the outer regions, the precise value of the power law index in the innermost galactic regions is still under debate. In particular it is still unclear whether the DM distributions present cuspy or shallow profiles in their innermost regions. The predic-

<table>
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<th>(\beta)</th>
<th>(\gamma)</th>
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<tr>
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<tr>
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<tr>
<td>Iso: Bergstrom et al.</td>
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**Figure 6.16**: The different parameters for the most widely used DM density profiles: Kravtsov et al. [232], Navarro, Frank and White (NFW) [233], Moore et al. [234] and Bergstrom et al. (modified isothermal) [235].
tions for the innermost slope in N-body simulations do not agree with the observation results. Normally, the simulation require a steeper profile than deduced from experimental measurements. Indeed there is an apparent contradiction between N-body simulations which favour cusped density profiles at small radii and observations which appear to favour significantly shallower density cores [239].

6.7 Interesting astrophysical sites for Indirect Dark Matter Searches

Since the $\gamma$-ray flux is proportional to the square of the DM density (see equation 6.10), a relevant question concerning the indirect search for DM annihilation products is where to look for hot DM spots in the sky.

6.7.1 The Galactic Center

In the past, the Galactic Center (GC), located at a distance of 8.5 kpc from Earth, was considered the best option for indirect DM searches. However, this is a very crowded region, which makes it difficult to discriminate between a possible $\gamma$-ray signal due to DM annihilation and that from other astrophysical sources. Moreover, the estimation for the DM annihilation flux from the GC is complicated by the large presence of baryons that could have modeled the DM distribution in a complicated way. While the well-studied motion of stars around the GC gives us important information on the average matter density, there is no clear indication on the density in the very central core, from where the most signal is expected.

Whipple [46], CANGAROO [47], MAGIC-I [49] and especially H.E.S.S. [48] have already carried out detailed observations of the GC. The measured H.E.S.S. spectrum, from 200 GeV to several TeV, was associated to the contribution from the several astrophysical emitters, while the study of the DM interpretation was performed by Profumo in [50]: only very massive neutralino of the order of 10 - 20 TeV could explain the H.E.S.S. results, for which the $\gamma$-ray yield is expected to be 2-3 orders of magnitude lower than the measured flux.

It is believed now that only very accurate and prolonged observation of the GC could unveil eventual spectral features that could be sign of DM, since, the expected DM signal it is orders of magnitude smaller than the astrophysical signal.

6.7.2 Dwarf Spheroidal Galaxies

Very promising targets with high DM density in relative proximity to the Earth (less than 100 kpc) are the dwarf Spheroidal (dSph) satellite galaxies of the Milky Way. These galaxies are believed to be the smallest (size $\sim$ kpc), faintest (luminosities $10^2$-$10^8$ $L_\odot$) astronomical objects whose dynamics are dominated by DM (see [240] and reference therein). Usually, their member stars show large circular velocities and velocity dispersions that, combined with their modest spatial extent, can be interpreted by the presence of a large DM halo of the order of $10^5$-$10^9$ $M_\odot$ with very high mass-to-light ratios (up to $\sim 10^3$ $M_\odot/L_\odot$). Moreover these objects are free from expected astrophysical $\gamma$-ray sources located in their vicinity.

The Sloan Digital Sky Survey (SDSS) [184] is leading to the discovery of a new population of MW satellites. This population of extremely low-luminosity galaxies can be very interesting for DM searches and for investigating galaxy formation at the lowest mass scales. The existence of a new class of ultra-faint MW satellites is also relevant because it provides a partial solution for the so-called missing “satellite problem” [231, 241] by partially filling the gap between the predicted and the measured number of galactic subhalos. These new galaxies
represent a population of extremely low luminosity objects, very interesting in the context of indirect DM searches and galaxy formation at the lowest mass scales. Some dSphs have already been observed in γ-rays by IACT experiments: Draco by Whipple [242], Canin Major [243] and Sagittarius [244] by H.E.S.S., Draco [245] and Willman 1 [246] by MAGIC-I, Draco, Ursa Minor, Bootes 1 and Willman 1 by VERITAS [247]. No significant signal from DM annihilations was found and only flux upper limits were estimated. Recently, the Fermi–LAT Collaboration published the results of the observations of 14 dwarf spheroidal galaxies taken during the first 11 months of survey mode operations [248]. No significant γ-ray emission was detected above 100 MeV from the candidate dwarf galaxies and upper limits to the γ-ray flux were provided.

Chapter 7 will be dedicated to the analysis of the observation carried out by the MAGIC-I telescope of Segue 1, an object discovered in 2006 by the SDSS Collaboration [249] and considered by many authors one of the most DM dominated dSphs so far known (with a mass–to-light ratio of the order of $10^3 M_\odot/L_\odot$).

### 6.7.3 Galaxy Clusters

Clusters of galaxies are the largest and most massive gravitationally bound systems in the Universe, with radii of the order of the Mpc and total masses around $10^{14}$–$10^{15} M_\odot$ [177]. These systems are thought to host enormous amounts of DM, which should gravitationally cluster at their center and present numerous local substructures and filaments [178] which could lead to a significant boost in the γ-ray flux. The large amount of DM should counteract their large distance compared to the already mentioned galactic targets. So far, no cluster has been firmly detected in γ-rays, but they are expected to be significant γ-ray emitters because are very active regions and therefore a substantial non-thermal emission is expected. Since they also contain large amounts of gas and strong magnetic fields, they are considered as natural place for CR acceleration. Three nearby clusters, Perseus, Coma and Virgo, are the most promising candidates for observation with MAGIC, not only for DM searches: indeed, a signal from the proton-induced pion decay emission could be very bright. The H.E.S.S. Collaboration observed Abell 496, Coma and Abell 85 clusters [250–252], VERITAS observed the Coma cluster [253], and MAGIC-I the Perseus cluster [254], in all cases without any significant γ-ray detection. Therefore only upper limits were provided. The Fermi–LAT Collaboration already reported their results achieved from the first 11 months of Fermi–LAT survey mode observations on DM searches in nearby clusters (AWM 7, Fornax, M49, NGC 4636, Centaurus and Coma), with no detection of DM γ-ray emission [255].

### 6.7.4 Intermediate Mass Black Holes

Another interesting DM target scenario is represented by the so-called intermediate mass black holes (IMBHs). The model described in [224] shows that studying the evolution of super massive black holes, a number of IMBHs do not suffer major merging and interaction with baryons along the evolution of the Universe. DM accretes on IMBH in a way that the final radial profile is spiky so that the IMBHs could be bright γ-ray emitters. These targets could be related to the unidentified EGRET and Fermi–LAT sources [22].
In view of indirect search for Dark Matter (DM) self annihilation signatures in $\gamma$-rays, MAGIC-I carried out the observation of the promising target Segue 1. This source is believed to be a dwarf spheroidal galaxy, satellite of the Milky Way, and it is located in the direction of the Leonis Constellation at a distance of 23 kpc. It was discovered in the catalog of the Sloan Digital Sky Survey (SDSS) together with other ultra-faint self-gravitating clusters. Nowadays, Segue 1 is considered one of the most DM dominated object known in the Universe with a huge mass-to-light ratio of the order of $10^3 M_\odot/L_\odot$, even if its nature is still under debate.

In this chapter, after a brief introduction to the source, the analysis of Segue 1 data recorded by the MAGIC-I telescope will be presented in detail. During the survey, the light of the star $\eta$-Leonis (apparent magnitude 3.5), was reflected in the inner part of the MAGIC-I camera, being the angular distance between Segue 1 and the star $\sim$0.7°. Special cares were thus necessary for minimizing the effects of the supplementary light of the star which resulted in a local loss of trigger efficiency for the lowest energy events. The related problems and the techniques used to get reliable results from the data, particularly for the lowest energy range (which is the most interesting one for the purpose of indirect DM searches) will be illustrated, as well as the tests performed to cross-check the consistency of the whole analysis chain.

No significant $\gamma$-ray emission coming from Segue 1 was found in roughly 30 hours of good quality data above an analysis energy threshold of 100 GeV. Upper limits on the $\gamma$-ray flux were therefore derived for different power law spectra and energy thresholds in view of a paper which is in preparation.

### 7.1 The satellite galaxy Segue 1

The dwarf spheroidal satellite galaxy Segue 1 was discovered by Belokurov et al. (SDSS Collaboration) [249] in 2006 in the SEGUE (Sloan Extension for Galactic Understanding and Exploration) campaign of SDSS data analysis. It is located at a heliocentric distance of $23\pm2$ kpc (28 kpc from the Galactic Center) at sky coordinates

$$RA = 10^h 07^m 04^s \quad DEC = 16^\circ 04' 55''$$
The nature of Segue 1 is still under debate: Belokurov et al. [249] and later on Niederste-Ostholt et al. [256] addressed the source as an extended globular cluster possibly belonging to the disrupted streams of Sagittarius. On the contrary, based on the Keck/DEIMOS [257] spectroscopy data, Geha et al. [258] claimed that Segue 1 belong to the class of the ultra-faint dwarf spheroidal satellite (dSph) galaxies. The uncertainties are connected to two basic facts: from one side, the data available are still scarce, and precise determination of astrophysical parameters are subject to assumptions. On the other hand, the estimated properties are just in between the class of globular clusters and that of dwarf galaxies. With a half-light radius of \( \sim 30 \) pc (the radius of a cylinder pointing to the Earth and containing half of the total luminosity), Segue 1 would be the largest globular cluster and the smallest dwarf known to date. The radial velocities measured by Geha et al. [258] on 24 members stars have a mean of 206 km/s and a velocity dispersion of \( \sim 4.3 \) km/s, which in the hydro-static equilibrium of the Jeans equation suggests a very large amount of DM \((\sim 10^3 \ M_\odot/L_\odot)^1\). Although Segue 1 spatially overlaps the leading arm of the Sagittarius stream, the authors calculated that the velocity is 100 km s\(^{-1}\) different than that predicted for Sagittarius tidal debris and they did not find strong kinematic evidence supporting tidal effects. The authors concluded thus that Segue 1 is the least luminous of the ultra-faint galaxies recently discovered around the Milky Way, and is thus the least luminous known galaxy. Finally, Xiang-Gruess et al. [259] presumed Segue 1 as a dwarf disk galaxy, i.e. stretched along one direction, which can partly mitigate the tidal disruption theorem.

Constraining the density profile of DM in Segue 1 is difficult being small and faint and with very few stars available to act as kinematic tracers of the gravitational potential. Only a more accurate understanding of the stellar composition and dynamics will allow to definitely qualify its nature. Several studies are currently ongoing in this direction and the results should be soon available [260].

### 7.2 MAGIC-I data

Segue 1 observation was carried out by MAGIC-I telescope during 35 dark nights between November 2008 and March 2009. The source was surveyed at Zenith angles between 13° and 34°, in order to guarantee the lowest trigger energy threshold. The data taking were performed in WOBBLE mode (see section 4.2.1) in which two opposite sky directions (hereafter W1 and W2) each 0.4° off the source, are tracked alternatively for 20 minutes each. The ON-source data (ON) are defined by calculating image parameters with respect to the source position, whereas background control OFF-source data (OFF) are obtained from the same data set, but with image parameters calculated with respect to the position on the opposite side of the camera, i.e. the anti-source position, which is always located at a nominal distance of about 0.8° away from the source. This technique allows, in good data taking conditions, a reliable estimation of the background with no need of extra observation time. As explained in section 4.2.1, the wobbling procedure permits to minimize the effect of the intrinsic trigger inhomogeneities at lower SIZES. Usually, a difference within \( \sim 15\% \) in the amount of observation time between W1 and W2 assures a good compatibility (outside the expected signal regions) between the ON and OFF distributions of the overall final sample. In Table 7.1, the basic daily pieces of information about the available MAGIC-I data set of Segue 1 are listed. The overall recorded subruns were 1253 (600 for W1 and 653 for W2 observations) corresponding to a total exposure time of roughly 43 hours. The data were mostly recorded

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1\ The mass and luminosity within 50 pc of Segue 1 are estimated to be respectively of the order of \( \sim 5 \times 10^3 \ M_\odot \) and \( \sim 340 \ L_\odot \) [258].
<table>
<thead>
<tr>
<th>Date</th>
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<th>Number of subruns</th>
<th>Zenith [°]</th>
<th>Azimuth [°]</th>
<th>Mean Trigger Rate [Hz]</th>
<th>PSF [mm]</th>
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<td>168.6 - 217.1</td>
<td>247±17</td>
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</tr>
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</table>

OVERALL 2589.8 (43.2 h) 1253 12.7 - 33.9 104.6 - 253.2 227±37 14.1±0.8

**Table 7.1:** Main daily pieces of information of Segue 1 data recorded by the MAGIC-I telescope between November 2008 and March 2009. The value of the PSF was calculated with the aid of selected muon events [80]. All data were taken during dark time.
7. OBSERVATION OF SEGUE 1 WITH THE MAGIC-I TELESCOPE

in similar MAGIC-I hardware conditions. Nevertheless, two problems affected the data:

- Between November 2008 and March 2009 the MAGIC-I telescope PSF worsened roughly constantly from \( \sim 13 \) mm to \( \sim 15 \) mm (see Table 7.1). This issue was faced by choosing a \( \gamma \)-ray Monte Carlo (MC-\( \gamma \)) sample with a PSF value of 14 mm, close to the mean value of the overall Segue 1 data (14.1 mm). The use of a unique MC-\( \gamma \) data sample was preferred to avoid the splitting of the overall Segue 1 data set in different subsamples which would have prevented the achievement of global results. In any case, no significant effects (quantifiable within few percents) correlated to the worsening of the PSF of the telescope were found, as it will be shown by the aid of a Crab Nebula data sample observed in the same period and used to optimize and cross-check the whole analysis chain.

