UNIVERSITY OF UDINE - ITALY

Department of Physics

Ph.D. Thesis

OBSERVATIONS OF GAMMA-RAY BURST AFTERGLOWS WITH THE MAGIC TELESCOPE

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Abstract

Very High Energy (VHE) photons are a powerful probe of fundamental physics under extreme conditions. During recent years, the observation of gamma-rays (γ-rays) is throwing light on the high energy Universe thanks to satellite-based detectors and ground-based telescopes. Ground-based γ-ray astronomy is part of a new field of fundamental research of Astroparticle Physics and recently made spectacular discoveries, studying sources that, already known in several other energy bands (i.e. γ-ray, X-ray, radio and/or optical frequencies), are expected to be VHE γ-rays emitters. Possible VHE emission detection from a particular class of γ-ray emitters, the Gamma-Ray Bursts (GRB) is an actual astrophysical challenge.

In this work I first introduce γ-rays and γ-astrophysics. In particular I describe the emerging VHE γ-ray sky reporting a description of the candidate γ emitters. Then I present the Imaging Atmospheric Cherenkov Technique (IACT) and the MAGIC Telescope, located at the Observatory of Roque de Los Muchachos, in the Canary Island of La Palma. MAGIC scientific activity is based on the detection of the Cherenkov photons produced in the Earth’s atmosphere by electromagnetic showers initiated by VHE cosmic γ-rays. The explored energy window is between 30 GeV and 10 TeV.

Later I review the current knowledge about GRBs, bright flashes of radiation with spectral energy distribution in the γ-ray band. The “prompt” emission lasts seconds and is followed by a fainter longer-lived “afterglow” (in X, radio, optical bands). I speak about GRB phenomenology and the mainstream theoretical framework in which GRB emission is understood as an expanding relativistic fireball. A special focus is given on the predictions of possible γ-ray emission at VHE. In the last part of this work I describe the GRB observation strategy with MAGIC, that is the most suitable current IACT instrument to perform observation of the prompt and afterglow GRB’s emission above 30 GeV. The MAGIC-I observation of GRB 080430 is reported and the results are used to evaluate the perspective for afterglow ob-
observation with ground based telescopes. Finally with a selected MAGIC-I GRBs data sample I try to infer general considerations on MAGIC GRB afterglow data. No signal was found in this analysis, but the obtained Upper Limits on flux, strongly emphasize the potential of continued GRB observation with MAGIC. If we observe a nearby GRB, we can constrain prompt/afterglow emission models. Performing a similar time-dependent stacked analysis, possible VHE features may appear in the light-curve, that may still help put constraints on emission models.

This work is organized as follows:

Chapter 1: A general introduction to $\gamma$-rays and $\gamma$-Ray astronomy.
Chapter 2: A description of IACT and of the MAGIC Telescope, in particular.
Chapter 3: A discussion of GRBs and their possible VHE emission.
Chapter 4: The MAGIC data analysis chain.
Chapter 5: GRB observation strategy with MAGIC.
Chapter 6: GRB 080430 afterglow observation, and related theoretical discussion.
Chapter 7: Analysis of MAGIC-I’s GRB database and emerging discussion.

I participate as analyzer in the article reported in the appendix, on the blazar 3c 454.3 and on the article on GRB 084030 (in chapter 6). A research note about the analysis presented in last chapter is in preparation. During my Ph.D I spent in total four data acquisition shifts at the Observatory. In two of them I was the shift leader and in one I worked for the commissioning of the MAGIC-II telescope.
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Gamma Rays: production, propagation, detection

Stable neutral particles, not deflected by electromagnetic fields (both galactic and intergalactic), point directly back to their original source. It’s easy to understand why they can be used as messengers across the Universe. Gamma-Ray Astronomy\(^1\) studies the radiation emitted by the highest energy sources in the Universe in order to refine theoretical models of high-energy particle emission and the nature itself of some elusive astrophysical sources. When we are dealing with $\gamma$-rays we are discussing about photons, so electromagnetic emission. On the other hand the Earth is constantly hit by a high flux of particles, collectively defined Cosmic Rays (CRs). CRs are much more frequent than $\gamma$-rays, which represent only a tiny fraction of the incoming flux: very roughly only 0.1% of CRs are $\gamma$-rays. Charged CRs are deflected by cosmic magnetic fields and reach the Earth almost isotropically, losing any information on their original arrival direction. In this context it’s easy to understand why $\gamma$-rays, that are always emitted where high-energy CRs are accelerated, are unique probes to study CRs accelerators.

\(^1\)Photons with energy bigger than 1 MeV are, for historical reasons, called gamma-rays ($\gamma$-rays and as said before, they are only a small fraction of the cosmic rays. In this work we classify gamma-rays in the following way:

- low–medium energy: 0.5 MeV–100 MeV
- High–Energy (HE): 100 MeV–30 GeV
- Very–High–Energy (VHE): 30 GeV–30 TeV
- Ultra–High–Energy (UHE): 30 TeV–30 PeV
- Extremely–High–Energy (EHE): 30 PeV—...
The aim of this chapter is to define what CRs and γ-rays are, how they can be produced, how they travel to the earth and how they can be detected. A description of the candidate emitters is reported in the last part of the chapter.

1.1 Cosmic rays

About one century ago, two pioneering works by the Austrian physicist Victor Hess and the Italian Domenico Pacini demonstrated, with different techniques, that ionizing radiation constantly reaches the Earth’s atmosphere [176]. The so-called CRs are considered messengers from the non-thermal part of the universe, since their energies far exceed the temperatures ordinary encountered in astrophysical objects. The origin, the composition and the propagation of the cosmic rays is still unclear and under debate.

The measured CR energy spectrum (Fig. 1.1) shows an energy range extending over more than 13 orders of magnitude. At 1 GeV, the intensity per unit solid angle per GeV in roughly 1000 particles per second per 1 m$^2$, but at $E > 6 \times 10^{19}$ eV, the probability (per unit solid angle and per EeV=10$^{18}$ eV) that a particle hits an area of 1000 km$^2$ is only about one per century!

Up to the energies $E \geq 1$ GeV the spectrum is unaffected by the propagation of solar wind particles: the differential flux $dN/dE$ is $dN/dE \propto E^{-\alpha}$, first with $\alpha \approx 2.7$ and then, for higher energies, with $\alpha \approx 3$ [149]. The power-law spectral shape suggest as possible processes responsible for the acceleration:

- shocks in planetary magnetospheres, interplanetary space, and the sun are believed to accelerate the low-energy CR

- shocks in expanding galactic supernova remnants (SNR) are thought to be responsible for higher energies ones, up to the $10^{15}$ eV

- shocks near active galactic nuclei (AGNs) may be responsible for the highest energy CRs.

The CR energy spectrum is not completely featureless and, by analyzing its dips and kinks, we can learn a great deal about the sources and the path these particles take on their trajectory to earth. Near $10^{15}$ eV, a “knee” occurs in the spectrum, presumably because this is
1.1. Cosmic rays

Figure 1.1: Differential energy spectrum of cosmic rays.

the highest energy that particles can reach in Galactic SNR. In the region near $10^{18} - 10^{20}$ eV, an additional structure in the spectrum, the “ankle”, has long been thought to be due to the propagation of the highest energy particles through intergalactic space. Up to the knee, the chemical composition is known: About 79% of the primary nucleons consist of protons, 70% of the rest consists of helium nuclei and only a small fraction of heavier nuclei [100]. Less than 1% of the cosmic rays consists of electrons-positrons, the spectrum of which is steeper than the one of protons and nuclei.

Physicists have struggled for decades to determine whether such ultrahigh-energy CR exist, where do they come from, what do they consist of, and how are they accelerated to energies
100 million times greater than (man-made) particle accelerators have reached.

In the past several years, answers started to appear, mainly from the Auger and HiRes experiments (see section 1.2 for the techniques description). In late 2007 The Pierre Auger Observatory collaboration provided evidence that the arrival directions of cosmic rays above $5.7 \times 10^{19}$ eV are correlated with the directions of nearby AGNs [15]. This fact is of utmost importance because it is the first experimental evidence that acceleration of ultrahigh-energy CRs (UHECR) is not isotropic but maybe associated to some emitters and that the active galactic nuclei are effective accelerators of particles.

In 2008, the two separate collaborations (HiRes and Auger), have published convincing evidence for the suppression of the UHECR due to their photopion interactions with the cosmic microwave background [7, 16]. This effect is called GZK after Greisen (1966) and Zatsepin Kuzmin (1966) predicted the effect [124, 245]. Although the HiRes and Auger observations of the GZK cutoff are extremely convincing, we cannot consider the issue of the origin and propagation of the UHECRs completely solved until we understand the following issues:

- the high energy interaction models
- the effects of heavier nuclear component in the CR beam
- the input to propagation models
- the acceleration mechanisms
- the nature of the sources

Because of interaction of the CRs with the interstellar medium, it is believed that the average injection spectrum is harder ($\gamma \approx 2.1$) than the locally measured CR spectrum [222].

The measured differential CR spectrum is power-law. When we deal with power-law spectra, non-thermal processes are likely involved: if the particle density is low, interactions are rare and may not lead to thermal equilibrium. Fermi-acceleration [92], can explain a power law spectrum. The popular model for particle shock acceleration is the diffuse shock acceleration model. Particles are accelerated when they repeatedly cross shock fronts. Magnetic field irregularities keep scattering the particles back and forth and the latter keep crossing the same shock. In the 1st-order Fermi acceleration the energy gain per acceleration cycle is proportional to the velocity $\beta$ of the shock, which carries the magnetic irregularity, and in
1.1. Cosmic rays

Figure 1.2: A sketch of the Fermi mechanism.

the 2nd-order it is proportional to $\beta^2$ [191].

1.1.1 Gamma–ray emission processes

The main mechanisms for gamma-rays emission from a power-law distribution of accelerated particles are the following:

**Bremsstrahlung:** A charged particle accelerated in an electric field emits Bremsstrahlung radiation. If we are dealing with a relativistic electron is in fact possible an interaction
between electron and Coulomb field without the capture of the electron. The radiation is emitted within a cone with opening angle of $\theta \simeq m_e c^2 / E_e = 1/\gamma$. The bremsstrahlung spectrum is flat up to roughly the electron kinetic energy $E_\gamma = (\gamma_e - 1) m_e c^2$ [82]. In case of a power law distribution energy spectrum of the hitting electrons, the bremsstrahlung spectrum follows the spectral index of the radiating particles’ spectrum.

**Synchrotron Radiation:** In the presence of magnetic fields, charged particles emit synchrotron radiation. Synchrotron radiation of accelerated electrons is one of the most important processes in the non-thermal Universe. If $E_e$ is the particle (electron) energy, the radiation is emitted mainly into an angle of $\theta \approx m_e c^2 / E_e \simeq 1/\gamma$ to the particle’s velocity vector. In the relativistic case electrons the energy of the emitted photons has a peak, whose position is proportional to the transverse component of the magnetic field $B$ and to the Lorentz factor of the electron:

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

(1.1)

It’s clear how, to push photon energies into the GeV-range, magnetic fields $B \geq 10^6$ G are needed. If the power-law distributions of particle energies is characterized by a spectral index $p > 0$, the spectral index of the corresponding synchrotron radiation is $(p - 1)/2$ [208]. The synchrotron radiation itself modifies the electron energy spectrum. The radiation “cools” the electrons with higher energy and produces a cooling break in the power-law distribution. The cooled portion of the spectrum is steeper by a half power, $\propto E^{-p/2}$

**Inverse Compton Scattering:** In the Compton scattering a “hot” photon gives some of its energy to the “cold” electron. In the Inverse Compton (IC) effect, the “cold” photon acquires some energy from the “hot” electron. $\gamma$-ray production can occur when relativistic electrons upscatter photons from interstellar optical, infrared or microwave radiation fields. The emission spectrum depends in particular on the incoming radiation energy and on the electron’s velocity distribution: the average energy of inverse Compton scattered photons, $E_\gamma$, can be calculated typically from the average energy of the scattering electrons $E_e$ [82]

$$E_\gamma \simeq 6.5 \cdot 10^3 \cdot \left( \frac{E_e}{\text{TeV}} \right)^2 \cdot \left( \frac{E_{\text{ph}}}{\text{eV}} \right) \text{ [GeV]},$$

(1.2)

for an ambient photon-field with typical energies of $E_{\text{ph}}$ such that $E_{\text{ph}} E_e \ll (m_e c^2)^2$ (Thomson-regime). As with this mechanism the original photon energy get multiplied by the square of treatment.
1.1. Cosmic rays

the Lorentz factor of the electron, photons can be upscattered to very high energies. For a power–law distribution of electrons with spectral index $p$, the resulting differential IC $\gamma$-ray spectrum is $\propto E^{(1-p)/2}$ for the non-relativistic regime ($E_{ph}E_e \ll (m_e c^2)^2$) [138] and $\propto E^{p+1-q}$ in the ultra-relativistic Klein-Nishina-regime, $E_{ph}E_e \gg (m_e c^2)^2$, where $q$ is the spectral index of the soft photons before being scattered [233].

As in the synchrotron radiation case, IC scattering of a population of “hot” electrons of a field of “cold” photons can cool the electrons and produce cooling-breaks in the $\gamma$-ray spectrum. IC scattering is important in regions of high photon densities. In the Thomson regime, the emitted photons follow the spectral shape of the seed photons. In the Klein-Nishina regime, the resulting spectrum has a sharp cut-off, which is determined by the maximum energy of the scattering electrons.

**Synchrotron Self–Compton (SSC) mechanism:** One basic leptonic scheme to explain HE and VHE $\gamma$-rays from astrophysical sources is the so called Synchrotron Self–Compton (SSC) Model. Ultra-relativistic electrons, accelerated in magnetic fields, emit synchrotron radiation in the form of a spectrum peaked in the infrared/X–ray range. Such photons interact with their own parent electron population, via Compton scattering. The Compton component can peak in GeV-TeV energy range and the two characteristic synchrotron and Compton peaks dominate the spectral energy distributions of typical non–thermal sources like blazars (Fig. 1.3)

**Neutral Pion Decay:** $\gamma$-rays can also be emitted by hadronic processes. Energetic collisions of protons $p$ with interstellar gas particles ($p_i$) or a radiation field ($\gamma_i$) will produce neutral pions:

$$p + p_i \rightarrow \pi^0 + k \cdot \pi^\pm + X \quad (1.3)$$
$$p + \gamma_i \rightarrow \Delta^+ \rightarrow p + \pi^0 \quad (1.4)$$

The $\pi^0$ ($\tau_{\pi^0} = 8.4 \cdot 10^{-17}$ s) decays almost immediately into two $\gamma$-rays with an energy distribution peaking at a broad maximum at about 70 MeV in the $\pi^0$ rest frame [150]. In the case of beamed annihilations, the resulting gamma ray energy will be Lorentz-boosted to higher energies. The $\pi^0$ decay is also the main responsible for the formation of secondary electromagnetic showers inside the atmospheric hadronic showers.
1.1.2 Propagation and absorption of $\gamma$-rays:

In astrophysics, we can distinguish two kinds of pair-production processes:

- Classical pair-production: a high energy $\gamma$-ray interacts with the electric field of an atomic nucleus or of an electron and produce the pair in a process $\gamma(\gamma) \rightarrow e^-e^+$ where the second photon is the virtual photon of the electron field.

- Photon-photon pair-production: a similar process, where the soft photons are available in the environment as stellar radiation from star, for example: $\gamma\gamma$(background) $\rightarrow e^-e^+$

The first process prevent high-energy $\gamma$-rays to cross the earth’s atmosphere and protects the life on the earth. The second one is responsible for the opacity of the Universe to $\gamma$-rays, due to the absorption of the latter in their travel through the intergalactic medium.

**Interaction with EBL and the $\gamma$-ray horizon** The Universe, excluding the galactic plane, is isotropically filled with soft photons at the so called Extragalactic Background Light (EBL) [119]. The EBL is composed by a contribution from stellar light summed up at different epochs plus a contribution coming from the dust-scattered light. The contributions are respectively at mid-IR and in the close UV wavelengths.

VHE photons have a large cross-section for pair-production with the EBL and for this reason the Universe is not completely transparent to VHE $\gamma$-rays. It’s so clear why far distant sources
cannot be observed, and a cosmological gamma-horizon can be defined. Along its travel from a source at redshift $z$ to the earth the probability for a photon to survive the absorption can be expressed in the form: $e^{-\tau(E,z)}$ where the coefficient $\tau(E,z)$ is called optical depth. The gamma-horizon is defined as the distance corresponding to the redshift $z$ for which $\tau(E,z) = 1$, that give an attenuation by a factor $1/e$.

The attenuation of the measured spectra in the TeV region can thus be used to derive constraints on the EBL density [219]. First limits on the EBL were obtained in [220], while recent determinations from the detection of distant VHE sources are reported in [19,157]. Present data on absorption seem not clearly compatible in the pure QED and other kinds of interaction can be speculated. Photons might interact with (possibly quintessential) very

![Gamma-ray horizon](image.png)

Figure 1.4: $\gamma$-ray horizon
light axion-like particles, which might change the absorption length [72–74]. In particular, in the DARMA model [74], such contribution might enhance the photon flux via a regeneration mechanism. Such an interaction would be mediated by the (intergalactic) magnetic fields. Mechanisms in which the absorption is changed through violation of the Lorentz invariance are also to be proved: such models are particularly fascinating within quantum gravity (QG) scenarios [42].

The AGNs detected at VHE (see Table 1.1) are almost uniformly distributed in the sky and they are very far from us. The spectral index measured is plotted in figure 1.5 and - within the small redshift range probed - show a substantial lack of correlation between measured TeV slopes and redshift. It will be important to increase the statistic in order to increase our knowledge.

The transparency of the Universe at γ−rays seems to be larger than expected from popular EBL models: MAGIC detected the Quasar 3C279 at redshift 0.536 [35]. This might also indicate the existence of new propagation physics.

![Figure 1.5: Observed spectral indeces of known VHE blazar are shown as filled circles (with error bars) [72]. Superposed is the predicted behaviour of the observed spectral index within two different scenarios: classical (light grey) and DARMA (dark grey).](image-url)
Table 1.1: TeV blazar data [184]. Col.1: source name. 
Col.2: source redshift. 
Col.3: observed 0.2-2 TeV photon spectral index, and associated statistical uncertainty. The corresponding systematic uncertainties are typically $\sim 0.1$ for H.E.S.S. and $\sim 0.2$ for MAGIC. Col.4: observed $>0.2$ TeV flux (in erg cm$^{-2}$ s$^{-1}$), and associated statistical uncertainty (from observed spectral normalization only). 
Col.5: Cherenkov telescope (CT) or array with which the data in col.3 have been collected: symbols stand for H=H.E.S.S., M=MAGIC, W=Whipple.

<table>
<thead>
<tr>
<th>Source</th>
<th>z</th>
<th>$\alpha_\gamma$</th>
<th>$F_\gamma$</th>
<th>CT</th>
</tr>
</thead>
<tbody>
<tr>
<td>Mrk 421</td>
<td>0.031</td>
<td>2.33 ± 0.08</td>
<td>1.0(±0.1)E-10</td>
<td>M</td>
</tr>
<tr>
<td>Mrk 501</td>
<td>0.034</td>
<td>2.28 ± 0.05</td>
<td>1.7(±0.1)E-11</td>
<td>M</td>
</tr>
<tr>
<td></td>
<td></td>
<td>2.45 ± 0.07</td>
<td>3.8(±1.0)E-12</td>
<td>M</td>
</tr>
<tr>
<td>PKS 2344+514</td>
<td>0.044</td>
<td>2.95 ± 0.12</td>
<td>1.2(±0.1)E-11</td>
<td>M</td>
</tr>
<tr>
<td>Mrk 180</td>
<td>0.045</td>
<td>3.30 ± 0.70</td>
<td>8.5(±3.4)E-12</td>
<td>M</td>
</tr>
<tr>
<td>1ES 1959+650</td>
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<td>2.72 ± 0.14</td>
<td>3.0(±0.4)E-11</td>
<td>M</td>
</tr>
<tr>
<td>BL Lacertae</td>
<td>0.069</td>
<td>3.60 ± 0.50</td>
<td>3.3(±0.3)E-12</td>
<td>M</td>
</tr>
<tr>
<td>PKS 2005-489</td>
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<td>4.00 ± 0.40</td>
<td>3.3(±0.5)E-12</td>
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</tr>
<tr>
<td>PKS 2155-304</td>
<td>0.116</td>
<td>3.37 ± 0.07</td>
<td>2.9(±0.2)E-11</td>
<td>H</td>
</tr>
<tr>
<td>1ES 1426+428</td>
<td>0.129</td>
<td>3.55 ± 0.46</td>
<td>2.5(±0.4)E-11</td>
<td>W</td>
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<tr>
<td>1ES 0229+200</td>
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<td>2.50 ± 0.19</td>
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<tr>
<td>H 2356-309</td>
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<td>2.6(±0.7)E-12</td>
<td>H</td>
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<tr>
<td>1ES 1218+304</td>
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<td>3.00 ± 0.40</td>
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<td>M</td>
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<tr>
<td>1ES 1101-232</td>
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<td>2.94 ± 0.20</td>
<td>4.5(±0.7)E-12</td>
<td>H</td>
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<tr>
<td>1ES 0347-121</td>
<td>0.188</td>
<td>3.10 ± 0.23</td>
<td>3.9(±0.7)E-12</td>
<td>H</td>
</tr>
<tr>
<td>1ES 1011+496</td>
<td>0.212</td>
<td>4.00 ± 0.50</td>
<td>6.4(±0.3)E-12</td>
<td>M</td>
</tr>
<tr>
<td>PG 1553+113</td>
<td>&gt;0.25</td>
<td>4.20 ± 0.30</td>
<td>5.2(±0.9)E-12</td>
<td>M,H</td>
</tr>
<tr>
<td>3C 279</td>
<td>0.536</td>
<td>4.1 ± 0.7</td>
<td>2.9(±0.5 )E-11</td>
<td>M</td>
</tr>
</tbody>
</table>
1.2 Detection techniques

γ-ray astronomy is the latest opened window on the exploration of the electromagnetic emission of the Universe. However, the window is wide: the low-energy limits can be placed at the electron mass, where the nuclear and electron-positron annihilation lines start to be observable. The high-energy limit is based on the observation of CRs beyond $10^{20}$ eV, but potentially there can be no upper limit. I previously defined the ranges in which we can split the γ-ray window, that roughly correspond to different acceleration and emission mechanism and are observable with different detectors. In Fig. 1.6 the transparency of the atmosphere for different photon energies and possible detection techniques are depicted.

The Imaging Atmospheric Cherenkov Technique (IACT) will be discussed in more detail in the next chapter. I’ll also summarize the main IACT and future developments.

Figure 1.6: Transparency of the atmosphere for different photon energies and possible detection techniques.
1.2.1 Particle detector experiments

Fluorescence experiments: With this class of detector it is possible to observe cosmic rays by imaging the extensive air shower (EAS) generated by a primary cosmic ray. At their passage, charged particles can ionize and excite the atmosphere’s molecules (mainly nitrogen and oxygen), which in turn deexcite and emit fluorescence light. This light, mostly of UV and optical frequencies, can be produced over a track of many kilometers in length. By focussing the light with the use of several small reflectors operate simultaneously and placed at a certain distance one from another, and measuring the intensity and the time separation of the events measured of the various telescopes, one is able to reconstruct the energy of the primary particle CR energy and its direction. Usually they are located at very high altitudes where the absorption is lower.

The HiRes experiment was operated on clear, moonless nights over a period of nine years (1997-2006) and was composed of two groups of telescopes at 12.6 km distance located at on the U.S. Army Dugway Proving Ground in Utah. HiRes has made the first statistically significant observation of the GZK suppression. The break energy was measured at $5.6 \times 10^{19}$ eV, with a significance of 5.3$\sigma$. No spatial association with any targeted was found [7].
1. **Gamma Rays: production, propagation, detection**

1.1 **Direct particle experiments:** Particle detector experiments directly detect the charged component of the atmospheric shower by measuring the arrival time of the shower front at several water tanks distributed over a large area. The direction of the primary CR can be calculated to about one degree accuracy.

The foremost representative of this technique is the Auger experiment, in Malargue, Argentina (http://www.auger.org). Auger is an hybrid detector, that combines a particle detector array with fluorescence telescopes. A large array of 1600 surface detectors (SDs), covers an area of 3000 km$^2$ and detects the particles at ground level by means of the Cherenkov radiation they produce in the water. At each of four sites on the periphery of the instrumental area, six inward-facing optical telescopes observe the sky on clear moonless nights, measuring the fluorescence light produced as an extensive air shower passes through the field of view. In the Auger experiments the surface detector technique and the fluorescence detectors (FDs) are complementary: The SDs measure the two-dimensional lateral structure of the shower at ground level, whereas FDs record the longitudinal profile and its development through the atmosphere. In 2007, the Pierre Auger Observatory collaboration established the extragalactic origin of the highest energy CR, and in 2008, for the first time it established the reality of the GZK effect. In fact one of the advantage of the Auger approach compared to HiRes is that by viewing the same event using both fluorescence and water Cherenkov technique, the energies can be calibrate by using one technique versus the other.

1.2.2 **γ-ray experiments**

**HE domain. Satellite-borne experiments:** 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, because γ-ray at high energies have both small fluxes and cannot be contained in a too short detector. Satellite HE gamma telescopes detect the primary photons at energies lower than ground-based telescopes. However constrained by the small effective area (of the order of 1 m$^2$) they cannot reach an high sensitivity. On the other hand they have the advantage of a large duty cycle (they are not constrained by night observation like ground-based ones) and a low background rate. They obviously have a large cost. The detectors are
1.2. Detection techniques

characterized by very large field of view (Fermi cover the entire sky in 3 hours) but quite small angular resolution (depending on the energy, but of the order of 1 deg). Due to the rapidly falling flux of $\gamma$-rays with energy, they are only marginally sensitive above the 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 (in the case of Fermi).

A famous representative experiment based on this technique was the EGRET experiment, on board of the CGRO satellite, which operated in 1991-2000. EGRET detected 271 sources at energy above 100 MeV of which about 170 remained unidentified.

The Fermi Gamma-ray Space Telescope (formerly known as GLAST), launched on june 11, 2008, is providing new insight into HE $\gamma$-ray astronomy. Fermi is composed by two instruments, the Large Area Telescope (LAT) and the GLAST Burst Monitor (GBM). The LAT is a pair-production telescope composed of $4 \times 4$ grid of towers: each tower consist of a silicon-strip detector and a tungsten-foil tracker-converter, mated with a hodoscopic cesium-iodide calorimeter. This grid of towers is covered with a segmented plastic–scintillator anti-coincidence detector.

The Fermi-LAT is sensitive to $\gamma$-rays with $20 \text{ MeV} < E < 300 \text{ GeV}$, and its on-axis effective area is about 8000 cm$^2$ for $E > 1 \text{ GeV}$. The Fermi-GBM includes two sets of detectors: 12 sodium iodide (NaI) scintillators, each 12.7 cm in diameter by 1.27 cm thick, and two cylindrical bismuth germanate (BGO) scintillators, each 12.7 cm in diameter and 12.7 cm in height. The NaI detectors are sensitive in the lower end of the energy range, few keV$ < E < 1 \text{ MeV}$ and provide burst triggers and locations. The BGO detectors cover the energy range 150 keV to 30 MeV, providing a good overlap with the NaI at the lower end and with the Fermi-LAT at the high end. The Fermi-GBM has a field of view (FoV) several times larger than the Fermi-LAT and provides spectral coverage of GRB that extends from the LAT lower limit down to 10 keV (http://glast.gsfc.nasa.gov/public/instruments.html).

In Fig. 1.8 an image of the emerging Fermi-sky is reproduced. I’ll discuss Fermi results in the next sections, trying to explain how and how much they have already changed our understanding of some types of $\gamma$-ray emitters. In particular for GRBs we can talk about a “Fermi revolution era”: this will be discussed in chapter 3.

