Observations of Transient Events with Very-High-Energy Gamma-Ray Telescopes

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Abstract

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Astrophysical events that evolve on short timescales (from milliseconds to years) are widely referred as transient events. In many cases, transient events are explosions or mergers of astrophysical objects that emit particles of all energies. This thesis focuses on very-high-energy (VHE; 100 GeV to 100 TeV) gamma rays, observed by the VERITAS telescope, to understand two types of transients, superluminous supernovae (SLSNe) and classical novae. In the first part, the background physics and technical approach of an imaging atmospheric Cherenkov technique deployed by VERITAS is reviewed in depth, including the analysis pipeline of VERITAS data, from camera data reduction to high level analysis output. In addition to supporting the ongoing work at VERITAS, the second part of this thesis describes the extensive effort to develop, commission and align the optical system of the prototype Schwarzschild-Couder telescope, also located at the VERITAS observatory. This new telescope provides an wider field of view and higher angular resolution compared with the conventional IACT design of current telescopes, and will join the next generation of VHE telescopes in the Cherenkov Telescope Array (CTA) project currently in development. The observation and analysis of two SLSNe is performed, on SN 2015bn and SN 2017egm, with both VERITAS and *Fermi*-LAT telescopes. The upper limit is reported on both events and is compared to a simple and a self-consistent model for parameter estimation. In addition, a population of sources were analyzed to estimate the future detection potential of new SLSNe with several gamma-ray observatories. Finally, the observation and

analysis of several novae were performed. These sources were observed with VERITAS and *Fermi*-LAT. The observation of Nova Herculis 2021 is noteworthy in that the overlap of both observations may constrain the cutoff energy of the nova spectral model, providing a estimate for the maximum energy of the accelerated particles in the nova's shock region.

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Dedication

To my family, for their unwavering support.

Chapter 1: Introduction and Background

The field of gamma-ray astrophysics is primarily concerned with the most energetic events in the universe. These events include active galactic nuclei, stellar explosions, mergers, turbulent jet structures and many others.

In many cases, the well-studied and confirmed gamma-ray sources are bright, constantly emitting sources with occasional flares. As the field developed, more sensitive and faster instruments have been able to resolve new time-dependent sources, known as transients, that appear in the sky without prior warning. Transients form the subfield of time-domain astrophysics, which is a main topic of this thesis.

Fast, high-energy, explosive events such as gamma ray bursts (GRBs) were discovered with gamma-ray detectors in the 1960's. Vela, the group of satellites meant to monitor nuclear detonations on Earth, inadvertently discovered gamma rays from the universe in 1967 [1]. This detection, and confirmation of the astrophysical nature, began the search for GRBs and other high-energy time-domain events, leading to the construction of several space-based gamma-ray detectors. Both BATSE of the Compton Gamma Ray Observatory (CGRO, [2]) and Beppo-Sax [3–5] followed up the detection of several thousands of GRBs, cataloging one of the most luminous source types in the universe at 10^{53} erg$\cdot$s isotropic luminosity [6].

In the meantime, the development of the imaging atmospheric Cherenkov technique (IACTs) grew with the detection of very high energy (VHE) (100 GeV to 100 TeV) gamma rays from the Crab Nebula [7] in 1989. This detection was pioneered by the Whipple 10-m telescope, which used a ground-based telescope that observed the Cherenkov effect in the extensive air showers created by gamma rays interacting with the atmosphere. HEGRA, an IACT array developed on the Spanish island La Palma, joined Whipple and the gamma-ray astronomy field in 1987, introducing an array design to exploit stereoscopic reconstruction of the incoming gamma rays [8].

The Very Energetic Radiation Imaging Telescope Array System (VERITAS) was proposed in 1995, broke ground in 2003, achieved first light in 2004, and completed the full array in 2007 [9]. VERITAS followed up on the success of Whipple, constructed on the same site, and is the main telescope used in this work (see chapter 2 for a technical overview). In addition, two more IACT observatories were constructed, forming the current generation of IACT arrays around the world. The High Energy Stereoscopic System (HESS) consists of five IACTs in Namibia, which achieved first light in 2004 and construction of HESS II in 2012 [10–12]. At the same site as HEGRA, the Major Atmospheric Gamma Imaging Cherenkov (MAGIC) Telescopes (later renamed to MAGIC Florian Goebel Telescopes) achieved first light in 2004, and construction of a second telescope MAGIC-II in 2009 [13, 14].

These ground-based IACT observatories sensitive to VHE gamma rays have been aided by the additional development of gamma-ray satellites sensitive to high-energy gamma-ray photons (HE; 30 MeV to 100 GeV). The BATSE detector on the CGRO satellite was accompanied by EGRET, an important surveying detector that cataloged 271 gamma-ray sources, mapping the entire gamma-ray sky from 1991 to 2000 [2, 15]. EGRET was followed by the Fermi Gamma-Ray Space Telescope, a still operational satellite containing the Large Area Telescope (LAT) and Gamma-ray Burst Monitor (GBM), launched in 2008 [16–18]. Fermi-LAT is a crucial instrument for the work in this thesis since it surveys the sky every three hours, enabling long term monitoring of various sources. Importantly, it includes a robust public data access system[1] with user friendly data reduction and analysis tools[2] provided by the Fermi Science Support Center (FSSC) [19].

The international collaboration developed over time by these observatories have led to many physics results and many more questions. An important, open question in the field is about the origin of cosmic rays. First discovered by Victor Hess in 1912 using balloon experiments, the flux of high-energy cosmic particles from the universe reaches the highest particle energies possible, up to 10^{21} eV [20].