- The star \( \eta \)-Leonis (mean apparent magnitude 3.5) is located at sky coordinates

\[
\text{RA} = 10^h \ 07^m \ 20^s \quad \text{DEC} = 16^\circ \ 45' \ 46''
\]

at \( \sim 0.68^\circ \) away from Segue 1 sky position, as sketched in Fig. 7.1(b). \( \eta \)-Leonis is a white supergiant with stellar classification A0Ib and it has the traditional Arabic name of Al Jabbah. With an absolute magnitude of -5.6, it is actually one of the brightest stars in the Leo Constellation. The Hipparcos astrometric data has estimated the distance of this star to be about \( 7 \times 10^2 \) pc (\( \sim 2.2 \times 10^2 \) ly) from the Earth [261]. Its apparent magnitude seems to be variable within a range between 3.4 and 3.6 and there is evidence that \( \eta \)-Leonis is part of a binary star system. Due to its quite high magnitude and being its angular distance from W1 and W2 pointing positions respectively \( \sim 0.76^\circ \) and \( \sim 0.81^\circ \), the light of the star, during the whole Segue 1 survey, was always present in the inner part of the camera (\( \sim 1^\circ \) in radius) where the trigger region is operating. This circumstance created a supplementary inefficiency in the trigger system (besides the intrinsic trigger inhomogeneities at lower SIZES) due to the fact that the Individual Pixel Rate (IPR) control (see section 3.2.5) dynamically enhanced the discrimination thresholds of the enlightened pixels, thus decreasing the trigger efficiency of the region of the camera nearby the spot of the star light. It is worth mentioning that the amplitude of this kind of inefficiency is smaller than the typical ones affecting the MAGIC-I camera at lower SIZES but it has a fundamental difference in nature: the position of the spot of \( \eta \)-Leonis light had not a fixed position, since the star is moving in the camera (due the Alt-Azimuth mount of the telescope). Therefore, the usual wobbling procedure (which is just performed to assure that the intrinsic fixed inhomogeneities of the camera are safely smoothed out), could not reduce the inefficiency due to the star. This issue made the analysis of Segue 1 data particularly delicate.

The analysis chain used for Segue 1 data reduction was basically the standard one applied to MAGIC-I data, outlined in chapter 4. The MC-\( \gamma \) events were chosen among the standard production, with simulated energies between 10 GeV and 30 TeV and spectral index of -2.6 (see section 4.2.3). Only events within the same Zenith range of Segue 1 data were considered. The PSF value was chosen to be 14 mm. For the optimization and the check of the whole analysis chain, \( \sim 15 \) hours of Crab Nebula data observed in the same period of Segue 1 survey were also taken into account.

The effects of the light of the star in the final selected data mainly resulted in a significant disagreement, at the lower energy range, between the ON and OFF distributions of the image
7.3 Data analysis

(a) Sketch of the Leo Constellation. (b) Sketch of Segue 1 and η-Leonis sky positions.

Figure 7.1: (a) Sketch of the Leo Constellation. The supergiant η-Leonis, with its absolute magnitude of -5.6, is one of the brightest stars of the Constellation. (b) Sketch of the sky nearby Segue 1. The star η-Leonis is located at an angular distance of ~0.68°. The telescope WOBBLE pointing positions W1 and W2 are also shown.

parameters. In particular, a residual mismatch in the signal region of the |ALPHA| parameter was present, as it will be shown. This dangerous mismatch did not allow a reliable estimation of the γ-ray excesses of the observation at the affected energies, and thus it was faced by conceiving a particular set of cuts based on the fact that the mutual positions between the nominal source and the star in the camera rely on certain particular symmetries. The tests performed on the MC-γ and Crab Nebula test samples for cross-checking the consistency of the method gave good results, as shown in the next section.

7.3 Data analysis

In the following subsections, each step of the analysis, together with the illustration of the problems related to the light of the star and the adopted countermeasures, are presented. The analysis description refers to Segue 1 data, but the MC-γ and Crab Nebula samples were processed as well through the same analysis chain.

7.3.1 Data preparation and calibration

As first step of the analysis, the program MERPP was used to translate the binary raw data coming out directly from the DAQ system, into the custom MARS file format and to merge them with all the available information stored in ASCII files coming from the telescope subsystems. All the pieces of information were afterward used for checking the quality of the observations and for selecting the final data samples of Segue 1 and Crab Nebula. Each subrun was then calibrated with the aid of the corresponding pedestal and calibration runs of each sequence, accordingly to the description given in section 4.3.2. In order to extract the signal and arrival time of each channel for each triggered event, the spline method method was used. The signal amplitudes were then converted to photoelectrons by using
the $F$-factor method. During this step, the pixels affected by hardware malfunctioning were automatically found out. The signals coming from these unsuitable pixels were replaced by the linear interpolation of the signals of the suitable neighbour pixels. Generally, the amount of unsuitable pixels per sequence for the Segue 1 data was of the order of 3% (10-20 pixels), a typical value for MAGIC-I data. Note that the light of the star had not significant effects on the determination of the bad pixels. In Fig. 7.2(a) and 7.2(b) the summary of the unsuitable pixels for Segue 1 data taking during the night of 2008/11/28 is reported as example.

![Diagram of unsuitable pixels](image)

(a) Unsuitable pixels. (b) Information for the unsuitable pixels.

**Figure 7.2:** (a) Positions of the unsuitable pixels in the MAGIC-I camera for Segue 1 data taking in 2008/11/28. (b) Information about the unsuitable pixels for Segue 1 data taking in 2008/11/28.

### 7.3.2 Identification of the $\eta$-Leonis presence and relative distances in the camera

The presence of $\eta$-Leonis light in Segue 1 data can be clearly identified in the fluctuations of the pedestal signals of the enlightened pixels\(^2\). In Fig. 7.3(a) and 7.3(c) the mean pedestal RMS values for the W1 and W2 observations carried out in 2008/11/28 are reported. In both cases the light of star produced a huge enhancement of the pedestal RMS (up to \(~4\) times the overall mean value). It is worth noting that the affected pixels are at the inner edge of the active trigger system region. Therefore, an effect on the trigger efficiency (related to the IPR control system) and consequently on the distribution of the image parameters of the recorded showers is expected. This effect is concentrated at the lower SIZE range (i.e. at the lower energy range) since events with high SIZE have a higher probability to overcome the local trigger inefficiency related to the star. Fig. 7.3(b) and 7.3(d) report the sketches of the important relative distances which defines the geometry of the problem for both the W1 and W2 particular cases discussed here. As shown, the nominal center of the area with highest pedestal fluctuations (which corresponds to the nominal star position) is always placed at an angular distance of about 0.68\(^\circ\) with respect to the nominal source position, as the angular distance between Segue 1 and $\eta$-Leonis is fixed in the sky. Instead, the angular distance between the star and the anti-source is slightly different in the two cases: for W1 data it has a value of \(~1.02\)^\(\circ\), whereas for W2 data \(~1.07\)^\(\circ\).\(^3\) The fact that these two distances are very

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\(^{2}\)Because of the AC-coupling of the PMT output, the NSB does not affect the pedestal value itself but the size of its fluctuations, which are characterized by the pedestal RMS.

\(^{3}\)Note that the nominal anti-source during W1 and W2 observations corresponds to two different positions in the sky.
close in value occurred just by chance. Indeed, the two distances would be nominally equal in value in case $\eta$-Leonis had exactly the same RA coordinates of Segue 1 (while the two RA values are close but not exactly the same). For the same reason, the relative distances between the star and the W1 and W2 positions (at the center of the camera) are also different: $\sim 0.76^\circ$ in case of W1 observation and $\sim 0.81^\circ$ in case of W2 observation.

The examples discussed so far for W1 and W2 observations are representative of a typical situation. For different subsamples, the locations in the camera of the nominal source and the star (whose presence always results in a spot of light in the inner camera with consequent local higher individual pixel mean pedestal RMS and discrimination threshold values) are in general different, since the nominal source position moves along the so called WOBBLE circle\textsuperscript{4}. Nevertheless, the important relative nominal angular distances which define the geometry of the problem are fixed in the whole data sample and are equal to those indicated respectively in Fig. 7.3(b) and 7.3(d).

7.3.3 Image cleaning and parameter reconstruction

After the calibration, the data subruns were processed, day by day, by the STAR program. The aim of this analysis step is to remove, for each event, pixels containing presumably only noise from the NSB and/or signals from small tracks far outside the triggered shower and subsequently to calculate the relative image parameters. During the process, a particular treatment based on the current of each individual channel of the camera was also applied.

For each event, pixels with large deviations (above $3.5\sigma$ with respect to the mean current value) were excluded for the determination of the image parametrization. This procedure was done to minimize possible deformations to the image parameters due to the extra light of the star. The free parameters of the algorithm were set with the standard values: $Q_{\text{core}} = 6$ phes and $Q_{\text{boundary}} = 3$ phes. Additional suppression of pixels containing only noise were achieved by requesting a narrow time coincidence between core pixel arrival times ($\sim 4.5$ ns) and between mean core pixel arrival time and boundary pixel ones ($\sim 1.5$ ns). The use of the arrival time constraints reduced the contribution of the incoherent light of $\eta$-Leonis on the image parameter reconstruction but it could not cure the related local inefficiency, as shown later. Higher values for $Q_{\text{core}}$ and $Q_{\text{boundary}}$ were also tested but did not succeeded to eliminate completely the inefficiency due to the light of the star without increasing the analysis energy threshold. For those reasons, the standard image cleaning was kept and the star related problems were treated afterward, by applying suitable cuts.

During this step of analysis, the information coming from the starguider subsystem (see section 3.2.1) were used for the correction of the reconstructed event origins from possible mispointing during the observation and, thus, for the accurate determination of the nominal source position in the camera.

7.3.4 Data selection

The final Segue 1 data sample was selected with the aid of the information coming from the different subsystem and from the image parametrizations of the events. For each subrun, the following criteria were applied:

- Mean L2 Trigger rate within 20% of the nominal expected value as a function of the

\textsuperscript{4}The WOBBLE circle is an ideal circle center in the origin of the camera coordinates with radius $0.4^\circ$, as shown in Fig 7.3(b) and 7.3(d) (solid black line circles). During WOBBLE observations the nominal source position (as well as the nominal anti-source one) always lies around the WOBBLE circle.
Figure 7.3: (a) Mean individual pixel pedestal RMS values for W1 data recorded during 2008/11/28. (b) Relative distances between the star position, the source, the anti-source and the center of the camera for W1 data recorded in 2008/11/28. (c) Mean individual pixel pedestal RMS values for W2 data recorded during 2008/11/28. (d) Relative distances between the star position, the source, the anti-source and the center of the camera for W2 data recorded in 2008/11/28.
Zenith angle $\theta$, given by
\[
\text{Rate [Hz]} = 250 \times \cos(\theta)^{0.5}
\] (7.1)

- Mean values of cloudiness lower than 40%.
- Arrival time for calibration and signal events of the inner and outer pixels within the FADC time slices reference windows (set respectively between 38 - 60 and 32 - 44 FADC slices).
- Successful correction of the nominal pointing position of the telescope during the observation through the starguider correction.
- Mean values of $\log_{10}(\text{SIZE})$, WIDTH, LENGTH, $|\text{ALPHA}|$, DIST, MEAN ARRIVAL TIME, NUMBER OF ISLANDS within 10% of the overall mean values calculated after the previous four selection criteria.