The Italian experiment Astro-rivelatore Gamma a Immagini LEggero (AGILE) was launched in 2007 and has a structure really similar to Fermi, but its effective area is about one order of
Figure 1.8: In order: pair-production in the typical structure of a satellite detector after an incident $\gamma$-ray enters the telescope (http://coss.csc.nasa.gov/docs/cgro/images/epo/gallery/glast/), the Large Area Telescope (LAT) (http://glast.gsfc.nasa.gov/public/instruments.html) and an image of the fermi-LAT first source catalogue sky.
1.2. Detection techniques

magnitude smaller (http://agile.asdc.asi.it/). Active since 2004, another satellite, Swift, has been really important for GRBs observations (http://heasarc.gsfc.nasa.gov/docs/swift/).

Swift is composed by three instruments:

- **BAT (5 - 150 keV):** with its large FoV (2 steradians) and high sensitivity, it detects about 100 GRBs per year, and computes burst positions onboard the satellite with arc-minute positional accuracy.

- **XRT (0.3 - 10 keV):** was built with existing hardware from JET-X. The XRT takes images and is able to obtain spectra of GRB afterglows during pointed follow-up observations. The images are used for higher accuracy position localizations, while the spectra are used to determine redshifts from X-ray absorption lines.

- **UVOT (170 - 650 nm):** takes images and obtains spectra of GRB afterglows during pointed follow-up observations. The images are used for 0.3 - 2.5 arc-second position localizations, while the spectra are used to determine redshifts and Lyman-alpha cut-offs.

A GRBs Coordinate Network (GCN) for GRBs alerts has been active since April 2005. Two kinds of circulars are distributed to the GRBs community. (1) the real-time (and near real-time) distribution of GRB locations detected by various spacecraft (Swift, HETE, INTEGRAL, IPN, etc), and (2) the distribution of follow-up observation reports submitted by the GRB community. MAGIC is a member of this network and performs follow-up observation of GRBs after satellite alerts, if the burst is observable at the MAGIC location during a clear dark night.

**UHE domain, Water Cherenkov experiments:** The water Cherenkov experiments reveal the Cherenkov light produced by the atmospheric shower’s 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 (Fig. 1.9). The FoV is therefore very large, and covers almost completely sky above, whereas the angular resolution is rather small. A representative of this technique is the Milagro experiment (www.lanl.gov/milagro/), situ-
ated at 2600 m above the sea level, in the the Jemez Mountains near Los Alamos, New Mexico, which operated until 2008. Essentially Milagro was a six million gallon pond of surface area $60 \times 80$ m$^2$ and 8 m in depth covered by a light-tight foil. More than 700 photomultiplier tubes arranged in a layer at the bottom and top of each detector detect the faint Cherenkov light produced by charged particles from EAS. A sparse array of smaller particle detector separation station surrounds the pond and enables efficient separation between air showers initiated by photons and those created by hadrons. Milagro measured $\gamma$-rays at $E \approx 15$ TeV. At the end of 2008 a puzzling results was reported from the collaboration [8]: unexpected localized fluxes of cosmic rays of unknown origin have been observed at $E \sim 10$ TeV on an angular scale of $10^\circ$. Both regions are inconsistent with pure $\gamma$-ray emission with high confidence and their energy spectra are moderately to strongly inconsistent with the spectrum of the isotropic CR flux. The nature of the hot spots is still unclear, but extended emission is currently not confirmed by IACT data.

Figure 1.9: The MILAGRO detector.
The successor of Milagro, called High Altitude Water Cherenkov Experiment (HAWC), will be located at the altitude of 4100 m in the Sierra Negra volcano, Mexico (http://hawk.umd.edu/). It will incorporate a new design solutions like placing the photomultipliers in isolated tanks, and adopting a larger spacing between them. HAWK will be at least an order of magnitude more sensitive than Milagro and is expected to perform a highly sensitive all-sky survey, monitoring known sources and discovering new types and classes of TeV sources.

1.3 The emerging VHE $\gamma$-ray sky

Thanks mostly to Cherenkov telescopes, a large amount of VHE sources has been detected and identified. In Fig. 1.10 we show a map of the known sources in the VHE sky: we know up to know 60 galactic sources and 35 extragalactic. In the next section I’ll give a brief description of the various types of $\gamma$-rays emitter, from Galactic (Supernova remnants, Pulsars, Bynaries systems and the Galactic center), to extragalactic (Active Galactic nuclei, Gamma-Ray Bursts). Finally, I’ll describe candidate sources for dark matter studies.

Figure 1.10: VHE-sky (http://www.mppmu.mpg.de/~rwagner/sources/).
1.3.1 Supernova Remnants

When a high mass star (final mass $\geq 3M_\odot$) collapses at the end of its life a supernova occurs. An enormous shock wave sweeps through the star at high speed, blasting away the various layers into space and leaving a neutron core and an expanding shell of matter known as a supernova remnant (SNR). The amount of energy released in such explosions in the form of kinetic energy of the ejecta is $\approx 10^{51}$ erg. According to the progenitor star mass, the neutron protostar becomes a spinning neutron star (Pulsar) or a black-hole (BH). Accordingly, SNRs are classified into shell-types (no conspicuous feature seen in the center), plerions (interior sub-structure indicating the presence of a pulsar in the center, emitting a “pulsar wind”) and composite (appearing like plerions or shell-type, depending on the observed wavelength). Galactic SNRs are extended-sources with a diameter of typically $0.1^\circ$.

When a supernova occurs about 99% of the initial gravitational energy is released into neutrinos ($\sim 10^{53}$ erg), 1% into kinetic energy of the remnants particles ($\sim 10^{51}$ erg) and only 0.01% into radiation ($\sim 10^{49}$ erg). Even with an efficiency of 0.01%, the supernova event prompts a brief burst of radiation that may temporarily outshine the entire host galaxy, before fading out over several weeks or months. During this short interval, a supernova can radiate as much energy as the Sun would emit over 200 million years! It is believed that SNRs are the sources for CRs up to the knee [22], mainly because the total power contained in CRs matches a few percent of the total mechanical energy released by supernovae in our Galaxy: the local density of CRs in the galaxy is $\rho_{CR} \approx 1 \text{ eV/cm}^3$ while the mean life time of cosmic rays in the galactic disk is $\tau_{CR} \approx 6 \cdot 10^6$ yrs. The needed power to generate all CRs in the Galaxy is thus: $L_{CR} = V_D \cdot \rho_{CR} / \tau_{CR} \approx 5 \cdot 10^{40} \text{erg/s}$ where $V_D$ is the Galaxy volume: $V_D = \pi R^2 d \approx \pi \cdot (15 \text{ kpc})^2 \cdot (200 \text{ pc}) \approx 4 \cdot 10^{66} \text{ cm}^3$. Supposing that a supernova has an acceleration efficiency $\epsilon_{SNR}$ of a few percent and a supernova rate of $r_{SN} \approx 0.03 \text{ yr}^{-1}$, the power provided by SNRs reads as: $L_{SNR} = \epsilon_{SNR} \cdot E_{SN} \cdot r_{SN} \approx \epsilon_{SNR} \cdot 10^{51} \text{ erg} \cdot 10^{-9} \text{s}^{-1} \approx \epsilon_{SNR} \cdot 10^{42} \text{ erg/s}$. Thus, even an efficiency of 0.01% is enough to cover the entire energy budget od GCRs [200]. Last but not least we have a convincing theory on particle acceleration, the diffusive (Fermi 1st-order) shock acceleration mechanism, which can describe the accelerations (almost) up to the knee fo the CR spectrum. Observing the secondary VHE $\gamma$-rays accompanied to the accelerate CRs we can study the acceleration mechanisms and try to discriminate between leptonic or hadronic acceleration from the SNR. However, it’s really difficult to disentangle
the hadronic VHE component (mainly produced by $\pi^0$ decay) from the leptonic one (produced by IC scattering of interstellar radiation field photons or cosmic microwave background photons of ultrarelativistic electrons): the resulting spectra are similar, at least in the relatively narrow range of energies sampled by IACTs.

In 2004, the HESS telescope array could resolve spatially the TeV-emission from the SNR shell of RX J1713.7-3946 [17] and correlate it with the X-ray emission detected by Rosat (Fig. 1.11). Deep HESS observation of the source were performed also later, reaching an unprecedented precision [21]. While this evidence seems to favor the leptonic acceleration, hadronic acceleration has still not been ruled out. The reason depend very much on the local presence of of gas densities and magnetic field which can boost or suppress one of the two contributes. MAGIC detected a new VHE $\gamma$-ray source located close to the Galactic Plane, namely MAGIC J0616+225 [28], which is spatially coincident with the SNR IC 443. The mea-
sured energy spectrum is well fitted by a power law with spectral index $3.1 \pm 0.3$. MAGIC J0616+225 is point-like for MAGIC spatial resolution, and appears displaced to the south of the center of the SNR shell. It is spatially correlated with a molecular cloud [225]. There is also an EGRET source centered in the SNR shell. The observed VHE radiation may be due to $\pi^0$ decays from interactions between cosmic rays accelerated in IC 443 and the nearly dense molecular cloud. However if the particle population were accelerated at the SNR shock front a harder spectrum could be expected. A possible distance of this cloud from IC 443 could explain the steepness of the measured VHE $\gamma$-ray spectrum as a diffusion effect. These facts hint to a possible hadronic origin of the VHE-$\gamma$ emission. A substantial improvement for the detection of SNRs requires a telescope with increased angular resolution (in order to perform morphological studies) and a sensitivity up to 100 TeV in order to explore the differences between acceleration mechanism channels.

The first object ever observed by a Cherenkov telescope was the plerion-type SNR Crab [239] which is also the strongest steady TeV gamma ray source in the sky. It is nowadays used as a calibration source for all IACTs. The SN in Crab exploded in 1054, and its light was so bright that it was seen in the sky for almost a month, even during daytime. The SN, at its peak, was as bright as 400,000,000 suns!

1.3.2 Pulsar and Pulsars Wind Nebulae

A pulsar is a highly magnetized rotating star (with periods between $\sim 1$ ms and $\sim 1$ s), formed after a SN explosion. The suggestion that pulsars were rotating neutron stars and could be observable at radio frequencies was put forth independently by Gold and by Pacini in 1968 [116,177], and was soon proven beyond reasonable doubt by the discovery of a pulsar with a very short (33 ms) pulse period in the Crab nebula. Pulsars have, subsequently, been found to emit also at visible, X-ray, and $\gamma$-ray frequencies. The pulsed emission is beamed along the magnetic field axis, which itself is mis-aligned with the spin axis. Therefore the emitting region rotates and sweeps the line-of-sight to the Earth at regular intervals. Magnetic field strengths $B \approx 10^{12}$ G, the highest deduced to exist in the Universe, are predicted for pulsar.

The rotation energy of the neutron star fuels the acceleration of CRs. The pulsar expels a relativistic wind of particles and magnetic fields which determine the formation of a syn-
1.3. The emerging VHE $\gamma$-ray sky

chrotron nebula in the surrounding of the source. The nebula is also called a pulsar wind nebula or plerion. Electron-positron pairs originate in the magnetosphere via interaction of the $\gamma$-rays with the magnetic field and they may escape through the polar cap regions in the form of a wind that eventually terminates in the surrounding interstellar medium. $\gamma$-rays can be produced through several radiation mechanisms in three different regions: magnetosphere, relativistic wind, surrounding nebula.

Up to few months ago, there was no wide consensus on the physical mechanism for magnetospheric emission and the most crucial issue about HE emission was to understand the location of the acceleration region: the polar cap [70] or in the outer gap [59,206]. (A sketch of the models in Fig. 1.12). In the first case, the synchrotron photons will interact with the strong magnetic field and create electron-positron pairs initiating electro-magnetic cascades. The electrons and positrons can in turn IC upscatter the synchrotron radiation field. In the other model, $\gamma$-rays of GeV energies are produced farther out in vacuum gaps in the outer magnetosphere where energetic electrons IC upscattered infrared/optical photons from outside the gap. Both models differ mainly in their predictions of the position of a cut-off in the energy spectrum: at tens of GeV in the polar cap model, at $\sim 100$ GeV in the outer gap model. Fermi results are deeply changing our knowledge about this and about pulsar physics in general. Fermi has identified 55 pulsars that radiate in the gamma range, in just one year of data collection. This can be compared to the 5 new pulsars discovered over nine years by Fermi's predecessor EGRET. 14 gamma-ray-only pulsars, 22 radio loud pulsars (already identified by radio telescopes), and 9 gamma-ray Milly Second Pulsars (MSPs). The main Fermi results, described in [14], about Pulsars can be summarized:

- in blind searches, Fermi found previously unknown pulsars that appear to radiate only $\gamma$-rays [11]. It may be that these pulsars emit in radio wavelengths, too, but their radio beams are simply not visible from Earth; follow-up observations should reveal if they radiate at other frequencies, including (faintly) in the radio;

- a large fraction of the local energetic pulsar are GeV emitters;

- 9 MSPs have been detected for the first time in $\gamma$-rays [9];

- for most of the pulsar, $\gamma$-ray emission appear to mainly come from the outer magnetosphere, while the polar-cap emission remains plausible for a remaining few.
Figure 1.12: Sketch of the pulsar models
The possible exclusion of the polar cap model, is a confirmation of an important MAGIC result. In 2008 MAGIC, thanks to a special trigger setup, the so-called MAGIC analogue sum trigger, for the first time in IACT history revealed the pulsed emission of the Crab Pulsar above 25 GeV with a statistical significance of $6.4\sigma$ [38]. As already mentioned, the Crab Nebula is still the brightest and steady emitter in the VHE sky, and is used as calibration candle. The $\gamma$-ray emission is dominated by the pulsed emission from the rotating neutron star below GeV energies, and by the the steady emission from the nebula above GeV energies. Its broad-band spectrum is reported in Fig. 1.13 in which two components are visible that are usually attributed one to synchrotron radiation and the other to the Compton upscattering of these same photons off the parent relativistic electrons (that emerge from the termination shock of the pulsar wind). The Crab is a prototype of the Pulsar Wind Nebulae (PWN), pulsars displaying a prominent nebular emission.

### 1.3.3 Binary systems

When a star is close to a compact object, like a neutron star, a pulsar or a black hole, it may start to orbit around it loosing an enormous amount of matter - via wind or Roche lobe overfilling - that collapse onto the dense object. The stream of collapsing particles is
accompanied by an associated X-ray emission and there is already evidence that particles are also accelerated to VHE. Like beamed emission in the AGN (see below), as result of the infalling material two jets step out in the direction orthogonal to the rotation plane of the compact object where the particles are directly accelerated. Due to this similarity with AGN, these objects are also called microquasars. Four TeV binaries have been detected so far: PSR B1259-63/SS2883, LS 5039, LSI+61 303, and CygX-1.

The pulsar-powered binary system PSR B1259-63/SS2883, composed of a 48 ms pulsar and a B2e star, was observed by HESS above 380 GeV [18]. The observation seems to confirm the IC scattering as mechanism for the acceleration, whereas hadronic emission is not ruled out. LSI+61 303, composed of a black-hole and a B0 star, was observed by MAGIC [27]. During a multiwavelength campaign set up during September 2007, performed by MAGIC, XMM-Newton and Swift XRT, a simultaneous X-ray and VHE-γ outburst was detected [30]. This correlation indicates a simultaneity of the emission processes and it favors a leptonic models. This fact, together with the measured photon indices, suggests that in LSI+61 303 the X-rays are the result of synchrotron radiation of the same electrons that produce VHE emission as a result of IC scattering of stellar photons.

The similar object LS 5039, observed with HESS, clearly show a periodicity of 3.9 days in γ-rays alone: the VHE spectrum may indicate photon-photon absorption or cascading. A hint of a signal from the well-established microquasar (MQ) Cygnus X-1 was found by MAGIC during a short-lived, intense flaring episode on September 24, 2006 [29]. This TeV flare was coincident with a historically high X-ray flux. The emission is point-like and excludes the nearby radio nebula powered by the relativistic jet. Cyg-X1 is the first stellar mass black hole, and hence the first established accreting binary, established as a VHE source.

1.3.4 The Galactic Center

The Milky Way center, crowded with many astrophysical sources, was observed by most IACTs: Whipple, CANGAROO, HESS and MAGIC. In particular the HESS and MAGIC results [20,26] are in agreement and measured a steady flux consistent with $\Gamma \approx 2.2$ up to $E \sim 20$ TeV with no apparent cutoff. The Galactic Center has three candidates γ-ray emitters: the shell type SNR Sgr A East, the closeby PWN G 359.95-0.04, and the supermassime black-hole Sgr A* itself.
1.3. The emerging VHE $\gamma$-ray sky

Possible radiation mechanisms includes IC scattering of energetic electrons, neutral pion decay in the interaction of hadrons with the interstellar medium or dense radiation fields, and curvature radiation of UHE protons close to Sgr A*. Dark-matter annihilation into VHE $\gamma$-rays, one of the very reason why the Galactic Center was observed, is currently disfavoured.

1.3.5 Active Galactic Nuclei

Active Galactic Nuclei (AGNs) are a class of galaxies that show evidence of hosting masses of $10^6$ to $10^9$ solar masses concentrated in regions of the size of the solar system. These central objects must be super-massive black-holes. In the unified scheme for AGNs [227] the central

Figure 1.14: The AGN NGC4261, as seen by the Hubble Space Telescope. The two jets, the dust torus and the accretion-disk could be spatially resolved.

Figure 1.15: AGN unified model sketch
BH is surrounded by a rotating luminous accretion-disk, heated up to temperatures in the optical and UV [196]. The plasma emits a thermal spectrum peaking in X-rays.

A thick dust torus, situated further outside, obscures the emission from the accretion disk and from clouds orbiting above the disk. By a yet not fully understood mechanism, the infall of material onto the BH turns on two jets of material and radiation, propagating in opposite directions perpendicular to the dish and torus plane. The jets can extend up to a thousand times the extension of the galaxy itself, up to Mpc scale.

Depending on the observation angle with respect to the jet axis, a rich phenomenology of AGNs can be observed leading to many classes and sub-classes of AGNs: Blazar, Quasar, Seyfert, Radiogalaxy. It’s generally believed that all AGNs show similar properties and that the surrounding medium only affects the acceleration of CRs at a secondary order. For $\gamma$-ray astronomy, the most important ones are those where a jet points directly in the direction of the Earth, i.e. Blazars.

The emission of blazar is amplified by relativistic Doppler effect. Blazars show strong flux variability at all frequencies. The emission can grow by up to two order of magnitude if the blazar is in “flaring state”. Otherwise, it is said to be in “quiet state”. The characteristics of the flares are very different: flares do not necessarily occur simultaneously at all frequencies, sometimes they last different times at different frequencies.

The mainstream interpretation of the blazars emission is within the SSC framework. A pure SSC model predicts a definite correlation between synchrotron and IC radiation. Evidence of neutrinos in a flare would provide smoking gun evidence for the presence of a hadronic emission (component). Simultaneous multi-wavelength observations are the best tool to understand and model the jet emission.

In the past several years MAGIC performed multiwavelength campaigns to observe various types of AGNs. Especially important are the two brightest blazars in $\gamma$-rays: Mrk 501 [193] and Mrk 421. (The latter was the first blazar ever observed with a Cherenkov Telescope [192].) As an example, let’s mention some of the latest MAGIC observations: the detection of the quasar 3C279, the most distant ever so far, the observations of BL Lac objects [32], the multiwavelength campaigns on the radiogalaxy M87 [34], on PG 1553+113 (with unknown redshift) [36], on 1ES 1959+650 [224] and on Mrk 421 [226]. A further aspect, is that the variability of the AGN in the VHE can provide information about possible Lorentz
invariance violation by means of the non trivial light dispersion relation expected in some quantum gravity models. The MAGIC data of Mrk 501 [31] showed a negative correlation between the arrival time of photons and their energies: assuming that the delay to be a quantum gravity effect, to a lower limit to the quantum gravity scale was deduced to be $\sim 0.03$ Planck mass.

Most TeV blazar were also detected by Fermi LAT. For a catalogue of the LAT Bright AGN Sample (LBAS), in the first three months of sky-survey operation, see [10]. Fermi discovered hundreds of new sources, proving that blazars dominate the extragalactic sky and it should be emphasized that:

- important spectral properties (correlation of photon index with blazar class, spectral breaks, relative constancy of photon index with flux) have been observed;

- variability time scales, ranging from sub-day to several months, were observed;

- many multifrequency studies have been triggered by Fermi observations, providing time-resolved SEDs and temporal interband (radio, optical, X-ray, TeV) correlation.

### 1.3.6 Production from self-annihilation of dark matter

In many Dark Matter (DM) scenarios, the annihilation of DM particles can produce $\gamma$-rays with a continuum spectrum that extends up to very high energies of the order of the electroweak symmetry breaking scale (hundreds of GeV). Astrophysical structures dynamically dominated by DM, such as dwarf Spheroidal Galaxies, Galaxy Clusters (the largest ones in the local Universe being mostly observable from the northern hemisphere) and Intermediate Mass Black Holes, can be considered as interesting targets to look at for DM annihilation with Imaging Atmospheric Cherenkov Telescopes (IACTs). Instead, the center of our Galaxy is strongly contaminated with astrophysical sources.

DM particle candidates should be weakly interacting with ordinary matter, otherwise they would have been already detected. The theoretically favoured ones are heavier than the proton, and are called Weakly Interacting Massive Particles (WIMPs). Experimental limits from accelerators indicate a minimum mass of 50 GeV. In supersymmetric models the lightest supersymmetric neutral particle, the neutralino, is predicted to be a Majorana particle, and thus a natural candidate for WIMP.
The most relevant neutralino interaction for the purposes of indirect DM searches is the self annihilation in fermion-antifermion pairs, gauge bosons pairs and final states containing Higgs bosons. The subsequent hadronization results in a $\gamma$-ray power-law spectrum with a sharp cutoff at the neutralino mass (expected to be between 50 GeV and several TeV).

Recently it was pointed out that the Internal Bremsstrahlung (IB) process may boost the $\gamma$-ray yield of the neutralino self-annihilation at the higher energies by up to four orders of magnitude, even for neutralino masses considerably below the TeV scale [58]. This discovery represents a very important issue for the indirect DM search, particularly for the IACTs which are sensitive to the energy range most affected by the $\gamma$-ray flux enhancement due to the IB process. Moreover, the IB introduces features in the $\gamma$-ray spectrum that potentially allow an easier discrimination between a DM source and the standard astrophysical sources located in the vicinity, whose spectrum is usually a featureless power law.

MAGIC has still observed two particularly promising candidate sources for dark matter annihilation signals, the nearby dwarf galaxies Draco [33] and Willman 1 [40]. No signal was found so far, but with MAGIC II and CTA the efforts to detect such a signal will continue.

1.3.7 Gamma-Ray Bursts

This phenomena will be discussed more extensively in chapter 3.
VHE-γ astrophysics was developed only in the last several years, because VHE photons’ detection is a complex experimental problem. The atmosphere shields us from photons of frequencies above the ultraviolet (UV). Photons in the X and soft γ range can be directly detected from space-borne instruments. VHE γ-ray detection would require very large effective collecting areas in space-borne detectors, because cosmic VHE γ-ray spectra are very steep ($\propto E^{-2.5}$ or steeper). So in practice VHE γ can only be detected by ground based detectors: with bigger effective areas and higher sensitivities this class of detector is sensitive to VHE and UHE photon fluxes. The sensitivities of current and incoming detectors are shown in Fig. 2.1. Ground-based VHE Extensive EAS such as Milagro and ARGO and the new generation of Cherenkov telescopes like CANGAROO, H.E.S.S., MAGIC and VERITAS detect the secondary particles and, respectively their Cherenkov light, of the atmospheric showers produced by the primary photons and the cosmic rays of energy higher than the primaries observed by satellites.

The MAGIC telescope is an Imaging Atmospheric Cherenkov Telescope (IACT) and its scientific activity is based on the detection of the Cherenkov photons produced in the Earth atmosphere by electromagnetic showers initiated by VHE cosmic γ-rays. The explored energy window is between $\sim 30$ GeV (with “sum”-trigger, see section 2.4.5) and $\sim 10$ TeV (also depending on the source spectrum). The aim of this chapter is to explain how and why the VHE gamma rays can be indirectly detected by MAGIC.
2. Imaging Air Cherenkov Technique and The MAGIC Telescope

Figure 2.1: Sensitivities of some present and future HE gamma detectors, measured as the minimum intensity source detectable at $5\sigma$. Data for EAS and satellite detector are based on one year of data taking; data for Cherenkov telescopes are based on 50 hours of data.

2.1 Extended Air Showers (EAS)

Cosmic rays hardly ever hit the ground, but instead collide with the nucleons, mainly nitrogen, present in the atmosphere. In such collisions, new particles are created which themselves interact with the atmospheric nuclei, leading to an air shower. Depending on whether the impinging particle is a hadronic (nucleus) or an electromagnetic (electron or $\gamma$-ray) particle, one makes the distinction between hadronic and electromagnetic air showers.

$\gamma$-rays above a critical energy $E_c \sim 20$ MeV lose energy in air primarily through pair creation, electrons and positrons above $E_e \approx 81$ MeV $[100]$ through bremsstrahlung, below through ionization $[52]$. The energy loss due to bremsstrahlung is proportional to the energy $E_e$ of the electron ($-dE_e/dx = E_e/X_0^\gamma$) characterized by the $X_0^\gamma$ (37 g/cm$^2$ in air). Analogously, the pair creation process by gamma rays can be characterized by a mean free path $X_0^\gamma = 9/7 \cdot X_0^e$ (47 g/cm$^2$ in air). The secondary particles created in these processes are again electrons, positrons and photons which can generate secondaries until they reach the respective critical energies. The number of created particles increases exponentially with the shower length until the shower maximum is reached (when the average energy of the cascade particles equals the critical energy) and the shower drops off. The longitudinal development of the number of electrons $N_e$ above a certain energy threshold with the slant depth $X$ in units of the radiation
2.1. Extended Air Showers (EAS)

length $X_0$ [122,207] (the ”Rossi Approximation B”) is:

$$< N_e(t, E_0) > \simeq \frac{0.31}{\sqrt{\ln(E_0/E_c)}} \cdot \exp(t - 1.5 t \ln s), \quad (2.1)$$

with:

$$t \equiv \frac{X}{X_0} \approx \frac{X_{air}}{X_0 \cdot \cos(\theta_{sh})} \cdot \exp(-\frac{H}{H_0}), \quad (2.2)$$

$$s \equiv \frac{3t}{t + 2 \ln(E_0/E_c)}, \quad (2.3)$$

$$\Delta N_e(s) \simeq \frac{9}{14} (s - 1 - 3 \ln s) \cdot N_e(s), \quad (2.4)$$

where $E_0$ is the primary particle’s energy, $X_{air} = 1013 \text{gcm}^{-2}$ is the column height of air at the ground, $H$ is the height from ground, $H_0 = 8 \text{ km}$ is the scale height of the atmospheric pressure. $\Delta N_e(s)$ is the increase of the number of photoelectrons at each new branching. $\gamma$-ray induced air showers in the energy range from 30 GeV to 30 TeV have their shower maximum well above 2200 m altitude. Most energetic showers are reaching deeper into the atmosphere. Below about 100 GeV, the showers die out completely before reaching the ground. Moreover, the distribution is asymmetric with respect to the shower maximum.

It is useful to invert equation 2.2:

$$H = H_0 \cdot \ln \left( \frac{R_{air}}{t} \right), \quad (2.5)$$
Figure 2.3: Longitudinal development of an electromagnetic shower: $N_e$ versus the slant depth $X$, measured in radiation lengths $t = X/X_0$. The lines are calculated for different values of $\ln(\frac{E_0}{E_c})$.

with $R_{\text{air}} = X_{\text{air}}/X_0 = 27.4$ to calculate typical heights $H$ of the shower maximum (Fig. 2.3.) For instance, at 100 GeV the shower maximum ranges from $\sim 12$ km to 10 km a.s.l.

There are large fluctuations from shower to shower, even for those of same energy. From eq. 2.4 follows that the fluctuations are smallest near the shower maximum$^1$. At 10 km height, a 100 GeV shower has then: $s = 1.38, \Delta N_e \approx 0.4N_e$ a 1 TeV shower: $s = 1.18, \Delta N_e \approx 0.2N_e$.