According to a recent review by Beatty and Westerhoff [21], the cosmic-ray spectrum has

[1] https://fermi.gsfc.nasa.gov/cgi-bin/ssc/LAT/LATDataQuery.cgi
[2] https://fermi.gsfc.nasa.gov/ssc/data/analysis/software/

Figure 1.1: The cosmic-ray spectrum, approximating a power-law spectral model with index 2.7.
The “knee" and “ankle" features represent the approximate energies of the breaks in the spectrum, which may be transition points when different physical processes dominate the emission [21].

distinct features at 10^{15} eV and 10^{18} eV, known colloquially as the “knee" and “ankle" for its resemblance to a human leg (annotated on Figure 1.1). The spectrum (de ning the $\frac{3\#}{3}$ per energy bin; see appendix C for further de nitions) approximates a power law with index 2.7, while the breaks may indicate the turnover between dominant physical processes within each energy range. The particles from this cosmic ray ux may represent highest energy particles in the universe and lead to some basic fundamental questions: what are their origins? Could gamma rays reveal the acceleration mechanism? Many studies focus on speci c sources of high-energy particles like active galactic nuclei (AGNs), supernova remnants, novae and GRBs to identify the physical mechanisms that accelerates particles up to 10^{21} eV and can explain the cosmic ray spectrum.

Most low-energy light from astrophysical sources consists of thermal photons, meaning that the energy radiated is proportional to the temperature of the source. This is known as blackbody radiation. For higher energy emission, the gravitational and rotational energies released by astrophysical sources power the relativistic acceleration of charged particles via processes such as shocks, reconnection, or turbulence. The relativistic, charged particles then may radiate gamma rays through various processes, both leptonic and hadronic.

The following sections are split into two parts: rst, the mechanisms that accelerate charged particles are reviewed in Section 1.1; and second, the radiative processes that emit gamma rays from these high energy charged particles are reviewed in Section 1.2.

1.1 Acceleration Processes

A major acceleration mechanism is due to shocks, which is the pressurized supersonic ow created in the interface between two uids of different velocities. These uids are made of standard gas and matter around astrophysical objects, like hydrogen or other star ejecta, thermal and non-thermal photons, or higher energy particles and nucleons. The target medium is generally ionized, so the magnetic eld of the charged particles is ampli ed and enhanced by self interaction. The particles in the shock are accelerated by the rst and second order Fermi acceleration, described in the following sections, as summarized by [22–24].

1.1.1 Second Order Fermi Acceleration

First proposed by Fermi [25] in 1949, second order or stochastic acceleration posits that randomly scattered galactic magnetic elds moving at velocity D deposit energy into charged particles, with the average gain

$$\frac{X}{} = \frac{8}{3}\frac{D}{2}^{2} \tag{1.1}$$

In this system, charged particles interact with gas clouds in the interstellar medium with irregularities in the magnetic eld. These particles are scattered inelastically, and since the cloud is moving with velocity D, individual particles gain energy. Head-on (or tail-on) collisions impart an energy gain that occurs just slightly more often than energy loss over time. The dependence on the second order 3 • 2 term leads to a smaller and less ef cient energy gain than rst order Fermi acceleration.

1.1.2 First Order Fermi Acceleration

The rst order, or diffusive shock acceleration, was also proposed by Fermi [26] in 1954, and posits that charged particles gain energy with successive crossings of a planar shock front [22, 23, 26]. The average energy gain of a particle crossing a shock front with velocity B

$$\frac{X}{} = \frac{4}{3}\frac{3_B}{2} \tag{1.2}$$

These collisions are effectively all head-on collisions, which impart a more regular energy gain than in second-order Fermi acceleration.

In both cases, the Fermi acceleration yields a similar energy distribution. If the fractional energy gain per interaction is X, a particle of initial energy $_0$ will reach = $_0$V= after = interactions; if the probability of remaining per scattering event in the accelerating region is ?, then after the same = interactions, only # = #$_0$?= of the original #$_0$ particles will remain. Combining the logarithm of both equations and cancelling out the number of interactions

$\ln {}^{1}\#^{1}\ {}^{o}\bullet\#_{0}{}^{o} = \ln {}^{1}\ \bullet\ {}_{0}{}^{o1}\ln ?\bullet\ln V^{o}$ and therefore

$$\frac{3\#^{1}\ {}^{o}}{3} / \frac{\#^{1}\ {}^{o}}{} \tag{1.3}$$

$$/ \quad {}^{1,\frac{\ln ?}{\ln V}} \tag{1.4}$$

This relationship yields $\frac{3\#}{3}$ that is a power law with index $\kappa = \ 1\ ,\ \ln ?\bullet\ln V$. For strong shocks, the Rankine-Hugeniot conditions describing the uid compression factor and mach number from the conservation of mass and energy between uids leads to a spectral index κ ' 2•5 [23]. This power-law outcome for both rst and second order Fermi accelerations is a tantalizing clue that shock environments may be a primary mechanism driving the acceleration of cosmic rays.

1.2 Radiative Processes

There are 4 main processes toward radiating photons described here, illustrated in Figure 1.2 adapted from [27]. The processes described here are synchrotron radiation, inverse Compton scattering, Brehmsstralung, and pion production. Unless otherwise speci ed, the derivations and discussion of these production process are given by the seminal reference text book by Longair [24].

1.2.1 Synchrotron

A basic source of the photon production is the radiation by a charged particle accelerating in a magnetic eld. In the simplest form, this is the spiraling of an electron within a uniform magnetic eld. The gyrofrequency of the electron in a magnetic eld moving relativistically with Lorentz factor $W = 1\bullet\sqrt{1\ \ V^{2}}$ is $a_{6} = \frac{4}{2cW<_{4}2}$, with $V = 3\bullet 2$. Using the Lorentz invariant radiation power $\%= \frac{@^{3}W^{4}}{6cn_{0}2^{3}}{}^{1}j0_{?}j^{2}\ ,\ W^{2}j0_{k}j^{2o}$, the radiation power becomes

$$\%= f\ )\ 2V^{2}W^{2*}\ \ - \tag{1.5}$$

Figure 1.2: Overview of four radiative processes that emit photons [27].

where $f_{)} = \frac{8c4^4}{3<^2_4 2^4}$ is the Thomson cross-section, and is the magnetic eld density of the region, averaged over the pitch angle between the particle velocity vector and B (from $2f_{)} 2V^2 W^2 * \sin^2 U$ for angle U).