Among the 1253 available Segue 1 subruns, 873 survived the quality selection (~ 70%, 404 for W1 and 469 for W2 observations). The rejection was mostly due to bad weather conditions. The final sample resulted in 29.4 hours of high quality data spanning the Zenith range between 12.7° and 33.0°.

All the selection criteria were also applied to the Crab Nebula sample: for this source, the overall amount of high quality data in the same Segue 1 Zenith range was about 9.8 hours.

### 7.3.5 $\gamma$-hadron separation and energy reconstruction

The calculation of the matrices used for the $\gamma$-hadron separation and the energy reconstruction was performed with the algorithms described in sections 4.3.5 and 4.3.6. For training the Random Forest method, a subsample (10%) of the final selected Segue 1 data was randomly chosen as hadron train sample, taking care that all the Zenith range was covered. A subsample of MC-$\gamma$ (30% of the overall used MC-$\gamma$) was taken as $\gamma$-ray train sample and not used in the final analysis steps. The relative contribution to the $\gamma$-hadron separation of the parameters used for the training is reported in Fig. 7.4(a) (Gini plot): the most important parameters are the (SIZE-rescaled) WIDTH and LENGTH, the source dependent parameter DIST and the time parameter TIME GRADIENT. Note that since we deal with WOBBLE data and since source dependent parameters were used for the training, set of matrices calculated with respect to the source as well as the anti-source were generated. Throughout the analysis only 1 OFF region (the nominal anti-source) was used\(^5\). During this analysis step, further cuts were applied to the events used for the training procedures:

- standard cuts for rejecting the spark events (which are spurious events produced by discharges close to the photocathodes)
- $\text{SIZE} > 80 \text{ phe}$
- $\text{LEAKAGE} < 0.2$
- $\text{NUMBER OF ISLANDS} \leq 2$

\(^5\) Usually, for $\text{SIZE}$ greater than ~300 phe and in case of point-like sources, 3 OFF regions can be safely considered. The background estimation can be thus improved, enhancing the significance of a given observation. Note, however, that the quantity $(\text{Excesses})/(\text{Background})^{0.5}$, which is normally used to infer the sensitivity of the instrument, does not significantly improve (in value) by the determination of the background from more than 1 OFF region. In any case, since Segue 1 is affected by the light of $\gamma$-Leonis, always only 1 OFF region was taken into account.
In Fig. 7.4(b) the Hadronness distributions for test subsamples of MC-\(\gamma\) and hadron events used for the calculation are displayed. A good overall separation above 80 phes is achieved for Hadronness values less than \(\sim 0.5\). The differential true energy distribution of the MC-\(\gamma\) events (with simulated spectral index \(-2.6\)) after all analysis cuts (see the next subsection), is shown in Fig. 7.5(a), whereas in Fig. 7.5(b) the energy resolution and its bias are reported. The energy threshold of the analysis is defined as the peak of the energy distribution (after all analysis cuts) and is placed at \(\sim 100\) GeV. An energy resolution of 15\%-20\% is achieved for events with energies \(\geq 200\) GeV. The energy resolution reduces down to 30\% below 70 GeV.

Figure 7.5: (a) Distribution of MC-\(\gamma\) after the optimized cuts of the analysis. The maximum of the distribution defines the energy threshold of the analysis (\(\sim 100\) GeV, dashed vertical line). (b) Quality of the energy reconstruction defined by the resolution (red points) and its bias (blue) points.

After the matrixes for \(\gamma\)-hadron separation and energy reconstruction were properly calculated, the program MELIBEA took care to apply them to Segue 1 data, as well as to the
Crab Nebula and to the test MC-γ samples, in order to get the values of the parameters Hadronness and Estimated Energy for each event.

### 7.3.6 Optimization of the cuts and analysis performance

The analysis of Segue 1 data was carried out following the ALPHA approach (see section 4.3.7). The Hadronness and |ALPHA| cuts were optimized taking into account the performance of the analysis upon a Crab Nebula data sample. Since we are interested in establishing a source not previously known to emit γ-rays, the cuts were optimized in order to get the best differential sensitivity (see section 4.3.12) in each defined logarithmic energy bin, with an additional requirement of a minimum efficiency for the MC-γ events after the cuts (typically 40-50%). Indeed, this procedure is usually applied when searching for new VHE γ-rays emitters whereas looser cuts are typically used to determine the energy spectrum and the light curve of already established γ-ray sources.

In view of the optimization of the cuts, the Crab Nebula selected sample was divided randomly in two subsamples, each one corresponding to roughly 4.9 hours of effective observation time. Then, the combined |ALPHA| and Hadronness cuts were scanned in order to maximize the differential sensitivity of the observation of one of the two subsamples (train sample). The optimization was performed in logarithmic energy bins (10 bins per decade) between $10^{1.8}$ and $10^4$ GeV. For energies outside the range of the optimization, fixed |ALPHA| cuts of 20° and Hadronness cuts of 0.9 were always used. Finally, the selected cuts were applied to the other statistically independent subsample of Crab Nebula data (test sample) in order to check the consistency of the optimization. In this process, a couple of further analysis cuts was applied: SIZE > 100 phes and NUMBER OF ISLANDS = 1. The following constraints were also imposed:

- A minimum overall γ-events efficiency of 50% was required.
- The minimum allowed Hadronness cut was set to 0.06 (this was done just to preserve a minimum statistics for the real events, since the MC-γ efficiency at higher energies is high even if the strongest Hadronness cuts are applied).
- The minimum allowed |ALPHA| cut was set to 4°.
- The number of excess events had to be at least 5% of the number of the background events.

In table 7.2 the final set of optimized Hadronness and |ALPHA| cuts are reported together with the noteworthy quantities extracted from the Crab Nebula test sample, with an effective observation time of $T_{eff} = (17481 \pm 29)$ s. The analysis results have a quite stable behavior between 100 GeV up to energies of the order of ~1-2 TeV. For higher energies large fluctuations are present, due to the reduction of the background statistics.

In Fig. 7.6 the integral sensitivity above 100 GeV computed from the test Crab Nebula sample and achieved with the optimized cuts in the considered estimated energy bins is shown (in percentage units of Crab Nebula flux). The best integral sensitivity is obtained at ~250 GeV and it has a value of about 2% of the Crab Nebula flux. Note that tighter cuts in Hadronness and higher values of SIZE are typically necessary for obtaining the best discovery sensitivity.

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6 Even though the energy threshold given by the MC-γ distribution is around 100 GeV, also lower energy bins were considered. Indeed, the results coming from bins below the energy threshold are necessary to get correct results from the unfolding algorithm.

7 This constraint was relaxed to 3% for energies below 100 GeV.
<table>
<thead>
<tr>
<th>Estimated Energy [GeV]</th>
<th>Hadronness</th>
<th></th>
<th>ALPHA</th>
<th>MC-γ events efficiency</th>
<th>Significance [σ_{LiMA}]</th>
<th>Excesses Rate [events/min]</th>
<th>Background Rate [events/min]</th>
<th>Background Suppression [%]</th>
<th>Diff. Sensitivity [% Crab flux]</th>
</tr>
</thead>
<tbody>
<tr>
<td>$10^{-3}$ - $10^{-2.9}$</td>
<td>&lt;0.46</td>
<td>&lt;12</td>
<td>0.57</td>
<td>1.21</td>
<td>0.316 ± 0.260</td>
<td>9.727 ± 0.183</td>
<td>95.60</td>
<td>90.16 ± 74.38</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.9}$ - $10^{-2.6}$</td>
<td>&lt;0.43</td>
<td>&lt;12</td>
<td>0.51</td>
<td>2.21</td>
<td>0.971 ± 0.439</td>
<td>27.558 ± 0.308</td>
<td>95.66</td>
<td>49.34 ± 22.29</td>
<td></td>
</tr>
<tr>
<td>$10^{-2}$ - $10^{-2.1}$</td>
<td>&lt;0.38</td>
<td>&lt;14</td>
<td>0.50</td>
<td>4.07</td>
<td>1.867 ± 0.459</td>
<td>29.792 ± 0.320</td>
<td>96.69</td>
<td>26.69 ± 6.57</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.1}$ - $10^{-2.2}$</td>
<td>&lt;0.29</td>
<td>&lt;14</td>
<td>0.51</td>
<td>4.74</td>
<td>1.352 ± 0.285</td>
<td>11.165 ± 0.196</td>
<td>98.64</td>
<td>22.56 ± 4.76</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.2}$ - $10^{-2.3}$</td>
<td>&lt;0.18</td>
<td>&lt;12</td>
<td>0.50</td>
<td>7.78</td>
<td>1.133 ± 0.146</td>
<td>2.540 ± 0.093</td>
<td>99.62</td>
<td>12.84 ± 1.68</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.3}$ - $10^{-2.4}$</td>
<td>&lt;0.13</td>
<td>&lt;8</td>
<td>0.50</td>
<td>12.13</td>
<td>1.143 ± 0.096</td>
<td>0.765 ± 0.051</td>
<td>99.81</td>
<td>6.99 ± 0.63</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.4}$ - $10^{-2.5}$</td>
<td>&lt;0.07</td>
<td>&lt;8</td>
<td>0.51</td>
<td>12.51</td>
<td>0.903 ± 0.074</td>
<td>0.354 ± 0.035</td>
<td>99.84</td>
<td>6.01 ± 0.58</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.5}$ - $10^{-2.6}$</td>
<td>&lt;0.06</td>
<td>&lt;6</td>
<td>0.53</td>
<td>11.76</td>
<td>0.642 ± 0.057</td>
<td>0.154 ± 0.023</td>
<td>99.89</td>
<td>5.59 ± 0.65</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.6}$ - $10^{-2.7}$</td>
<td>&lt;0.06</td>
<td>&lt;6</td>
<td>0.61</td>
<td>10.11</td>
<td>0.511 ± 0.052</td>
<td>0.148 ± 0.023</td>
<td>99.85</td>
<td>6.86 ± 0.88</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.7}$ - $10^{-2.8}$</td>
<td>&lt;0.06</td>
<td>&lt;4</td>
<td>0.56</td>
<td>11.42</td>
<td>0.443 ± 0.042</td>
<td>0.038 ± 0.011</td>
<td>99.95</td>
<td>4.01 ± 0.71</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.8}$ - $10^{-2.9}$</td>
<td>&lt;0.06</td>
<td>&lt;4</td>
<td>0.62</td>
<td>8.55</td>
<td>0.296 ± 0.037</td>
<td>0.048 ± 0.013</td>
<td>99.91</td>
<td>6.78 ± 1.24</td>
<td></td>
</tr>
<tr>
<td>$10^{-2.9}$ - $10^{-3}$</td>
<td>&lt;0.08</td>
<td>&lt;6</td>
<td>0.77</td>
<td>11.30</td>
<td>0.395 ± 0.039</td>
<td>0.021 ± 0.008</td>
<td>99.95</td>
<td>3.32 ± 0.75</td>
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</tr>
<tr>
<td>$10^{-3}$ - $10^{-3.1}$</td>
<td>&lt;0.06</td>
<td>&lt;4</td>
<td>0.69</td>
<td>9.02</td>
<td>0.244 ± 0.030</td>
<td>0.010 ± 0.006</td>
<td>99.97</td>
<td>3.80 ± 1.19</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.1}$ - $10^{-3.2}$</td>
<td>&lt;0.08</td>
<td>&lt;8</td>
<td>0.81</td>
<td>8.93</td>
<td>0.216 ± 0.028</td>
<td>0.003 ± 0.003</td>
<td>99.98</td>
<td>2.47 ± 1.28</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.2}$ - $10^{-3.3}$</td>
<td>&lt;0.06</td>
<td>&lt;4</td>
<td>0.70</td>
<td>6.79</td>
<td>0.130 ± 0.022</td>
<td>0.003 ± 0.003</td>
<td>99.98</td>
<td>4.10 ± 2.16</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.3}$ - $10^{-3.4}$</td>
<td>&lt;0.06</td>
<td>&lt;4</td>
<td>0.77</td>
<td>5.49</td>
<td>0.089 ± 0.018</td>
<td>0.003 ± 0.003</td>
<td>99.97</td>
<td>5.99 ± 3.24</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.4}$ - $10^{-3.5}$</td>
<td>&lt;0.08</td>
<td>&lt;6</td>
<td>0.86</td>
<td>3.70</td>
<td>0.062 ± 0.018</td>
<td>0.014 ± 0.007</td>
<td>99.87</td>
<td>17.31 ± 6.54</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.5}$ - $10^{-3.6}$</td>
<td>&lt;0.09</td>
<td>&lt;6</td>
<td>0.89</td>
<td>2.62</td>
<td>0.031 ± 0.012</td>
<td>0.007 ± 0.005</td>
<td>99.93</td>
<td>24.48 ± 13.08</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.6}$ - $10^{-3.7}$</td>
<td>&lt;0.30</td>
<td>&lt;8</td>
<td>0.96</td>
<td>3.68</td>
<td>0.051 ± 0.015</td>
<td>0.007 ± 0.005</td>
<td>99.93</td>
<td>14.69 ± 6.72</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.7}$ - $10^{-3.8}$</td>
<td>&lt;0.34</td>
<td>&lt;10</td>
<td>0.95</td>
<td>2.68</td>
<td>0.031 ± 0.012</td>
<td>0.007 ± 0.005</td>
<td>99.90</td>
<td>24.48 ± 13.08</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.8}$ - $10^{-3.9}$</td>
<td>&lt;0.42</td>
<td>&lt;10</td>
<td>0.96</td>
<td>0.00</td>
<td>0.048 ± 0.013</td>
<td>0</td>
<td>100</td>
<td>not defined</td>
<td></td>
</tr>
<tr>
<td>$10^{-3.9}$ - $10^{-4}$</td>
<td>&lt;0.44</td>
<td>&lt;12</td>
<td>0.92</td>
<td>0.00</td>
<td>0.007 ± 0.011</td>
<td>0</td>
<td>100</td>
<td>not defined</td>
<td></td>
</tr>
</tbody>
</table>