Multiple scattering of the electrons and positrons deflects the secondary particles away from the primary $\gamma$-ray direction, the shower axis. The transverse extension of an electromagnetic shower can be parameterized by the Molière radius $R_M$ [52]:

$$R_M = 21.2 \text{ MeV} \cdot \frac{X_0}{E_c} \approx 9.3 \text{ g/cm}^2 \text{ (in air)}$$

(2.6)

On average, only 10% of the lost energy lies outside a cylinder with radius $R_M$ ($\approx 78$ m at sea level, 190 m at 10 km height) and about 99% is contained within $3.5R_M$. Traditionally, the lateral spread can be parameterized by the Nishimura-Kamata-Greisen (NKG)-formula [123, 173], although only strictly valid between $s = 1.0$ to 1.4:

$$\rho_e(r) = K \cdot \frac{N_e}{R_M^2} \cdot \left( \frac{r}{R_M} \right)^{s-2} \cdot \left( 1 + \frac{r}{R_M} \right)^{s-4.5},$$

(2.7)

where $K$ is some normalization constant.

$^1$However, eq. 2.4 does not include development fluctuations, i.e. the dependence of $s$ with height.
Hadronic Air Showers  A hadronic shower is produced in the collision of a nucleus with another (atmospheric) nucleus, creating mainly pions and Kaons and further nucleons. The last ones and the possible fragments of the original nucleus form part of the hadronic core of the cascade. The cascade continues to suffer collisions with the air, similar to the first one, until its energy per nucleon falls below the pion production threshold of about 1 GeV.

Hadronic showers have various components:

- *Hadronic component:* The core of energetic hadrons, consisting of nucleons and mesons. As these hadrons are heavy, they can transfer significant transverse momentum to their decay products.

- *Electromagnetic component:* Consists of electrons, positrons and photons from electromagnetic sub-cascades, initiated by the decay of mesons, mainly the neutral pions \( \pi^0 \). About one third of the collision energy is transferred to the electromagnetic component in each hadronic interaction. This process transfers continuously energy from the hadronic to the electromagnetic component which ends up being the dominant one at the shower tail. If occurring at a high transverse momentum, the pure electromagnetic component can resemble a stand-alone electromagnetic shower and produces an irreducible background for Cherenkov Telescopes.

- *Muonic component:* Some of the charged pions and Kaons decay before they interact, most of them into muons and neutrinos\(^3\). The muons themselves interact almost exclusively by ionization and usually reach the ground before decaying. Single muons with large transversal momenta can travel far from the shower core and are recognized by their clean Cherenkov rings (see e.g. Fig. 2.9 right).

An interaction length can be defined also for hadronic showers: \(-dE/dx = E/X_h\), characterized by the hadronic radiation length \( X_h \) (83 g/cm\(^2\) in air for protons and 107 g/cm\(^2\) for pions). The average shower maximum can be expressed as [99]:

\[
    t_{\text{max}} = \ln\left( \frac{E_0}{AE_c} \right),
\]

\(^2\)The \( \pi^0 \) has a decay time of only \( \tau = 8.4 \cdot 10^{-17} \) s, while the \( \pi^\pm \) particles decay much slower: \( \tau = 2.6 \cdot 10^{-8} \) s [183].

\(^3\)The neutrino component of the air shower is not visible for Air Cherenkov Telescopes and is further neglected here.
with \( A \) being the mass number of the primary nucleus. At the maximum of shower development, there are approximately 2/3 particles per GeV of primary energy \([100]\). For protons, the characteristic interaction length is more than twice of the corresponding electromagnetic interaction lengths, hadronic showers penetrate thus deeper into the atmosphere than electromagnetic ones of the same energy. Moreover, hadronic showers show larger fluctuations from shower to shower, than compared to the electromagnetic case.

### 2.2 Cherenkov Light and the Imaging Air Cherenkov Technique

Observations of \( \gamma \)-rays using the subsequently produced electromagnetic showers in the atmosphere is possible by detecting the electromagnetic radiation of the secondary shower particles. Charged particles moving through a dielectric medium of refraction index \( n \) with a velocity \( \beta c \) larger than the local phase velocity of light \( (c/n) \) generate a photonic shock-wave: the Cherenkov emission. The threshold condition for this type of light emission is:

\[
\beta \cdot n > 1 \\
E > m_0 c^2 \cdot \frac{n}{\sqrt{n^2 - 1}}
\]

(2.9)

with \( m_0 \) being the particle rest mass and \( E \) the particle energy. It’s possible to translate this condition into an energy threshold for electrons and positrons of about 20 MeV in dry air at 2200 m a.s.l. \((n = 1.0004 \text{ at } 400 \text{ nm})\). In principle, there is also a wavelength-dependency of the refraction index in air. The refractivity \( (n - 1) \) varies by less than 1% over the visible spectrum and has been neglected. Cherenkov light is emitted on a narrow cone around the direction of the emitting particle with an opening angle \( \theta_c \) of:

\[
\cos(\theta_c, \lambda) = \frac{1}{\beta n(\lambda)}
\]

(2.10)

The Cherenkov light cone opened by gamma ray induced air showers has an opening angle of typically 1.2 deg at 2200 m a.s.l., 0.75 deg and 0.36 deg at 10 km and 20 km altitude, respectively. From everywhere in the particle’s track, the Cherenkov light hits the ground forming an ellipse on the ground. As there are many charged particles produced in an air shower, the individual Cherenkov cones overlap, and the sum of all Cherenkov light photons illuminate
uniformly an area of typically 50000 square meters at 2200 m a.s.l.: the Cherenkov light pool. This number hardly depends on whether the light has been emitted at 20 km or at 10 km altitude.

In an electromagnetic air shower, about 500 Cherenkov photons are created per GeV of incident gamma ray energy in the spectral range between 300 nm and 600 nm. A considerable fraction of these photons is lost due to scattering and absorption in the air. Since the involved scattering angles are large compared to the Cherenkov angle, scattered photons are usually lost (as well as absorbed photons) for a Cherenkov telescope. The cross section of Rayleigh-scattering is $\propto \lambda^{-4}$, thus the UV part of the spectrum is especially affected by the Rayleigh-scattering, which is the dominant contribution to the Cherenkov light attenuation under perfect weather conditions. Mie-scattering which occurs to particles larger than the photon wavelength (water droplets, dust and calima $^4$) has cross section is $\propto \lambda^{-(1-1.5)}$: it prevails under poorer weather conditions. At altitudes above 10 km, absorption by ozone may play a significant role, for $\lambda < 400$ nm. Absorption by H$_2$O and CO$_2$ molecules in the air affect $\lambda > 800$ nm. As the absorption affects especially short wavelengths below 300 nm, it introduces a zenith-angle dependency of the Cherenkov light spectrum: Showers from particles incident at higher zenith angles have spectra shifted toward larger wavelengths (redder

Figure 2.4: Sketch of Cherenkov light production

$^4$Calima is the name of the ultra-thin Saharan sand transported by the wind form Africa, which for some days a year fills the entire sky
Fig. 2.5 shows the simulated average Cherenkov light spectrum for electromagnetic showers seen from different angles with respect to vertical incidence: the effect of the wavelength shift can be seen. There is also an energy-dependency of the absorption and scattering: low-energy showers lose their energy at higher altitudes, their Cherenkov light needs to cross a larger amount of air and is more affected by absorption than showers of higher energy.

Fig. 2.6 shows the average radial Cherenkov light photon density for vertically incident 100 GeV gamma ray and 400 GeV proton induced showers. The values have been averaged over all azimuth angles. One can see that the radial light density is almost constant up to about 130 m impact distance for the electromagnetic shower, with a slight enhancement at about 80–120 m and a consecutive steep fall beyond 125 m. There is even light emission beyond 200 m impact distance. With the mirror area of the MAGIC telescope, shown as a dashed line, effective collection areas of $\pi \cdot (180 \degree)^2 = 10^5 \text{ m}^2$ should thus be reached.

Fig. 2.7 shows the locations of Cherenkov photons hitting the ground at 2200 m a.s.l. from simulated 300 GeV and 100 GeV $\gamma$-ray showers. While the 300 GeV Cherenkov light pool is
2.2. Cherenkov Light and the Imaging Air Cherenkov Technique

Figure 2.6: Lateral distributions of Cherenkov photon densities for 100 GeV gamma ray and 400 GeV proton induced showers at an altitude of 2200 m a.s.l. for vertical incidence. Figure from [48].

Figure 2.7: Examples of the lateral distribution of Cherenkov light from simulated vertically incident gamma showers on ground at 2200 m a.s.l. The area displayed covers 400×400 m with the shower core at the center. Atmospheric extinction has not been taken into account. Left: 300 GeV gamma ray, right: 100 GeV gamma ray. Figures from [51].
filled out uniformly, at 100 GeV the individual ring structures become visible.
Proton showers show their brightest average emission at 20 m impact parameter, steeply falling down beyond. However, local fluctuations are much bigger for hadronic showers, due to the Cherenkov emission of penetrating single pions and Kaons, and the muons produced in the respective decays. For the same reason, light can be spread to larger distances from the shower core. While the hadronic cascades are isotropic in arrival directions, a Cherenkov telescope searches for an enhancement of electromagnetic showers from a hypothetical source direction. Only with the invention of the Imaging Atmospheric Cherenkov Technique (IACT) by the Whipple collaboration in 1989 [239], the hadronic background could be reduced to such an extent that the first TeV gamma rays from the Crab Nebula could be detected.
An IACT uses statistically different shapes of hadronic and electromagnetic showers to differentiate between both. A sketch of the IACT technique can be seen in Fig. 2.8. For a γ-initiated shower, whose axis is parallel to the telescope axis, photons emitted higher in the atmosphere are reflected close to the camera center, while the shower tail image extends towards the camera edges. The quantity of the collected light is then correlated in first order to the number of particles producing Cherenkov light of the EAS and, at the second order to the primary particle energy. Since electromagnetic showers are narrower and better confined than hadronic showers, their images can be reduced to a couple of statistical parameters (the so-called Hillas parameters [130]) which in turn are used to suppress more than 99% of the unwanted hadronic background. Fig. 2.9 shows three typical examples of such shower images in the camera of MAGIC. Two other differences between gamma ray and hadron induced showers concern the distributions of arrival times of the Cherenkov photons on ground:
gamma-ray showers: 1 - 3 nsec
hadron showers: 3 - 7 nsec.

2.3 Current IACT experiments

Present instruments in the field of ground-based gamma-ray astronomy are sensitive to photons with energies above 50-100 GeV. The most sensitive of these instruments are IACTs. The pioneering experiment in this field was the Whipple telescope, which in 1989 detected the first γ-ray signal from the Crab Nebula [238]. Significant improvements were then achieved by HEGRA (http://www.mpi-hd.mpg.de/hfm/CT/CT.htm) through the introduction of stere-
2.3. Current IACT experiments

Figure 2.8: Schematic view of IACT Technique. The image in the camera has an elliptical shape, the so called head and tail of the shower are here labelled as A and C.
Figure 2.9: Example of the three different types of shower images recorded by the MAGIC camera: Left: gamma-like shower, center: hadronic shower, right: single muon.

...oscropy, i.e. the simultaneous imaging, from different angles of the same shower with multiple Cherenkov telescopes. The HEGRA experiment does not exist anymore, and the Whipple telescope was upgraded with a more effective camera [139].

At the moment, there are four big IACT experiments running worldwide. The CANGAROO-III (Collaboration of Australia and Nippon (Japan) for a Gamma Ray Observatory in the Outback) experiment (http://icrhp9.icrr.u-tokyo.ac.jp/) located in South Australia consists of four telescopes with a diameter of 10 m. The VERITAS (Very Energetic Radiation Imaging Telescope Array System) experiment (http://veritas.sao.arizona.edu/) has an array of four 12 m optical reflectors, based on the design of the existing 10 m telescope of the Whipple Observatory. Since 2003, the H.E.S.S. (High Energy Stereoscopic System) experiment (http://www.mpi-hd.mpg.de/hfm/HESS/) located in Namibia is in operation consisting of four single telescopes. This array will be upgraded in 2010 with a fifth, and larger, instrument. The experiment using the world’s largest single telescope is the MAGIC (Major Atmospheric Gamma-ray Imaging Cherenkov) experiment located in the North-Atlantic island of La Palma, Spain, and running since 2004 by an international collaboration of 23 institutes in 8 European countries and one in the USA. The aim of this project was to achieve the lowest energy threshold among all IACTs of the third generation (down to 30 GeV) and the possibility to move the telescope to any sky position within few tens of seconds in order to quickly react to GRB alerts. The elements of the MAGIC telescope are described in the next section. A second MAGIC telescope, on the same site, with further improvements towards low energy, is now operational, after a commissioning phase which finished recently (end 2009). A brief
discussion of the differences introduced in the second telescope will be discussed in the last section.

2.4 The MAGIC Telescope

The Major Atmospheric Gamma-ray Imaging Cherenkov (MAGIC) telescope is the largest operating single telescope of the world. It is located at the Roque de los Muchachos (2200 m a.s.l.) on the Canary Island of La Palma. Fig. 2.10 shows a picture of the MAGIC telescope.

Figure 2.10: View of the MAGIC Telescope.

2.4.1 Frame and Reflector

One of the aims of the MAGIC project is to detect GRBs. The mirrors are mounted on a light weight space frame structure made of carbon fiber reinforced plastic tubes (5 tons). Due to this light weight construction, the MAGIC telescope can be repositioned in any direction within less than 30 seconds. The reflector itself consists of a 17 m diameter tassellated parabolic shaped mirror, fragmented in 964 spherical (all aluminum) mirrors with an individual size of $49.5 \times 49.5 \text{ cm}^2$. This leads to a total reflecting surface of 236 m$^2$. The surface
of the mirrors is coated with a protective quartz layer and has a high reflectivity of more than 80% in the wavelength range between 250 nm and 750 nm, see Fig. 2.11. Because of the parabolic shape of the dish, the time spread of Cherenkov light reflected from the dish into the camera is minimized and the time window from which the signal is extracted can be reduced. This leads to less noise integration and thus to a better signal to noise ratio, i.e. sensitivity. Another advantage is the very good separation between hadronic and electromagnetic showers because of their time information. The mirrors are mounted in groups of four (at the edge of three) on one panel. For compensation of degradations of the optical performance due to the deformations of the telescope structure depending on the pointing position, each panel can be re-adjusted during the telescope operation by the Active Mirror Control (AMC) [2]. The alignment of the mirrors is performed with an artificial light source at a distance of 980 m (the Roque lamp), with a displaced camera to match the focal length for this procedure. In the center of each mirror panel, a red laser (685 nm) points towards the camera. The position of the laser spot of each mirror on the camera is recorded as a result of the alignment procedure with the Roque lamp. The reflector is usually focused to 10 km because this corresponds to the typical height of the shower maximum for 100 GeV gamma ray induced showers at low zenith angle observations (see Fig. 2.3 and Eq. 2.5). From geometrical arguments follows that a mirror dish of diameter $D$ “sees” the shower until a maximum height of $H > (d_\perp - D/2)/\tan(\theta_c)$, where $d_\perp$ is the maximum possible impact parameter $d_\perp \approx 65$ m. Even in the “worst case” – a shower occurring at the maximum impact.
2.4. The MAGIC Telescope

parameter – the shower is imaged up to a height of \( H \approx 3.3 \text{ km} \) from the telescope. This number corresponds to about 14 electromagnetic radiation lengths for a vertically incident \( \gamma \)-ray. From Fig. 2.3 it follows that showers below 100 GeV are practically always fully contained in the camera image. Until the highest imaged gamma ray energies of 20 TeV, at least the shower maximum is always fully contained in the camera image.

2.4.2 Camera

The Cherenkov photons are reflected and focused onto a 577 pixel photomultiplier camera. In the inner part of the camera are mounted 397 hexagonal pixels with a diameter of 30 mm = 0.1 deg. 325 of these pixels are used to give the trigger. The FoV if the inner part of the camera is 2.1 - 2.3 deg. In the outer part of the camera are placed 180 hexagonal pixels with a diameter of 60 mm = 0.2 deg. The total FoV of the camera is 3.5 deg. Every PMT is equipped with a light guide on the top (Winston Cones). This is a hollow, hexagonal-shaped, non-imaging light concentrator to compensate the dead space between the round PMTs. The hemispherical entrance window of the PMTs is coated with a milky lacquer doped with a wavelength shifter. All this results in an increase of the effective quantum efficiency of the PMTs of around 20% compared to flat window PMTs. The photomultipliers were selected by their quantum efficiencies in the blue part of the spectrum in order to maximize their response to gamma ray showers of small energies. Temperature and humidity are monitored at three points in the camera and can be controlled using a water-based cooling system. In front of the light guides, the MAGIC camera is protected by a plexiglass with uniform transmission over all wavelengths down to 300 nm where it drops down.

![Diagram of Light Collector and PMT](image)
2.4.3 Star-guider

In order to monitor the tracking system online, a sensitive CCD-camera (0.0003 lx) of type Watec WAT902H has been installed in the center of the mirror dish. Is has a $4.6^\circ \times 4.6^\circ$ FoV and image the sky in the telescope pointing direction as well as part of the PMT camera. It delivers 25 frames per second which are getting integrated over 5 seconds to achieve a better signal-to-noise ratio [202]. Six reference points (red LEDs) on the camera frame indicate the position of camera while individual stars get recognized by a dedicated analysis software and compared to starfield catalogs. With this information, the real pointing position of the telescope can be retrieved.

2.4.4 Signal Processing from the Camera to the Counting House

The readout and the trigger electronics are external of the telescope camera. In this way the camera has less weight which leads to a better mobility of the telescope and a better stability of the camera position. The electronics are situated in a place (Counting house) with generous dimensions for easy maintance. The electrical signal of the PMTs is converted into an optical one. Then, along the distance of 162 m to the Counting house, the fast analog signals are transmitted over optical fibers, driven by Vertical Cavity Surface Emitting Laser Drivers (VCSELs). The use of optical fibers reduces cable weight and allows electrical decoupling, noise immunity as well as signal transmission with only weak attenuation and pulse deformation. In the Counting house, the optical signal is received and converted back into an electrical one with a fast Gallium Arsenide (GaAs) PiN photo diode. The signal is then split into two branches. One branch is again split into two, by which the limited dynamical range of the 300 MHz 8 - bit Flash Analog-to-Digital-Converter (FADC) is increased. So a time resolution of 3.3 ns is reached. This FADC system was replaced in February 2007 by a multiplex 2 GSample/s FADC system with 10 bit amplitude resolution and a 0.5 ns time resolution. One half of the split signal is amplified by a factor of ten (high gain) whereas the other half of the signal no (low gain). If the high gain exceeds a certain amplitude, the 50 ns delayed low gain signal is digitized as well, but before this, the whole signal has to be stretched with a time constant of 6 ns. Then the FADC continuously writes the digitized amplitude information into a ring buffer. The other branch leads the signal to the trigger. In case of a trigger, the digitization stops and the corresponding part of the ring buffer is
2.4. The MAGIC Telescope

written to disk.

2.4.5 Trigger

The trigger makes the decision if an event should be read out from the camera. The inner part of the MAGIC camera is segmented into 19 overlapping macro cells of 37 pixels each. As the images of the air showers are extended, it is not necessary to include all 577 pixels into the trigger. The MAGIC trigger system has three decision levels.

Figure 2.13: The 19 active trigger regions in the inner camera.

- Level-0 (Discriminator): If the analog signal of one camera pixel exceeds a preset threshold, which can remotely be controlled by the central control PC, a comparator gives an approximately 6 nsec long logic output signal.

- Level-1: gives a trigger if close packed 4NN (4 next neighbour) pixels are “firing”, e.g.
passing level-0 trigger, in a time window of 2-5 nsec. Also this trigger condition can be changed by the central control PC.

- Level-2: allows a further online discrimination based on the topology of an event that already passed level-1.

**Sum trigger**

A new trigger system was mounted in the MAGIC Telescope during 2007, and has hallowed MAGIC to lower the trigger threshold to only 25 GeV for some dedicated observations, like the detection of the Crab Pulsar [38].

### 2.4.6 Calibration system

In order to know the shower energy is important to reconstruct the exact number of the phonons which materially produced the triggered event. A relative calibration of the camera is done firing a special combination of LEDs onto the camera at a frequency of 50 Hz. With the calibration light it is possible to artificially flat-field the camera by multiplying each and every PMT signal for a corrective factor. The best-know method to do this is the F-factor method, which used an intrinsic PMT parameter. A second aim of the calibration system is to absolute calibrate the the camera, i.e. to find the conversion factor between FADC units and photoelectrons (or photons) hitting the PMTs. This is done with the already mentioned F-factor method with the use of filtered pixels which observe the single photoelectrons, and with the use of a PIN-diode which measures the direct light from the calibration pulser and confronts it with the signal from a radioactive source [108].

### 2.4.7 Data flow

The raw data are stored by the Data Acquisition system (DAQ). Each run-file is limited to 2 GB. All files are transferred via internet or they are taped and shipped from La Palma to the MAGIC data centers Port d’Informacio Cientifica (PIC) in Barcelona and to the data center in Wurzburg where they are calibrated afterwards.
2.4.8 New solutions in MAGIC II

The structure of the MAGIC II Telescope clone, namely the foundations, and the telescope chassis with the space frame structure, are basically a perfect repetition of MAGIC I. They were assembled and mounted in La Palma already in 2006-07. On the other hand, all the other subsystem were substantially improved. This overall enhancement should substantially increase MAGIC II performance, compared to MAGIC I, and cure some of its design weaknesses. Going over the subsystem, principals new solutions are:

- **Camera:** A major effort was made to improve the MAGIC camera photon detection system. First of all, on the market there were already photomultipliers with quite larger quantum efficiency compared to MAGIC I EMI-Thomson devices [167,168]. After a market selection Hamamatsu PMTs were chosen. Their photon conversion efficiency is 32% at 350 nm without lacquer coating. All the PMTs have the same size of 0.1°. The number of the PMTs has increased from 577 to 1039 in total. Also the triggers have been increased and now comprise 559 PMTs in the central camera. The design of the PMT housing has also been completely renewed. The PMTs are grouped into clusters of 7. Each cluster can be easily removed in case of problems. Inside the cluster, the HV is produced, a sampling signal can be injected to test electronics after the PMT. The PMT is read via Lemo cables and controlled with Ethernet communication. The new camera will bring an increased effective area, and an increased signal to noise ratio. For the future, it is also foreseen the replacement of the entire camera with an avalanche photon detector (HPD) Hamamatsu R9792U-40 [127,166,172]. This device is composed of a GaAsp photocathode, where the photon is converted into a photoelectron, which is subsequently accelerated directly by an intense voltage drop of 8 kV. The electron hits an avalanche photon detector (APD) where the signal is amplified with a total gain of around 5,000. The HPDs have a much higher quantum efficiency than PMT, better single-photon resolution and are almost free of after-pulses.

- **Electronics:** The signal from the PMTs is transmitted via optical cable to the counting house, where receiver boards convert the signal back into charge, as for MAGIC I. The signal, after some trigger logic, is digitized with Domino III chips, which sample the signal at 2 GHz. The L1 trigger board (topological trigger) was adapted from MAGIC
I for an increased trigger area. For the moment, the topological next-neighbors logic has been implemented, whereas different logic could be adapted in the future. The calibration is performed with an ultra-bright laser mounted on the main reflector and controlled remotely. A diffusing sphere is mounted in front of the laser to illuminate homogeneously the MAGIC II camera.

- Mirrors: The MAGIC II mirrors reflector is composed of two types of mirrors: 143 full-aluminium mirrors similar to MAGIC I mirrors but with a larger area of 1 m² and improved design and 104 cold-slumped glass-aluminium mirrors sandwich developed at INAF-Milano.
Gamma-Ray Bursts (GRBs) are among the most fascinating phenomena in the Universe. They are bright flashes of radiation with spectral energy distributions peaking in the $\gamma$-ray band. The “prompt” emission (peaked at a typical energy of the order of $\sim 100$ keV) lasts seconds and is likely to produce direct flows of relativistic matter with kinetic (isotropic) luminosities exceeding $10^{53}$ erg s$^{-1}$, making them the most luminous events known in the Universe since the Big Bang. It is followed by a fainter longer-lived “afterglow” (in X, radio, optical bands). All evidence points to a gravitational power source associated with the cataclysmic formation of a relativistic star remnant or to a precursor stage whose inevitable end point is a stellar mass BH.

While interesting on their own, GRBs are now becoming powerful tools to study detailed properties of their host galaxies and of the spacetime of the Universe. Their apparent association with massive star formation and their brilliant luminosities make them unique probes of the high-redshift universe and galaxy evolution. Absorption spectroscopy of GRB afterglows is being used to study the Inter Stellar Medium (ISM) in evolving galaxies. One of the most interesting use of GRBs in cosmology is as probes of the early phases of star and galaxy formation, and the resulting reionization of the Universe at $z \approx 6$-20.

GRBs are bright enough to be detectable (in principle) much farther out than the most luminous quasar or galaxies detected up to now. Promptly localized GRBs could provide information about earlier epochs in the history of the Universe than currently probed galaxies and quasars, shining through the pregalactic gas.

In this chapter I'll review current knowledge about GRBs: their phenomenology and the mainstream theoretical framework in which GRB emission is understood as an expanding
relativistic fireball \[162,194,214\], with the beamed radiation due to internal/external shocks (prompt/afterglow phase, respectively). Finally I’ll outline how, in the framework of the fireball model, GeV-TeV emission in the prompt and afterglow phase is predicted by several authors and how the detection of VHE emission from the GRB will allow us to discriminate among competing theoretical models.

3.1 History

GRBs were discovered in the late 1960s by the U.S. Vela satellites, which were built to detect $\gamma$ radiation pulses emitted by nuclear weapons tested in space. On July 2, 1967, at 14:19 UTC, the Vela 4 and Vela 3 satellites detected a flash of $\gamma$ radiation unlike any known nuclear weapons signature. The team at the Los Alamos Scientific Laboratory, led by Ray Klebesadel, filed the data away for investigation. An additional Vela satellites were launched with better instruments, the Los Alamos team continued to find inexplicable GRBs in their data.

By analyzing the different arrival times of the bursts as detected by different satellites, the team was able to determine rough estimates for the sky positions of sixteen bursts and definitively rule out a terrestrial or solar origin. The discovery was declassified and published in 1973 as an Astrophysical Journal article titled “Observations of Gamma-Ray Bursts of Cosmic Origin” \[140\].

Many theories were advanced to explain these bursts, most of which posited nearby sources within the Milky Way Galaxy. Little progress was made until the 1991 launch of the Compton Gamma Ray Observatory (CGRO) and its Burst and Transient Source Explorer (BATSE) instrument, an extremely sensitive $\gamma$-ray detector. fig:batsedistr BATSE measured more than 2700 events revealing that about two visible bursts occur somewhere in the Universe on a typical day. This instrument provided crucial data indicating that the distribution of GRBs in the sky is isotropic \[158\] (see Fig. 3.1.)

Because of the flattened shape of the Milky Way Galaxy, sources within our own galaxy would be strongly concentrated in or near the Galactic plane. The absence of any such pattern in the case of GRBs provided strong evidence that GRB must come from beyond the Milky Way \[185\]. However, some Milky Way models were still consistent with an isotropic distribution \[146\].

For decades after the discovery of GRBs, a counterpart was searched but any astronomical
3.1. History

Figure 3.1: Positions on the sky of all GRBs detected during the BATSE mission. The distribution is isotropic, with no concentration towards the plane of the Milky Way, which runs horizontally through the center of the image. (credit: CGRO BATSE team)

object in positional coincidence with a recently observed burst was found. Many distinct classes of objects, including white dwarfs, pulsars, supernovae, globular clusters, quasars, Seyfert galaxies, and BL Lac objects [134] were considered. All such searches were unsuccessful, [A notable exception is the 5 March event of 1979, an extremely bright burst that was successfully localized to supernova remnant N49 in the Large Magellanic Cloud. This event is now interpreted as a magnetar giant flare, more related to SGR flares than “true” GRBs] and in a few cases particularly well-localized bursts (those whose positions were determined with what was then a high degree of accuracy) could be clearly shown to have no bright objects of any nature consistent with the position derived from the detecting satellites.