The synchrotron power is $W^2 * {}_B \bullet <^2$, which implies strong dependence on the energy of the spiraling electron and inverse dependence on the charged particle mass squared. This mass relationship implies that electrons will play the largest role in synchrotron power emission when compared to proton-based synchrotron emission in the same environment.

The most important result of this radiation is the combined effect on an electron sea that yields a distinguishable spectrum of the synchrotron radiation. In an extended feat of algebra by Lightman and Rybicki [28], the energy emitted by wavelength is calculated to get the emissivity of the electron as $9^1 I^{\,o} / a^{1\bullet 2} \exp^1 a\bullet a_2{}^o$, where $a_2 = \frac{3}{2} W^2 a_6 \sin U$, where U is the angle between the magnetic eld and particle direction of motion. For an electron eld with power-law distribution, $\frac{3\#_4}{3} / \quad ?$, where ? is the power-law index, the resulting emissivity is proportional to $a^{o} / {}^1$constant$^{o\ 1?, 1^o\bullet 2} a^{1\ ?\ 1^o\bullet 2}$. In this case, the synchrotron spectral index $0 = {}^1? \quad 1^o\bullet 2$. This result shows that an electron distribution with a steep power law will lead to a power-law emission spectrum.

1.2.2 Inverse Compton Scattering

In inverse Compton (IC) scattering, photons gain energy from a relativistic electron, that is:

$$4_{,} W_{low\ energy} ! \quad 4_{low\ energy,} W_{VHE} \tag{1.6}$$

The typical environments contain plenty of low energy photons, whether from the cosmic microwave background (CMB), the local thermal emission or nearby bright sources – e.g. optical or UV from nearby stars or IR from dust clouds. The resulting photon is increased by the relativistic

factor W of the electron:

$$\frac{0}{W} / W^2 \quad W \tag{1.7}$$

The Compton power is $W^{2*} \bullet < \frac{2}{4}$, implying strong dependence on the energy of the target electron and linear dependence on the local radiation eld density $\frac{*}{rad}$. The full expression, in the classical Thomson regime where $\frac{}{W} < {}_4 2^2$,

$$\% = \frac{4}{3} f_{\,)} 2 V^2 W^{2*}{}_{rad} \tag{1.8}$$

where the Compton cross section is frequency independent, $f_{\,)} = \frac{8c4^4}{3<\frac{2}{4}2^4} \quad 6\bullet 65 \quad 10^{\ 25}\,cm^2$. In the high energy regime $_W < {}_4 2^2$, the quantum relativistic scattering cross-section is the Klein-Nishima formula,

$$f_{KN} = \frac{3f_{\,)}}{8G} \quad 1 \quad \frac{2^1 G, \ 1^o}{G^2} \quad \ln{}^1 2G, \ 1^o \ , \ \frac{1}{2} \ , \ \frac{4}{G} \quad \frac{1}{2^1 2G, \ 1^{o2}} \tag{1.9}$$

where $G = \ _W \bullet < {}_4 2^2$, with Y the energy of incident photon. This reduces to the Thomson regime when $_W \ddot{Y}\ddot{Y} < {}_4 2^2$, and the cross section decreases $\sim 1 \bullet G$ for $G \quad 1$.

The general result is that the frequency of a photon is up-scattered by $\frac{0}{W} V^2$ a. For an electron eld with power-law distribution, $\frac{3\#_4}{3} / \quad ?$, where ? is the power-law index, it follows from a similar argument made with synchrotron radiation that the resulting IC spectral index $0 = 1 ? \ 1^o \bullet 2$. Identifying the IC component of a source spectrum is an important part of nding the electron spectrum of the underlying source environment, providing crucial properties that will help understand the nature of these systems.

1.2.3 Bremsstrahlung

In the presence of the electric eld of a nucleus, a charged particle experiences Bremsstrahlung, or "braking radiation", emitting photons that slows down the charged particle. This effect is par-

ticularly important in the production of extensive air showers, described in section 2.1.1, but also in the production of photons in dense, metal rich astrophysical environments.

The power of the nonrelativistic Bremsstrahlung is $\propto Z^2 n_e n_i$, which implies strong dependence on the metallicity and density of the environment. This is the effective behavior in astrophysical environments with relatively slow moving electrons. Relativistically (above the electron rest mass energy), the Bremsstrahlung power is $\propto E$, which is implies an exponential energy function rather than the polynomial relationship of the non-relativistic case. In this regime, the position dependence also enables the calculation of the radiation length, which is the distance necessary for $1/e$ energy loss. In the air, the critical energy is 83 MeV, which is when the ionization losses equal the Bremsstrahlung losses. At this energy, the radiation length is $X_0 = 36.62$ gcm^{-2}.

1.2.4 Pions

The production of pions is a critical pathway toward radiating gamma rays. Charged and neutral pions are produced in equal amounts by the interaction between hadronic cosmic rays and protons, nucleons and photons in astrophysical environments.