Table 7.2: Hadronness and |ALPHA| cuts, for each considered energy bin, determined by the optimization of the differential sensitivity on the Crab Nebula train sample. Other quantities, such as the significance, the excesses rate, the background rate and suppression, and the differential sensitivity, are also reported and were calculated by the aid of the statistical independent Crab Nebula test sample. Note that above ∼1-2 TeV the analysis suffers large fluctuations due to the small amount of survived background events. In particular, in the last two energy bins no background events survived the cuts, preventing the determination of the sensitivity in that energy range. The energies below 100 GeV (separated by a black horizontal line) are below the analysis energy threshold but they were anyway considered in the optimization in view of the subsequent unfolding calculation of the Crab Nebula energy spectrum.
7.3. Data analysis

of the telescope, which is about 1.6% of the Crab Nebula flux at around 250 GeV [137].

The spectrum of the Crab Nebula (test sample) was also calculated and subsequently un-

folded with three different methods: Tikhonov [146], Bertero [147], Schmelling [148]. Good
agreement between all methods was found. In Fig. 7.7 the result of Tikhonov method is
shown. The resulting differential energy spectrum can be well described by a power law
function between 100 GeV and 3 TeV (black solid line in Fig. 7.7) with:

$$dF/dE = (4.8 \pm 0.2_{\text{stat}}) \times 10^{-10} \times (E/300\text{GeV})^{-2.36\pm0.05_{\text{stat}}} \left[\text{TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}\right]$$  (7.2)

The result is in good agreement with the Crab Nebula energy spectrum power law fit published
by the MAGIC Collaboration in 2008 [42], as shown by the dashed red line in Fig. 7.7:

$$dF/dE = (5.7 \pm 0.2_{\text{stat}}) \times 10^{-10} \times (E/300\text{GeV})^{-2.48\pm0.03_{\text{stat}}\pm0.2_{\text{syst}}} \left[\text{TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}\right]$$  (7.3)

The flux normalization factor of this parametrization is 15% higher than in the present anal-
ysis, which is within the systematic uncertainties (∼30%, see section 4.3.13). The value
obtained for the spectral index is also compatible within the errors. The results obtained by
the HEGRA [153] and H.E.S.S. [154] Collaborations (blue and yellow dashed lines in Fig. 7.7,
respectively) are also shown and agree well with the present result.

Energy spectra obtained with the other unfolding methods can also be well described by
a power law fit. The fit parameters are summarized in Table 7.3. The achieved results and
the values of $\chi^2$/NDF demonstrate the good agreement between different unfolding methods.
The systematic error related to the unfolding procedure can be estimated from the power law
fits of the different methods and it turns out to be much smaller and negligible with respect

![Figure 7.6: Integral sensitivity computed from the Crab Nebula test sample. The cuts used in each estimated energy bin are those reported in table 7.2. The best integral sensitivity of this analysis is about 2% of the Crab Nebula flux in the energy range between 250 GeV and 800 GeV.](image)
Figure 7.7: Differential energy spectrum of the Crab Nebula (test sample) unfolded using the **Tikhonov** method. The black solid line is the fit by a power law to the data points. The red, blue dashed lines and the yellow solid line are the parametrizations published respectively by the MAGIC Collaboration [42], by the HEGRA Collaboration [153] and by the H.E.S.S. Collaboration [154].

<table>
<thead>
<tr>
<th>Method</th>
<th>$f_0[10^{-10}$ TeV$^{-1}$ cm$^{-2}$ s$^{-1}$]</th>
<th>$\Gamma$</th>
<th>$\chi^2$/NDF</th>
</tr>
</thead>
<tbody>
<tr>
<td><strong>Tikhonov</strong></td>
<td>4.8±0.3</td>
<td>-2.36±0.05</td>
<td>5.42/8</td>
</tr>
<tr>
<td><strong>Bertero</strong></td>
<td>4.8±0.2</td>
<td>-2.34±0.05</td>
<td>12.27/8</td>
</tr>
<tr>
<td><strong>Schmelling</strong></td>
<td>4.9±0.2</td>
<td>-2.36±0.05</td>
<td>8.87/8</td>
</tr>
</tbody>
</table>

Table 7.3: Parameter values from a power law fit of form $dF/dE = f_0(E/300 \text{ GeV})^\Gamma$ to the results of different unfolding methods for the energy spectrum of Crab Nebula between 100 GeV and 3 TeV. The quality of the fits is given by the $\chi^2$ divided by the number of degrees of freedom (NDF).
to other systematic effects of the analysis (see section 4.3.13).

The results reported in this subsection show that the cuts are well defined as well as the criteria used for selecting the data. Also, the γ-hadron separation is working properly, demonstrating that the presence of the supplementary inefficiency due to the star in Segue 1 data (which actually were used as hadron train sample) does not affect the determination of the Hadronness parameter. Moreover, the problem related to the worsening of the PSF does not seem to affect the analysis results (less than few percents), since the Crab Nebula flux (extracted from data recorded in the same period of Segue 1 observation) is in good agreement with previous published results. In addition, as reported in Fig. 7.8, the monthly light curve shows an almost constant behavior and the estimated fluxes above 100 GeV are consistent with the expectations.

In the following data analysis steps, the Hadronness cuts summarized in table 7.2 were always applied in the corresponding energy bins.

7.3.7 Determination of the signal coming from Segue 1

Once the optimized cuts of the analysis were determined and successfully tested, they were applied on Segue 1 data for determining the significance of the observation above the energy threshold. In Fig. 7.9(a) the Alphaplot above 100 GeV for Segue 1 final data set is shown. The used Hadronness cuts are those listed in table 7.2, whereas a common cut in $|\text{ALPHA}|$ of 12° was applied to infer the number of excesses and the significance. Unfortunately, a clear mismatch between ON and OFF distributions is present, preventing a reliable determination of the significance of the observation. Hence, it is evident that the applied optimized Hadronness cuts could not remove completely the local inefficiency due to the light of the star. The fact that the ON distribution is systematically below the OFF one (for $|\text{ALPHA}| \leq 60°$) basically occurs because the distance between the star and the source ($\sim 0.68°$) is smaller than the distance between the star and the anti-source ($\sim 1.07°$ in case of W1 observations and $\sim 1.02°$ for W2 ones), making the acceptance of the camera (which depends on the parameter
DIST) slightly smaller with respect to the source position rather than with respect to the anti-source position.

In order to better quantify the mismatch, the distribution of the significances in each Alphaplot bin between 20° and 80° (the so called tail region, see section 4.3.7) was calculated\(^8\). In case of good quality WOBBLE data, the ON and OFF distribution should match in that region within the statistic errors, as no γ signal is expected there\(^9\). Moreover, no extra normalization (a part from the geometrical ones given by the number of considered OFF positions, which is equal to 1 in our case) should be necessary for guaranteeing the matching of the two distributions. Therefore, a possible good check consists in monitoring the distribution of the significances \((\sigma_{\text{LiMa}})\) calculated in each Alphaplot bin of the tail region. A good behavior of the ON and OFF distributions should result in a mean value of the significances around zero with RMS close to 1. In Fig. 7.9(b), the distribution of the significances calculated in each Alphaplot bin of the tail region are shown (blue empty crosses). The distributions in the first half of the region (20° - 50°, red empty crosses) and the second half (50° - 80°, green empty crosses) were also calculated. The solid vertical lines represent the mean values of the distributions. As expected, the significances do not distribute around zero, especially in the first part of the tail region, being the mean values respectively -0.941±0.909 (whole region), -1.459±0.894 (first half), -0.423±0.565 (second half).

For constraining the range of energy affected by a significant mismatch, the distributions of the significances of each Alphaplot bin of the tail region were monitored separately in different energy bins above 100 GeV. To cross-check the scan, a subsample of NGC 1275 data was also considered. This source was observed by the MAGIC-I telescope during November and December 2008 under the same hardware and observational conditions of Segue 1 data. No significant excesses were found above an energy threshold of 100 GeV \([254]\). Since no extra light due to bright stars in the field of view was present during NGC 1275 survey, these data can be considered as a good guideline for comparing the characteristics which Segue 1 data should have in the |ALPHA| tail region in absence of significance star related effects. NGC 1275 data were analyzed exactly in the same way Segue 1 data were processed (from calibration and data selection to hadron separation and energy reconstruction). Moreover, only NGC 1275 data taking in those days in which both the sources were observed were considered for the check, namely the 28th, 29th of November and the 1st and 2nd of December 2008. Also, the Zenith range of Segue 1 and NGC 1275 observations were very similar (~13° to ~34°). The final amount of NGC 1275 data available for the test was roughly 10 hours. The scan was performed by using exactly the same cuts in each energy bin for both the data samples (see table 7.2). All the bins with energies greater than ~400 GeV were stacked together in a single Alphaplot. Tab. 7.4 and 7.5 summarize the achieved results. In case of Segue 1 data, the distributions of the significances start to show an expected behavior only for energies above \(10^{2.3} \sim 200 \text{ GeV}\) (hereafter we refer to this energy simply as 200 GeV), with mean values compatible to those calculated for the unbiased NGC 1275 data. For energies below 200 GeV significant disagreements between the two data sets are present: particularly in the first half of the tail region, the mean values for Segue 1 data are all greater (in module)

\(^8\)As shown in Fig 7.9(a), the number of bins used for the Alphaplots were always set to 45, corresponding to \(2°\) of |ALPHA| parameter each.

\(^9\)Note that a good agreement in the tail region for WOBBLE data is achieved if also no huge γ-ray signal is present in the source position (as in the Crab Nebula data for energies above ~200 GeV): in fact, a huge number of excesses located in the ON signal region (i.e. at low |ALPHA| values) would be spread in the OFF |ALPHA| distribution at middle values of |ALPHA|, jeopardizing the good agreement of the distributions in the tail region. Nevertheless, since no huge signal is expected from Segue 1 data, a good agreement in the |ALPHA| tail region can be used as a good test for checking a fair behavior of the ON and OFF |ALPHA| distributions. For more detail about this issue see [137].
7.3. Data analysis

Figure 7.9: (a) Alphaplot above 100 GeV for Segue 1 final sample. A residual mismatch between ON distribution (black histogram) and OFF distribution (red histogram) is evident. The green histogram is the excess distribution. (b) Distributions of the significances calculated respectively in the whole tail region ($20^\circ < |\alpha| < 80^\circ$, blue crosses), in the first half ($20^\circ < |\alpha| < 50^\circ$, red crosses) and in the second half ($50^\circ < |\alpha| < 80^\circ$, green crosses). The corresponding mean values are indicated by the solid vertical lines.
than the maximum values found for NGC 1275 data.