This suggested an origin of either very faint stars or extremely distant galaxies [135]. Even the most accurate positions contained numerous faint stars and galaxies, and it was widely agreed that final resolution of the origins of cosmic GRB would require both new satellites and faster communication [93]. Several models for the origin of GRBs postulated [179] that the initial burst of $\gamma$-rays should be followed by slowly fading emission at longer wavelengths created by collisions between the burst ejecta and interstellar gas. Early searches for this “afterglow” were unsuccessful, largely due to the difficulties in observing a burst’s position at longer wavelengths immediately after the initial burst.
The smoking gun came in February 1997 when the satellite BeppoSAX detected tGRB 970228\(^1\) and when the X-ray camera was pointed towards the direction from which the burst had originated, it detected fading X-ray emission [63]. Ground-based telescopes later identified a fading optical counterpart as well [229]. Once the GRB faded, deep imaging was able to identify a faint, distant host galaxy at the location of the GRB as pinpointed by the optical afterglow. Because of the very faint luminosity of this galaxy, its exact distance was not measured for several years. Well before then, another major breakthrough occurred with the next event registered by BeppoSAX, GRB 970508. This event was localized within four hours of its discovery, allowing research teams to begin making observations much sooner than any previous burst. The spectrum of the object revealed a redshift of \(z = 0.835\), placing the burst at a distance of roughly 6 billion light years from Earth. This was the first accurate determination of the distance to a GRB, and together with the discovery of the host galaxy of 970228 proved that GRBs occur in extremely distant galaxies [54].

Within a few months, the controversy about the distance scale ended: GRBs were extragalactic events originating within faint galaxies at cosmological distances. The following year, GRB 980425 was followed by a bright supernova (SN 1998bw), indicating a clear connection between GRBs and the deaths of very massive stars [137]. This burst provided the first strong clue about the nature of the systems that produce GRBs. BeppoSAX functioned until 2002 and CGRO (with BATSE) was deorbited in 2000. However, the revolution in the study of GRBs motivated the development of a number of additional instruments designed specifically to explore the nature of GRBs, especially in the earliest moments following the explosion.

The first such mission, HETE-2 [201], launched in 2000 and operating until 2006, providing most of the major discoveries during this period. One of the most successful space missions to date, Swift, was launched in 2004 and is currently still operational [110]. Swift is equipped with a very sensitive gamma ray detector as well as on-board X-ray and optical telescopes, which can be rapidly and automatically slewed to observe afterglow emission following a burst. More recently, the \textit{Fermi} mission was launched carrying the GBM, which detects bursts at a rate of several hundred per year, some of which are bright enough to be observed at extremely

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\(^1\)The naming convention of a GRB follows the rule: GRB YYMMDD. E.g. GRB970228 has to be read as: The GRB occurring on the 28\(^{th}\) of February, 1997.
3.2 Phenomenology

While most astronomical transient sources have simple and consistent time structures (typically a rapid brightening followed by gradual fading, as in a nova or supernova), the light curves of GRBs are extremely diverse and complex. No two GRB light curves are identical, important differences being observed in almost every property: the duration of observable emission can vary from milliseconds to tens of minutes, there can be a single peak or several individual subpulses, and individual peaks can be either symmetric or with a fast brightening and a very slow fading. Some bursts are preceded by a “precursor” event, a weak burst that is then followed (after seconds to minutes of no emission at all) by the much more intense “true” bursting episode. The light curves of some events have extremely chaotic and complicated profiles with almost no discernible patterns. Although some light curves can be roughly reproduced using certain simplified models, little progress has been made in understanding the full diversity observed.

Many classification schemes have been proposed, but these are often based solely on differences in the appearance of light curves and may not always reflect a true physical difference in the progenitors of the explosions. Fig. 3.2 shows a couple of typical light curves taken in the 100 keV energy regime by BATSE. One can see immediately that GRBs can exhibit complicated light curves and time structures $\delta T$ much smaller than $T$, their total duration. The burst morphology can be subdivided into four groups [95]: single-pulsed events, smooth events with well-defined peaks, bursts with distinct, well-separated episodes of emission and

high energies with Fermi’s Large Area Telescope.

Meanwhile, on the ground, numerous telescopes have been built or modified to incorporate robotic control software that responds immediately to signals sent through the GRB Coordinates Network (GCN). This allows the telescopes to rapidly re-point towards a GRB, often within seconds of receiving the signal and while the gamma-ray emission itself is still ongoing. New developments over the past few years include the recognition of short GRBs as a separate class (likely due to merging neutron stars and not associated with supernovae), the discovery of extended, erratic flaring activity at X-ray wavelengths lasting for many minutes after most GRBs, and the discovery of the most luminous (GRB 080319B) [55] and the most distant (GRB 090423) [210] objects in the universe.
Figure 3.2: Four typical light curves of GRBs as detected by BATSE. Top left: burst with distinct, well-separated episodes of emission, top right: chaotic, spiky burst, bottom left: smooth event, multi-peaked, bottom right: single-pulsed (here double-pulsed). All times in seconds.
finally erratic, chaotic and spiky bursts. The shortest rise times recorded are roughly equal
to the shortest structures within time histories, namely \( \delta T \approx 0.2 \, \text{ms} \). This number constrains severely the size and the speed of the emitting region, since \([118, 178]\):

\[
R_{\text{emission}} < \Gamma^2 \cdot c \cdot \delta T_{\text{min}} \approx \Gamma^2 \cdot 60 \text{ km}
\]

where \( R_{\text{emission}} \) is the path along which the emission takes place, \( \Gamma \) is the Lorentz factor of the emitting region moving towards the observer and \( c \) the speed of light. Observed GRB durations span at least five orders of magnitude (see Fig. 3.4) following a bimodal \([143]\) distribution with peaks at 0.5 and 34 seconds \([94]\). GRBs can be divided into those with durations smaller than and those greater than 2 seconds: the so-called short-bursts and long-bursts.

Short bursts have a tendency to display harder spectra \([143]\) and to have only a few major pulse structures. The prompt emission spectra of GRBs are simple (no absorption or emission lines) and range out to GeV energies \([136]\). In general, both the time-integrated spectra and
Figure 3.4: Distribution of burst durations obtained with the BATSE detector

Figure 3.5: Distribution of the low-energy spectral indices $\alpha$ (left) and high-energy spectral indices $\beta$ (right), obtained from fits of GRB spectra from the 4th BATSE catalog to the "Band function" (eq. 3.2). On the right side, all bursts which could not be fitted beyond the break energy or with fit results $\beta < -4$ are included in the first bin. Figure from [188].
spectra of shorter intervals within a burst, can be well fit with a smooth double power law with break energy \( E_b \) (the so-called “Band-spectrum” [46]):

\[
N(E) = N_0 \cdot \begin{cases} 
E^\alpha \exp(-E/E_b) & \text{for } E < (\alpha - \beta)E_b \\
((\alpha - \beta)E_b)^{\alpha-\beta} E^\beta \exp(\beta - \alpha) & \text{for } E > (\alpha - \beta)E_b
\end{cases}
\] (3.2)

In this model, the peak energy in the \( \nu F_\nu \) representation relates to the break energy like: \( E_{\text{peak}} = (\alpha + 2)E_{\text{break}} \). Most bursts have a high-energy spectral index close to \( \beta \approx -2.25 \) (see figure 3.5). Spectral softening is usually observed throughout a GRB and in sub-pulses within a burst (see figure 3.3 and [189]). On the low-energy side, the spectral indices \( \alpha \) are most often centered around \( \alpha \approx -1 \). The break energies \( E_b \) of BATSE bursts are distributed very closely around 200 keV.

Afterglows at longer wavelengths have been observed since the launch of Bepposax in 1996. As the afterglow light curves decline with time faster than \( t^{-1} \), they do not contribute significantly to the total energy budget of the emission [76, 101]. Nevertheless, the XRT detector on Swift discovered bright X-ray flares [85] in the first two hours after the onset of burst, following by several less energetic flares in the same energy range up to two days after the burst [71]. Swift discovered flaring activity appearing well after the prompt phase in the 50% of X-ray afterglows [60, 84] (for more details see section 3.5.)

Observations of breaks in the afterglow light curves suggest that the \( \gamma \)-ray emission is beamed into solid angles of \( \Omega \approx 0.1 \text{ sr} \) [144], where the break happens at the transition from a regime dominated by relativistic beaming to the one dominated by the intrinsic beaming of the jet itself. With the measured time at which a jet break occurs and general assumptions about the circumburst medium density, the beaming angle \( \theta \) can be derived [213]. The ensuing collimation-corrected bolometric energy of the bursts \( E_\gamma = (1 - \cos \theta) \cdot E_{\gamma,iso} \), correlated with the peak energy in the GRB rest frame: \( E_{\text{peak}}^{\text{obs}} (1+z) \propto E_\gamma^{0.7} \). This relation will allow to GRBs to be used as standard candles at cosmological distances [111].

### 3.3 GRB modeling

#### 3.3.1 Progenitors

Because of the large distances of most GRBs, identification of the progenitors, the systems that produce these explosions, is particularly challenging. All hypotheses presented here have
Figure 3.6:
Example 0.2-10 keV band X-ray flare light curve, observed by XRT on SWIFT [85].
Figure 3.7: Ghirlanda relation: In black: $E_{\gamma,iso}$ without beaming correction; red and green: collimation-corrected bolometric energy ($E_\gamma$); blue arrows: upper/lower limits to $E_\gamma$ due to upper/lower limits on their jet break time. Figure from [111], update from [112].
in common that a GRB comes together with the creation of a stellar-sized BH, surrounded by an accretion disk. With an almost maximally rotating BH, the geometry of the system has two favored directions, identified with the rotation axis of the BH, along which emission jets form.

The most widely accepted mechanism for the origin of long-duration GRBs is the collapsar model [154, 242, 243], in which the core of an extremely massive, low-metallicity, rapidly-rotating star collapses into a BH in the final stages of its evolution. Matter near the star’s core rains down towards the center and swirls into a high-density accretion disk. The infall of this material into a black hole drives a pair of relativistic jets out along the rotational axis, which funnel through the stellar envelope and eventually break through the stellar surface and radiate as $\gamma$-rays. Alternatively some models replace the BH with a newly-formed magnetar [165], although most other aspects of the model (the collapse of the core of a massive star and the formation of relativistic jets) are the same.

A variant of the collapsar model (Vietri Stella in 1998 [230]): the supranova model predicts that a SN explosion initially results in the formation of a comparatively massive, very fast rotating NS. This NS is supramassive but for its fast rotation, it would immediately collapse into a BH. Subsequently, the NS loses rotational momentum through a pulsar wind and gravitational waves until it finally collapses to a BH and triggers the GRB.

A first hint to a close connection between GRBs and SN came with the observation of GRB 980425 [102], the closest burst observed so far ($z = 0.0085$), spatially coincident with the SN 1998bw. However, GRB980425 was a very unusual burst, showing a total energy release of orders of magnitude lower than typical GRBs at higher redshifts. With GRB 030329, a very close and bright burst could be intensely followed by telescopes at all wavelengths. After $\sim 10$ days since the burst, the optical spectra clearly showed features of a SN [132], thus confirming the association for at least some long bursts [160]. The closest Galactic analogs of the stars producing long GRBs are likely the Wolf-Rayet stars, which are extremely hot and massive stars which have shed most or all of their outer hydrogen layers due to radiation pressure.

The massive star model probably does not explain all types of GRB. There is strong evidence that some short GRB occur in systems with no ongoing star formation, such as galaxy halos and intergalactic space [190].
The favored picture for the origin of most short gamma-ray bursts is the NS-merger scenario (Paczyński Goodman [118,170,178,197]) which assumes two NSs or a NS + BH or two BHs orbiting around each another. Due to loss of orbital momentum through the emission of gravitational waves, the two compact remnant spiral together and coalesce. At that moment, a black-hole is formed and a GRB occur (see Fig. 3.8). GRBs in this models should therefore occur mainly outside galaxy cores or star forming regions: the shedding of a big part of the star’s envelope from a point displaced from the center of gravity of the binary system, must transfer momentum to the binary system containing the neutron star remainder, the so-called “momentum kick”. Due to the fact that these bursts occur in much less dense regions of a galaxy, their optical afterglows are predicted to be about a factor 1000 dimmer than the ones obtained from supernova models. These predictions could be checked on July 9th, 2005.

Figure 3.8: Sketch of NS-NS merger. Figure from [133].

\footnote{Neutron star or black hole binary systems have life times of typically $10^9$ yr and travel 1-100 kpc before merging [232].}
when the Hete-II satellite quickly provided the position of a short burst [231], and follow-up observations provided data on its optical afterglow, the first time for a short burst [133].

Several other models have been proposed for short gamma-ray bursts, including the accretion-induced collapse of a neutron star [53] or the evaporation of primordial black holes [61].

### 3.3.2 Fireball Model

Matter ejection at relativistic speed is a necessary requirement for producing GRB emission. GRBs vary on such short timescales that the size of the emitting region must be very small, or else the time delay due to the finite speed of light would “smear” the emission out in time, wiping out any short-timescale behavior. At the energies involved in a typical GRB, so much energy stored into such a small space would make the system opaque to $\gamma$-rays by photon-photon pair production, making the burst far less luminous and also giving it a very different spectrum from what is observed.

However, if the emitting system is moving towards Earth at relativistic velocities, the burst is compressed in time (as seen by an Earth observer, due to the relativistic Doppler effect) and the emitting region inferred from the finite speed of light becomes much smaller than the true size of the GRB. A related constraint is imposed by the relative timescales seen in some bursts between the short-timescale variability and the total length of the GRB. Often this variability timescale is far shorter than the total burst length. For example, in bursts as long as 100 seconds, the majority of the energy can be released in short episodes less than 1 second long. If the GRB were due to matter moving towards Earth (as the relativistic motion argument enforces), it is hard to understand why it would release its energy in such brief interludes.

The generally accepted explanation for this is that these bursts involve the collision of multiple shells traveling at slightly different velocities; so-called “internal shocks”. The collision of two thin shells flash-heats the matter, converting kinetic energy into random motion of particles, greatly amplifying the energy release.

Which physical mechanisms are at play in producing the observed photons is still an being debated, but the most likely candidates appear to be synchrotron radiation and inverse Compton scattering. The Fireball Model (Mészáros & Rees [198, 199], Narayan & Piran [159, 170]) is based on the relativistic generalization of SNR models. It assumes that a central object
produces a highly variable ultra-relativistic ($\Gamma \gg 100$) outflow of an optically thick plasma shell (the fireball), containing mostly electrons, positrons and $\gamma$-rays.

This energy is transported via bulk motion out to $\sim 10^{13} - 10^{15}$ cm before the plasma becomes optically thin and radiates the GRB. The baryonic mass of the outflow needs to be below $10^{-4}$ solar masses to allow these highly relativistic expansion speeds.

In order to reconvert the kinetic energy efficiently into radiation, relativistic shocks ($\Gamma > 100$) are required [47,197], accelerating the electrons-positrons at very short time scales, compared to the shock expansion timescale [76]. Large-scale turbulences distribute the energy dissipated in the shock over the shocked gas.

Possibly, various such outflows can collide with each other – producing relativistic “internal-shocks” [199] at about $10^{14} - 10^{16}$ cm from the triggering event – and emit the prompt $\gamma$-rays. The observed radiation is then a convolution of contributions from regions of the blast wave emitting at different times and moving with different Lorentz factors and directions with respect to the observer.

The internal shocks were shown to be able to produce the complex time-profiles of some bursts [50,68]. At a later stage, the shell impacts on the circumburst-medium, a relativistic “external-shock” [198] is produced which radiates the afterglow and, depending on the density of the circumburst medium, the smooth GRB time profiles. In both cases, a fraction of electrons is accelerated by repeated diffusion across the shock front, undergoing first order Fermi acceleration. The electrons spiral in the turbulent magnetic fields and radiate by synchrotron mechanism. (Given the observational constraints, a fireball is always in the Klein-Nishina-regime of IC scattering [211]).

At the same time, the electrons cool by the their own synchrotron emission. If the primary electron Lorentz-factors are distributed like $f(\gamma) \propto \gamma^{-p}$ in an interval of ($\gamma_{\text{min}} < \gamma < \gamma_{\text{max}}$), then the resulting photon spectra consist of three parts: $N(E) \propto E^{-2/3}$ before $E = \gamma_{\text{min}}^2 \cdot h\nu_B$ ($\nu_B = eB_\perp/(2\pi m_e c)$ is the gyrofrequency), then $N(E) \propto E^{-(p+1)/2}$ where cooling effects are not yet important and last $N(E) \propto E^{-(p+2)/2}$ after the so-called “cooling break”. The hardest possible spectral index of this model if therefore $\alpha \approx -1.5$, in contradiction to some of the spectral fit results, shown in figure 3.5 (see however [148] and [80] for possible solutions to this problem).

Because the fireball expands relativistically, the observed radiation is blue-shifted. Due to
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Figure 3.9: A sketch of the fireball model: a central object (here a black hole created from a collapsing massive star) produces a highly relativistic and variable jet of electron-positron plasma. Different shells of the outflow collide with each other and form internal shocks. Further away, an external shock is created when the outflow runs into the interstellar medium. At a distance of $10^{12}$ cm the plasma becomes optically thin to $\gamma$-rays and the GRB is radiated. Subsequently, the plasma loses opacity at lower wavelengths and radiates the afterglow. In the case of an accompanying supernova explosion, an additional SNR has to be taken into account.
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Figure 3.10: Sketch of the various kind of shocks.

relativistic beaming, only a small fraction of the expanding shock is visible at the beginning and thus the time profile of the arriving gamma rays is not smoothed out by the simultaneous emission from different points in the expanding shell [182, 211]. As long as the jet has an opening angle of $\theta \gg \Gamma^{-1}$, the observed geometry is determined by relativistic beaming. As $\Gamma$ decreases, the beaming angle opens up and the expansion is seen as if it were isotropic. Once these effects appear the jet fades very rapidly, an effect that is visible as a power-law “break” in the afterglow light curve. This is the so-called “jet break” that has been seen in some events and is often cited as evidence for the consensus view of GRBs as jets. The initial interaction of the fireball ejecta with the surrounding gas produces also a “reverse-shock”. The twice-shocked material can produce a bright optical/UV flash, which has been seen in a few GRBs [23]. There can be “refreshed-shocks” if the central engine continues to release fast-moving matter in small amounts even out to late times, these new shocks will catch up with the external shock to produce something like a late-time internal shock. This explanation has been invoked to explain the frequent flares seen in X-rays and at other wavelengths in many bursts, though some theorists are uncomfortable with the apparent demand that the progenitor continues to remain active for very long.

The strength of the fireball model consisted in predicting very exactly the behaviour of the afterglow before these could be verified experimentally, assuming only synchrotron emission (yielding a power-law spectrum with three breaks [161]). Later, the optical signature of
3. Gamma-Ray Bursts

A reverse-shock predicted in 1999 [212], was verified one month later with the ROTSE telescope [24]. The currently most used formulation of the – nevertheless complex – fireball model can be found in [103, 240]. Although the fireball model has become widespread and largely accepted, there is still no agreement about its details [76]. Currently, additional thermal components of the spectrum are discussed (see e.g. models by Felix Ryde [209]), dominating the emission at the beginning of the GRB.

3.3.3 Other Models

The Cannonball Model (CB) (DeRújula & Dar [66, 67, 69, 81]) assumes first a core-collapse supernova and the creation of a neutron star. About ten days later, the neutron star collapses and bipolar jets of hydrogen plasma clouds – the so-called cannonballs (CBs) – are ejected. These CBs have Lorentz factors of the order to $\Gamma \sim 10^3$ and consist of baryons and electrons in equal parts. A CB is emitted, when a part of the accretion disk falls abruptly onto a compact central object (as observed e.g. in microquasars).

When the CB crosses the SNR with its large Lorentz factor, its surface becomes heated to keV temperatures due to collisions with the shell particles. It emits thermal bremsstrahlung radiation as it reaches the transparent outskirts of the shells. This radiation, in turn, is boosted and collimated by the CB’s motion and seen like a single pulse in a GRB. Due to relativistic beaming, GRBs are observable only if the angle $\theta$ subtended by the CB velocity vector and the line of sight to the observer is small: $\theta \sim (1/\Gamma_0)$.

In the Compton Drag Model [113] a massive star collapses and creates a fireball which traverses a funnel inside the (not yet blown away) massive star envelope. The fireball IC scatters soft ambient photons radiated by the walls of the funnel (the so-called Compton Drag) and produces thus the GRB.

Another model (Stern & Poutanen [221]) suggests that an SSC mechanism at rather low shell velocities ($\Gamma \approx 30 - 100$) can reproduce the gamma ray and X-ray spectra if the electrons get continuously re-accelerated and heated. In contradiction to the fireball model, particle acceleration is then not dominated by Fermi acceleration, but by other mechanisms which are not further discussed here: Plasma instabilities behind the shock front [98] or magnetic field energy dissipation [152, 228]. The strength of this model consists in explaining very

\footnote{\textsuperscript{3}However, calculated within the framework of a $\Lambda = 0$ universe.}
Figure 3.11: Sketch of the CB model: a core-collapse SN results in a compact object and a fast-rotating accretion disk of non-ejected fallen-back material. Matter accreting into the central object produces a narrowly collimated beam of CBs. As these CBs pierce the SNR, they heat up and re-emit photons. They also scatter light from the shell. Both emissions are Lorentz-boosted and collimated by the CBs’ ultra-relativistic motion. Figure from [69].
naturally the hard-to-soft evolution of spectra, as shown in figure 3.3, and the possibility to reach rather hard values of \( \alpha > -1.5 \), obtained from Band-function fits (equation 3.2) of some GRB spectra.

Finally in the Poynting flux model (Usov [153, 228]) a highly magnetized, fastly-rotating NS, created in a collapsar, develops a strongly magnetized wind that flows with a high Lorentz factors. Magnetic reconnection can then transfer magnetic energy to electron-positron pairs in the wind. Later, large-scale electro-magnetic waves can efficiently accelerate parts of the electrons and positrons. The resulting spectra will be again a combination of thermal spectra with non-thermal synchrotron emission [209].

### 3.4 HE and VHE emission from GRBs

In the past, several attempts were made to observe GRBs in the MeV-GeV energy range, The EGRET detector could observe seven GRBs emitting gamma rays with 100 MeV < \( E < 18 \) GeV [136, 217]. Figure 3.12 shows the averaged spectrum of the four brightest bursts, observed by EGRET: GRB 910503, GRB 930131, GRB 940217 and GRB 940301. The continuation of the power law is visible up to the GeV-range, without any apparent cut-off.

Especially one burst, GRB940217 [136], showed an event at \( E > 18 \) GeV 90 minutes af-

![Figure 3.12: Averaged spectra of the four brightest GRBs observed by EGRET: GRB 910503, GRB 930131, GRB 940217 and GRB 940301. Figure from [83].](image)
3.4. HE and VHE emission from GRBs

Figure 3.13: The light curve of GRB940217, as observed by the ULYSSES detector (black lines) and EGRET (red points). In the last case, each point corresponds to one single gamma ray event and the vertical scale is the photon energy (right scale in red). During the time period marked with “earth occultation”, the burst was not visible for EGRET. The probability to observe the 10 events after earth occultation from background fluctuations is $6 \cdot 10^{-4}$, the probability to detect one single event with energy $E > 18$ GeV from background is $5 \cdot 10^{-6}$. Figure and numbers from [136].
ter the prompt low-energy emission of the burst (see figure 3.13). Results from the TASC shower counter, part of the EGRET detector, jointly fit with BATSE data, indicate that the spectrum of GRB 941017 contain a high energy emission component in the addition to the Band function [117]. The resolved spectral component, presumably due to ultra-relativistic hadrons, has differential photon flux spectral index of $\alpha = -1$ with no cut-off up to the TASC detector energy limit at 200 MeV.

More recently, MeV-GeV detections of GRBs have been also obtained by AGILE and Fermi. AGILE Gamma-Ray Imaging Detector (GRID), sensitive in the 30 MeV-30 GeV range, firmly detected three GRBs (GRB 080514B, GRB 090401B and GRB 090510) [155]. GRB 090510 provides the first case of a short burst with delayed $\gamma$-ray emission [115]. Fermi-LAT as of october 2009 detected already the emission from 9 GRBs and HE emission at GeV and a long-lived HE emission were observed in both long and short LAT GRBs. For a brief summary of Fermi GRB physics result I refer to section 3.5.

On the contrary, in spite of continuous attempts, no convincing detections at higher energies (TeV range) have so far been obtained. The Milagrito experiment claimed to have observed emission at about 0.1 TeV for GRB970417A [43] at a $\sim 3\sigma$ significance; however, Milagrito’s successor, Milagro, did not detect any significant signal in more than 50 GRBs observed. Null detections so fare have also been reported by various Imaging Atmospheric Cerenkov Telescopes (IACTs) as HESS, VERITAS and in particular MAGIC which is characterized by the lowest energy threshold which has been brought down to 25 GeV under favourable conditions in the 2008 upgrade and is now becoming more sensitive with the advent of the MAGIC-II telescope.

### 3.4.1 Theoretical expectations: emission models

It is very likely that particles are accelerated to very high energies close to the emission site in GRBs. This could either be in shock fronts, where the Fermi I mechanism or other plasma instabilities can act, or in magnetic reconnection sites. In the general fireball model scenario, possible HE and VHE emission mechanisms are predicted for both leptonic and hadronic models [45]. In the leptonic models, synchrotron emission by relativistic electrons can explain the 10 keV -1 MeV spectrum in $\sim 2/3$ of bursts (e.g., see Preece et al. 1998 [189]), and inverse Compton (IC) scattering of low energy seed photons generally results in GeV band
3.4. HE and VHE emission from GRBs

Emission. These processes could operate in both internal and external shock regions (see, e.g., Zhang & Mészáros 2001 [249]), with the relativistic electrons in one region scattering the “soft” photons from another region (Fragile et al. 2004 [96]; Fan et al. 2005 [89]; Mészáros et al. 1994 [164]; Waxman 1997 [236]; Panaitescu et al. 1998 [182]). Correlated HE and LE emission is expected if the same electrons radiate synchrotron photons and IC scatter soft photons. In Synchrotron Self-Compton (SSC) models the electrons’ synchrotron photons are the soft photons and thus the HE and LE components should have correlated variability (Guetta & Granot 2003 [120]; Galli & Guetta 2008 [105]).

Alternatively, photospheric thermal emission might dominate the soft keV - MeV range during the early part of the prompt phase (Rees & Mészáros 2005 [195]; Ryde 2004 [209]). Such a component is expected when the outflow becomes optically thin, and would explain low energy spectra that are too hard for conventional synchrotron models (Crider et al. 1997 [64]; Preece et al. 1998 [189], 2002 [187]). An additional power law component might underlie this thermal component and extend to high energy; this component might be synchrotron emission or IC scattering of the thermal photons by relativistic electrons.

In a hadronic environment relativistic protons scatter inelastically off the ~100 keV burst photons producing (among other possible products) neutral pions, that decay into γ-rays. Similarly, if neutrons in the outflow decouple from protons, inelastic collisions between neutrons and protons can produce pions and subsequent high energy emission (Derishev et al. 2000 [75]; Bahcall & Mészáros 2000 [44]). High energy neutrinos that may be observable are also emitted in these interactions (Waxman & Bahcall 1997 [237]).

Many variants of hadronic cascade models have been proposed: high energy emission from proton-neutron inelastic collisions early in the evolution of the fireball (Bahcall & Mészáros 2000 [44]); proton-synchrotron and photo-meson cascade emission in internal shocks (e.g., Totani 1998 [226]; Zhang & Meszaros 2001 [249]; Fragile et al. 2004 [96]; Gupta & Zhang 2007 [125]); and proton synchrotron emission in external shocks (Bottcher & Dermer 1998 [56]).