The production of pions in the interaction with high-energy protons ($\sim 10^{16}$ eV) and optical photons also yields neutrons $\sim \sqrt{W} = \sqrt{\epsilon_\gamma \epsilon_{photon} \sin\theta}$, where the center of momentum energy of the photon photon interaction. In a more common interaction, both in the atmosphere and a dense astrophysical environment, the interaction of high-energy protons with a nucleon creates a shower of protons, charged and neutral pions (strange particles and antinucleons may also be created in this shower). The resulting pions decay in the following common modes:

$$\pi^0 \rightarrow \gamma + \gamma \tag{1.10}$$

$$\pi^+ \rightarrow \mu^+ + \nu_\mu \tag{1.11}$$

$$\pi^- \rightarrow \mu^- + \bar{\nu}_\mu \tag{1.12}$$

The gamma rays are the most common decay mode of the neutral pion decay (equation 1.10),

occurring more than 98% of the time with very small lifetimes (10^{-9} s) [29]. Due to the direct decay of particles in these chains, a power-law distribution of the primary protons will also yield a power-law distribution of gamma rays produced.

1.3 Thesis Overview

This thesis is structured into three main parts. The first part focuses on the background physics and technical approach. In chapter 1, the basic acceleration and radiative processes are introduced. These processes are referenced throughout and are fundamental to the field of gamma-ray astrophysics. In chapter 2, the VERITAS technical approach is described in detail. This includes both the physics and processes for developing IACTs, and the data analysis pipeline of VERITAS data, from camera data reduction to high-level analysis output.

In the second part, the development of the prototype Schwarzschild-Couder telescope is discussed in chapter 3. The primary effort of this thesis was the construction, commissioning and alignment of the optical system. This effort took place in parallel with analytical and scientific work of chapters 4 and 5.

In chapter 4, the observation and analysis of superluminous supernovae (SLSNe) is discussed. This work was submitted to The Astrophysical Journal and is reproduced here verbatim. Two SLSNe were analyzed, SN 2015bn and SN 2017egm, with both VERITAS and Fermi-LAT telescopes. In addition, a population of sources were analyzed to estimate the future detection potential of new SLSNe with several gamma-ray observatories.

In chapter 5, the observation and analysis of several novae were discussed. These novae were observed with VERITAS and Fermi-LAT. The observation of Nova Herculis 2021 is noteworthy in that the overlap of both observations may constrain the cutoff energy of the nova spectral model, providing a estimate for the maximum energy of the accelerated particles in the nova shock region.

Chapter 2: VERITAS Technical Approach

2.1 Imaging Atmospheric Cherenkov Technique

Gamma rays of all energies reach the Earth, but different observational methods are necessary to detect them accurately. At very high energy (VHE, 100 GeV to 100 TeV), the atmosphere interacts directly with incoming gamma-ray photons, and while that atmospheric opacity makes it impossible to directly detect the VHE gamma ray with a ground-based detector, it provides a new opportunity for an indirect detection. The incoming photon is transformed into an extensive air shower that can be imaged by high resolution, fast triggering, ground-based optical telescopes via the imaging atmosphere Cherenkov technique (IACT), described here. In this chapter, the physics behind this technique is introduced, followed by a description of the VERITAS telescope and the analysis software necessary to process the incoming gamma-ray data.

2.1.1 Shower Cascade

A suf ciently high-energy incoming particle that interacts with the atmosphere will create a secondary particle that in turn will also interact with the atmosphere. This cascade process continues through many generations until the secondary is absorbed by the atmosphere. In this process, an extensive air shower is formed from the secondary particles.

In the presence of a nucleus, a gamma ray with a minimum energy of $2 \times m_e c^2$ (1.022 MeV) can pair produce [illegible]. If the gamma-ray energy is high enough, the resulting leptons are highly relativistic and will radiate through Bremsstrahlung in the presence of the nuclei in the atmosphere, producing further VHE photons that can also pair produce (the minimum energy required for the Bremsstrahlung process is 84 MeV per particle). The radiation length X_0 de nes the distance required for a photon to travel and lose $1/e$ of its original energy due to radiation. In

Figure 2.1: An example of the gamma-ray shower cascade from [31].

air, with Bremsstrahlung $\lambda_0 = 36.62\ \mathrm{g\,cm^{-2}}$, which is 7.9 of the pair production mean free path [29, 30]. This implies that the length a photon travels before pair production is a similar length to an electron undergoing Bremsstrahlung. The typical atmospheric depth is $1033\ \mathrm{g\,cm^{-2}}$ (i.e. the height of the atmosphere), so the interaction must happen near the top of the atmosphere as shown in Figure 2.1.

The number of particles grows exponentially, but at each generation energy is lost from ionization and Bremsstrahlung ($84\ \mathrm{MeV}$ per particle), so the process ceases when the kinetic energy is depleted. At this critical point, the critical energy $\varepsilon_c = 84\ \mathrm{MeV}$, the shower has the maximum number of particles [30]. This basic model was developed by [32], which is simplified by the existence of only photons, electrons and positrons. In the air, the point at which ionization losses dominate for incoming gamma rays of energy $100\ \mathrm{GeV}$ to $1\ \mathrm{TeV}$ occurs at around $10\ \mathrm{km}$ above sea level [33].

There are several features to this air shower that can be directly exploited by an IACT. The only particles produced are photons, electrons and positrons, which simplifies the shower cascade (and

any subsequent simulation effort). The very high energy of the incoming photon, along with the lack of nuclear interactions, translates into most of the particle production occurring longitudinally along the path of the original gamma ray. However, due to Coulomb scattering, transverse broadening of the shower occurs by as much as 30 m. This relatively tight shower helps with the imaging and location of the shower morphology. The height of the shower maximum is proportional to the logarithm of the energy of the incoming photon, while the number of particles created is directly proportional to that energy. These two features can be mapped directly onto the image of the air shower to enable reconstruction of the true energy of the original particle.