In Fig. 7.10, the \textit{Alphaplot} for energies above 200 GeV is reported together with the distrib-

\begin{table}[h]
\centering
\begin{tabular}{cccccc}
\hline
Energy [GeV] & Entries on/off & \langle \sigma_{LiMa} \rangle (RMS) & \langle \sigma_{LiMa} \rangle (RMS) & \langle \sigma_{LiMa} \rangle (RMS) & \langle \sigma_{LiMa} \rangle (RMS) \\
\hline
\multirow{4}{*}{\textit{10}^{2} - \textit{10}^{2.1}} & 166005(168042) & -0.642(0.986) & -0.712(1.022) & -0.573(0.945) & -0.753(0.945) \\
\multirow{4}{*}{\textit{10}^{2.1} - \textit{10}^{2.2}} & 58912(60315) & -0.722(1.306) & -1.281(1.186) & -0.163(1.175) & -0.163(1.175) \\
\multirow{4}{*}{\textit{10}^{2.2} - \textit{10}^{2.3}} & 16854(17034) & -0.170(1.126) & -0.647(1.140) & 0.307(0.883) & 0.307(0.883) \\
\multirow{4}{*}{\textit{10}^{2.3} - \textit{10}^{2.4}} & 8614(8516) & 0.130(1.077) & 0.255(0.954) & 0.006(1.041) & 0.006(1.041) \\
\multirow{4}{*}{\textit{10}^{2.4} - \textit{10}^{2.5}} & 3648(3613) & 0.136(0.831) & -0.125(0.892) & 0.295(1.012) & 0.295(1.012) \\
\multirow{4}{*}{\textit{10}^{2.5} - \textit{10}^{2.6}} & 2188(2185) & 0.345(0.913) & 0.395(0.839) & 0.295(1.012) & 0.295(1.012) \\
\multirow{4}{*}{\textit{10}^{2.6} - \textit{10}^{2.7}} & 4391(4487) & -0.183(0.824) & -0.271(0.689) & -0.094(0.932) & -0.094(0.932) \\
\hline
\end{tabular}
\caption{Mean value ($\langle \sigma_{LiMa} \rangle$) and RMS (in brackets) of the distribution of the significances calculated from Segue 1 data in each \textit{Alphaplot} bin of the tail region for different energy ranges. The corresponding quantities calculated in the first and second part of the tail region are also reported.}
\end{table}

\begin{table}[h]
\centering
\begin{tabular}{cccccc}
\hline
Energy [GeV] & Entries on/off & \langle \sigma_{LiMa} \rangle (RMS) & \langle \sigma_{LiMa} \rangle (RMS) & \langle \sigma_{LiMa} \rangle (RMS) & \langle \sigma_{LiMa} \rangle (RMS) \\
\hline
\multirow{4}{*}{\textit{10}^{2} - \textit{10}^{2.1}} & 61154(61619) & -0.243(0.832) & -0.037(0.710) & -0.450(0.891) & -0.450(0.891) \\
\multirow{4}{*}{\textit{10}^{2.1} - \textit{10}^{2.2}} & 21076(21473) & -0.356(0.947) & -0.231(0.927) & -0.482(0.950) & -0.482(0.950) \\
\multirow{4}{*}{\textit{10}^{2.2} - \textit{10}^{2.3}} & 6410(6196) & 0.345(0.913) & 0.395(0.799) & 0.295(1.012) & 0.295(1.012) \\
\multirow{4}{*}{\textit{10}^{2.3} - \textit{10}^{2.4}} & 3327(3269) & 0.136(0.831) & -0.125(0.892) & 0.397(0.670) & 0.397(0.670) \\
\multirow{4}{*}{\textit{10}^{2.4} - \textit{10}^{2.5}} & 1313(1251) & 0.226(0.828) & -0.112(0.857) & 0.563(0.639) & 0.563(0.639) \\
\multirow{4}{*}{\textit{10}^{2.5} - \textit{10}^{2.6}} & 721(769) & -0.273(1.328) & 0.040(1.257) & -0.586(1.323) & -0.586(1.323) \\
\multirow{4}{*}{\textit{10}^{2.6} - \textit{10}^{2.7}} & 1661(1680) & -0.060(0.7870 & -0.118(0.792) & -0.002(0.778) & -0.002(0.778) \\
\hline
\end{tabular}
\caption{Mean value ($\langle \sigma_{LiMa} \rangle$) and RMS (in brackets) of the distribution of the significances calculated from NGC 1275 data in each \textit{Alphaplot} bin of the tail region for different energy ranges. The corresponding quantities calculated in the first and second part of the tail region are also reported.}
\end{table}

As resulting from these tests, Segue 1 observation can be considered not affected by significant star related inhomogeneities for energies above 200 GeV. A possible way to proceed can...
### 7.3. Data analysis

#### Figure 7.10:

(a) *Alphaplot* above 200 GeV for Segue 1 final sample. No significant mismatch between ON distribution (black histogram) and OFF distribution (red histogram) is present, as well as no hint of signal. The green histogram is the excess distribution.

(b) Distributions of the significances calculated in each *Alphaplot* bin respectively of the whole tail region (*20° < |ALPHA| < 80°*, blue crosses), the first half (*20° < |ALPHA| < 50°*, red crosses) and the second half (*50° < |ALPHA| < 80°*, green crosses) for energies above 200 GeV. The corresponding mean values are indicated by the solid vertical lines.
7. OBSERVATION OF SEGUE 1 WITH THE MAGIC-I TELESCOPE

Figure 7.11: (a) Distributions of the DIST parameter for estimated energies between 100 GeV and 200 GeV calculated with respect to the source (ON data, black histogram) and with respect to the anti-source (OFF data, red histogram). In the region delimited within vertical dashed red lines the entries of the ON distributions are systematically less than the entries of the OFF distribution. (b) Distributions of the DIST parameter for estimated energies above 200 GeV calculated with respect to the source (ON data, black histogram) and with respect to the anti-source (OFF data, red histogram). The two distributions are well matching within the errors.
thus be the application of an energy cut at 200 GeV (hence losing the possibility to extract some results in the lowest energy range) or a drastic cut in DIST between \(\sim 0.6^\circ\) and \(\sim 1^\circ\), for energies below 200 GeV (which, anyway, should have the consequence to spoil almost completely the achievable results). Therefore, since the low energy regions are particularly interesting for indirect DM search purposes, the symmetries of the star and source positions were taken into account with the aim of finding out a set of cuts below 200 GeV which could eliminate the problems related to the star inefficiency (rejecting as less events as possible) and allow the extraction of physical results also between 100 and 200 GeV. The basic idea was to remove the events lying in a certain part of the camera near the spot of the star light. Of course, since the data were taken in WOBBLE mode, special care had to be addressed in order to maintain the basic symmetries between source and anti-source. To better understand this approach it is necessary to introduce a set of positions in the camera which we must deal with (see Fig. 7.12):

- the nominal source position \((x_S, y_S)\) and the nominal anti-source position \((x_{AS}, y_{AS})\)
- the nominal star position \((x_\eta, y_\eta)\)
- the nominal position of the star as seen by the anti-source \((x_\eta W, y_\eta W)\)
- the nominal position symmetric to \((x_\eta W, y_\eta W)\) with respect to the axis passing through the anti-source position and perpendicular to the axis joining the nominal source and anti-source positions, tagged as \((x_\eta WP, y_\eta WP)\)

In Fig. 7.12 the positions listed above are drawn for one chosen data subrun taken as example. Of course, for different subruns, the positions in the camera are in general different since the geometry is constrained by the source position which is always moving along the WOBBLE circle (but the relative distances remain unchanged). To make possible the calculation of the relative distances in the camera with respect to the positions drawn in Fig. 7.12, new simple classes were defined in MARS (see section 4.1). First of all, given the RA and the DEC coordinates of \(-Leonis\) it was straightforward to calculate the coordinates \((x_S, y_S)\) in the camera by using the already implemented MARS classes [143], which are exactly those used to calculate the nominal source position in the camera giving the RA and the DEC of Segue 1. Then, once the coordinates of the star in the camera were available for each event, the other positions of Fig. 7.12 were deduced just applying basic geometrical relations between the axes joining the different positions, leading to:

\[
\begin{align*}
    y_\eta W &= -y_S - y_S - y_\eta + y_S - y_S + x_S \\
    x_\eta W &= x_\eta + x_S,y_S \cdot (y_\eta W - y_\eta)
\end{align*}
\]

\(^{10}\)W stands for WOBBL ED
\(^{11}\)P stands for PARITY
Figure 7.12: Positions in the camera listed in the text which were used to define the star-cuts.
and

\[ x_{\eta WP} = \frac{-g - \sqrt{g^2 - 4fh}}{2g} \]

\[ y_{\eta WP} = y_{\eta W} + \frac{y_{AS}}{x_{AS}} \left( x_{\eta WP} - x_{\eta W} \right) \]

with

\[ f = 1 + \frac{y^2_{AS}}{x^2_{AS}} \]

\[ g = -2x_{\eta W} \cdot \left( 1 + \frac{y^2_{AS}}{x^2_{AS}} \right) \]

\[ h = x^2_{\eta W} \cdot \left( 1 + \frac{y^2_{AS}}{x^2_{AS}} \right) - 4 \left[ (-x_{\eta W} + i)^2 + (-y_{\eta W} + l)^2 \right] \]

\[ l = \frac{y_{AS} - x_{\eta W} \cdot \frac{x_{AS}}{y_{AS}} + y_{\eta W} \cdot \frac{x^2_{AS}}{y_{AS}} + \frac{x^2_{AS}}{y_{AS}}}{1 + \frac{x^2_{AS}}{y^2_{AS}}} \]

\[ i = x_{\eta W} + (l - y_{\eta W}) \cdot \frac{x^2_{AS}}{y_{AS}} \]

After all the coordinates in the camera of the positions sketched in Fig. 7.12 were available for each event, the following cut was considered (only for events with estimated energy below 200 GeV): the events with CoG (i.e., the centroid of the image of each parametrized event) lying at a distance from the source between 0.6° and 1° and at the same time within a circle of radius 0.8° centered in the star position \((x, y)\) were removed. In order to keep the symmetry with respect to the anti-source, a similar cut was applied also for the OFF distributions by replacing the coordinates of the star \((x, y)\) with \((x_{WP}, y_{WP})\). Hereafter these cuts will be called star-cuts. The value 0.8° was chosen as the minimum value which guarantees a good agreement between ON and OFF DIST distributions between 0.6° and 1°, as shown in Fig. 7.13. The location of the events removed by the star-cuts, once applied to the same subrun used to sketch the main positions in Fig. 7.12, is shown in Fig. 7.14 respectively for the OFF distribution (left plot) and for the ON distribution (right plot) of the CoGs. Note that those plots must be considered as examples since for different subruns the locations of the rejected events due to the star-cuts are in general different, since the source as well as the star always rotate in the camera.

In Fig. 7.15, the Alphaplot between 100 GeV and 200 GeV is shown before and after the star-cuts. A nice agreement between the ON and OFF distribution in the tail region is achieved after removing the (star position dependent) parts of the camera affected by the star related inefficiency. The number of events rejected by the star-cuts are of the order of 20%.

In order to be able to apply the star-cuts on the MC-\(\gamma\) data (which is essential for calculating in the correct way the final Effective Collection Area) as well as to check those cuts on the

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12To be correct, another cut should be performed. Indeed, also the anti-source could be affected by the loss of efficiency due to the star around \(~1.05°\). Anyway, at those distances the acceptance at low energies is much smaller compared to the distances which are affected by a clear mismatch, as shown in Fig. 7.11(a). No further cuts were therefore considered.
Figure 7.13: Distributions of the DIST parameter for estimated energies between 100 GeV and 200 GeV calculated with respect to the source (ON data, black histogram) and with respect to the anti-source (OFF data, red histogram) after the application of the star-cuts. A good agreement between 0.6° and 1° is achieved within the errors. The rejected events due to the star-cuts are of the order of 20% (see Fig. 7.11(a)).

Figure 7.14: CoG plots after the application of the star-cuts for the ON data (left plot) and for the OFF data (right plot), for the same subrun used as example for illustrating the relevant positions in the camera which allow the definition of the star-cuts (see Fig. 7.12). Note that in order to have a good statistic from a single subrun only the star-cuts were applied for producing these plots.
7.3. Data analysis

| α | 0 10 20 30 40 50 60 70 80 90 |
|---|---|---|---|---|---|---|---|---|---|---|
| # counts / 2 | 0 | 1000 | 2000 | 3000 | 4000 | 5000 | 6000 | 7000 | 8000 |

\[ T_{\text{ON}} = 29.42 \text{ h} \]
\[ N_{\text{ON}} = 61482 \]
\[ N_{\text{OFF}} = 63533 \pm 252.1 \]
\[ N_{\text{EXCESS}} = -2051 \pm 353.6 \]
\[ \sigma_{\text{LiMa}} = -5.80 \sigma \]

$\alpha$ Alphaplot between 100 GeV and 200 GeV before the star-cuts.

$\alpha$ Alphaplot between 100 GeV and 200 GeV after the star-cuts.