A hadronic model was invoked to explain the additional spectral component observed in GRB 941017 (Dermer & Atoyan 2004 [79]). The emission in these models is predicted to peak in the MeV to GeV band (Bottcher & Dermer 1998 [56]; Gupta & Zhang 2007 [125]), and thus would produce a clear signal in the LAT energy band. However, photon-meson interactions would result from a radiatively inefficient fireball (Gupta & Zhang 2007 [125]), which is in contrast
with the high radiative efficiency that is suggested by Swift observations [121, 174]. Thus, the hadronic mechanisms for γ-ray production are many, but the \textit{Fermi} measurements of the temporal evolution of the highest energy photons will provide strong constraints on these models, and moreover discern the existence or otherwise of distinct GeV-band components.

### 3.4.2 Delayed GeV Emission

The observations of GRB 940217 [136] demonstrated the existence of GeV-band emission long after the 100 keV “prompt” phase in at least some bursts. With the multiplicity of shock fronts and with synchrotron and IC components emitted at each front, many models for this lingering high energy emission are possible. In combination with the prompt emission observations and afterglow observations by \textit{Swift}, ground-based telescopes, and \textit{Fermi} LAT observations may detect spectral and temporal signatures to distinguish between the different models.

These models include [45]: Synchrotron Self-Compton (SSC) emission in late internal shocks (LIS) (Zhang & Mészáros 2002 [250]; Wang et al. 2006 [235]; Fan et al. 2008 [88]; Galli & Guetta 2008 [105]); external IC (EIC) scattering of LIS photons by the forward shock electrons that radiate the afterglow (Wang et al. 2006 [235]); IC emission in the external reverse shock (RS) (Wang et al. 2001 [234]; Granot & Guetta 2003 [120]; Kobayashi et al. 2007 [142]); SSC emission in forward external shocks (Mészáros & Rees 1994 [163]; Dermer et al. 2000 [78]; Zhang & Mészáros 2001 [249]; Dermer 2007 [77]; Galli & Piro 2007 [106]).

A high energy IC component may be delayed and have broader time structures relative to lower energy components because the scattering may occur in a different region from where the soft photons are emitted (Wang et al. 2006 [235]). The correlation of GeV emission with X-ray afterglow flares observed by \textit{Swift} would be a diagnostic for different models (Wang et al. 2006 [235]; Galli & Piro 2007 [106]; Galli & Guetta 2008 [105]).

### 3.4.3 HE and VHE predictions from other models

The CB model predicts narrow GeV emission flares from pion decay, each associated with one of the CBs and arriving about 1 second earlier than the GRB emission. The Compton drag model predicts no GeV γ radiation at all because the plasma is so dense that it opaque to the energetic photons (getting self-absorbed via \(γ + γ \rightarrow e^+ + e^-\)) when the GRB is radiated.
Similarly, the Stern and Poutanen model predicts no measurable $\gamma$-ray emission above about 20 GeV.

### Universe transparency at HE and VHE

![Cumulative fraction of the GRBs as a function of redshift for 44 pre-swift bursts (blue) and 108 Swift bursts (red). Overplotted is a simple model for the expectations of the redshift distribution of GRBs (see Fig. 3.14).](image)

It’s important to remember that the universe is not transparent to VHE photons and that unless the event is relatively nearby ($z < 1$) and/or the emission suitably strong, the EBL attenuation could prevent the detection of GRB emission at VHE. In figure 3.14 is reported the distribution of the redshift of the swift burst updated to november 2009. The mean value is $z = 2.25$ but bursts at low redshift are also presents and a possible detection from IACTs instruments, if VHE emissions are at play, is possible.

### 3.5 GRBs in Swift and *Fermi* eras: results and open questions

**Swift era**:

Swift was designed to investigate the GRB afterglows by filling the temporal gap between observations of the prompt emission and the afterglow. The combined power of BAT and XRT
has revealed that prompt X-ray emission smoothly transitions into the decaying afterglow. One of the most striking result is that many of the early X-ray afterglows show a canonical behaviour: the light curve broadly consist of three distinct power law segments (Fig. 3.15).

A bright rapidly falling $(t^{-\alpha}$ where $\alpha > 3)$ afterglow immediately after the prompt emission [223] is followed by a steep-to-shallow transition, which is usually accompanied by a change in the spectral power-law index. This is consistent with an interpretation [175, 248] in which the first break occurs when the slowly decaying emission from the forward shock becomes dominant over the steeply decaying tail emission of the prompt $\gamma$-rays as seen form large angles [145]. The presence of the shallow decay phase implies that most of the energy in the afterglow shock was either injected at late times after the prompt $\gamma$-ray emission was over or was originally in slow material that would not have contributed to the prompt $\gamma$-ray emission.

Most of Swift X-ray light curves are broadly consistent with this basic temporal description, although in most cases we don’t see all three power-law segments, either because not all are present or because of limited temporal coverage. Swift has also discovered the flaring behaviour appearing well after the prompt phase in 50% of X-ray afterglows. the rapid rise and decay, multiple flares in the same burst, and cases of fluences comparable to the prompt emission suggest that the flares are due to the same mechanism responsible for the prompt emission, which is usually attributed to the activity of the central powerhouse.

Swift also detected the first GRB confirmed to be bright enough to be seen with the naked eye, GRB 080319B at redshift $z = 0.937$ [55]. Figure 3.16 shows the comparison between GBR 080319B, quasars, and one of the most energetic supernova recorded, SN 2006gy [216]. While evolutionary effects in all three populations are surely to be important at some level, in the context of probing the high-redshift universe, the overall impression is clear: for 30 minutes in the rest frame (what would be $\sim 4$ hr in the observer’s frame at $z = 7$), GRB 080319B would have been brighter than the brightest known QSO in the universe. Swift discovered also the most distant known object in the universe, GRB 090423 at a reshift $z = 8.1$ [210]. The properties of this burst, similar to those of GRBs observed at loser redshifts, suggest that the meachanism and progenitors that give rise to this burst are not strongly different from those producing GRBs about $10,000,000,000$ years later.
Figure 3.15: A synthetic cartoon of the X-ray lightcurve based on the observational data from the Swift XRT. The phase “0” denotes the prompt emission. Four power law lightcurve segments together with a flaring component are identified in the afterglow phase. The segments I and III are most common, and they are marked with solid lines. Other three components are only observed in a fraction of bursts, so they are marked as dashed lines. Typical temporal indices in the four segments are indicated in the figure. The spectral indices remain unchanged for segments II, III and IV, with a typical value of $\beta_X \sim 1$ ($F_X \propto \nu^{-\beta_X}$). The segment I sometimes has a softer spectrum (e.g. $\beta_X \sim 1.5$), but in some other cases it has a similar spectral index as the other three segments. The flares (segment V) have similar spectra as the segment I, and time evolution of the spectral index during the flares has been observed in some bursts (e.g. GRB 050502B).
Figure 3.16: Rest frame comparison of the most luminous optical-IR probes of the distant universe, showing the absolute magnitude versus time of GRB 080319B (red circles) and SN 2006gy (blue triangles). Transformed light curves of GRB 990123 (yellow stars) and GRB 050904 (green squares) are also shown. For reference, the most luminous known QSO is shown with a dashed horizontal line; the distribution of SDSS QSO magnitudes is shown as horizontal banding. Figure from [55]
3.5. GRBs in Swift and Fermi eras: results and open questions

**Fermi era: have we seen what we expected to see?**

The detection of the long extremely energetic GRB 080916C [13], over a broad energy range spanning about 7 decades of γ-ray energy opened the possibility to investigate the onset delay between LE and HE emission. GRB 080916C shows a single spectral form from 8 keV to 13.2 GeV consistent with a single Band function, suggestive with a single dominant emission mechanism in the observed energy range.

The redshift of this burst, $z = 4.35$ means that this burst has the largest reported apparent isotropic γ-ray energy release: $E_{\text{ISO}} \approx 8.8 \times 10^{54}$ ergs. It also sets a stringent lower limit on the GRB outflow Lorentz factor, $\Gamma_{\text{min}} \approx 890$, and limits on the quantum gravity mass scale, $M_{\text{QG}} > 1.3 \times 10^{18}$ GeV/c$^2$.

Delayed HE emission were detected in both classes of long and short bursts. In the case of GRB 080825C the late time HE emission (the highest energy photon was detected after 28 s when the low energy emission was already down to a low level) has harder photon spectrum than the earlier HE emission, which is consistent with an external shock origin. Let’s underline that the converse is true for GRB 080916C, that (with the highest photon was detected after 17 s, while the LE emission was still very active) favours a somewhat different model such an external compton scattering of late time X-ray flare photons by forward shock electrons.

GRB 090902B is one of the brightest burst detected by the GBM and LAT instruments [12]. It clearly shows LE and HE excess emission during the prompt phase, requiring a hard power-law component in addition to the usual Band function in order to fit the data. The origin of this component is not understood, and its presence in this burst poses real challenges to the theoretical models.

Like the two others bright bursts detected by the LAT, GRB 080916C and GRB 090510, GRB 090902B appears to possess a very high Lorentz factor for the bulk outflow, about 1000, and has some suggestion of the delayed onset of the emission above 100 MeV. The 33.4 GeV photon, the highest energy yet detected from GRB, and the $z = 1.822$ redshift of this burst have allowed to place significant constraints on some EBL models.

This puzzling results from Fermi are a challenge for theoreticians, that are searching to discriminate between a large variety of models. To clarify this recently opened challenge I conclude this section referring to the interpretations of the data on 4 LAT bursts, given in [65]. Dado and Dar, starting from the fact that in four LAT GRBs (GRB 080916C, GRB 0909002B,
GRB 090510, GRB090510) the arrival times of the HE photons did not coincide with the times of the brightest X-rays’ and MeV peaks, and the HE emission lasted longer than the prompt KeV–MeV emission, explain how this could be a problem for standard fireball model, that predict simultaneous emission at all energy bands.

In the CB model, indeed, this seems a natural consequence of the model: the dominant leptonic and hadronic emission mechanism are IC of synchrotron radiation of Fermi accelerated electrons (in the collision of the jet of highly relativistic CBs with the wind/ejecta blown form the progenitor or copanion star long before the GRB) and the neutral pion decay produced in hadrons collisions between the CB nuclei and the nuclei of the hadronic environment the jet passes into. The main predictions of the model for the early HE emission (for a detailed description see [65]) are consistent with the present available data on HE emission form GRBs obtained from the satellites...

With Fermi (and Swift still in orbit) a new era for GRB studies has been recently opened and it’s clear how, increasing the statistic will be possible to improve the understanding of the prompt GRB emission mechanism and the physical properties of the emission region, as well as the understanding of the early afterglow and the (delayed) HE emissions. It’s also clear how observations in the GeV-TeV band would be really useful in order to solve the ”GRB puzzle”.
This chapter gives an overview of the main analysis concepts, before these will be applied to data in the next chapter. The standard data transfer procedure foresees to send data from La Palma to Europe, either via tape or via a 2 MByte/s Ethernet link. These data consist of ON data, whenever the telescope was pointing at the presumed source, and OFF data, with the telescope pointing at a different part of the sky, at the same zenith angle, observational conditions and average star light intensity as the one obtained with the ON pointing. Also simulated gamma-ray showers are available from the diverse institutes of the collaboration, for different telescope pointings and point-spread functions of the mirrors. Additionally, auxiliary data from the individual subsystems is gathered. All three main data types can contain three different run types which are combined in an observation sequence:

**Pedestal Run:** This run contains usually 1000 events, taken with random trigger and used to calculate the pedestal offsets, due to Night Sky backgorund (NSB), for the calibration run.

**Calibration Run:** This run contains usually 4096 events, taken with the standard calibration light pulses from 10 UV Leds and the calibration trigger. It is used to determine the conversion factors of the signal from FADC counts to photo electrons (phe) produced by the photo effect inside the PMT. Interleaved calibration events are also taken, with a rate of 50 Hz.

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1There is also data taken in “wobble”-mode, i.e. the telescope pointing 0.4° off the presumed source, alternating the offset at fixed time intervals. With steady, point-like sources, that observation mode yields higher sensitivities since no time is lost to take OFF data. Because of the dead-time introduced when changing between the Wobble-positions, we never use the “wobble”-mode for Gamma Ray Burst observations.
Cosmics Runs: These consist of as many data runs as the source was observed. Each run contains usually 10075 (or less) events (for the new FADC system), taken with the level-1 majority trigger, which requires 4 neighboring channels to exceed a pre-set discriminator threshold (DT). The DT’s depend on the amount of light of night sky and are usually adjusted such that a Galactic source is observed with a 25% higher threshold condition than an extra-galactic source. This configuration yields an average event rate of typically 200 Hz.

These runs are merged into sequences. One sequence consist of one pedestal, one calibration and several data runs. Subsequently, all three data types are passed through the same analysis chain:

Run selection procedure: Eliminates too short runs (having less than 10 events), or runs taken under obviously bad observation conditions, e.g. runs explicitely marked as test runs, very low event rates or strong mis-pointing of the telescope.

Signal extraction from FADC slices: Calculates the charges and arrival times from the FADC slices of each channel and each event.

Calibration: Converts the extracted charges to equivalent numbers of photo-electrons, using the calibration light pulses, applies the timing offsets and excludes mal-functioning pixels.

Image cleaning: Recognizes and eliminates pixels which contain no signal from a Cherenkov light flash.

Pre-selection: Based upon very basic quality criteria, like the number of remaining channels or the total number of accumulated photo-electrons after cleaning, a first event selection is made, removing typically 3–5% of the events.

Calculation of quality parameters: Mainly the Hillas-parameters [131] and a couple of further quality criteria are calculated. These are then getting combined to one single quality parameter, called "hadronness".

Application of quality cuts: Only those events are getting selected which have a low "hadronness"-parameter and sometimes fulfill further requirements, based on the to-
tal number of photo-electrons or further image parameters which were not included in
the calculation of “hadronness”.

**Calculation of significances:** Finally, the reconstructed angle between the incidence di-
rection of the cascade and the telescope pointing direction, expressed in the so-called
ALPHA-parameter, is used to determine whether an excess of cascades from the ob-
served source position can be seen. If so, the significance of the number of excess events
is calculated.

**Calculation of upper limits:** In case of too low significances (usually taken as $5 \sigma$), an
upper limit to the number of gamma rays is calculated for the observation period.

Most of the functionality is available as executable programs within the Magic Reconstruc-
tion and Analysis Software (MARS):

- **callisto** $^2$: performs the signal extraction and calibration steps.
- **star** $^3$: performs the image cleaning and calculation of quality parameters.
- **osteria**$^4$: perform the Random Forest statistical learning.
- **melibea**$^5$: performs the calculation of the “hadronness” and offers the possibility to apply
  first quality cuts.
- **flux**: calculates spectra and effective areas.

**4.0.1 Calibration**

In a first step, the calibration parameters (mean number of photo-electrons, conversion factors,
and time offsets) are calculated from a dedicated calibration file which itself uses pedestal offsets
calculated from a previously taken pedestal file. This calibration procedure determines which
pixels have to be excluded, calculates the conversion factors from extracted signal to equivalent
photo-electrons and the arrival time offsets. These numbers are then used to calibrate the first
10 seconds of cosmosics data. The conversion factors are subsequently getting updated with the

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$^2$ stands for: **CALibrate LIght Signals and Time Offsets**

$^3$ stands for: **STandard Analysis and image Reconstruction**

$^4$ stands for: **Optimize STandard Energy Reconstruction and Image Anelysis**

$^5$ stands for: **MErge and Link Image parameters Before Energy Analysis**
results of every 500 interlaced calibration pulses, while the mean number of photo-electrons is updated every 5000 pulses.

The calibration of a Gamma Ray Burst data set is slightly different from this standard calibration because there is no time to take pedestal and calibration runs before the source is observed. Instead, the interlaced calibration events from the observation of the previous source are used which calibrate directly the first 10 seconds of the GRB data. All following data is then calibrated with the interlaced calibration events of the same GRB data set.

### 4.0.2 Image Cleaning

The image cleaning removes pixels which apparently do not form part of the shower image. As the later shower reconstruction relies on second moments of the image [131], any disturbing pixels far from the rest of the shower, can severely mis-lead the image reconstruction. It is therefore mandatory not to add pixels to the image which contain only noise. On the other hand, any unnecessary tightening of the exclusion limits will lead to signal loss at low gamma-ray energies. For this reason, clever algorithms are needed which have to be adjusted such as to provide a low energy threshold and ensure that the vast majority of the images are un-affected by noise.

#### Absolute Image Cleaning

First, all pixels with a number of phe above a predefined threshold are selected. These are the so-called "core pixels". If they have at least one direct neighbored core pixel, they survive the first cleaning step. Then, in direct vicinity of the core pixels, the so-called "boundary pixels" are searched for. They have to contain a signal above a predefined threshold that is lower than the first one and they have to exhibit an already cleaned next neighbor. With this cleaning method, an easy analysis is possible, but a low energy threshold cannot be reached: if we define a threshold to low, we’ll include in the cleaned image pixels just illuminated by NSB.

#### Time Image Cleaning

Since 2008 the standard image cleaning is the Time Image Cleaning [39]. To reduce the number of the pixels surviving the cleaning that only contain a NSB signal, an additional
coincidence in time of the extracted signals between adjacent pixels is requested. The time
distribution of the photons coming from the NSB is diffuse, unlike the photons of shower
events that arrive at the camera in a well defined time window. Timing is used to constrain
the selection of core and boundary pixels in the Image Cleaning algorithm. In a first step,
Absolute Image Cleaning with low threshold is applied. In the second and third step, pixels
are rejected by comparison of their arrival times. The arrival time of the signal is given in
terms of time slices (t.s.) i.e. in terms of digitation samples units (1 t.s. = 0.5 ns for the new
FADCs). Usually, a time coincidence window of 1.5 ns or 2.0 ns is used. In Fig. 4.1 different
cleaned events are shown.

4.0.3 Quality Parameters

Those \( N \) pixels which survived the image cleaning are then used to calculate the classical
Hillas parameters [131].

**SIZE** The **SIZE** of the image is the sum of the pixel charges, expressed in equivalent-phes.
It is proportional to the total integrated light content of the shower. In the case of the MAGIC
camera, \( \eta_{comm} \) has been chosen as the average photo-detection efficiency of the inner pixels.
There, the equivalent number of photo-electrons equals the number of actually registered
photo-electrons on average. Outer pixels, in turn, have on average a factor \( \eta_{inner}/\eta_{outer} \approx 1.5 \)
higher numbers of equivalent photo-electrons than actual numbers of photo-electrons. This
construction has the advantage to avoid discontinuities of the reconstructed image parameters
at the edge between areas of different pixel types. On the other hand, the simple assumption
that the variance of **SIZE** equals the same **SIZE**, because of Poissonian fluctuations, is only
true for images entirely comprised in the inner part of the camera.

As the number of created electrons in an electro-magnetic shower is proportional to the
energy of the shower and the emitted number of Cherenkov photons per unit path length of an
electron is almost constant, the sampled part of the emitted shower light reflects already well
the shower energy and **SIZE** can be used therefore as a first measure of the shower energy.

**WIDTH, LENGTH, DIST and ALPHA** Further (classical) Hillas parameters are
sketched in figure 4.2 and contain the parameters [90,131]:

Figure 4.1: Illustrative $\gamma$-event images. First row: raw data; second row: Absolute cleaned event (different levels); third row: image obtained with the Time Image Cleaning.
Figure 4.2: Sketch of image parameters: The red areas show those pixels having survived the image cleaning procedure. The black dotted line show the reconstructed ellipse, with the parameters $\text{LENGTH}$, $\text{WIDTH}$, $\text{DIST}$, $\text{ALPHA}$ and $\text{LEAKAGE}$. Head ($H$) and tail ($T$) of the shower are displayed as well as three additional “islands”.

$\text{WIDTH}$: The RMS spread of light along the minor axis of the image, a measure of the lateral development of the cascade.

$\text{LENGTH}$: The RMS spread of light along the major axis of the image, a measure of the vertical development of the cascade.

$\text{DIST}$: The distance from the centroid of the image to the source position in the camera.

$\text{ALPHA}$: The angle between the major axis of the image and the radius drawn from the center of the camera through the center of the image.

In order to calculate these parameters, the first and second moments of the image are calculated from the positions $x$ and $y$ (measured in degrees) of each pixel in the camera which has survived the image cleaning procedure:

\[
< x > = \frac{\sum_{i=0}^{N} x_i w_i}{\sum_{i=0}^{N} w_i}, \quad < y > = \frac{\sum_{i=0}^{N} y_i w_i}{\sum_{i=0}^{N} w_i},
\]

\[
< x^2 > = \frac{\sum_{i=0}^{N} x_i^2 w_i}{\sum_{i=0}^{N} w_i}, \quad < y^2 > = \frac{\sum_{i=0}^{N} y_i^2 w_i}{\sum_{i=0}^{N} w_i},
\]

\[
< xy > = \frac{\sum_{i=0}^{N} x_i y_i w_i}{\sum_{i=0}^{N} w_i}
\]  

(4.1)
where the weights $w_i$ are typically a power $\alpha$ of the pixel size in equivalent photo-electrons: $w_i = \text{SIZE}_i^\alpha$.

The first moments can be combined into a parameter describing the distance of the center of the ellipse from the center of the camera\footnote{In case the source position is not the camera center (e.g. when observing in Wobble mode), the first moments have to be calculated with respect to the new source position.}:

$$\text{DIST} = \sqrt{< x >^2 + < y >^2} \quad (4.2)$$

In the following, we set $\tilde{\text{var}}(x) := < x^2 > - < x >^2$, $\tilde{\text{var}}(y) := < y^2 > - < y >^2$ and $\tilde{\text{cov}}(x, y) := < xy > - < x >< y >$, where the hats indicate that the expressions yield true variances only in the case of $\alpha = 0$ (all weights are 1). The second moments get then combined in the "covariance" matrix $M$:

$$M = \begin{pmatrix} \tilde{\text{var}}(x) & \tilde{\text{cov}}(x, y) \\ \tilde{\text{cov}}(x, y) & \tilde{\text{var}}(y) \end{pmatrix}, \quad (4.3)$$

which is then rotated by an angle $\delta$ such that $M^{\text{rot}}$ becomes diagonal in the new coordinate system:

$$M^{\text{rot}} = \begin{pmatrix} \cos \delta & \sin \delta \\ -\sin \delta & \cos \delta \end{pmatrix} \cdot M \cdot \begin{pmatrix} \cos \delta & -\sin \delta \\ \sin \delta & \cos \delta \end{pmatrix} = \begin{pmatrix} \lambda_1 & 0 \\ 0 & \lambda_2 \end{pmatrix} \quad (4.4)$$

The solution of eq. 4.4 and the requirement $\lambda_1 \geq \lambda_2$ yields the rotation angle and the diagonalized matrix:

$$\tan \delta = \frac{\tilde{\text{var}}(y) - \tilde{\text{var}}(x) + \sqrt{(\tilde{\text{var}}(y) - \tilde{\text{var}}(x))^2 + 4 \cdot \tilde{\text{cov}}(x, y)^2}}{4 \cdot \tilde{\text{cov}}(x, y)^2}$$

$$\text{LENGTH}^2 := \lambda_1 = \frac{\tilde{\text{var}}(x) + 2 \cdot \tan \delta \cdot \tilde{\text{cov}}(x, y) + \tan^2 \delta \cdot \tilde{\text{var}}(y)}{1 + \tan^2 \delta}$$

$$\text{WIDTH}^2 := \lambda_2 = \frac{\tilde{\text{var}}(x) - 2 \cdot \tan \delta \cdot \tilde{\text{cov}}(x, y) + \tan^2 \delta \cdot \tilde{\text{var}}(y)}{1 + \tan^2 \delta}, \quad (4.5)$$

yielding the ellipse’s major ($\text{LENGTH}$) and minor ($\text{WIDTH}$) half axes. With the above definitions, a unit vector along the main axis of the ellipse is obtained:

$$\vec{a} = \frac{1}{1 + \tan^2 \delta} \begin{pmatrix} 1 \\ \tan \delta \end{pmatrix} \quad (4.6)$$
and such the angle $ALPHA$ can be calculated (see fig 4.2)\(^7\):

$$ALPHA = \arccos \left( \frac{< x > + \tan \delta \cdot < y >}{DIST \cdot \sqrt{1 + \tan^2 \delta}} \right),$$

(4.7)

with $ALPHA$ ranging from $-90^\circ$ to $+90^\circ$, depending on whether the shower center is found in the left or right part of the camera\(^8\).

The reason to define parameters based on the second moments of the image lies in the fact that electromagnetic showers tend to produce “cleaner” reconstructed ellipses, i.e. showing on average a smaller ratio $WIDTH/LENGTH$, than images from hadronic showers which tend to look more fragmented and yield “rounder” reconstructed ellipses. On the contrary, the $ALPHA$ parameter gives an estimate of the arrival direction of the shower. Especially showers traveling parallel to the telescope axis yield $ALPHA$-values concentrated around zero.

While cosmics rays (hadronic showers) do not show any preferred arrival direction, gamma rays from astro-physical sources do so and show thus excesses at low $ALPHA$-values.

**CONC, CONC4, LEAKAGE, NUMBER ISLANDS** Further (non-classical) image parameters are:

- **CONC**: Fraction of equivalent photo-electrons, contained in the two brightest pixels, per total image $SIZE$.

- **CONC4**: Fraction of equivalent photo-electrons, contained in the four brightest pixels, per total image $SIZE$.

- **LEAKAGE**: Fraction of equivalent photo-electrons, contained in the outmost ring of the camera, per total image $SIZE$.

- **NUMBER ISLANDS**: Number of separated areas with signal, after the image cleaning.

While the concentration parameter $CONC$ is used mainly to identify and reject non-shower events, the leakage parameter $LEAKAGE$ is an approximate measure of the part of a shower

\(^7\)Equation 4.7 does not include possible other reference points than the center of the camera, nor the inclusion of the “head-tail” information, i.e. knowledge about the direction in which the shower was moving. For more detailed information about these possibilities, see [215, 241].

\(^8\)There is also a different definition of $ALPHA$ ranging from $-180^\circ$ to $+180^\circ$, which makes sense only if the head-tail information is known, i.e. the time-flow direction of the shower. In this work, only the absolute value of $ALPHA$, obtained from the classical definition between $-90^\circ$ to $+90^\circ$, is used. See also [241].
which lies outside of the sensitive area of the camera. Images with large values of \textit{LEAKAGE} lack necessarily precision in the determination of the image parameters. Finally, the number of separated islands has turned out to be a useful quality parameter to distinguish hadronic (more than one island) from electromagnetic showers (only one island).

4.0.4 Multivariate Classification

A common problem in experimental physics is the separation of signal from background where quality parameter distributions overlap and the separability of both is low. In such a situation, probability density functions (p.d.f.) should be constructed, which in the ideal case incorporate the full description of the physical models of all processes at play and the entire information captured by the detector. A p.d.f. for electromagnetic showers and one for hadronic showers can be constructed, depending on the telescope parameters, which yield directly both signal and background probabilities. Unfortunately, especially the hadronic shower development is hard to model statistically, and many necessary parameters cannot be included in a straight-forward way, e.g. atmospheric conditions. Nevertheless, efforts are made in that direction, without being completed so far [156].

In this analysis only reduced information, based on the image parameters, is used to determine the p.d.f. for a given data set. In the case of one only parameter, signal and background acceptance can be readily visualized as a distribution along that variable. In the case of multiple variables – which are in general correlated – this is not possible any more and sophisticated multivariate have been invented to extract the acceptances from a set of parameters and construct a one-dimensional p.d.f. Many of these methods employ random techniques used to construct the probability density model on training and to determine their efficiencies on test. For a detailed description and comparison of these methods, see [203].