2.1.2 Hadronic Showers

In addition to the gamma-ray extensive air shower, a signi cant number of hadronic cosmic-ray showers are created in the atmosphere, made up of charged, highly relativistic protons and other nuclei. In most cases, the hadrons interacting with air nuclei break off into charged and neutral pions:

$$c^0 \; ! \quad W, \; W \tag{2.1}$$

$$c^{\flat} \; ! \quad `^{\flat} \; , \; a^{`} \tag{2.2}$$

$$c \; ! \quad ` \; , \; a^{`} \tag{2.3}$$

Hadronic cosmic-ray showers that initially decay into neutral pions quickly form into predominantly gamma-ray showers, which make them nearly indistinguishable from pure gamma-ray showers. On the other hand, showers decaying into charged pions form muons that reach the ground before decaying into the leptonic particles, (a_4, $a^{`}$, a_4 and $4^{\flat}$) due to time dilation. These muons lose energy through Cherenkov radiation. These strong interactions with heavy nuclei in hadronic showers produce transverse momentum, increasing the lateral length much higher than the Coulomb broadening induced in the electromagnetic showers.

The large number of decay channels complicate a general model for the hadronic cosmic-ray

Figure 2.2: An example of the hadron cosmic-ray shower cascade from [31].

shower, which make it dif cult to distinguish hadronic cosmic ray showers from pure gamma-ray incoming photons, as shown in Figure 2.2. Electron or positron cosmic rays will also induce electromagnetic showers, further complicating the true source of the incoming particle. The similarities between hadronic or leptonic cosmic-ray showers and gamma-ray showers introduce a nearly irreducible background of events that must be accounted for in the analysis.

2.1.3 Cherenkov Radiation

Once the charged particles are produced by the extensive air shower, Cherenkov radiation is the primary process converting that energy to optical light that can be detected by the IACT on the ground. In Cherenkov radiation, charged particles with phase velocity larger than the speed of light in a medium with index of refraction n will constructively polarize particles in that medium, which quickly and simultaneously radiate when the particles return to the ground state. The Cherenkov light emitted from these particles will form a cone around the trajectory of the charged particle with angle\ ,

Figure 2.3: *Top*: Distribution of shower extent for a pure electromagnetic shower (from a 300 GeV gamma ray; left) and hadrons (a 1 TeV proton; right). *Bottom*: The distribution of Cherenkov light on the ground from these simulations. Both images are from Monte Carlo simulations by S. Fegan in Aharonian, Buckley, Kifune, et al. [34].

Figure 2.4: A simple Cherenkov radiation diagram. Figure from Wikimedia [35].

$$\cos\theta = \frac{1}{\beta n(\lambda)} \tag{2.4}$$

The Cherenkov angle is dependent on the height of the interaction because the index of refraction depends on the density of the local medium, such that the angle decreases with increasing height. At sea level, this angle is about $1.3°$, while $\theta \approx 0.74°$ at 8 km to $0.66°$ (12 mrad) at 10 km. This height dependence introduces a de-focusing effect as the altitude drops, while Coulomb broadening of the electrons in the cascade fill in the light emission turning it into a pool of light on the ground. The resulting Cherenkov light emitted at 10 km has radius of about $r \approx 10\,\text{km} \times 12\,\text{mrad} \approx 120\,\text{m}$ at ground level. VERITAS exploits this approximate shower size by placing IACTs telescopes within a baseline 100 m so that nearby telescopes can see the same shower and far enough away for stereo vision to improve direction reconstruction.

The spectrum of the resulting Cherenkov radiation is dependent on $n(\omega)$, see appendix A.1 and A.2 for discussion on the wavelength dependence. An oscillating particle forms a dielectric constant $\varepsilon(\omega)$ in equation A.10, which defines the index of refraction of the medium. This constant is then used in the application of Maxwell's equations to a moving particle in A.1. The Frank-

Tamm equation A.9 is derived in Fourier transform as a function of the frequency, and is rewritten here as a function of the wavelength:

$$\frac{3^2\#}{3G3\_} = \frac{2cUI^2}{\_^2} \quad 1 \quad \frac{1}{V^2=^1\_^{o2}} \tag{2.5}$$

where the critical angle is when the phase velocity 3_{phase} i $2 \bullet =$, and $\cos\backslash = \frac{1}{V=^1F^o}$, where $V = 3 \bullet 2$. Examples of the Cherenkov light pool generated by primary gamma rays and protons are shown in Figure 2.3.

2.2 VERITAS

The Very Energetic Radiation Imaging Telescope Array System (VERITAS) is an IACT observatory located in Southern Arizona, USA (see Figure 2.5). VERITAS has four 12 m Davies-Cotton (DC) design single-mirror telescopes, each with 350 individual mirror facets, a 499 pixel camera with 3•5° eld of view. VERITAS had two major upgrades that altered the instrument sensitivity: in 2009, the T1 telescope was moved across the site, increasing the baseline distance between telescopes to ò 100 m and improving the ground area distribution for stereo reconstruction; and in 2012, the photo-multiplier tubes (PMTs) were replaced by higher quantum ef ciency PMTs to lower the energy threshold and improve the overall sensitivity. VERITAS is sensitive to gamma rays within the energy range 100 GeV to v 30 TeV, has an angular resolution (68% containment) of v 0.1° at 1 TeV, and an energy resolution of 15% at 1 TeV.

The many optical photons from the Cherenkov shower pass through the optical system of the IACT to a camera for conversion from analog to digital. The PMT converts light to electronic signals through the photoelectric principle, in which photons eject electrons from the surface of a photo-cathode, triggering a cascade of electrons in the presence of large electric eld, and producing an output current. The electronic noise and night sky background (NSB) produce both positive and negative signal uctuations in the output current, so an arti cial signal is injected so the PMT

Figure 2.5: A photograph of VERITAS at Fred Lawrence Whipple Observatory.

signal never crosses zero. The NSB is stellar and man-made light scattered in the atmosphere that dominates over the electronic noise. The injected arti cial signal is known as the pedestal, and the variation on the pedestal (the “pedvar") is directly correlated to the camera noise parameters.