**Figure 7.15:** (a) ON and OFF $|\text{ALPHA}|$ distributions between 100 GeV and 200 GeV for Segue 1 final sample. The mismatch between ON distribution (black histogram) and OFF distribution (red histogram) is evident. The green histogram is the excess distribution. (b) ON and OFF $|\text{ALPHA}|$ distributions between 100 GeV and 200 GeV for Segue 1 final sample, after the application of the star-cuts (see text for details). The mismatch between ON distribution (black histogram) and OFF distribution (red histogram) is completely removed. The green histogram is the excess distribution.
Crab Nebula data, a fake $\eta$-Leonis like position in the camera was determined for those two data samples. For MC-$\gamma$ data, since the source position is always placed in the camera at coordinates $x=0.4^\circ$ and $y=0^\circ$, the position of the fake star was simply calculated as the mean value of the $\eta$-Leonis positions (in the camera) of those events of Segue 1 data having nominal source position around $x=0.4^\circ$ and $y=0^\circ$. In Fig 7.16 the location of MC-$\gamma$ events (with estimated energy between 100 GeV and 200 GeV) removed by the star-cuts is shown. Note that for MC-$\gamma$ events only the star-cuts with respect to the nominal source position must be applied. In case of Crab Nebula data, the RA and DEC coordinates of the fake star were calculated by replacing Segue 1 coordinates with the Crab Nebula ones and by determining the position of the fake star as to simulate exactly the relative position between Segue 1 and $\eta$-Leonis. Then, all the other coordinates in the camera (needed to define the star-cuts), were calculated just following the same procedure adopted for Segue 1 data.

Once the fake star was simulated in the MC-$\gamma$ and Crab Nebula data samples, some checks were performed. The differential true energy distribution of the MC-$\gamma$ events after all analysis cuts and the star-cuts (applied only for events with estimated energy below 200 GeV) is shown in Fig. 7.17(a), whereas in Fig. 7.17(b) the energy resolution and bias are reported. The energy threshold of the analysis is not affected by the star-cuts, as the energy peak is still around 100 GeV. The percentage of $\gamma$-ray events rejected by the star-cuts is of the order of 20%. Also the energy resolution and the bias have values similar to those found without the application of the star-cuts. In Fig. 7.18, the Effective Collection Area before and after the analysis cuts and after the analysis cuts plus the star-cuts is shown. The effect is not negligible up to $\sim$300 GeV in true MC-$\gamma$ energies.

In order to check the correct behavior of the analysis after the introduction below 200 GeV of the star-cuts, the spectrum of the Crab Nebula (test sample) was recalculated (after the application of the analysis cuts listed in Tab. 7.2 and the star-cuts on both the Crab Nebula and MC-$\gamma$ data) and then unfolded as done in subsection 7.3.6. In Fig. 7.19 the result of Tikhonov method is shown: the resulting parametrization (black solid line in Fig. 7.19) given by a power law fit of the data between 100 GeV and 3 TeV is still compatible with the results.
7.3. Data analysis

Figure 7.17: (a) Distribution of MC-γ after the application of the star-cuts for estimated energy below 200 GeV. The maximum of the distribution defines the energy threshold of the analysis (~100 GeV, dashed vertical line). (b) Quality of the energy reconstruction defined by the resolution (red points) and the bias (blue) points, after the star-cuts.

Figure 7.18: Effective Collection Area before and after the analysis cuts (blue and green points) and after the application of the analysis cuts and of the star-cuts (red points). The effect of the star-cuts is not negligible up to ~300 GeV.
7. OBSERVATION OF SEGUE 1 WITH THE MAGIC-I TELESCOPE

<table>
<thead>
<tr>
<th>Method</th>
<th>$f_0 \times 10^{-10}$ Tev$^{-1}$ cm$^{-2}$ s$^{-1}$</th>
<th>$\Gamma$</th>
<th>$\chi^2$/NDF</th>
</tr>
</thead>
<tbody>
<tr>
<td>Tikhonov</td>
<td>4.6±0.2</td>
<td>-2.35±0.05</td>
<td>5.36/8</td>
</tr>
<tr>
<td>Bertero</td>
<td>4.7±0.2</td>
<td>-2.33±0.05</td>
<td>10.99/8</td>
</tr>
<tr>
<td>Schmelling</td>
<td>4.7±0.1</td>
<td>-2.35±0.05</td>
<td>6.44/8</td>
</tr>
</tbody>
</table>

Table 7.6: Parameter values from a power law fit of form $dF/dE = f_0 (E/300 \text{ GeV})^\Gamma$ to the results of different unfolding methods for the energy spectrum of Crab Nebula after the application of the star-cuts. The fits were performed between 100 GeV and 3 TeV. The quality of the fits is given by the $\chi^2$ divided by the number of degrees of freedom (NDF).

Achieved without star-cuts (see subsection 7.3.6):

$$dF/dE = (4.6 \pm 0.2_{stat}) \times 10^{-10} \times (E/300 \text{ GeV})^{-2.35\pm0.05_{stat}} [\text{TeV}^{-1}\text{cm}^{-2}\text{s}^{-1}] \quad (7.6)$$

In Table 7.6 the results obtained by the different unfolded methods are listed. A good agreement between different methods and previous results (i.e., without the application of the star-cuts) is achieved.

These last results obtained after the application of the star-cuts, show that the analysis of Segue 1 data can be performed also below 200 GeV. Indeed, the symmetric rejection of the events which lie nearby the part of the camera affected by the light of $\eta$-Leonis succeeded in removing the significant mismatches between ON and OFF distributions and gave reliable
results once applied to Crab Nebula test sample. Note that since the star-cuts were applied also to MC-\(\gamma\) data, the Effective Collection Area is properly calculated. This is particularly important for deriving correct values for the upper limit of the \(\gamma\)-ray flux coming from Segue 1 source.

### 7.4 Results

In Fig. 7.20, the Alphaplot above 100 GeV of Segue 1 final sample after the analysis cuts and the star-cuts is shown. For the computation of the excesses a common |\(\alpha\)| cut of 12\(^\circ\) was applied. No hint of signal is present, being the overall significance equal to -0.69.

Since no signal was found, the upper limits on the flux of Segue 1 above 100 GeV were computed from the Alphaplots, considering separately each energy bin defined in Table 7.2. For energies below 200 GeV, the star-cuts were always applied. Details on the upper limit calculation can be found in section 4.3.10. In Tab. 7.7 the differential upper limits on the flux achieved from Segue 1 observation are listed for different energy ranges and for a general power law spectrum with index -1.8. In Fig. 7.21, the corresponding Alphaplots are shown, whereas in Fig. 7.22 the achieved differential upper limits summarized in Table 7.7 are graphically displayed.

In Tables 7.8 to 7.13, the integral flux upper limits from Segue 1 observation are reported for generic power law spectra with index \(\Gamma\) respectively -1, -1.5, -1.8, -2, -2.2 and -2.4 and for rising energy threshold between \(10^2\) GeV and \(10^3\) GeV (with logarithmic step of 0.1). In Fig. 7.23 the achieved integral upper limits summarized in Tables 7.8 to 7.13 are graphically displayed.
7. OBSERVATION OF SEGUE 1 WITH THE MAGIC-I TELESCOPE

(a) Alphaplot for $10^2 < E \text{[GeV]} < 10^{2.5}$.

(b) Alphaplot for $10^{2.5} < E \text{[GeV]} < 10^3$.

(c) Alphaplot for $10^3 < E \text{[GeV]} < 10^{3.5}$.

(d) Alphaplot for $10^{3.5} < E \text{[GeV]} < 10^4$.

Figure 7.21: Segue 1 final sample Alphaplots between $10^2$ GeV and $10^{2.5}$ GeV (a), $10^{2.5}$ GeV and $10^3$ GeV (b), $10^3$ GeV and $10^{3.5}$ GeV (c) and $10^{3.5}$ GeV and $10^4$ GeV (d). The black histograms are the ON distributions and the red histograms are the OFF distributions.
7.4. Results

Figure 7.22: Differential flux upper limits summarized in Table 7.7.

Figure 7.23: Plot of the integral flux upper limits ($\Phi^{UL}$) summarized in Tables 7.8 to 7.13 for generic power laws with spectral indices -1, -1.5, -1.8, -2, -2.2 and -2.4 and rising energy threshold between $10^2$ GeV and $10^3$ GeV (with logarithmic step of 0.1).
### Differential Flux Upper Limits, Spectral index $\Gamma = -1.8$

<table>
<thead>
<tr>
<th>Energy range [GeV]</th>
<th>$N_{on}$</th>
<th>$N_{off}$</th>
<th>$N_{U.L.}$</th>
<th>$\sigma_{LiMa}$</th>
<th>$A_{eff}$</th>
<th>$E_s$</th>
<th>$\phi_{U.L.}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$10^2 - 10^{2.5}$</td>
<td>51871</td>
<td>52271</td>
<td>398.8</td>
<td>-1.24</td>
<td>1.453 x 10^9</td>
<td>202.2</td>
<td>5.506 x 10^{-11}</td>
</tr>
<tr>
<td>$10^{2.5} - 10^3$</td>
<td>696</td>
<td>657</td>
<td>155.9</td>
<td>1.06</td>
<td>2.881 x 10^9</td>
<td>602.5</td>
<td>3.610 x 10^{-12}</td>
</tr>
<tr>
<td>$10^3 - 10^{3.5}$</td>
<td>99</td>
<td>77</td>
<td>71.8</td>
<td>1.66</td>
<td>4.049 x 10^9</td>
<td>1834.8</td>
<td>3.888 x 10^{-13}</td>
</tr>
<tr>
<td>$10^{3.5} - 10^4$</td>
<td>69</td>
<td>57</td>
<td>48.1</td>
<td>1.07</td>
<td>5.129 x 10^9</td>
<td>5937.0</td>
<td>6.352 x 10^{-14}</td>
</tr>
</tbody>
</table>

**Table 7.7:** Differential flux upper limits ($\phi_{U.L.}$) from Segue 1 observation in four energy bins between $10^2$ and $10^4$ GeV, for a general power law spectrum with index -1.8. The value $E_s$ represents the mean value of the energy in a particular energy range and is given by equation 4.17.

### Integral Flux Upper Limits, Spectral index $\Gamma = -1$

<table>
<thead>
<tr>
<th>Energy [GeV]</th>
<th>$N_{on}$</th>
<th>$N_{off}$</th>
<th>$N_{U.L.}$</th>
<th>$\sigma_{LiMa}$</th>
<th>$A_{eff}$</th>
<th>$\phi_{U.L.}$</th>
<th>$\Phi_{U.L.}$</th>
</tr>
</thead>
<tbody>
<tr>
<td>$&gt;10^2$</td>
<td>52978</td>
<td>53301</td>
<td>453.4</td>
<td>-0.99</td>
<td>1.706 x 10^{10}</td>
<td>1.614 x 10^{-13}</td>
<td>6.339 x 10^{-12}</td>
</tr>
<tr>
<td>$&gt;10^{2.1}$</td>
<td>18835</td>
<td>19233</td>
<td>174.1</td>
<td>-2.04</td>
<td>1.686 x 10^{10}</td>
<td>6.197 x 10^{-14}</td>
<td>2.357 x 10^{-12}</td>
</tr>
<tr>
<td>$&gt;10^{2.2}$</td>
<td>6122</td>
<td>6374</td>
<td>92.9</td>
<td>-2.25</td>
<td>1.662 x 10^{10}</td>
<td>3.307 x 10^{-14}</td>
<td>1.215 x 10^{-12}</td>
</tr>
<tr>
<td>$&gt;10^{2.3}$</td>
<td>3012</td>
<td>3088</td>
<td>110.2</td>
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<td>1.716 x 10^{-12}</td>
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<td>$&gt;10^{2.7}$</td>
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<td>1.460 x 10^{10}</td>
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<td>123.8</td>
<td>0.84</td>
<td>1.412 x 10^{10}</td>
<td>4.447 x 10^{-14}</td>
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<td>145.7</td>
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<td>1.352 x 10^{10}</td>
<td>5.255 x 10^{-14}</td>
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</table>

**Table 7.8:** Integral flux upper limits ($\Phi_{U.L.}$) from Segue 1 observation for a general power law spectrum with index -1 as function of different energy thresholds between $10^2$ GeV to $10^3$ GeV.