Random Forest The analysis presented in this thesis, is based on an implementation of the Random Forest technique, developed by Leo Breiman [57] and implemented in the MAGIC analysis software by Thomas Hengstebeck [129]. As other classification methods, it is based on the construction of decision trees: If every image parameter (out of \( N \)) is thought of as a vector in an \( N \)-dimensional parameter space, all signal and background events (e.g. from the training sample) are contained in an \( N \)-dimensional hypercube. The tree starts at one corner
of the hypercube, called the “root node”. Subsequently one image parameter (dimension of the hyperspace) is selected randomly and a ‘cut value’ on that parameter chosen such that the gini-index of the two samples gets minimized. The Gini-index can be expressed in terms of the number of events on the left or right side of the cut value:

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

(4.8)

The next step chooses another parameter randomly, starting from the obtained cut value (the first node). This partitioning of the hypercube in smaller sub-cubes is performed until the node is “pure”, i.e. the Gini index is zero or the number of remaining events in the sub-hypercube is less than a certain threshold, usually 1–10 events. The corresponding node is then called a “terminal node”.

In order to avoid overfitting, various (usually 50 to 100) random trees are grown, starting from different points of the hypercube on a random subsample of the training sample. 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. Depending on whether the event lies inside or outside the corresponding terminal hyper-cube, a value of 0 (for gamma-like terminal nodes) or 1 (for hadron-like terminal nodes) is assigned to the event. As there is a whole forest, each random tree assigns a value \( l \) to the event and the parameter \( \text{HADRONNESS} \) can be calculated:

\[ \text{HADRONNESS} = \frac{\sum_{i=0}^{N_{trees}} l_i}{N_{trees}} \]  

(4.9)

Per definition, the \( \text{HADRONNESS} \) ranges from 0 to 1 and yields values distributed close to 1 for hadron-like and close to 0 for gamma-like event, if the two event types are separable within the chosen set of parameters. Otherwise, values distributed around 0.5 are expected for both types of events.

A separate feature of the Random Forest algorithm has been used frequently in the collaboration: By adding a scale parameter, like the number of equiv. photo-electrons, to the list of quality parameters, the classifier can be brought to optimize the separation of signal from background for every bin of the scale parameter. To function properly, this procedure requires that the two original distributions of the scale parameter are brought to match, either
by removing events from the background training sample or from the signal training sample. The parameter with the highest discrimination power, \textit{ALPHA}, is not included in the list of parameters for the random forest training, since parts of the distribution of \textit{ALPHA} are needed to estimate the background in the signal region from an unbiased estimate of the background from the pure background region of that distribution (see the analysis part of the next chapter). This capability is lost if \textit{ALPHA} is included in the calculation of \textit{HADRONNESS} and later cuts on \textit{HADRONNESS} are applied.

\textbf{Energy Estimation}  
Already the number of photo-electrons per shower image, expressed in the variable \textit{SIZE}, is correlated with the shower energy and can serve thus for an energy estimate \textit{ENERGY}. Especially at high energies, this approach works fine, although the conversion factor between \textit{SIZE} and energy has to be retrieved for every telescope pointing zenith angle. Moreover, the total amount of light detected from a shower reflects not only an energy dependence but also a distance dependence.

A more precise approach uses a slightly modified version of the Random Forest algorithm: Instead of separating signal from background distributions, random forest is getting trained to separate the population of events with matching (simulated) energy from those having energy outside a corresponding energy bin. Then, a loop over all energy bins is performed to train the estimator. 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 getting selected. This approach yields energy resolutions of the order of $\sim 25\%$ at higher energies and $\sim 30\%$ around the energy threshold.

\textbf{4.0.5  Determination of Significances and Cuts}  

After the calculation of the \textit{HADRONNESS} parameter, the analyzer remains with simulated MC gamma data sets, an ON and various OFF data sets. In the following, efficient cuts on \textit{HADRONNESS} have to chosen, separately for every bin in \textit{ENERGY}, and finally an \textit{ALPHA-plot} is made showing the distribution of the absolute value of \textit{ALPHA} ($|\textit{ALPHA}|$) for ON data and for $\kappa$·OFF data, where $\kappa$ is a normalization constant obtained by demanding that the histogram integral from $|\textit{ALPHA}| = 30^\circ$ to $|\textit{ALPHA}| = 80^\circ$ be the same for ON and normalized OFF data. That range in the distribution of $|\textit{ALPHA}|$ is assumed to be almost
signal-free [90]. Figure 4.3 gives an example of an \textit{ALPHA}-plot, if the observed source (in this case the Crab Nebula) is visible. Later, the number of excess events $N_{\text{ex}}$ between ON and OFF data in the range from $|\text{ALPHA}| = 0$ to $|\text{ALPHA}| = \text{ALPHA}_{\text{cut}}$ is counted. The value $\text{ALPHA}_{\text{cut}}$ has to be determined beforehand, otherwise the calculated significance of the number of excess events would be biased. Following a maximum likelihood calculation by Ti-Pei Li and Yu-Qian Ma [147], the significance can then be obtained from the above numbers:

$$S = \sqrt{2} \cdot \left[ N_{\text{ON}} \cdot \ln \left( \frac{1+\kappa}{\kappa} \cdot \frac{N_{\text{ON}}}{N_{\text{ON}} + N_{\text{OFF}}} \right) + N_{\text{OFF}} \cdot \ln \left( 1 + \kappa \cdot \frac{N_{\text{OFF}}}{N_{\text{ON}} + N_{\text{OFF}}} \right) \right]^{1/2}. \quad (4.10)$$

![Figure 4.3: a) schematic of the \textit{ALPHA} b) Example of an \textit{ALPHA}-plot.](image)

4.0.6 Calculation of Effective Areas

By lack of an adequate calibrated high-energy gamma-ray source in the sky, the effective area $A_{\text{eff}}$ of a Cherenkov telescope has to be derived by means of MC simulation. $A_{\text{eff}}$ depends
on the energy $E$ of the incident gamma-ray, on the telescope pointing zenith angle $\theta$ (at low energies also on the azimuth angle) and the atmospheric conditions. It can be defined as:

$$A_{\text{eff}}(E, \theta) = \lim_{N_{\text{sim}}(E) \to \infty, A_{\text{sim}} \to \infty} A_{\text{sim}} \cdot \frac{N_{\text{analysis}}(E, \theta)}{N_{\text{sim}}(E)},$$  \hspace{1cm} (4.11)$$

where $N_{\text{sim}}(E, \theta)$ is the number of simulated gamma-rays with energy $E$, $A_{\text{sim}}$ the simulated incidence area and $N_{\text{analysis}}(E, \theta)$ the number of remaining events after correct simulation of the telescope zenith angle, trigger and the application of all analysis cuts. In practice, the limit is replaced by a sufficiently large numbers $A_{\text{sim}}$ and $N_{\text{sim}}(E)$.

In case one works with energy bins (e.g. in the calculation of upper limits), $N_{\text{analysis}}(E, \theta)$ will also contain cuts on the reconstructed energy. Figure 4.4 shows a typical distributions of $A(E)$: Without cuts on ENERGY and cutting out an energy bin. One can see that the true (simulated) energy extends will beyond the limits of reconstructed energy, because of the limited energy resolution.

![Effective area](image)

Figure 4.4: Distributions of effective areas for a telescope pointing zenith angle $\sim 38^\circ$ before and after typical analysis cuts.

### 4.0.7 Calculation of Upper Limits

Knowing a gamma-ray flux $dN_\gamma/dE dA dt$, observed with a Cherenkov telescope of effective area $A(E)$, during a time interval $[T_{\text{min}}, T_{\text{max}}]$ with an effective on-time function $\epsilon(t)$, one
could predict the number of events registered by the telescope as follows:

\[ N_{\text{obs}} = \int_0^\infty \int_{T_{\text{min}}}^{T_{\text{max}}} \frac{dN_\gamma}{dE \, dA \, dt} \cdot A(E) \cdot \epsilon(t) \, dE \, dt , \quad (4.12) \]

or, dividing the observed number of excess events in bins of reconstructed energy \( E_r \):

\[ N_{\text{obs}}(E_r^{\text{min}}, E_r^{\text{max}}) = \int_0^\infty \int_{T_{\text{min}}}^{T_{\text{max}}} \frac{dN_\gamma}{dE \, dA \, dt} \cdot A(E, E_r^{\text{min}}, E_r^{\text{max}}) \cdot \epsilon(t) \, dE \, dt , \quad (4.13) \]

where \( E_r^{\text{min}} \) and \( E_r^{\text{max}} \) are the bin limits of reconstructed energy.

In case of no signal, the telescope measures on average zero events \( N_{\text{obs}} = 0 \), but also different (positive and negative) numbers are possible since \( N_{\text{obs}} \) follows a statistical distribution. From the number of observed events, the analyzer derives first an upper limit on the number of observable events, typically giving a confidence level of 95%: \( N_{>95\%} \). In a next step, eq. 4.13 is translated to an un-equality:

\[ N_{>95\%}(E_r^{\text{min}}, E_r^{\text{max}}) > \int_0^\infty \int_{T_{\text{min}}}^{T_{\text{max}}} \frac{dN_\gamma}{dE \, dA \, dt} \cdot A(E, E_r^{\text{min}}, E_r^{\text{max}}) \cdot \epsilon(t) \, dE \, dt , \quad (4.14) \]

which has to be transformed into an upper limit for \( dN_\gamma/dEdAdt \).

**Upper Limits for Number of Observable Events**

Typically, the method of the ALPHA-plot yields the following numbers:

\( N_{\text{obs}}^{\text{ON}} \): The number of observed events from ON data within the signal region (e.g. from \( 0 < \text{ALPHA} < \text{ALPHA}_{\text{cut}} \)).

\( N_{\text{obs}}^{\text{OFF}} \): The number of observed events from OFF data within the signal region

\( \Delta N_{\text{OFF}} \): The statistical uncertainty on \( N_{\text{OFF}} \)

\( \Delta \epsilon \): A global systematic uncertainty of the efficiency of the detector

Based on these number, the observed number of excess events is calculated:

\[ N_{\text{EX}}^{\text{obs}} = N_{\text{obs}}^{\text{ON}} - N_{\text{obs}}^{\text{OFF}} . \quad (4.15) \]

An upper limit on \( N_{\text{EX}} \) is then constructed in two steps: Construction of a probability density function (pdf) for \( N_{\text{EX}} \), given a hypothesis on the mean physical number \( \mu_{\text{EX}} \), and inversion of the original p.d.f. to yield a p.d.f. for \( \mu_{\text{EX}} \), given all observed numbers.
Construction of p.d.f.  Given a hypothesis on the true mean number of excess events $\mu_{EX}$ and background events $\mu_{OFF}$ (usually not an integer numbers), the probability distribution for observing values of $N_{ON}$ is constructed: $P(N_{ON} \mid \mu_{EX}; \mu_{OFF}; \Delta \mu_{OFF})$. E.g. for the (unrealistic) case of pure Poissonian fluctuations for $N_{ON}$ and a perfectly known mean background $\mu_{OFF}$ without uncertainty ($\Delta \mu_{OFF} = 0$), one gets

$$P(N_{ON} \mid \mu_{EX}; \mu_{OFF}) = \frac{(\mu_{EX} + \mu_{OFF})^{N_{ON}} \cdot e^{-(\mu_{EX} + \mu_{OFF})}}{N_{ON}!}.$$  (4.16)

The confidence interval defines a range $[N_{ON}^{low}, N_{ON}^{up}]$ outside which would lie the results $N_{ON}$ of only a percentage $\alpha$ of experiments (carried out under the same conditions):

$$P\left(N_{ON} \in [N_{ON}^{low}, N_{ON}^{up}] \mid \mu_{EX}; \mu_{OFF}; \Delta \mu_{OFF}\right) = 1 - \alpha.$$  (4.17)

In case of upper limits, the confidence interval counts the percentage of experiments yielding $N_{ON} > N_{ON}^{up}$.

Inversion of p.d.f.  Starting from probability distributions $P(N_{ON} \mid \mu_{EX}; \mu_{OFF})$, a second distribution $P(\mu_{EX} \mid N_{ON}^{obs}, N_{OFF}^{obs})$ has to be retrieved, predicting the probability for the true mean number of excess events, given the observations $N_{ON}^{obs}$ and $N_{OFF}^{obs}$. To solve this (non-trivial) problem, two branches of statistics have been developed: a bayesian and a frequentist approach. The first uses Bayes’ theorem:

$$P(\mu_{EX} \mid N_{ON}^{obs}, \mu_{OFF}; \Delta \mu_{OFF}) = P(N_{ON} \mid \mu_{EX}; \mu_{OFF}; \Delta \mu_{OFF}) \cdot P(\mu_{EX} \mid \mu_{OFF}; \Delta \mu_{OFF}),$$  (4.18)

where $P(\mu_{EX} \mid \mu_{OFF}; \Delta \mu_{OFF})$ are called “prior probabilities” for $\mu_{EX}$. The strength of this method lies in the fact that un-physical regions or other prior knowledge about $\mu_{EX}$ can be included in $P(\mu_{EX} \mid \mu_{OFF}; \Delta \mu_{OFF})$. One possible choice for the prior probability was made by O. Helene [128], namely $P(\mu_{EX} \mid \mu_{OFF}; \Delta \mu_{OFF}) = N$, where $N$ is simply a normalization constant. That choice assumes thus a uniform probability for the entire signal range from $[0, \infty]$. An upper limit is then obtained via the condition:

$$\int_{N_{ON} > 95\%}^{\infty} P(\mu_{EX} \mid N_{ON}^{obs}, \mu_{OFF}; \Delta \mu_{OFF}) d\mu_{EX} = \alpha.$$  (4.19)

The confidence level (CL) is then defined as:

$$CL = (1 - \alpha) \cdot 100\%.$$  (4.20)
By convenience, confidence levels are typically chosen to be 95% \(^9\). The frequentist approach proceeds to construct a confidence belt according to the prescription of Neyman [171]: For each hypothesis \(\mu_{EX}\), a horizontal confidence interval satisfying eq. 4.17 is drawn (see figure 4.5 for the example p.d.f. of eq. 4.18, with background \(\mu_{OFF} = 3\)). Given a measured value \(N_{ON}^{obs}\), one can then invert the confidence interval to obtain limits \([\mu_{EX}^{low}, \mu_{EX}^{up}]\) (depending on \(N_{ON}^{obs}\)) which satisfy a confidence interval for \(\mu_{EX}\):

\[
P(\mu_{EX} \in [\mu_{EX}^{low}, \mu_{EX}^{up}] \mid N_{ON}; \mu_{OFF}; \Delta \mu_{OFF}) = \alpha \tag{4.21}
\]

or:

\[
P(\mu_{EX} < \mu_{EX}^{up} \mid N_{ON}; \mu_{OFF}; \Delta \mu_{OFF}) = \alpha \tag{4.22}
\]

whereby the first is said to yield a central confidence interval and the second an upper confidence limit,\(^{10}\). As above, the choice of \(\alpha\) is free, but usually 0.05 is taken (\(\equiv 95\%\) CL). The prescription to construct the limits \(\mu_{EX}^{low}\) and \(\mu_{EX}^{up}\) is then said to yield the correct coverage if in a series of experiments the true (unknown) value \(\mu_{EX}^{true}\) is found in between the limits in \((1 - \alpha)\) of the cases. The Bayesian construction eq. 4.18 does in general not yield the correct

\(^9\)Other choices sometimes seen are: 90\% or 99\%.

\(^{10}\)There is a separate problem related to the choice of both methods, known as the “flip-flop” problem which is not further explained here.
coverage. The prescription shown in figure 4.5 however does so by construction.

After the commonly used prescription by Gary Feldman and Robert Cousins [91], the construction of a confidence interval for the case of more than one measured parameter or more than one possibility for $\mu_{EX}$ to yield a same limit $N_{ON}^{up}$ or $N_{ON}^{low}$, those values of $\mu_{EX}$ are chosen which have the highest relative probability and are physically allowed.

The original prescription of [91] was enlarged by Wolfgang Rolke and Angel López [204] to a confidence belt construction which includes a probability model for the background

$$[\mu_{EX}^{low}, \mu_{EX}^{up}] = [\mu_{EX}^{low}, \mu_{EX}^{up}](N_{ON}, N_{OFF}, \Delta N_{OFF})$$

and later [205] including a probability model of the overall efficiency $\epsilon$ and its p.d.f.:

$$[\mu_{EX}^{low}, \mu_{EX}^{up}](N_{ON}, N_{OFF}, \Delta N_{OFF}, P(\epsilon))$$

Although traditionally upper limits in Cherenkov telescope have been published using the Bayesian approach of O. Helene [128], that method can result in considerable under-coverage and does not give any prescription of how to include systematic uncertainties on the efficiency.

The (GRB working group of the) MAGIC collaboration has published so far upper limits using the enlarged method by Rolke et al. [205] (e.g. in [25]) to obtain upper limits $\mu_{EX}^{up} \equiv N_{ON}^{95\%}$.

One of the authors of [205] has provided an implementation of their method [62] for seven typical experimental situations. One of these models (model #3) was found to match most of the situations found in typical analyses of MAGIC data: $N_{ON}$ distributed Poissonian, the down-scaled number of OFF events $\kappa \cdot N_{OFF}$ in first approximation Gaussian with a width $\kappa \cdot \sqrt{N_{OFF}}$ and the efficiency Gaussian, centered around 1. The rather complicated formulae used in [62] cannot be presented here, but are fully described in [205].

### Particle Flux Upper Limits

Since in the case of no signal neither the energy spectrum of the source $dN_\gamma/dE$, nor the time evolution $dN_\gamma/dt$ are unknown, some reasonable assumptions have to be made: Typically in these cases, power-law spectra can be assumed or combinations of power-law spectra and

---

11 Note that for the case $\kappa > 1$ and $N_{OFF} < 10$, this assumption is not valid any more and model #2 should be used, instead.
exponential cut-offs:

\[
\frac{dN_\gamma}{dE \, dA \, dt} = f_0 \cdot \left( \frac{E}{E_0} \right)^{-\alpha} \cdot \exp \left( -\frac{E - E_0}{E_b} \right),
\]

(4.25)

where \(E_0\) is the mean energy at which the limit is calculated and \(\alpha\) the hypothetical spectral index. \(E_b\) is the break energy of the exponential cut-off and \(A_0\) a reference area. The light curve \(dN_\gamma/dt\) has to be assumed approximately constant, otherwise the observation time window \([T_{\text{min}}, T_{\text{max}}]\) would need to be split into smaller parts.

Making the above assumption, eq. 4.13 can be re-written to yield an expression for the flux limit:

\[
f_0 < \frac{N_{\text{\%95}}(E_{\text{r min}}, E_{\text{r max}})}{\int_0^{\infty} \left( \frac{E}{E_0} \right)^{-\alpha} \cdot A(E, E_{\text{r min}}, E_{\text{r max}}) \, dE \cdot \int_{T_{\text{min}}}^{T_{\text{max}}} \epsilon(t) \, dt},
\]

(4.26)

with only measured values on the right side, except for the spectral index \(\alpha\). The integral \(\int_{T_{\text{min}}}^{T_{\text{max}}} \epsilon(t) \, dt\) is also called the effective ON time \(T_{\text{eff}}\) and the integral over the effective area:

\[
\int_0^{\infty} \left( \frac{E}{E_0} \right)^{-\alpha} \cdot A(E, E_{\text{r min}}, E_{\text{r max}}) \, dE
\]

can be called the weighted effective area average \(< A_{\text{eff}} >\).

In its short form, equation 4.26 is then written:

\[
f_0 < \frac{N_{\text{\%95}}(E_{\text{r min}}, E_{\text{r max}})}{< A_{\text{eff}}(E_{\text{r min}}, E_{\text{r max}}, \alpha)> \cdot T_{\text{eff}}},
\]

(4.27)

\(f_0\) has conveniently the units: [photons/cm\(^2\)/s/TeV].

**Particle Fluence Upper Limits**

In case a limit on the fluence from a hypothetical gamma-ray source is desired, the test spectra eq. 4.25 have to be converted to fluences:

\[
\frac{dN_\gamma}{dE \, dA} = F_0 \cdot \left( \frac{E}{E_0} \right)^{-\alpha} \cdot \exp \left( -\frac{E - E_0}{E_b} \right),
\]

(4.28)

and limits on the fluence \(F_0\) are derived:

\[
F_0 < \frac{N_{\text{\%95}}(E_{\text{r min}}, E_{\text{r max}})}{< A_{\text{eff}}(E_{\text{r min}}, E_{\text{r max}}, \alpha)>}.
\]

(4.29)

\(^{12}\)Only differential upper limits will be treated here. Since at poor energy resolutions, the integration boundaries are not well defined, integral limits do not make much sense and are skipped in the following further discussions.
Expression 4.29 makes only sense, if the information is provided that the source had been observed during an effective ON time $T_{eg}$ in the time interval from $T_{min}$ to $T_{max}$. Especially for highly variable sources like GRBs, limits on the fluence are preferred since they are independent from the light-curve of the hypothetical emission. $F_0$ has conveniently the units: [photons/cm$^2$/TeV].

**Spectral Energy Density Upper Limits**

Very often, the spectral energy density SED $E^2 \cdot dN_\gamma / dEdAdt$ at a given energy is a preferred parameter over the particle flux since many physical processes emit approximately the same power in different energy ranges. In order to place an upper limit on the SED, the test spectra eq. 4.25 are transformed in energy density spectra:

$$
E^2 \frac{dN_\gamma}{dE dA dt} = P_0 \cdot \left( \frac{E}{E_0} \right)^{-\alpha+2} \\
E^2 \frac{dN_\gamma}{dE dA dt} = P_0 \cdot \left( \frac{E}{E_0} \right)^{-\alpha+2} \cdot \exp \left( -\frac{E - E_0}{E_b} \right), \tag{4.30}
$$

It is straightforward to see that:

$$
P_0 < \frac{N_{>95\%} < E(\alpha)^2 >}{< A_{eff}(E_{r}^{min}, E_{r}^{max}, \alpha + 2) > \cdot T_{eff}}, \tag{4.31}
$$

where the mean square energy $< E(\alpha^2) >$ has to estimated from MC simulations of the corresponding test spectra and application of all cuts. $P_0$ is mostly given in units of [erg/cm$^2$].
MAGIC GRBs observation strategy

MAGIC is currently the most suitable instrument to perform observation of the prompt emission and early afterglow emission from GRBs above 25 GeV. The instrument is designed to have the lowest possible energy threshold and fastest reaction time to GRB Coordinate Network (GCN) alerts. The MAGIC I telescope started to follow-up GCN alerts at the end of 2004. Since then, more than 50 candidates were observed. Just now MAGIC II has started to operate and the with the stereo system will have a factor of three better sensitivity. In this chapter I summarize the characteristics of the MAGIC I “GRB observation mode” and the strategy we’ll follow with the stereo system.

5.1 Observation of GRBs with ground based experiments

GRBs occur at random times in unpredictable directions. In order to get a significant enough detection rate, dedicated satellite experiments for GRB surveys need to observe large portions of the sky. Large FoV results in poor spatial resolution and crowded field at the detector. Therefore new satellite experiments, like SWIFT for example, are equipped with additional detectors, which allow studying the source with higher precision and in a broader wavelength range. These special detectors have smaller FOVs and need therefore to be aligned to the new sky position. This is done by slewing the satellite and performing follow-up observations with some delay after the GRB onset. Ground based experiments are larger than the satellite detectors and therefore much more sensitive. In order to allow follow-up observations of GRB events by ground based experiments, the GCN was invented.
5.1.1 GRB coordinates network

The GRB coordinates network was established in 1994. Its aim was to distribute the GRB sky coordinates from the BATSE detector – on board of the CGRO – to ground based experiments, while the GRB was still ongoing. After the de-orbiting of CGRO, the previously called BATSE COordinates DIstribution NEtwork (BACODINE) was renamed by the more general name GCN\(^1\). Three satellite experiments provided in the past GRB locations to the GCN in real-time: SWIFT, HETE-2 and INTEGRAL. Since April 2007 the HETE-2 satellite is no more fully operational. Currently SWIFT, INTEGRAL, AGILE and FERMI activate the follow-up observations. The main task of the GCN is to distribute the GRB coordinates to registered clients around the world (see Fig. 5.1). In addition, it allows the GRB community to share quickly their observation results over mailing lists (GCN circulars). The data flow between the satellite and the GCN is different from experiment to experiment. The time delay in the distribution of coordinates is the sum of several factors. First of all there is a delay onboard the satellite. This is variable and depends on the instrument and of course on the intensity and the time profile of the GRB event. The average delay to accumulate a 5\(\sigma\) detection is about 5 s, but it can last much longer. Before the alert package is transmitted, conversion to

\(^1\)http://gcn.gsfc.nasa.gov
sky coordinates and a comparison with a list of known variable sources is performed. This procedure depends again on the detector and lies in the range of $\sim 2$ s. The signal propagation delay to the ground station is small ($\sim 0.6$ s) and the distribution through TCP/IP protocol to the various ground based instruments takes less than 1 s.

## 5.2 MAGIC alert system and observation criteria

The GCN provides GRB alerts over different ways (e-mail, pager, phone, fax etc.) to its clients. The shortest delay between validation of signal at ground and acceptance of the alert at the client of $0.1 - 2.0$ s is warranted with the TCP/IP internet socket connection, connecting a computer at the instrument site directly with the GCN computer over internet. The time delay of the propagation of the coordinates package varies with the distance between the two computers. The internet socket connection method is used in the MAGIC project. A *Gamma Sources POinting Trigger* (gspot) demon program \cite{104} performs a full-time survey of the alerts provided by GCN and validates them with the predefined observability criteria. The satellite experiments distribute different kinds of information over GCN. In the first stage, gspot filters out alert packages from selected satellites, which contain all mandatory information. In the second step, the remaining packages are validated based on the following criteria:

- the Sun is a zenith angle of $> 103^\circ$, part of the astronomical twilight is thus explicitly included in the allowed observation time;

- the angular distance from the GRB to the Moon is $> 30^\circ$;

- the zenith angle for the GRB observation is $< 60^\circ$. In case of moon shine the maximal zenith angle is reduced to $55^\circ$;

- if the uncertainty of the GRB coordinate is $> 1.5^\circ$, the telescope observe 15 minutes and aborts observation unless a more precise position update arrives.

Nominal observation duration are from the start of observability until $T_0 + 4$ hours. As the redshift of the source is normally only known a few hours to days later, one is obliged to observe all candidates, although a later redshift measurement can classify the observation as
useless. As the energy threshold of an IACT depends sensitively on the observation zenith angle, most of GRBs are observed with higher threshold than the lowest possible.

5.3 pointing performances

The lightweight design of MAGIC supporting cradle together with the optimized steering system allows MAGIC I to slew very fast to any position in the sky. During the design phase of MAGIC II improvements of the drive system were made. The new steering electronics and optimization of the control loop parameters made the system more robust and allowed to increase the repositioning speed. Consequently MAGIC I drive system was recently upgraded and allows to reach the opposite location of the sky (180° movement in azimuth) now in only 20 s, while it was 54 s before the upgrade.

5.4 GRBs observed by MAGIC I

For the first years of operation the majority of the GRB observations by MAGIC I were triggered by the Swift-BAT detector. Since the end of 2008, the Fermi GBM detector started to dominate the alert frequency. Within the IV observation cycle (May 2008 - June 2009) in total 251 Fermi, 155 Swift, 12 INTEGRAL and 6 AGILE GCN alerts were received by MAGIC. Out of these, 14 Swift, 12 Glast and one INTEGRAL alerts fulfilled the observation criteria. An always update list of the MAGIC follow-up observation can be found in the webpage of the MAGIC GRB working group (http://mojorojo.magic.iac.es/grb). For a more detailed information about the GRB analyzed in this work I’ll remind to the last chapter.
MAGIC observation of GRB080430 afterglow

The MAGIC observation started at 21:12:14 UTC and ended at 23:52:30 UTC. The observation was disturbed by clouds so that 40 min of the data, starting from 21:12:14 UTC had to be rejected. The begin of the observation was at $T_0 + 4753$ s, well after the end of the prompt emission phase. The redshift of the burst was measured to be $z = 0.767$. The observation with MAGIC started at a zenith angle of $Z_d = 23^\circ$, reaching $Z_d = 35^\circ$ at the end. The energy threshold at the beginning was $E_{\text{thr}} = 112$ GeV, increasing slightly during the second half of the observation. No significant excess was found and upper limits are presented in this chapter. In the last part results of MAGIC observation of this bursts are used to evaluate the perspective for late-afterglow observation with ground based telescopes.