The ash of light from the Cherenkov event is integrated over 2 ns samples within a custom-built ash analog-to-digital-converter (FADC) at the sampling rate of 500 MHz and stored temporarily in a 64 µs buffer. The digital signal has an 8-bit (256 digital counts per channel) dynamic range. A 32 ns window of the FADC readout is stored in a memory buffer to be sent to the data acquisition for recording, timed to compensate for the shower arrival time from each telescope. Since not all data is needed, there are hardware triggers to monitor and lter the readout of the electrical signal to the data acquisition system. The three triggers are: L1, the pixel-level trigger; L2, the telescope-level trigger; and L3 the array-level trigger. The array-level event trigger will signal to re the memory buffer to the data acquisition [36, 37].

Speci c hardware con guration is designed for each trigger level. Each PMT is connected to a constant fraction discriminator (CFD), which nds the maximum voltage of the PMT pulse. Whenever the running maximum exceeds an experimentally set threshold (generally 45 mV), the pixel generates an L1 trigger. These pixel-level triggers can come from the NSB, which is made up of uncorrelated light of varying brightness. In order to lter out these individual ashes, L2 triggers when 3 neighboring PMTs are triggered within 5 ns of each other [38]. Further, single-telescope

events like muon rings (that cause L2 triggers) need to be filtered out. L3 triggers when 2 or more telescopes fire L2 triggers within 50 ns (experimentally optimized). A signal is sent to the FADCs to read out and record the 32 ns (16 samples) [39]. However, at this point, data is not taken during the readout period, which introduces a deadtime to the entire array. A typical dead-time percentage for the VERITAS array is around 10% at an L3 trigger rate of 300 Hz.

Observations by VERITAS are designed to optimize the data quality recorded based on the hardware considerations. The full use of the camera field of view is necessary to reconstruct events near the source position, since the shower image must be large enough to calculate image moments for each event. Additionally, observations of a background region must be taken to help set a baseline set of properties to discriminate between gamma-ray and non-gamma-ray events and to estimate the residual rate of background events at the location of interest. To accommodate this, observations are taken in a few standard ways: On/Off Tracking and wobble mode.

Each observation is taken in 30 min segments called runs. In On/Off tracking, a dedicated run is taken for the source in the center of the field of view, immediately followed by a second run taken 30 minutes (7.5°) offset from the original coordinates. This strategy allows for the direct observation of a background region in very similar and nearby sky conditions. In the wobble mode, the background and source observations are taken simultaneously by offsetting the tracking by 0.5° from the center of the camera. This offset is applied in the cardinal directions (N, E, S, W), so that the target source is always inside the camera field of view. The offline analysis of this data takes care of various observation strategies in the computation of the background, described in section 2.3.3.

2.3 VEGAS

The large amount of data generated by the array is analyzed using a clear software paradigm. In the context the next generation gamma-ray observation, the Cherenkov Telescope Array (CTA), a data organization and management was developed in anticipation of the multi-wavelength collaboration of gamma-ray data analysis [40, 41]. Although the VERITAS software predates this

Table 2.1: Mapping of data analysis stage to the responsible VEGAS tool and associated data level [40, 41].

VEGAS Tool	Data Level		Analysis stage	Description
vaStage1	DL0	DL1	Calibration	Calculation of T0 and pedestal level.
vaStage2	DL1		Image Analysis	Trace integration. Single telescope image analysis. A deprecated vaStage3 used to perform image moment calculation, but this functionality was moved to vaStage2.
vaStage4	DL1	DL2	Event Reconstruction	Stereo reconstruction. Energy and core position calculation. Lookup tables required.
vaStage5	DL2	DL3	Background Rejection	Gamma-hadron cuts applied.
vaStage6	DL3	DL4	Result Extraction	Spectral analysis, sky map creation, lightcurve analysis, source detection. Effective area tables required.

structure, it is a useful guide to understanding the data reduction strategy used here. This strategy created the gamma-ray analysis data format (GADF), which outlines the elds and le formats for each data level (DL) created from DL1 to DL5. Each data level refers to a speci c responsibility: DL0 is the reduced raw camera data; DL1 is the calibrated camera data; DL2 is the reconstructed shower parameters (energy, direction, etc.) per event; DL3 are the selected gamma-like events and instrument response functions (IRFs); DL4 are the compiled spectra, skymaps, light curves from the data reduction; and DL5 are the observatory catalog or survey sky maps [40, 41].

This data management structure helps to explain the VERITAS data analysis strategy. Asserting cross-checking and validation, VERITAS maintains two independent data analysis packages to process data from DL0 to DL4: EventDisplay [42] and VEGAS [43]. In this thesis, VEGAS is used as the analysis package for all VERITAS analysis and all scienti c results are con rmed by a collaborator using EventDisplay.. Each tool in VEGAS is responsible for calculations that can be mapped onto a speci c data level, as outlined in Table 2.1. VEGAS is written with, relying on the CERN ROOT library for both computational methods and the le formatting [44].

Table 2.2: VEGAS image parameters calculated in stage 2.

Parameter	De nition
width	The RMS spread of light parallel to the shower image major axis.
length	The RMS spread of light perpendicular to the shower image minor axis.
distance	The angular distance of the image centroid from the center of the camera FOV.
miss	The perpendicular angular distance of the image axis from the center of the camera FOV.
azimuth-width	The RMS image width relative to an axis drawn from the image centroid to the center of the camera FOV.
size	The total number of digital counts from all the pixels making up the shower image.
nTubes	The total number of pixels that make up the shower image.