### Integral Flux Upper Limits, Spectral index $\Gamma = -1.5$

<table>
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<tr>
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<th>$N_{on}$</th>
<th>$N_{off}$</th>
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<th>$\sigma_{LiMa}$</th>
<th>$A_{eff}$</th>
<th>$\phi_{U.L.}$</th>
<th>$\Phi_{U.L.}$</th>
</tr>
</thead>
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<td>53301</td>
<td>453.4</td>
<td>-0.99</td>
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<td>2.914 x 10^{-13}</td>
<td>1.090 x 10^{-11}</td>
</tr>
<tr>
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<td>19233</td>
<td>174.1</td>
<td>-2.04</td>
<td>1.684 x 10^{10}</td>
<td>1.097 x 10^{-13}</td>
<td>3.855 x 10^{-12}</td>
</tr>
<tr>
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<td>6122</td>
<td>6374</td>
<td>92.9</td>
<td>-2.25</td>
<td>1.661 x 10^{10}</td>
<td>5.727 x 10^{-14}</td>
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<td>3088</td>
<td>110.2</td>
<td>-0.97</td>
<td>1.636 x 10^{10}</td>
<td>6.655 x 10^{-14}</td>
<td>2.047 x 10^{-12}</td>
</tr>
<tr>
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<td>1654</td>
<td>194.2</td>
<td>0.57</td>
<td>1.600 x 10^{10}</td>
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<td>3.390 x 10^{-12}</td>
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<td>249.7</td>
<td>1.67</td>
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<td>761</td>
<td>147.4</td>
<td>0.79</td>
<td>1.513 x 10^{10}</td>
<td>8.183 x 10^{-14}</td>
<td>2.309 x 10^{-12}</td>
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<td>$&gt;10^{2.7}$</td>
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<td>139.7</td>
<td>0.96</td>
<td>1.459 x 10^{10}</td>
<td>7.517 x 10^{-14}</td>
<td>2.076 x 10^{-12}</td>
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<tr>
<td>$&gt;10^{2.8}$</td>
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<td>509</td>
<td>123.8</td>
<td>0.84</td>
<td>1.410 x 10^{10}</td>
<td>6.494 x 10^{-14}</td>
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<td>1.350 x 10^{10}</td>
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<td>134.5</td>
<td>1.36</td>
<td>1.271 x 10^{10}</td>
<td>6.653 x 10^{-14}</td>
<td>1.797 x 10^{-12}</td>
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</tbody>
</table>

**Table 7.9:** Integral flux upper limits ($\Phi_{U.L.}$) from Segue 1 observation for a general power law spectrum with index -1.5 as function of different energy between thresholds $10^2$ GeV to $10^3$ GeV.
### 7.4. Results

#### Integral Flux Upper Limits, Spectral index $\Gamma = -1.8$

<table>
<thead>
<tr>
<th>Energy [GeV]</th>
<th>$N_{on}$</th>
<th>$N_{off}$</th>
<th>$N^{*}$</th>
<th>$\sigma_{LiMa}$ [cm$^2$]</th>
<th>Aeff [cm$^2$]</th>
<th>$\Phi^{U.L.}$ [TeV$^{-1}$ cm$^{-2}$ s$^{-1}$]</th>
<th>$\Phi^{U.L.}$ [cm$^{-2}$ s$^{-1}$]</th>
</tr>
</thead>
<tbody>
<tr>
<td>$&gt;10^2$</td>
<td>52978</td>
<td>53301</td>
<td>453.4</td>
<td>-0.99</td>
<td>1.704·10$^{10}$</td>
<td>5.743·10$^{-13}$</td>
<td>1.505·10$^{-11}$</td>
</tr>
<tr>
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<td>18835</td>
<td>19233</td>
<td>174.1</td>
<td>-2.04</td>
<td>1.684·10$^{10}$</td>
<td>2.087·10$^{-13}$</td>
<td>5.168·10$^{-12}$</td>
</tr>
<tr>
<td>$&gt;10^{2.2}$</td>
<td>6122</td>
<td>6374</td>
<td>92.9</td>
<td>-2.25</td>
<td>1.660·10$^{10}$</td>
<td>1.053·10$^{-13}$</td>
<td>2.439·10$^{-12}$</td>
</tr>
<tr>
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<td>3088</td>
<td>110.2</td>
<td>-0.97</td>
<td>1.635·10$^{10}$</td>
<td>1.184·10$^{-13}$</td>
<td>2.557·10$^{-12}$</td>
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<tr>
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<td>1654</td>
<td>194.2</td>
<td>0.57</td>
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<td>1.950·10$^{-13}$</td>
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<td>1.67</td>
<td>1.559·10$^{10}$</td>
<td>2.343·10$^{-13}$</td>
<td>4.959·10$^{-12}$</td>
</tr>
<tr>
<td>$&gt;10^{2.6}$</td>
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<td>0.79</td>
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<td>1.291·10$^{-13}$</td>
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<td>0.96</td>
<td>1.458·10$^{10}$</td>
<td>1.136·10$^{-13}$</td>
<td>2.417·10$^{-12}$</td>
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<tr>
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<td>123.8</td>
<td>0.84</td>
<td>1.410·10$^{10}$</td>
<td>9.493·10$^{-14}$</td>
<td>2.023·10$^{-12}$</td>
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<tr>
<td>$&gt;10^{2.9}$</td>
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<td>1.350·10$^{10}$</td>
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<td>1.270·10$^{10}$</td>
<td>9.002·10$^{-14}$</td>
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</table>

**Table 7.10:** Integral flux upper limits ($\Phi^{U.L.}$) from Segue 1 observation for a general power law spectrum with index -1.8 as function of different energy thresholds between $10^2$ GeV to $10^3$ GeV.

#### Integral Flux Upper Limits, Spectral index $\Gamma = -2$

<table>
<thead>
<tr>
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<th>$N_{on}$</th>
<th>$N_{off}$</th>
<th>$N^{*}$</th>
<th>$\sigma_{LiMa}$ [cm$^2$]</th>
<th>Aeff [cm$^2$]</th>
<th>$\Phi^{U.L.}$ [TeV$^{-1}$ cm$^{-2}$ s$^{-1}$]</th>
<th>$\Phi^{U.L.}$ [cm$^{-2}$ s$^{-1}$]</th>
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<tbody>
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<td>$&gt;10^2$</td>
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<td>53301</td>
<td>453.4</td>
<td>-0.99</td>
<td>1.703·10$^{10}$</td>
<td>1.004·10$^{-12}$</td>
<td>1.837·10$^{-11}$</td>
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<td>-2.04</td>
<td>1.683·10$^{10}$</td>
<td>3.528·10$^{-13}$</td>
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<td>6374</td>
<td>92.9</td>
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<td>1.722·10$^{-13}$</td>
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<td>-0.97</td>
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<td>0.57</td>
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<td>0.79</td>
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<td>580</td>
<td>139.7</td>
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<td>123.8</td>
<td>0.84</td>
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<td>1.278·10$^{-13}$</td>
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<td>1.34</td>
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<td>1.372·10$^{-13}$</td>
<td>2.504·10$^{-12}$</td>
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<tr>
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<td>373</td>
<td>134.5</td>
<td>1.36</td>
<td>1.270·10$^{10}$</td>
<td>1.140·10$^{-13}$</td>
<td>2.213·10$^{-12}$</td>
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**Table 7.11:** Integral flux upper limits ($\Phi^{U.L.}$) from Segue 1 observation for a general power law spectrum with index -2 as function of different energy thresholds between $10^2$ GeV to $10^3$ GeV.
### Integral Flux Upper Limits, Spectral index $\Gamma = -2.2$

<table>
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<th>Energy [GeV]</th>
<th>$N_{on}$</th>
<th>$N_{off}$</th>
<th>$N^{U.L.}$</th>
<th>$\sigma_{LiMa}$</th>
<th>$A_{eff}$</th>
<th>$\phi^{U.L.}$</th>
<th>$\Phi^{U.L.}$</th>
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<td>19233</td>
<td>174.1</td>
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<td>7.379 $\cdot 10^{-12}$</td>
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<tr>
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<td>6374</td>
<td>92.9</td>
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<td>5.259 $\cdot 10^{-13}$</td>
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<td>123.8</td>
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<td>1.348 $\cdot 10^{10}$</td>
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<td>134.5</td>
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<td>1.468 $\cdot 10^{-13}$</td>
<td>2.392 $\cdot 10^{-12}$</td>
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</table>

**Table 7.12:** Integral flux upper limits ($\Phi^{U.L.}$) from Segue 1 observation for a general power law spectrum with index -2.2 as function of different energy thresholds between $10^2$ GeV to $10^3$ GeV.

### Integral Flux Upper Limits, Spectral index $\Gamma = -2.4$

<table>
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<tr>
<th>Energy [GeV]</th>
<th>$N_{on}$</th>
<th>$N_{off}$</th>
<th>$N^{U.L.}$</th>
<th>$\sigma_{LiMa}$</th>
<th>$A_{eff}$</th>
<th>$\phi^{U.L.}$</th>
<th>$\Phi^{U.L.}$</th>
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<td>-0.99</td>
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<td>19233</td>
<td>174.1</td>
<td>-2.04</td>
<td>1.682 $\cdot 10^{10}$</td>
<td>1.081 $\cdot 10^{-12}$</td>
<td>8.693 $\cdot 10^{-12}$</td>
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<td>$&gt;10^{2.2}$</td>
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<td>6374</td>
<td>92.9</td>
<td>-2.25</td>
<td>1.658 $\cdot 10^{10}$</td>
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<td>110.2</td>
<td>-0.97</td>
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<td>5.005 $\cdot 10^{-13}$</td>
<td>3.796 $\cdot 10^{-12}$</td>
</tr>
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<td>$&gt;10^{2.4}$</td>
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<td>1654</td>
<td>194.2</td>
<td>0.57</td>
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<td>249.7</td>
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</tr>
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<td>139.7</td>
<td>0.96</td>
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<td>3.124 $\cdot 10^{-13}$</td>
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</tr>
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<td>$&gt;10^{2.8}$</td>
<td>536</td>
<td>509</td>
<td>123.8</td>
<td>0.84</td>
<td>1.408 $\cdot 10^{10}$</td>
<td>2.401 $\cdot 10^{-13}$</td>
<td>2.621 $\cdot 10^{-12}$</td>
</tr>
<tr>
<td>$&gt;10^{2.9}$</td>
<td>486</td>
<td>445</td>
<td>145.7</td>
<td>1.34</td>
<td>1.348 $\cdot 10^{10}$</td>
<td>2.435 $\cdot 10^{-13}$</td>
<td>2.953 $\cdot 10^{-12}$</td>
</tr>
<tr>
<td>$&gt;10^3$</td>
<td>411</td>
<td>373</td>
<td>134.5</td>
<td>1.36</td>
<td>1.269 $\cdot 10^{10}$</td>
<td>1.901 $\cdot 10^{-13}$</td>
<td>2.578 $\cdot 10^{-12}$</td>
</tr>
</tbody>
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**Table 7.13:** Integral flux upper limits ($\Phi^{U.L.}$) from Segue 1 observation for a general power law spectrum with index -2.4 as function of different energy thresholds between $10^2$ GeV to $10^3$ GeV.
7.5 Discussion

This chapter has been dedicated to the analysis of the promising source Segue 1 observed by the MAGIC-I telescope in view of indirect DM searches in γ-rays. The presence of the star η-Leonis made the analysis particularly delicate for the lowest energy range (100 to 200 GeV). Nevertheless, suitable cuts (star-cuts) have been conceived and it has been demonstrated, with the aid of a Crab Nebula sample observed in the same period of the Segue 1 survey, that reliable results could be achieved also below 200 GeV, where the star related effects were problematic.

The derived upper limits for Segue 1 reported in section 7.4 are of the same order of those achieved from the observation by MAGIC-I of the dwarf satellite galaxies Draco [245] and Willman 1 [246]. In those papers it was found that the achieved flux upper limits required flux boost factors of three orders of magnitude even in the most optimistic scenarios to match the predictions of typical mSUGRA benchmark models [262]:

- **Draco observation**

  Draco is a dSph galaxy accompanying the Milky Way at a galactocentric distance of about 82 kpc. From a kinematical analysis of a sample of 200 stellar line–of–sight velocities it was possible to infer the DM profile: the result of the fit, assuming a Navarro-Frenk-White (NFW) smooth profile [233], indicates a virial mass of the order of $10^9 \, M_\odot$, with a corresponding mass–to–light ratio of $M/L \sim 200 \, M_\odot/L_\odot$. With 7.8 hours of observation performed by MAGIC-I in 2007 a 2σ upper flux limit on steady emission of $1.1 \times 10^{-11}$ photons cm$^{-2}$ s$^{-1}$ was found, under the assumption of a generic annihilation spectrum without cutoff and a spectral index of -1.5 for photon energies above 140 GeV. For different mSUGRA model parameters using the benchmark points defined by Battaglia et al. [263] and for other models, the γ-ray spectrum expected from neutralino annihilations was computed. Assuming these underlying spectra and a smooth DM density profile as suggested in [264], the upper limits on the integrated flux above 140 GeV were calculated and compared to the experimental ones. As one can see from Fig. 7.24, the resulting upper limits are at least three orders of magnitude higher than expected by mSUGRA without enhancements.