6.1 GRB 080430 MAGIC analysis

This analysis give an example of the standard analysis performed on MAGIC GRB data. When a GRB is observed, a report, similar to this is produced, and lead to a publication of a GCN Circular with the results.
Trigger:
Satellite: BAT
Time: 19:53:02 UTC
Trigger: 310613
Coordinates: RA=11ʰ01ᵐ14.66ˢ Dec=+51ᵈ41ᵐ084ˢ (J2000), given by UVOT
Type of burst: long
Duration: $T_{90} = 16.2 \pm 2.4$ s
Mean fluence 15−150 keV: $(1.2 \pm 0.1) \cdot 10^{-6}$ erg cm$^{-2}$ s$^{-1}$
Spectral index: $\alpha_{\text{BAT}} = 1.74 \pm 0.09$
GCN Circular: 7647 (Guidorzi et al.)

6.1.1 Results from multi-wavelength observations

**BAT GCN Circulars**

GRB080430 was triggered by the BAT detector [?]. The BAT light curve is shown in figure 6.1. The light curve shows a single FRED peak starting at $T_0 - 0.7$ s, peaking at $T_0 + 1.5$ s, and ending at $T_0 + 60$ s, before it returns to baseline. The duration of GRB080430 was $T_{90} = 16.2 \pm 2.4$ s, thus it can be classified as a long burst.

The time averaged spectrum from $T_0 - 0.3$ s to $T_0 + 21.3$ s is best fit by a single power-law model. The power-law index of the time-averaged spectrum is $\alpha_{\text{BAT}} = 1.73 \pm 0.09$. The fluence in the 15−150 keV band is $(1.2 \pm 0.1) \cdot 10^{-6}$ erg cm$^{-2}$. The 1 s peak photon flux measured from $T_0 + 1.70$ s is the 15−150 keV band is $(2.6 \pm 0.2)$ ph cm$^{-2}$ s$^{-1}$. All the quoted errors are at the 90% confidence level.

The SWIFT XRT began the follow-up observation at $T_0 + 55$ s [? and continued until $T_0 + 3.5 \cdot 10^6$ s. The XRT light curve is shown in figure 6.2. The XRT light curve can be modeled with a double broken power-law with slopes of $\alpha_{\text{XRT}} = 2.35 \pm 0.14$ up to a break at $T_0 + 294 \pm 40$ s, which is followed by a shallow decay with a slope of $\alpha_{\text{XRT}} = 0.45 \pm 0.04$ until $T_0 + 32 \pm 5$ ks, and finally joins $\alpha_{\text{XRT}} = 1.15 \pm 0.05$.

The spectrum formed from all the WT data (from 55 to 138 s) can be modeled with
an absorbed power-law of photon index $\gamma = 2.42 \pm 0.3$ and a column density of $NH = 4.6 \cdot 10^{20} \text{ cm}^{-2}$, which is in excess of the average Galactic column density in this direction ($9.6 \cdot 10^{19} \text{ cm}^{-2}$). The spectrum formed from the PC data, spanning from 5.6 to 24.7 ks, is fit with a higher absorbing column of $NH = (2.0 \pm 0.4) \cdot 10^{21} \text{ cm}^{-2}$.

No flares could be seen in the XRT data.

The source was also detected by the UVOT [?, ?] starting with the first exposure at $T_0 + 58 \text{ s}$. The optical transient was visible in all seven filters. The initial brightness had magnitude 17 in the white filter. The temporal slope in this filter band out of 30400 s was approximately $\alpha = 0.23$.

Optical observations

Various optical telescopes performed follow-up observations of this event and found an afterglow counterpart. We mention here the observation from the 2.2 m Calar Alto telescope which measured a redshift of $z = 0.75$. This redshift was confirmed by various other optical telescopes and refined to $z = 0.767$. 
6.1.2 Observation delays

- Burst onset time $T_0$: 19:53:02 UTC
- Alert received: 19:53:15 UTC
- Alert delay: 13 s
- Start of the observation: 21:12:14 UTC
- End of the observation: 23:52:30 UTC
- Delay of MAGIC observation to $T_0$: 4753 s
- Total duration of the MAGIC observation: 9616 s
- Zenith angle at the begin: 23.3 deg.

The GCN trigger was received by *gspot* 13 s after the burst onset $T_0$. The burst onset was before the astronomical dark time. Therefore, at the the time of the alert, no data taking with the MAGIC telescope was ongoing yet.

The shift crew accepted the alert at 21:12:14 UTC, 4739 s after the alert income.
6.2 Description of the Data used for analysis

The observational conditions were not optimal during the data taking. Especially starting from 22:56:32 UTC claudiness values > 50%, measured with the pyrometer and representing the cloud coverage in the FoV of MAGIC, required the rejection of 40 min of the data. The analysis and results are split therefore into two data sets respectively:

From 21:12:14 UTC (RUN NUMBER 349193, $\theta = 23^\circ$) to 22:56:32 UTC (RUN NUMBER 349266, $\theta = 27^\circ$), with the corresponding OFF data:

<table>
<thead>
<tr>
<th>source name</th>
<th>runs</th>
<th>date</th>
</tr>
</thead>
<tbody>
<tr>
<td>OffAgi-Igr-1</td>
<td>314373-314413</td>
<td>20071215</td>
</tr>
<tr>
<td>1ES2344+514-W0.40+000</td>
<td>315820-315822</td>
<td>20071227</td>
</tr>
<tr>
<td>1ES2344+514-W0.40+180</td>
<td>315823-315834</td>
<td>20071227</td>
</tr>
<tr>
<td>1ES2344+514-W0.40+000</td>
<td>315835-315858</td>
<td>20071227</td>
</tr>
<tr>
<td>1ES2344+514-W0.40+180</td>
<td>315859-315866</td>
<td>20071227</td>
</tr>
<tr>
<td>B3-1009+427-W0.40+180</td>
<td>327923-327941</td>
<td>20080114</td>
</tr>
<tr>
<td>B3-1009+427-W0.40+000</td>
<td>327946-327949</td>
<td>20080114</td>
</tr>
</tbody>
</table>

and from 23:34:54 UTC (RUN NUMBER 349289, $\theta = 31^\circ$) to 23:52:30 UTC (RUN NUMBER 349300, $\theta = 34^\circ$) with the corresponding OFF data:
The muon ring analysis yielded a PSF of $\sigma = 10.3\, \text{mm}$ for all ON data. Gamma MC files, simulated with a PSF of $\sigma = 10.7\, \text{mm}$ were used.

### 6.2.1 Observation conditions during data taking

Fig. 6.3 and 6.4 show the event rate (after melibea) for the individual analysis periods. Fig. 6.5 and 6.6 show a couple of parameters to characterize the GRB observation conditions. Because the discriminator threshold were set higher than the usual dark night values on all data sets, the following efficiency corrections had to be applied:

- From 21:12:14 UT to 22:56:32 UT: $\epsilon = 0.90$
- From 23:34:54 UT to 23:52:30 UT: $\epsilon = 0.90$

Using the relations following relations for the corrections for the different discriminator thresholds and zenith angles,

$$E_{\text{thr}}(\theta) = E_{\text{thr}}(\text{zenith}) \cdot (\cos \theta)^{-2.7}$$
$$E_{\text{thr}}(DT) = E_{\text{thr}}(DT = 14) + 0.41 \cdot (< DT > -14)$$

we expect the following average energy thresholds:

- From 21:12:14 UT to 22:56:32 UT: $E_{\text{thr}} = 112.5\, \text{GeV}$
- From 23:34:54 UT to 23:52:30 UT: $E_{\text{thr}} = 143.0\, \text{GeV}$
Unfortunately, the average cloudiness parameter was sometimes high also in the remaining two analysis periods ($<\text{Cloudiness}>=44.0$). We do not yet know how to correct for the effective absorption in the atmosphere!

Figure 6.3: Event rates (after melibe) for the first analysis time slot from 21:12:14 UTC to 22:56:32 UTC. The arrow indicates the time when the telescope was pointing to the correct position.

### 6.3 Analysis of the MAGIC data set

The data was taken in ON mode. The Random Forest method was used to separate the $\gamma$ from hadron events, using Time Image Cleaning algorithm and time parameters. Because the high cloudiness value in the middle of the observation period, the data sample had to be split into two parts: from 21:12:14 UTC ($\theta = 23^\circ$) to 22:56:32 UTC ($\theta = 27^\circ$) and from 23:34:54 UTC ($\theta = 31^\circ$) to 23:52:30 UTC ($\theta = 34^\circ$). The data quality check rejected therefore 40 min of the data set.

#### 6.3.1 Filter cuts for background reduction

The following precuts on Hillas parameters were applied in osteria:
Figure 6.4: Event rates (after melibea) for the second analysis time slot from 23:34:54 UT to 23:52:30 UT.

- **Precuts:**

  \[ \text{LEAKAGE} < 0.2 \]
  \[ \text{COREPIXELS} > 2 \]
  \[ \text{ISLANDS} < 4 \]
  \[ \text{SIZE} > 60 \text{ ph.e.} \]

- **Spark cut:**

  \[ 1.5 - 4 \cdot \log(\text{CONC}) > \log(\text{SIZE}) \]

- **Car flashes cut:**

  \[ \log \left( \frac{\text{WIDTH} \cdot \text{LENGTH}}{297 \cdot \text{DIST}} \right) > -0.3 \]

Finally, after the RF training, the following filter cuts were chosen:

- **Cut on the DIST parameter:**

  \[ 0.2^\circ < \text{DIST} < 1.15^\circ \]
6.3. Analysis of the MAGIC data set

Figure 6.5: Observation conditions for the first analysis time slot from 21:12:14 UTC to 22:56:32 UTC. The upper left plot shows the mean discriminator threshold settings (the reference line the default setting for dark nights). The upper right plot shows the nominal and current telescope pointing zenith angle as a function of time. The lower left plot shows the nominal and current telescope pointing azimuth angle as a function of observation time. The arrows indicate the time when the telescope was pointing to the correct position. One can see that at the beginning of the observation, the telescope was not yet pointing to the source. The lower right plot shows the cloudiness parameter of the pyrometer (the reference line shows a typical acceptable value).

- **Further cuts:**

  \[ ISLANDS < 2 \]

  Because the re-zenithing is still not yet as trustable as we would like to have it, osteria applied only a re-sizing of the hadron distribution. Later, no explicit zenith angle was used as a training parameter for random forest. However, the MC was chosen such that the zenith angle ranges match well.
Figure 6.6: Observation conditions for the second analysis time slot from 23:34:54 UTC to 23:52:30 UTC. The upper left plot shows the mean discriminator threshold settings (the reference line the default setting for dark nights). The upper right plot shows the nominal and current telescope pointing zenith angle as a function of time. The lower left plot shows the nominal and current telescope pointing azimuth angle as a function of observation time. The arrows indicate the time when the telescope was pointing to the correct position. One can see that at the beginning of the observation, the telescope was not yet pointing to the source. The lower right plot shows the cloudiness parameter of the pyrometer (the reference line a typical acceptable value).

6.3.2 Results from the ALPHA plot analysis

An ALPHA plot analysis was performed for the two observation time sets. No significant signal could be found in these time periods. Fig. 6.7 and 6.8 show the corresponding ALPHA plots for the different energy bins. The upper limits were calculated using the Rolke method with 95% confidence level and 30% systematic uncertainty on the telescope efficiency. Because the discriminator threshold were set higher than the usual dark night values on all data sets, the following efficiency corrections had to be applied:
6.3. Analysis of the MAGIC data set

From 21:12:14 UT to 22:56:32 UT: $\epsilon = 0.90$
From 23:34:54 UT to 23:52:30 UT: $\epsilon = 0.90$

Upper limits on the gamma-ray emission during the observed time period were derived. The results are shown in tables 6.1 and 6.2.

6.3.3 Results from the MAGIC Light Curve analysis

Light curves of 10 min bins were calculated for different energy bins. Because of the short duration of the second time bin, light curves for the first time bin are only shown here. Also this attempt did not reveal any significant excess over background. Figure 6.9 shows the
Figure 6.8: ALPHA plots for the second analysis time slot from 23:34:54 UTC to 23:52:30 UTC for different energy bins.

Table 6.1: Upper limits on the VHE γ-ray emission from GRB080430. The upper limits correspond to 6257 s of the MAGIC observation from 21:12:00 UTC to 22:56:00 UTC.
### 6.3. Analysis of the MAGIC data set

<table>
<thead>
<tr>
<th>$E$  [GeV]</th>
<th>$\langle E \rangle$ [GeV]</th>
<th>HADRONNESS cut</th>
<th>ALPHA cut</th>
<th>Upper Limits [ph cm$^{-2}$ keV s]</th>
<th>[erg cm$^{-2}$]</th>
<th>C.U.</th>
</tr>
</thead>
<tbody>
<tr>
<td>125-175</td>
<td>141.4</td>
<td>0.40</td>
<td>10</td>
<td>7.5e-19</td>
<td>2.5e-08</td>
<td>0.18</td>
</tr>
<tr>
<td>175-300</td>
<td>223.2</td>
<td>0.20</td>
<td>7</td>
<td>3.5e-19</td>
<td>3e-08</td>
<td>0.27</td>
</tr>
<tr>
<td>300-1000</td>
<td>484.8</td>
<td>0.15</td>
<td>5</td>
<td>9.4e-21</td>
<td>3.7e-09</td>
<td>0.05</td>
</tr>
</tbody>
</table>

Table 6.2: Upper limits on the VHE $\gamma$-ray emission from GRB080430. The upper limits correspond to 1056 s of the MAGIC observation from 23:34:01 UTC to 23:52:00 UTC.

Figure 6.9: Light curves for the first analysis time slot from 21:12:14 UTC to 22:56:32 UTC for different energy bins. In red the number of excess events, in blue the evolution of the background is shown.
6.3.4 Search for a peaked emission phase in the MAGIC data set

In addition to the standard analysis method the data was divided into 72 equal time intervals of 100 s each. By sliding the intervals we searched for a possible delayed VHE short emission component. The remaining ON data was used for background subtraction. The same procedure was done with a 50 s phase shift. Fig. 6.10 and 6.11 show number of excesses vs. time, for different energy bins and the 2 analysis time periods, without phase shift, while Fig. 6.12 through 6.13 show the corresponding distributions of significances from both original and phase shifted distributions. All significances spread around zero with a width compatible with one. From each individual time slot, an upper limit on the γ-ray emission during that period can be derived. The worst of these limits yields the results shown in tables 6.3 through 6.4

<table>
<thead>
<tr>
<th>$E$ [GeV]</th>
<th>$\langle E \rangle$ [GeV]</th>
<th>HADRONNESS</th>
<th>ALPHA cut</th>
<th>Upper Limits</th>
</tr>
</thead>
<tbody>
<tr>
<td>80-125</td>
<td>95.2</td>
<td>0.50</td>
<td>15</td>
<td>5.3e-17</td>
</tr>
<tr>
<td>125-175</td>
<td>141.4</td>
<td>0.40</td>
<td>10</td>
<td>2.1e-17</td>
</tr>
<tr>
<td>175-300</td>
<td>223.2</td>
<td>0.20</td>
<td>7</td>
<td>3.1e-18</td>
</tr>
<tr>
<td>300-1000</td>
<td>484.8</td>
<td>0.15</td>
<td>5</td>
<td>2.4e-19</td>
</tr>
</tbody>
</table>

Table 6.3: Upper limits on the VHE γ-ray emission from GRB080430. The upper limits correspond to any 100 s time interval in the MAGIC observation from 21:12:00 UTC to 22:56:00 UTC.

<table>
<thead>
<tr>
<th>$E$ [GeV]</th>
<th>$\langle E \rangle$ [GeV]</th>
<th>HADRONNESS</th>
<th>ALPHA cut</th>
<th>Upper Limits</th>
</tr>
</thead>
<tbody>
<tr>
<td>125-175</td>
<td>141.4</td>
<td>0.40</td>
<td>10</td>
<td>1e-17</td>
</tr>
<tr>
<td>175-300</td>
<td>223.2</td>
<td>0.20</td>
<td>7</td>
<td>1.5e-18</td>
</tr>
<tr>
<td>300-1000</td>
<td>484.8</td>
<td>0.15</td>
<td>5</td>
<td>1.8e-19</td>
</tr>
</tbody>
</table>

Table 6.4: Upper limits on the VHE γ-ray emission from GRB080430. The upper limits correspond to any 100 s time interval in the MAGIC observation from 23:34:01 UTC to 23:52:00 UTC.
Figure 6.10: Excess events vs. observation time for the first analysis time slot from 21:12:14 UTC to 22:56:32 UTC for different energy bins.
Figure 6.11: Excess events vs. observation time for the second analysis time slot from 23:34:54 UTC to 23:52:30 UTC for different energy bins.
Figure 6.12: Distribution of significances for the first analysis time slot from 21:12:14 UTC to 22:56:32 UTC for different energy bins.
Figure 6.13: Distribution of significances for the second analysis time slot from 23:34:54 UTC to 23:52:30 UTC for different energy bins.
6.4 Perspective for the late afterglow observation

The results for the first time interval, giving the lower energy upper limit, were used in [37] to evaluate perspective for late-afterglow observation. The article is reported in the next sections.

6.4.1 Afterglow light-curve and spectral energy distribution

Preliminary results shows this afterglow can not be satisfactory described within any commonly referred scenario. In particular the early afterglow is puzzling likely requiring two distinct components, one begin the regular afterglow, with separated time evolution. However, at the epoch of the MAGIC observations (about 8 ksec from the high-energy event), the afterglow seems to have entered a rather stable phase. Analysis of the spectral (from the optical to X-rays) information shows that the afterglow can be well described as due to the interaction of a relativistic outflow with the circumburst medium surrounding the progenitor [186, 247]. The outflow is relativistic and shocks form with consequent particle acceleration. Details of the acceleration process are not known and it is usually assumed that electrons follow a power-law distribution in energy with a slope $p$. The late-afterglow of GRB 080430 can be characterized by a constant circumburst density environment with typical number density $n \sim 1 \, \text{cm}^{-3}$. The electron distribution index turns out to be $p \sim 2.1$. Given the afterglow spectral properties, it is possible to predict the time decay which, in the optical, is well consistent with the predictions. On the contrary, X-ray data\footnote{http://swift.gsfc.nasa.gov/docs/swift/archive/grb table.html} show a much milder decay than expected. It is difficult to attribute this behavior to a specific physical ingredients. Common additions to the reference model [246] which may or may not modify the spectrum involve late energy injection, structured jets, flares, circumburst density variations, etc. (see e.g. [181, 248], for comprehensive discussions about these factors). We model the VHE emission assuming the afterglow can be described in the context of the standard afterglow model [186, 247]. Then in the discussion we comment possible modifications induced by additional phenomena which in general can even increase the expected VHE flux.

In order to characterize the afterglow spectrum we must compute the synchrotron injection, $\nu_m$, and cooling, $\nu_c$, frequency values. The injection frequency is wheremost of the synchrotron emission occurs and the cooling frequency identifies where electrons cool effectively. In case
of constant circumburst medium [86,244] we have:

\[ \nu_m \approx 4.3 \times 10^{14} C_p^{-1/2} \epsilon_e^{2/2} \epsilon_B^{1/2} E_k^{1/2} t_3^{-3/2} (1 + z)^{1/2} \text{Hz} \] (6.2)

\[ \nu_c \approx 1.8 \times 10^{16} \epsilon_B^{-3/2} n^{-1} E_k^{-1/2} t_3^{-1/2} (1 + z)^{-1/2} \text{Hz} \] (6.3)

where \( z \) is the redshift of the source, \( n \) the medium particle density, \( E_k \) the kinetic energy of the outflow, \( t \) the time delay after the GRB and \( C_p = 13(p - 2)/[3(p - 1)] \). We will assume for the micro-physical parameters \( \epsilon_e \), the fraction of total energy going to electrons, and \( \epsilon_B \), the fraction of total energy going to magnetic fields, the values 0.1 and 0.01, respectively. These are typical values measured during late-time afterglow and in agreement with the result analysis of GRB 080430. The relation for the cooling frequency is indeed approximate since we are neglecting the possible role played by additional inverse Compton (IC) cooling.

We will consider again this issue in next section. The total energy can be derived from the burst isotropic energy \( E_{iso} \) with some assumptions about the spectrum and by correcting it for the fireball radiative efficiency \( \eta \). We estimate \( E_{iso} \) as the integral of the burst spectral model [218] in the \( 1-10^4 \) keV band [41], energy range covering most of the prompt emission of GRBs. In this energy band the spectrum of a burst is typically described by a Band function. According with this method we estimate \( E_{peak} = 39 \) keV and \( E_{iso} = 3 \times 10^{51} \) erg. If we then assume an radiative efficiency \( \eta \) of 10\%, we find the total kinetic energy going to the outflow \( E_{k,iso} = 3 \times 10^{52} \) erg. The prompt phase radiative efficiency can indeed vary among individual bursts [246]. Still lacking of a satisfactory treatment of the prompt emission phase emission process we choose 10\% higher for events characterized by a shallow decay phase [174] at the X-rays as it might be the case for GRB 080430.

6.4.2 SSC during the afterglow

For GRB 080430 the MAGIC observations were performed with a delay of 4753 s from the \( T_0 \), well after the burst onset. In this case we rule out any residual prompt emission and, any dominant flaring activity at the observation epoch. Assuming to be in this simple situation, and once the parameters of the lower-energy synchrotron emission are known, it’s possible to predict the SSC component. We followed the recipe by Fan and Piran [87]. We can assume to know all the required parameters for a modeling of the high energy emission from the afterglow: \( E_{iso} \sim 3 \times 10^{52} \) erg, \( \epsilon_e \sim 0.1 \), \( \epsilon_B \sim 0.01 \), \( p \sim 2.1 \) (assuming some degree of
6.4. Perspective for the late afterglow observation

energy injection), the circumburst density profile \( n \sim 1 \) and the redshift \( z \sim 0.76 \). In the SSC process a new spectral component, superposed to the underlying synchrotron spectrum, with the same global shape up to a cut-off frequency:

\[
\nu_{m,SSC} \sim \frac{\Gamma^2 m_e^2 c^4}{h^2 \nu_c}
\]  

(6.4)

where \( \nu_c \) is the synchrotron cooling frequency, \( \Gamma \) the fireball bulk motion Lorentz factor, \( m_e \) the electron mass, \( c \) the speed of light and \( h \) the Planck constant. Above this frequency the SSC emission is no more in the Thomson regime and becomes much weaker (Klein-Nishina regime). For typical bulk motion Lorentz factors [169] the SSC emission of the afterglow is however in the Thomson regime.

Assuming we are in a constant density circumburst environment, the predicted SSC spectrum is characterized by the cooling frequency \( (\nu_c) \) and the synchrotron typical frequency \( (\nu_m) \):

\[
\nu_{m,SSC} \approx 6.2 \times 10^{21} \text{Hz} C_p^4 \nu_{e,-1}^{1/2} E_{k,53}^{3/4} t_3^{-9/4} (1 + z)^{5/4}
\]  

(6.5)

\[
\nu_{c,SSC} \approx 4 \times 10^{24} \text{Hz} (1 + z)^{-4} \nu_{e,-2}^{-9/4} E_{k,53}^{-5/4} t_3^{-1/4} (1 + z)^{-3/4}
\]  

(6.6)

where \( C_p = 13(p - 2)/[3(p - 1)] \) and \( Y_{SSC} = U_{syn}'/U'_B \) is the rest frame synchrotron to magnetic field energy density ratio and \( z \) the redshift. It can be shown that:

\[
Y_{SSC} \simeq - \frac{1 + \sqrt{1 + 4 \eta \epsilon_e / \epsilon_B}}{2}
\]  

(6.7)

where \( \eta \) is defined \( \eta = (\nu_m/\nu_c)_{(p-2)/2} \).

The peak synchrotron frequency to cooling frequency ratio is:

\[
\nu_m/\nu_c \simeq 0.024 (1 + z) C_p^2 \nu_{e,-1}^2 E_{k,53} t_3^{-1}
\]  

(6.8)

With the assumed parameters \( C_p = 0.39 \), \( \nu_m/\nu_c = 0.00025 \) and \( Y_{SSC} \simeq 2.1 \). The equations 6.5 and 6.6 become \( \nu_{m,SSC} \approx 1.1 \times 10^{18} \text{ Hz} \) (\( \simeq 5 \text{ keV} \)) and \( \nu_{c,SSC} \approx 7.4 \times 10^{22} \text{ Hz} \) (\( \simeq 310 \text{ MeV} \)).

The energy band covered by the MAGIC observation \( (E_{\text{MAGIC}} \sim 90 \text{ GeV}) \) are clearly much higher than the cooling SSC frequency. We are therefore in the spectral range where the SSC spectrum is softer, following a power-law \( \nu^{p/2} \simeq \nu - 1.05 \).

The flux density at the typical SSC frequency is:

\[
F_{\nu_{m,SSC}} \simeq 10^{-11} \text{erg cm}^2 \text{s}^{-1} \text{MeV}^{-1} 0.07 n^{5/4} \nu_{e,-1}^{1/2} E_{k,53}^{5/4} t_3^{1/4} \left( \frac{1 + z}{2} \right)^{3/4} D_{L,28.34}^{-2}
\]  

(6.9)
where $D_L \sim 4.8$ Gpc ($\sim 1.5 \times 10^{28}$ cm). With our parameters, $F_{\nu_m,SSC} \simeq 5.1 \times 10^{-13}$ erg cm$^{-2}$ s$^{-1}$ MeV$^{-1}$.