2.3.1 Calibration

The primary function of vaStage1 is the calibration of the original camera data. This includes determining the timing of the digital signal[45], the gain, and the noise level of the uctuating NSB [46]. It also includes the identi cation of bad pixels within the camera. In conjunction with science data taking, every night calibration data is taken by tracking a dark patch of sky to measure the camera response to the asher, which uniformly illuminates the camera at a xed, high frequency rate [47].

Once the camera data from both data and asher runs are calibrated and combined, vaStage2 applies the asher calibration and is used to calculate the resulting image moments in the camera. Image cleaning is performed to remove bright pixels, any pixels with NSB level brightness, and truncated images [46]. A full description of this lower level analysis can be found in [48]. In brief, the Hillas parameters [49] are computed, which are summarized in Table 2.2 and shown in the diagram in Figure 2.6. In an earlier state of VEGAS, vaStage3 calculated these image moments for each event, but that functionality was merged into vaStage2 to improve the speed and ef ciency of the analysis.

An important development in the image analysis is the use of templates to determine the shower parameters from the camera image. The image template method (ITM), used extensively in this

Figure 2.6: A diagram of the shower properties projected on the IACT camera from [50].

thesis, was developed by simulating gamma-ray showers in a well-tested atmosphere and detector model (incorporating the mirror model and camera electronics) [51–53]. This simulation allows the construction of a pro le likelihood for every pixel in the camera, which can be t to the data to nd optimal gamma-ray shower [54]. One important cause of the improvement provided by this method is the reconstruction of truncated images at the edge of the camera, which recovers events previously thrown away in the standard method. This method has been extensively validated, including in conjunction with a BDT analysis in vaStage5 [54, 55].

2.3.2 Shower Reconstruction and Selection

In vaStage4, the stereo reconstruction of the shower using multiple telescopes is performed to calculate the direction, energy and shower core location (the hypothetical position of the light pool impact on the ground). Figure 2.7 shows this stereoscopic projection onto the ground reference frame. This stereo reconstruction enables the cutting of events with too few telescopes, or core locations that are too distant from the array.

Figure 2.7: Example of the direction reconstruction using stereoscopic projection. Figure from [8].

The calculated image parameters from stage 2 do not yet connect to real showers. Using simulations, which enable the knowledge of true parameters such as the energy, to determine what a shower should look like on the camera can help create a lookup table to relate the calculated image parameters to the physical shower parameters. From these lookup tables, measured values for the length and width parameters are scaled by the lookup table values in the following form, as mean scaled `parameter' (MSP)

$$"(\% = \frac{1}{\#_{tel}} \sum_{8=1}^{\#_{tel}} \frac{?_8}{?_{sim}(\zeta_{zen}, (_8, A_8)} \tag{2.6}$$

where $?_8$ is the parameter of interest (length or width), and $?_{sim}$ is the mean simulation value for that parameter given the shower zenith angle ζ_{zen}, image size $(_8$ and the impact distance A_8. Using this general equation for $"(\%$, the speci c value for mean scaled length (MSL) or mean scaled width (MSW) is calculated.

In vaStage5, the remaining cuts are applied to select the best gamma-ray candidate events. The parameter scaling is used so that uniform cuts may be applied to the entire data set to separate

Figure 2.8: A sample distribution of parameters used for the gamma-hadron selection cuts: mean-scaled length, mean-scaled width, shower max height, mean-scaled length RMS, mean-scaled width RMS, and energy RMS for event energies above 1 TeV.

gamma rays from hadrons. The standard gamma-hadron box cuts may look like MSW 0•05 1•1 and 0•05 MSL 1•3, which can be applied to the distributions in Figure 2.8.

In vaStage4, additional cuts were also applied in conjunction with the calculation of the scaled parameters. These include the minimum number of pixels in the image,) D1 4B 5 and image distance to center of camera 8BC 1•43. Finally, the size (sum of charge deposited on camera) is cut in three ways, corresponding to the standard cuts used in VERITAS - soft, medium, hard - which are optimized for analysis outcomes corresponding to well-known spectral behavior.

The box cuts on scaled parameters used in vaStage5 are optimized on the analysis of known sources. However, while these cuts are easy and fast to process, there are limitations to using box cuts due to the non-linear relationships between parameters. Non-linear algorithms such as BDTs are currently being explored for VEGAS [55], while it has been robustly veri ed in EventDisplay [56], and known to be a component of the H.E.S.S analysis [57].

2.3.3 Analysis Summation

Background Estimation

Each candidate gamma-ray event reconstructed by the preceding analysis may originate from a different location on the sky, some of which may originate from strong gamma-ray sources. Given the uncertainty in the event reconstruction precision (the gamma-ness of an event), the creation of a large and well-observed background region affords another opportunity to statistically discriminate gamma-rays and non-gamma-rays in the data. By comparing the counts of the signal and the background regions, it is possible to estimate the difference in counts to nd the source gamma-ray emission rate.

Two common background construction methods are used within both of the IACT analysis software in VERITAS: the ring background method (RBM), and the wobble or re ected region (RR) method. The observation of a source is performed by explicitly offsetting the telescope tracking position from the source position, pointed toward standard cardinal directions (i.e. wobbling around the source). This offset is tuned to balance the simultaneous measurement of the background and source positions in the same eld of view while also staying within the well-behaving radius of the acceptance of the camera. The reliability of gamma-ness from event reconstruction weakens radially from the center of camera, as the Cherenkov showers are truncated from the eld of view and become more dif cult to analyze. Both the RBM and RR methods bene t from these wobbled observations due to the increased background observation regions and well-performing acceptance at these positions.

In the analysis of a point source (which is the primary type of source analyzed this thesis), a small circle is used to specify the region from which to count events from the target source position. This circle generally has the radius comparable in size to the point spread function (PSF) for reconstructed gamma rays, around 0.1°, and is called the on region. The size of this circle is given by angular radius θ from the target position, which becomes an important cut in the gamma/hadron discrimination (this cut is generally referred by the θ^2 value, since the radius

squared is linearly related to the count rate within an area and is easier to model on a histogram). For both RBM and RR methods, the counts in the on region are compared to those of off region counts. Regions of the map which include bright stars, known sources, or the observational target itself, are excluded from the background estimation.