- **Willman 1 observation**

  Willman 1 dSph galaxy is located at a distance of 38 kpc in the Ursa Major constellation. It represents one of the least massive satellite galaxies known to date ($M \sim 5 \times 10^5 \, M_\odot$) with a very high mass–to–light ratio, $M/L \sim 500 – 700 \, M_\odot/L_\odot$, making it one of the most DM dominated objects in the Universe. Following the work of Strigari et al. [265], its DM halo was parametrized with a NFW profile.

  The observation by MAGIC-I of Willman 1 took place in 2008 for a total amount of 15.5 hours. No significant γ-ray emission was found above 100 GeV, corresponding to 2σ upper flux limits on steady emission of the order of $10^{-11}$ photons cm$^{-2}$ s$^{-1}$, taking into account a subset of four slightly modified Battaglia mSUGRA benchmark models, as defined by Bringmann et al. [266] (see table 7.14). These benchmark models represent each a different interesting region of the mSUGRA parameter space, namely the bulk ($I'$), the coannihilation ($J'$), the funnel ($K'$) and the focus ($F^*$) point regions, and they include for the first time the contribution of IB process in the computation of the cross sections and spectra.

  A comparison with the measured flux upper limit and the fluxes predicted assuming the underlying mSUGRA benchmark spectra and the chosen Willman 1 density profile was computed. The results are summarized in table 7.15. Although the boost factor
upper limits seem to show that a DM detection could still be far (the most promising scenario, $K'$, being three orders of magnitude below the sensitivity of the telescope), it is important to keep in mind the large uncertainties in the DM profile and particle physics modeling that may play a crucial role in detectability. In particular the possible presence of substructures in the dwarf, which is theoretically well motivated, may increase the astrophysical factor and therefore the flux of more than one order of magnitude. Furthermore, since the parameter space was not fully scanned, it is likely that there are models of neutralino with higher $\Phi^{PP}$.

<table>
<thead>
<tr>
<th>BM</th>
<th>$m_{1/2}$ [GeV]</th>
<th>$m_0$ [GeV]</th>
<th>$\tan\beta$</th>
<th>$A_0$</th>
<th>$\text{sign}(\mu)$</th>
<th>$m_\chi$ [GeV]</th>
<th>$\langle\sigma v_{\chi\chi}\rangle$ [cm$^3$/s]</th>
<th>$\Phi^{PP}(&gt;100)$ [cm$^2$ GeV$^{-2}$ s$^{-1}$]</th>
</tr>
</thead>
<tbody>
<tr>
<td>$I'$</td>
<td>350</td>
<td>181</td>
<td>35</td>
<td>0</td>
<td>+</td>
<td>141</td>
<td>$3.62 \times 10^{-27}$</td>
<td>$7.55 \times 10^{-34}$</td>
</tr>
<tr>
<td>$J'$</td>
<td>750</td>
<td>299</td>
<td>35</td>
<td>0</td>
<td>+</td>
<td>316</td>
<td>$3.19 \times 10^{-28}$</td>
<td>$1.23 \times 10^{-34}$</td>
</tr>
<tr>
<td>$K'$</td>
<td>1300</td>
<td>1001</td>
<td>46</td>
<td>0</td>
<td>−</td>
<td>565</td>
<td>$2.59 \times 10^{-26}$</td>
<td>$6.33 \times 10^{-33}$</td>
</tr>
<tr>
<td>$F^*$</td>
<td>7792</td>
<td>22100</td>
<td>24.1</td>
<td>17.7</td>
<td>+</td>
<td>1926</td>
<td>$2.57 \times 10^{-27}$</td>
<td>$5.98 \times 10^{-34}$</td>
</tr>
</tbody>
</table>

Table 7.14: Willman 1 observation: definition of benchmark models as in Bringmann et al. [266] and computation of the particle physics factor $\Phi^{PP}$ above 100 GeV.

The possible constraints to the mSUGRA parameter space which can be derived from the upper limits calculated from the MAGIC-I observation of Segue 1 should not drastically differ from what found for the cases of Draco and Willman 1 observations. Nevertheless, it is worth noting that the astrophysical factor for Segue 1 should be larger than those derived for Draco and Willman 1, making the possible constraints more stringent.

A paper dedicated to the MAGIC-I observation of Segue 1, based on the analysis reported in this chapter, is in preparation. In the paper it is foreseen to derived possible constraints...
Table 7.15: Willman 1 observation: comparison of estimated integral flux above 100 GeV for the chosen benchmark models and the upper limit in the integral flux $\Phi^{U.L.}$ above 100 GeV coming from MAGIC-I data in units of photons cm$^{-2}$ s$^{-1}$. On the rightmost column, the corresponding upper limit on the boost factor $B^{U.L.}$ required to match the two fluxes is calculated.

to the recent DM interpretations (see e.g. [193,194]) of the electron and positron fluxes from PAMELA [160], Fermi–LAT [162], and H.E.S.S. [163].

<table>
<thead>
<tr>
<th>BM</th>
<th>$\Phi^{model}$</th>
<th>$\Phi^{U.L.}$</th>
<th>$B^{U.L.}$</th>
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<tr>
<td>$I'$</td>
<td>2.64$\times 10^{-16}$</td>
<td>9.87$\times 10^{-12}$</td>
<td>3.7$\times 10^4$</td>
</tr>
<tr>
<td>$J'$</td>
<td>4.29$\times 10^{-17}$</td>
<td>5.69$\times 10^{-12}$</td>
<td>1.3$\times 10^5$</td>
</tr>
<tr>
<td>$K'$</td>
<td>2.32$\times 10^{-15}$</td>
<td>6.83$\times 10^{-12}$</td>
<td>2.9$\times 10^3$</td>
</tr>
<tr>
<td>$F^*$</td>
<td>2.09$\times 10^{-16}$</td>
<td>7.13$\times 10^{-12}$</td>
<td>3.4$\times 10^4$</td>
</tr>
</tbody>
</table>
Conclusions

The work presented in this thesis was carried out in the ambit of the MAGIC experiment, an array of two telescopes based on the Imaging Atmospheric Cherenkov technique, which can detect gamma-rays in the energy range between \( \sim 50 \) GeV and \( \sim 10 \) TeV. The main topics which were reported dealt with the implementation in the software of the experiment of the Azimuth dependence in the Effective Collection Area calculation finalized to the stereoscopic observations, and the analysis of the data taken with the MAGIC-I telescope of the source Segue 1, so far considered to be a dwarf spheroidal galaxy satellite of the Milky Way, characterized by a huge mass–to–light ratio of the order of \( 10^3 \) M\(_\odot\)/L\(_\odot\) and thus very interesting for indirect Dark Matter (DM) searches.

The main features of the two telescopes were introduced. The new telescope MAGIC-II is structurally a clone of the first telescope, although many hardware improvements have been implemented, especially for the PMT camera and the readout system. From the data analysis point of view, the software of the experiment was upgraded in order to allow the analysis of the stereoscopic observations of the gamma-ray sources. The current performance of the stereoscopic system was reported: to date, the MAGIC telescopes are the more sensitive world-wide ground-based detector for gamma-rays in the energy range between \( \sim 50 \) and \( \sim 150 \) GeV. This is due to the enhancement of the sensitivity (by a factor \( \simeq 1.5-2 \), and even larger for energies below \( \sim 150 \) GeV), of the angular and energy resolution, of the suppression of the hadronic background and to the lowering of the analysis energy threshold to \( \sim 50 \) GeV, when the two telescopes are operated in stereoscopic mode. These features open great potentialities for new discoveries and for a better understanding of the gamma-ray emission of several astrophysical objects in synergy with Fermi–LAT, a new generation gamma-ray detector on-board of the Fermi satellite which detects gamma-rays between 20 MeV and 300 GeV. Nowadays, these two instruments can be considered as complementary experiments and can give precise measurements of many gamma-ray emitters within a wide range of energies between tens of MeV and several TeV.

A major effort is ongoing inside the MAGIC Collaboration to further improve the performance of the stereoscopic system. The upgrade of the MAGIC-I telescope (foreseen for summer 2011) will give a decisive contribution in this sense.
Concerning the upgrade of the software of the experiment, I actively contributed to the introduction of the Azimuth dependence in the Effective Collection Area calculation in view of the stereoscopic observations. In fact, the Geomagnetic field effects on the development of the showers and the fixed configuration of the MAGIC telescopes’ array combine in a nontrivial way. Therefore the Azimuth dependence of the stereoscopic observation was systematically taken into account. This dependence is expected to be particularly relevant for observations at middle and high Zenith angles and for the lowest energy range. This work required a deep modification and updating of some basic classes and executables of the software of the experiment. It was shown, thanks to the aid of stereo MC gamma data, that the main contribution to the Azimuth dependence of the Effective Collection Area is given by the Geomagnetic field effect for Zenith angles below 45°, particularly at the energies close to the analysis threshold and for increasing Zenith angle ranges. On the other hand, it was not possible to clearly identify the telescopes’ array configuration effect: this effect could play an important role for stereoscopic observations carried out at Zenith angles above 45°. More studies are needed in order to verify this hypothesis and they will be done when stereo MC gamma and stereoscopic observations of the Crab Nebula at higher Zenith angles will be available. In any case, the implemented changes for the calculation of the Effective Collection Area allow to take into account the global resulting effect, obtained by the combination of the two contributions (i.e., the Geomagnetic field effect and the telescopes’ configuration effect), which would be otherwise difficult to disentangle. Finally, thanks to some tests performed on a sample of Crab Nebula data taken with MAGIC-I, it was shown that the implementation of the Azimuth dependence of the Effective Collection Area was indeed successfully accomplished and it is right now ready to be used in the stereo analysis chain. So far, the Azimuth parameter (unlike the Zenith one) is not used in the gamma-hadron separation and energy reconstruction algorithms. Further studies are foreseen in order to verify if the Azimuth parameter introduced in those algorithms can give some improvements in terms of sensitivity and smaller systematic errors.

The source Segue 1 is so far considered one of the most DM dominated dwarf spheroidal galaxies satellite of the Milky Way and thus represents one of the best target for indirect DM searches. Regarding the MAGIC-I observation of the Segue 1 source, its analysis was reported in detail. The data were affected by the light of the star η-Leonis, a white supergiant with apparent magnitude 3.5 and located at ~0.7° far away from the sky position of Segue 1. In fact, during the whole Segue 1 survey, the light of this star was always present in the inner part of the MAGIC-I camera (which has a radius of ~1°), where the trigger system is operating. The main effect was a local inefficiency due to the dynamic enhancement by the Individual Pixel Rate control of the discrimination thresholds of the pixels enlightened by the star. This local inefficiency resulted in a mismatch of the image parameter distributions for energies below 200 GeV which prevented the extraction of reliable results below that energy. Nevertheless, since the low energy ranges are particularly interesting for indirect DM search purposes, suitable cuts were conceived and trustworthy results were achieved also for energies between 100 and 200 GeV, thanks to the consideration of the geometrical properties of the star related problem. The techniques used to obtain such results were tested on a sample of Crab Nebula data observed in the same period of the Segue 1 survey. No significant gamma-ray emission from Segue 1 was found above an energy threshold of 100 GeV and upper limits on the flux emission derived from different assumed power law spectra were calculated. The achieved integral upper limits were of the same order (~10^{-11} - ~10^{-12} [ph cm^{-2} s^{-1}], depending on the assumed power law spectrum and energy threshold) of those obtained from the observations carried out by MAGIC-I of the dwarf satellite galaxies Draco and Willman 1 which required
flux boost factors of three orders of magnitude even in the most optimistic scenarios to match the predictions of typical mSUGRA benchmark models. A paper dedicated to the MAGIC-I observation of Segue 1, based on the results reported in this thesis, is in preparation. In the paper, besides other phenomenological considerations, it is also foreseen to derived possible constraints to the recent DM interpretations of the electron and positron fluxes measured by PAMELA, Fermi–LAT, and H.E.S.S. experiments.
Bibliography


[27] http://veritas.sao.arizona.edu/.


[34] http://top.gae.ucm.es/.


[58] http://swift.gsfc.nasa.gov/docs/swift/swiftdc.html.


(available at http://wwwmagic.mppmu.mpg.de/publications/theses/).


[121] O. Blanch and A. Moralejo. How to use the Camera simulation program 0.7. MAGIC-TDAS internal note 04-07, 2004.


(available at http://wwwmagic.mppmu.mpg.de/publications/theses/).

(available at http://wwwmagic.mppmu.mpg.de/publications/theses/).


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