The expected flux density at the MAGIC observation energy can be calculated by extrapolating the spectrum as:

$$F_{90 \text{GeV}} \sim F_{\nu_m,SSC} \left( \frac{\nu_c,SSC}{\nu_m,SSC} \right)^{-(p-1)/2} \left( \frac{\nu}{\nu_c,SSC} \right)^{-p/2}$$  \hspace{1cm} (6.10)

and, with our parameters, $F_{90 \text{GeV}} \sim 2.9 \times 10^{-18}$ erg cm$^{-2}$ s$^{-1}$ MeV$^{-1}$. The flux integrated in the MAGIC band, the parameter to be compared to the reported upper limits, can be well approximated by $\nu F\nu$ at about 90 GeV, and we have $F_{MAGIC} \sim 2.6 \times 10^{-13}$ erg cm$^{-2}$ s$^{-1}$ at the epoch of the MAGIC observation, $t \sim 8$ ks form the prompt emission. Any uncertainty in the underlying afterglow parameters of course affects the VHE predictions. Some of these uncertainties have however a rather limited (considering the present observational limits) impact because one of the relevant factors, the ratio between the injection and cooling synchrotron frequency, is constrained by the afterglow SED and uncertainties from micro-physical parameters should still keep the ratio close to the observed value. The $\nu_m/\nu_c$ ratio drives the importance of the IC component and the position of the cooling SSC frequency, i.e. where the VHE flux begins to decrease steeply going to higher energies. The total energy on the contrary is estimated assuming an efficiency for the GRB prompt emission process. This is a weakly known factor given that at present no satisfactory description of the GRB prompt emission process exists [151]. It is therefore possible [247] that the efficiency be substantially higher increasing the total energy and therefore the expected flux. Circumburst matter density has an important effect on the expected SSC flux and with the present afterglow data can essentially be estimated only coupled to the micro-physical parameters. A higher density would make the SSC component more important and possibly detectable at lower energies (see e.g. [126]). However, in most cases with data allowing a detailed modeling of the afterglow the derived circumburst density is consistent with the value we report for GRB 080430 [180]. The milder than expected temporal decay at the X-rays together, however, with the consistency of the observed SED with the reference afterglow model prediction, raises some concern about the reliability of the adopted theoretical scenario.
Figure 6.14: Predictions at different time delays from the HE event for the SSC emission during the afterglow of GRB 080430. Black triangles are the 95% CL upper limit derived by MAGIC at various energies. Lines of a same color show the same SSC modle, but different EBL absorption model. The blue lines to correspond to the MAGIC observation window.
6.4.3 EBL attenuation and discussion

Gamma-rays in the GeV energy regime are absorbed through pair-production processes with the EBL. The precise light content of the EBL is strongly debated. We can rely on many different models and predictions of which at $z \sim 1$ span a wide range of optical depths, from less than 1 up to 6 [87]. Moreover, the MAGIC collaboration published a striking observational result [35] suggesting that the EBL attenuation could be much lower than previously assumed. Thus at the redshift of GRB 080430 ($z \sim 0.76$) and at the MAGIC energy ($E \sim 90$ GeV) an optical depth $\tau$ not far from unity is possible. We included four representative models from Kneiske et al. (2004) [141], Franceschini et al. (2008) [97] and Gilmore et al. (2009a) [114] and show the range of possible absorbed spectra in Fig. 6.14. The blue lines correspond to the MAGIC observation delay, the other lines show the spectrum at earlier observation times, in principle easily accessible to IACTs. On average, we can assume an attenuation of the received flux from the afterglow of GRB080430 of the order a factor 3 or even less, allowing us to estimate $F_{\text{MAGIC}} \sim 2.6 \times 10^{-13}$ erg cm$^{-2}$ s$^{-1}$ as the predicted flux in the MAGIC band. As a matter of fact, our choice is possibly very conservative as [114] described models, in agreement with the observations reported in [35], with an optical depth as low as $\tau \sim 0.4$ at the same conditions of these MAGIC observations. The results appear to be well below the reported upper limits. Furthermore, our assumed low opacity for the EBL is in agreement with current observations (see also [114]). At any rate, this pilot case shows fairly interesting perspectives for a late-afterglow detection at high energies. In general, to increase the flux expected from a GRB afterglow (for SSC) it is mandatory to try to decrease the observation energy which is also very important for the minimization of the EBL attenuation. If the telescope sum trigger hardware upgrade had already been implemented before the observations, a limit above an energy of 45 GeV would have been obtained [109]. At these energies, the strong effect of the EBL could probably be neglected and the low energy threshold together with the expected performances of MAGIC-II will undoubtedly increase the chances of positive detections. As a matter of fact, GRB 080430 was an average event in terms of energetics. More energetic GRBs are indeed relatively common, and due to the mild positive dependence on the isotropic energy of a GRB, much higher fluxes than in the present case can be foreseen. This is also true if we consider the uncertainty in the present total energy determination which is based on an average value for the prompt emission efficiency. The time delay of the observation
from the GRB has a clear impact essentially because the observed SSC component is strictly related to the underlying synchrotron component which rapidly decays in intensity with time, depending on the specific environment and micro-physical parameters. Had MAGIC been able to start observations right at the start of the late afterglow phase (e.g. at $T_0 + 1$ ks), the flux predictions would have increased by more than an order of magnitude. The time delay of about two hours, coupled with the mediocre observing conditions, were more than enough to depress the observed flux and raise the reported upper limits.
Upper limits on MAGIC-I GRBs population emission

In no one of the 57 MAGIC-I GRB follow-up observation up to now, significant emission of VHE $\gamma$-rays could be detected. The idea of a stacking analysis is to merge together the information coming from individual GRB observations and try to infer general considerations on MAGIC GRB afterglow data. For sure this is a complicated issue because the uniqueness of each GRB emission is to be taken into account. Nevertheless, using events observed by MAGIC with substantially the same hardware performances during the afterglow phase, some general discussion about the early afterglow emission can be done for the first time.

All the discussion in this chapter will start from the assumption that at the epochs of the MAGIC observations the afterglow emission was in a late phase, so that we can model the VHE emission in the context of the standard afterglow model. Exactly like in the previous chapter we consider as the possible emission process the SSC. In this case, summing up many observation we cannot choose a unique set of parameters to model the general afterglow, but we have to try with many possibilities and compare the different prediction (always from the Fan & Piran 2008 [87]) with the analysis results.

7.1 The selected MAGIC-I GRBs data sample

Generally all observation for 2009 were carried out with the sum trigger 2.4.5. However only GRB090102 had precise coordinates and a measured redshift such that meaningful upper limits could be obtained, with an analysis threshold energy of only $E_{\text{thr}} = 32\text{GeV}$ [109]. We rejected the data sets (the ones not reported in Tab. 7.1) of 13 GRBs in cycle III and IV, due
to a bad quality data, either in the pointing accuracy or the weather conditions. All GBM alerts have generally too large errors to allow physical interpretation of the observations. Fig. 7.1 show the final localization errors from all Fermi-GBM triggers up to May 2009: only in the 6% of the cases the alert coordinates have a precision better than 2 deg. The Fermi-LAT detector is able to pinpoint the GRB sky coordinates with much higher precision, however in this case the more precise coordinates were sent after the end of the MAGIC-I observation and lay outside the FoV of the the camera. A summary of the GRBs observed by MAGIC-I in cycle III and IV, and for which was possible to perform a meaningful analysis, are reported in Tab. 7.1. For all these bursts an analysis similar to the one presented for GRB 080430 had been done and the results already published in GCN circulars [3–6]. I restricted the analysis to only this 8 GRBs because of similar telescope hardware performance, in fact:

- I selected GRBs observed with a similar telescope PSF in order to use the same MC sample in the calculation of the upper limits: PSF=10.7 mm.
- I divided my analysis into 5 sub-analyses, depending on the different zenith angle (ZA) of the observations.

### 7.2 Data analysis and flux upper limit

GRB data used for the sub-analysis and the observation time are:

1. ZA 0-33 : GRB 080430 (7418 s) + GRB 090102 (8191 s) + GRB 090113 (8682 s)
7.2. Data analysis and flux upper limit

<table>
<thead>
<tr>
<th>GRB</th>
<th>alert (sat.)</th>
<th>z</th>
<th>$T_{90}$ (s)</th>
<th>$\Delta \text{Az}$ (deg)</th>
<th>rep. time (s)</th>
<th>ZA (deg)</th>
<th>start obs. (s)</th>
<th>obs. time (s)</th>
</tr>
</thead>
<tbody>
<tr>
<td>080315</td>
<td>BAT</td>
<td>-</td>
<td>65</td>
<td>251</td>
<td>69</td>
<td>32</td>
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<td>290</td>
<td>1736</td>
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<td>BAT</td>
<td>1.51</td>
<td>61</td>
<td>129</td>
<td>-</td>
<td>48</td>
<td>91</td>
<td>1754</td>
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<tr>
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<td>0.758</td>
<td>16</td>
<td>-</td>
<td>-</td>
<td>23</td>
<td>4753</td>
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<td>080603B</td>
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<td>60</td>
<td>-</td>
<td>-</td>
<td>41</td>
<td>5578</td>
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<td>-</td>
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<td>8430</td>
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</tr>
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<td>-</td>
<td>5</td>
<td>4603</td>
<td>9405</td>
</tr>
</tbody>
</table>

Table 7.1: Summary of GRB events observed by the MAGIC telescope in the III and IV observation cycle and used in this work. The columns from left: burst name, triggering satellite, redshift, bursts duration, repositioning distance in azimuth direction, required repositioning time, zenith distance at the beginning of the MAGIC observation, begin of the MAGIC observation respect to the onset $T_0$ and the total observation time with MAGIC.

2. ZA 34-40: GRB 090102 (1395 s) + GRB 080319A (1736 s) + GRB 080315 (3025 s)

3. ZA 41-47: GRB 090102 (1325 s) + GRB 080603 (5619 s) + GRB 080903 (2978 s)

4. ZA 48-52: GRB 090102 (1366 s) + GRB 080330 (999 s) + GRB 080903 (1345 s)

5. ZA 52-60: GRB 080330 (769 s) + GRB 080903 (1570 s)

We have to remember that also after this selection, the telescope performances were not exactly the same and that we are summing up different GRBs, at different $z$, observed at different delay times from the prompt.

I weighted the number of ON and OFF (Non and Noff) depending from the source redshift (a factor $z^2$). For the bursts with unknown $z$ I used the mean value of the known ones (1.55). The calculated new ON and OFF numbers and upper limits results reported in tab. 7.2, 7.3. I considered the standard analysis energy ranges: 80-125 GeV, 125-175 GeV, 175-300 GeV and 300-1000 GeV. In Fig. 7.2 the (95% confidence level) flux upper limit are shown. The
Table 7.2: Resulting $N_{ON}$, $N_{OFF}$ and signal significance for the different sub-analysis.

upper limit are calculated with the assumption of steady emission and a 30% of systematic uncertainty on the absolute detector efficiency.
7.2. Data analysis and flux upper limit

Figure 7.2: 95% CL upper limits for the different ZA. The limit is reported in the mean energy of the bin (calculated from MC data). The considered energy bin are (depending from the zenith range): 80-125 GeV, 125-175 GeV, 175-300 GeV and 300-1000 GeV.
<table>
<thead>
<tr>
<th>Zd range (deg)</th>
<th>E (GeV)</th>
<th>mean E (GeV)</th>
<th>UL (counts)</th>
<th>Flux UL (erg cm(^{-2}) s(^{-1}))</th>
<th>Fluence UL (erg cm(^{-2}))</th>
</tr>
</thead>
<tbody>
<tr>
<td>0-33</td>
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<td>0-33</td>
<td>175-300</td>
<td>220</td>
<td>199</td>
<td>(1.2 \times 10^{-15})</td>
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</tr>
<tr>
<td>34-40</td>
<td>125-175</td>
<td>143</td>
<td>0.02</td>
<td>(1.4 \times 10^{-15})</td>
<td>(4.5 \times 10^{-11})</td>
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<tr>
<td>34-40</td>
<td>175-300</td>
<td>231</td>
<td>0.005</td>
<td>(1.2 \times 10^{-15})</td>
<td>(1.3 \times 10^{-11})</td>
</tr>
<tr>
<td>40-47</td>
<td>125-175</td>
<td>141</td>
<td>0.0003</td>
<td>(6.4 \times 10^{-17})</td>
<td>(7.4 \times 10^{-13})</td>
</tr>
<tr>
<td>40-47</td>
<td>175-300</td>
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<td>0.22</td>
<td>(7.8 \times 10^{-14})</td>
<td>(1.8 \times 10^{-11})</td>
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<td>48-52</td>
<td>175-300</td>
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<td>0.00005</td>
<td>(2.1 \times 10^{-17})</td>
<td>(1.2 \times 10^{-13})</td>
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<tr>
<td>48-52</td>
<td>300-1000</td>
<td>498</td>
<td>22</td>
<td>(7 \times 10^{-12})</td>
<td>(5.6 \times 10^{-8})</td>
</tr>
<tr>
<td>53-60</td>
<td>300-1000</td>
<td>539</td>
<td>0.06</td>
<td>(2.5 \times 10^{-14})</td>
<td>(1.4 \times 10^{-10})</td>
</tr>
</tbody>
</table>

Table 7.3: Resulting 95% CL upper limits for the different sub-analysis.
7.3 Discussion

For discussion sake, with the above obtained flux ULs I implement the model by Fan & Piran [87] we used for the GRB 080430 to more broadly explore the models parameters space. Predictions for different time delays from the high energy event for the SSC emission during a classical afterglow are reported for different combinations of the modeling parameters of $E_{\text{iso}}$, and $z$ ($\epsilon_e=0.1$, $\epsilon_B=0.01$, $p=2.3$, $n = 1$ cm$^{-3}$) in Fig. 7.3, 7.4, 7.5 (predictions at 90 GeV) and 7.6, 7.7, 7.8 (predictions at 30 GeV).

Furthermore at redshift $z > 1$, high EBL absorption dampens the incoming VHE flux. In this context the obtained low UL results are only confirming we were likely observing OFF VHE-$\gamma$ regions (and confirm the unknown redshifts of some analyzed GRB were high). In any case we want to stress the importance to have the possibility to perform, with the stereo system, observations with a lower energy threshold. We can confirm the really strong importance to continue GRB observation with MAGIC. If at the end we’ll have the possibility to catch a near GRB, observing it at a relatively low zenith angle (e.g GRB 080319B that was observed by MAGIC only in twilight), we can for sure say something about SSC emission afterglow models.

Last but not least we can suggest the possibility to made a similar time-dependent analysis, in order to have a real comparison with the model’s full predictions. Possible VHE time-depending features, can be amplified with a stacked analysis of data for GRB afterglows of similar temporal decay index and appear in the light-curve. A work like this, was not possible, because of the low number of observations at similar time delay.
Figure 7.3: Predictions of SSC emission at 90 GeV at different time delays. $z = 1$, $E_{\text{iso}}$: 1) $E_{\text{iso}} = 0.1 \times 10^{53}$ erg; 2) $E_{\text{iso}} = 0.5 \times 10^{53}$ erg; 3) $E_{\text{iso}} = 1 \times 10^{53}$ erg.
Figure 7.4: Predictions of SSC emission at 90 GeV at different time delays. $z = 1.55$, $E_{\text{iso}}$.
1) $E_{\text{iso}} = 1 \times 10^{53}$ erg; 2) $E_{\text{iso}} = 0.5 \times 10^{53}$ erg; 3) $E_{\text{iso}} = 0.1 \times 10^{53}$ erg;
Figure 7.5: Predictions of SSC emission at 90 GeV at different time delays. $z = 0.5$, $E_{\text{iso}}$:
1) $E_{\text{iso}} = 0.1 \times 10^{53}$ erg; 2) $E_{\text{iso}} = 0.5 \times 10^{53}$ erg; 3) $E_{\text{iso}} = 1 \times 10^{54}$ erg;
7.3. Discussion

Figure 7.6: Predictions of SSC emission at 30 GeV at different time delays. $z = 1$, $E_{\text{iso}}$: 1) $E_{\text{iso}} = 0.1 \times 10^{53}$ erg; 2) $E_{\text{iso}} = 0.5 \times 10^{53}$ erg; 3) $E_{\text{iso}} = 1 \times 10^{53}$ erg;
Figure 7.7: Predictions of SSC emission at 30 GeV at different time delays. $z = 1.55$, $E_{\text{iso}}$:
1) $E_{\text{iso}} = 0.1 \times 10^{53}$ erg; 2) $E_{\text{iso}} = 0.5 \times 10^{53}$ erg; 3) $E_{\text{iso}} = 1 \times 10^{53}$ erg;
Figure 7.8: Predictions of SSC emission at 30 GeV at different time delays. $z = 0.5$, $E_{\text{iso}}$:

1) $E_{\text{iso}} = 0.1 \times 10^{53}$ erg; 2) $E_{\text{iso}} = 0.5 \times 10^{53}$ erg; 3) $E_{\text{iso}} = 1 \times 10^{53}$ erg;
During my PhD work I was also one of the responsible analyzers for the article: MAGIC collaboration, *MAGIC upper limits to the VHE gamma-ray flux of 3C454.3 in high emission rate*, Astron. Astrophys., Vol. 498, pages 83-87.

In the article are reported the upper limits to the Very High Energy emission ($E > 100$ GeV) of the flat spectrum radio quasar 3C454.3 ($z = 0.859$) derived by MAGIC during the high states of July/August and November/December 2007. The upper limits derived in July are compared with the available quasi-simultaneous MeV-GeV emission observed by the AGILE γ-ray satellite and an interpretation of the observational results in the context of the leptonic emission models is done.

The source was observed during the active phases of July-August 2007 and November-December 2007 and the data were analysed with the standard methods. For the period around the end of July, characterized by with the most complete multifrequency coverage, we constructed the spectral energy distribution using our data together with nearly simultaneous optical, UV, X-ray and GeV data. Only upper limits could be derived from the MAGIC data.

The upper limits of July, once corrected for the expected absorption by the extragalactic background light, together with nearly simultaneous optical, UV, X-ray and GeV data, allow us to constrain the VHE emission of 3C454.3. The data are consistent with the expectation for the model based on the inverse Compton scattering of the ambient photons from the broad line region by relativistic electrons, which robustly predicts a sharp cut-off above 20-30 GeV.

An adapted shorter version of the article is reported in this appendix.
A.1 Introduction

3C454.3 is a well known FSRQ, detected several times in the $\gamma$-ray band by the EGRET telescope onboard CGRO, with an average photon index of $\Gamma = 2.2$ (Hartmann et al. 1999). In 2005 it underwent a very active phase in optical and X-ray bands, triggering intensive observations in the radio, optical and X-ray (Swift, Chandra, INTEGRAL) bands (Villata et al. 2006, Giommi et al. 2006, Pian et al. 2006). Unfortunately no $\gamma$-ray satellite was operating at that time and no information in GeV band was obtained. During the summer of 2007, 3C454.3 was active again, reaching a level of the optical emission comparable to 2005. Several observations in the optical, X-ray and $\gamma$-ray band were activated (optical: KVA, optical-UV: UVOT onboard Swift, X-ray: XRT onboard Swift, GeV band: AGILE). The AGILE satellite (Tavani et al. 2008), still in the commissioning phase, detected intense emission from 3C454.3 (Vercellone et al. 2008).

Triggered by these observations, the MAGIC Cherenkov telescope started observations of 3C454.3 on July 18 and observations were prolonged until July 21. Then another intensive set of observations with MAGIC was performed in August. Another $\gamma$-ray active phase was recorded by AGILE in November-December 2007 (Chen et al. 2007, Pucella et al. 2007), which triggered another set of observations with MAGIC during that period. In all these observations, the source was not detected and only upper limits can thus be derived. In the following (Section A.2) we describe the MAGIC observations and the analysis procedure. In Sect. A.3 we interpret the results in the framework of the widely assumed External Compton (e.g. Sikora et al. 1994) model. Since only the AGILE measurements taken in July are public we limit the discussion to that period.

A.2 MAGIC observations and data analysis

The observation was carried out in ON/OFF mode, in which, the source was observed on axis while observations from a region in the sky from where no gamma rays are expected to come were used to estimate the background, during July 2007 (total 9.3 hours) and August 2007 (total 9.6 hours). Later, in November and December 2007, new observations were performed in the false-source tracking (wobble) mode (Fomin et al. 1994) in which the telescope pointed alternatively for 20 minutes to two opposite sky positions at 0.4° off the source, in November
and December 2007 (total 6.8 hours). The zenith angle of all these observations ranged from 12 to 30 degrees.

The analysis was performed using the standard MAGIC analysis software (Bretz et al. 2005). After calibration and two levels of image cleaning tail cuts (for image core and boundary pixels, see Fegan 1997), the camera images were parameterized by the so-called Hillas image parameters (Hillas 1985). Two additional parameters, namely the time gradient along the main shower axis and the time spread of the shower pixels, were computed (Tescaro et al. 2007). Hadronic background suppression was achieved using the Random Forest (RF) method (Breiman 2001), in which for each event the so-called Hadronness is computed, based on the Hillas and the time parameters. The "Hadronness" parameter can be calculated for every event, which is a measure of the probability that the event is not γ like. The RF method was also used for the energy estimation. The Crab Nebula data from the same periods and zenith angle distributions were studied using the same analysis chain to check the validity of the results.

Since there was no significant signal found, upper limits (95% CL) were calculated (Rolke et al. 2005) taking into account 30% of systematic error. Table A.1 shows the results in July-August observations, which we compare to the AGILE observation in this work. Also the derived upper limits for November-December observation are given (Table A.2).

<table>
<thead>
<tr>
<th>⟨E⟩ [GeV]</th>
<th>U.L. July [erg cm⁻² s⁻¹]</th>
<th>U.L. August [erg cm⁻² s⁻¹]</th>
</tr>
</thead>
<tbody>
<tr>
<td>83</td>
<td>0.04 0.78 × 10⁻¹¹</td>
<td>0.02 0.3 × 10⁻¹¹</td>
</tr>
<tr>
<td>186</td>
<td>0.05 0.62 × 10⁻¹¹</td>
<td>0.03 0.3 × 10⁻¹¹</td>
</tr>
<tr>
<td>476</td>
<td>0.03 0.169 × 10⁻¹¹</td>
<td>0.01 0.09 × 10⁻¹¹</td>
</tr>
</tbody>
</table>

Table A.1: Derived upper limits on flux for July’s and August’s data. The columns represent respectively: the average true energy, the flux upper limit in Crab Unit (C.U.) and [erg cm⁻² s⁻¹].
Table A.2: Derived upper limits on flux for wobble data. The columns represent respectively: the average true energy, the flux upper limit in Crab Unit (C.U.) and [erg cm$^{-2}$ s$^{-1}$].

<table>
<thead>
<tr>
<th>$\langle E \rangle$ [GeV]</th>
<th>U.L. Nov. &amp; Dec. [C.U.]</th>
<th>[erg cm$^{-2}$ s$^{-1}$]</th>
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<tr>
<td>113</td>
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</tr>
<tr>
<td>235</td>
<td>0.09</td>
<td>$0.9 \times 10^{-11}$</td>
</tr>
</tbody>
</table>

A.3 Discussion

The SED of 3C454.3 around the epoch of the July MAGIC observations, assembled with the available data, is shown in Figure A.1. We report the nearly simultaneous data in the optical (KVA), optical-UV (UVOT onboard Swift), X-ray (XRT onboard Swift) and $\gamma$-ray (AGILE) band. For comparison we also show (open circles) historical data (see the Figure caption for references).

In the period July 24-30 AGILE observed an almost constant emission with an average flux above 100 MeV of $F(>100\text{ MeV}) = (280 \pm 40) \times 10^{-8}$ ph cm$^{-2}$ s$^{-1}$ (Vercellone at al. 2008). The slope of the spectrum measured by AGILE is still unpublished (Vercellone et al., in prep). We report the flux at 1 GeV (filled circle) assuming a spectral slope of $\Gamma = 2.5$ with the errorbar indicating the values for slopes in the range $\Gamma = 2 - 3$.

In the same figure, upper limits from observations with MAGIC are shown in triangles (observed: empty; deabsorbed: filled) (see Table A.1). For the deabsorption we used the LowSFR model of Kneiske et al. (2004) which predicts a low level of the EBL close to what is presently inferred from observations (Aharonian et al. 2006b, Mazin & Raue 2007).

One can see from Fig.A.1 that if the $\gamma$-ray (100 MeV-100 GeV) spectrum is relatively hard ($\Gamma < 2.5$) the MAGIC (absorption corrected) upper limit at $\sim$100 GeV is inconsistent with the extrapolation from GeV energies, thus indicating either that the spectrum is very soft ($\Gamma > 2.5$) or there is break (or a cut off) of the emission between the GeV and the 100 GeV band. As discussed below, this is consistent with the expectations from the simplest leptonic model.

Emission from blazars is dominated by the non-thermal continuum of a relativistic jet closely aligned toward the observer. The SED of FSRQs is widely interpreted in terms of synchrotron
A.3. Discussion

Figure A.1: SED of 3C454.3 assembled with multifrequency information available for the period close to the MAGIC observation at the end of July 2007 (optical: KVA, optical-UV: UVOT onboard Swift, X-ray: XRT onboard Swift, GeV band: AGILE). For AGILE we report the flux at 1 GeV (filled circle) assuming a spectral slope of $\Gamma = 2.5$ with the errorbar indicating the values for slopes in the range $\Gamma = 2 - 3$. Triangles report the observed (empty) and the deabsorbed (filled) upper limits of MAGIC in three different bands. For comparison we also report (open circles) historical data (Kuhr et al. 1981, NED, Gear et al. 1994, Stevens et al 1994, Impey & Neugebauer 1988, Smith et al. 1988 for radio and optical; Tavecchio et al. 2007 for X-rays from Chandra). The open circle and the bow-tye in the MeV-GeV region indicates the average EGRET spectrum (Hartman et al. 1999). Solid and long dashed lines report the results of the modelling with the synchrotron-inverse Compton model (see text for details and model parameters). The dotted line shows the emission from the accretion disk, while the spikes (solid line) around $10^{15}$ Hz shows the emission lines produced by the broad line region, used as soft photons for the inverse Compton process (see Tavecchio & Ghisellini 2008 for details).
and inverse Compton emission from high-energy electrons (leptonic models). The latter component is probably dominated by the scattering of the external photons (originating in the disk and/or in the broad line region [BLR], Sikora et al. 1994), though the synchrotron self-Compton emission (Maraschi et al. 1992) can significantly contribute in the X-ray band.

The SED of 3C454.3, including optical, X-rays and GeV measurements around the end of July, has been already discussed and modelled by Ghisellini et al. (2007). Here (solid line in Figure 1) we report a similar model. However, given the focus on the VHE emission we used a more refined calculation, including the full Klein-Nishina (KN) cross section for the IC scattering, and also considering the absorption of $\gamma$-ray photons through pair production within the BLR. Moreover, the external radiation field (assumed to be isotropic in the frame of the black hole), usually approximated by a black body peaking in the UV region, has been calculated using the photoionization code CLOUDY (Ferland et al. 1998). Details on the emission model can be found in Maraschi & Tavecchio (2003), while the description of the calculation of the external radiation field is reported in Tavecchio & Ghisellini (2008). Briefly, we assume that the emission is produced within a spherical region of radius $R = 5 \times 10^{16}$ cm, in motion with bulk Lorentz factor $\Gamma = 18$ at an angle $\theta = 3.2$ deg with respect to the line of sight. The tangled magnetic field has an intensity $B = 4.2$ G. The emitting particles, with total density $n = 1.7 \times 10^4$ cm$^{-3}$, follow a (steady state) broken-power law energy distribution extending from $\gamma_1 = 1.5$ to $\gamma_2 = 5 \times 10^3$, with indices $n_1 = 1.7$ and $n_2 = 3.25$ below and above the break at $\gamma_b = 16$. This purely phenomenological distribution has been assumed to reproduce the observed shape of the blazar SEDs, without any specific assumption on the acceleration/cooling mechanism acting on the particles. With this choice we are allowed to assume extreme low-energy slopes ($n_1 < 2$) such as those required for 3C454.3, which cannot be obtained under standard conditions. It is conceivable that, at least in these cases, the electron distribution derives from two (continuously operating) different acceleration mechanisms (see e.g. Sikora et al. 2002). We also neglect the effects related to the cooling of particles in the KN regime, discussed by Moderski et al. (2005). We note, however, that these effects should produce a bump in the optical-UV synchrotron emission which is not apparent in the available data, though the poor coverage does not allow a firm conclusion. We model the external radiation field assuming that the disk emission (dotted line in figure), with a total luminosity of $L_{\text{disk}} = 5 \times 10^{46}$ erg/s, is reprocessed by clouds of the BLR, a sphere
with radius $3 \times 10^{17}$ cm (we assume that clouds are characterized by standard values of the density $n_{BLR} = 10^{11}$ cm$^{-3}$ and hydrogen column density, $N_H = 10^{23}$ cm$^{-2}$). The model (with parameters similar to those obtained using the self-consistent particle distribution of Ghisellini et al. 2007) allows us to reproduce reasonably well the multiwavelength SED of 3C454.3.

The rapid decrease of the emission above few tens of GeV is related to two effects: i) the decrease of the scattering cross section and ii) the absorption of the produced $\gamma$-rays through pair production. The energy above which the KN effects become important can be roughly expressed as: $E_{KN} \simeq 22.5\nu_{\alpha,15}^{-1}$ GeV, where $\nu_{\alpha,15}$ is the frequency of the external photons (in units of $10^{15}$ Hz). The emission including only the KN effects, neglecting the absorption, is shown by the long-dashed line. The frequency above which the absorption of $\gamma$-rays become effective can be roughly expressed as: $E_{\gamma\gamma} \simeq 60\nu_{t,15}^{-1}$ GeV, where $\nu_{t,15}$ is the frequency of the target photons (in units of $10^{15}$ Hz). Therefore, as shown by the solid line in Fig.A.1 (calculated including both effects), the expected emission above 20-30 GeV is rather small, consistently with the observed upper limits. Note that, although the limit set by KN effects is a characteristic feature of leptonic models, absorption of $\gamma$-rays by soft photons can also be relevant for hadronic models (e.g. Reimer 2007).

Our observations, the first of a FSRQs at VHE nearly simultaneous with observation in the MeV-GeV band, show that even upper limits can be rather useful to test current emission models.

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A.3. Discussion


A. 3c454.3 FSRQ analysis
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