For RBM, a ring is created around the target position with upper and lower radii, forming an annulus of a given size. The top of Figure 2.9 shows the ring in light green at a random test position near the target region. This method enables the collection of a large amount of data, but with a varying acceptance that must be accounted for in the calculation. For the RR method, the bottom of Figure 2.9, a series of adjacent off regions are rotated around the center of the camera in each wobble observation.

Li & Ma Signi cance

The fundamental statistical process used in this eld is the use of the optimized likelihood ratio test, formed on the assumption that count rates in the signal and background regions can be modeled (with some analytical form such as a Poisson distribution), and that the measurement of the background region gives an accurate estimate of the background rate within the signal region. This estimation is improved by increasing the background region so that the statistical uncertainty of the background is negligible compared to the uncertainty of the signal region.

$$\#_B = U\#_o \tag{2.7}$$

$$\#_S = \#_{on} \quad U\#_o \tag{2.8}$$

$$U = \frac{{}_{on}C_{on}}{{}_o\ )_o} \tag{2.9}$$

$$5^1\#_{on} - \#_o\ ^o = 5^1\#_{on} - \#_o\ j\#_S - \#_B{}^o \tag{2.10}$$

$$= 5^1\#_{on} j\ h\#_{on} i^o\ 5^1\#_o\ j\ h\#_o\ i^o \tag{2.11}$$

The event counts in the on ($\#_{on}$) or off ($\#_o$) regions have background counts ($\#_B$) and signal

Figure 2.9: A diagram showing the RB (top, a) and re ected region (bottom, b) methods for background estimation. For both methods, the target region is wobbled to the west (right) in pink, a teston region is toward the top right in dark green, with its background in light green, while two exclusion regions from stars are also included. In addition to the calculating the background and excess counts for target region, the RB method is performed on test bins throughout the entire eld of view, each with their own ring background, to generate a distribution of signi cances for the entire observation and enable map generation.

counts ($\#_S$) from the target source. To increase the background modeling accuracy, the off region is observed with larger exposure which is scaled by U times the on region. The U factor is the ratio of the areas or exposures between the on and off regions; the reciprocal number of off regions is the U parameter in the simplest cases. The probability distribution is modeled by the Poisson distribution of both the signal and background counts.

The strategy outlined by Li and Ma [58] assumes this probability distribution in the likelihood L to form the log-likelihood ratio test, noting that Wilke's Theorem [59] reduces the log-likelihood to a χ^2 distribution with 1 degree of freedom when

$$\lambda = \frac{L_0}{L} \tag{2.12}$$

$$\zeta = \sqrt{-2\ln\lambda} \tag{2.13}$$

$$= \sqrt{2}\left[\#_{on}\ln\left[\frac{1+U}{U}\frac{\#_{on}}{\#_{on}+\#_o}\right] + \#_o\ln\left[(1+U)\frac{\#_o}{\#_{on}+\#_o}\right]\right]^{1/2} \tag{2.14}$$

This derivation assumes that the signal counts are equal to the background counts in the null likelihood case L_0, and also assumes that the number of background counts is equal to $\#_B = \frac{U}{1+U}(\#_{on}+\#_o)$.

This fundamental equation 2.14 is used throughout this work and is sometimes referred to as Equation 17 of Li & Ma [58].

Time Dependence

In the conventional analysis of sources observed by IACTs, background and signal regions are determined a priori so that an excess count can be calculated that characterizes the source count rate during an observation period. Following the criteria established, the statistical significance is calculated using Eq. 17 of Li & Ma [58]. However, in this formulation, both the source and background rates are assumed to be time independent. In the case of GRBs and other strongly time-dependent sources, the excess counts may be dominant in certain time bins within the observation and negligible in others, such that the excess counts summed over all time bins are not significant.

To solve this problem, the time dependence of the source count rate is assumed in a parameterized form $B(t)$ and the test statistic is re-derived following the strategy outlined by Li and Ma [58] [60]. The maximum likelihood estimate is used to optimize the signal and background count rates, collected within the On and Off observation regions de ned as $N_{on} = (\mu + B(t))\,T_{on}$ and $N_{off} = (\tau\mu)\,T_{off}$ for observation times T_{on} and T_{off}, respectively. Again assuming a Poisson distribution of counts, the general form of the test statistic is a ratio of the likelihoods for the null and alternate hypotheses L_0 and L, rewritten here:

$$\frac{L_0}{L} = \frac{\mu_0^{N_{on}}\,\tau^{N_{off}}}{(\mu + B)^{N_{on}}\,\mu^{N_{off}}}. \tag{2.15}$$

where μ and B are the time average background and signal count rates, μ_0 is the time averaged rate when all counts observed are background.

To apply the time dependence, the T_{on} observation time is split into N bins of Δt, where the probability of the number of counts for each bin is independent N and $\Delta t = T_{on}/N$. For a parameterized $B(t)$, at the limit of $N \to \infty$, $\Delta t \to 0$, the likelihood ratio reduces to:

$$\frac{L_0}{L} = \frac{\mu_0^{N_{on}}\,\tau^{N_{off}}}{\prod_{i=1}^{N}(\mu + B_i)\,\mu^{N_{off}}} \tag{2.16}$$

where $B(t) = \lambda\,t^{-1}$. The time dependence of the signal was assumed to be 1/t since it is most similar to many GRBs light curves detected by Fermi-LAT after the burst [61], where λ is a normalization constant for the time-dependent count rate.

This modi ed likelihood ratio was validated with a Monte Carlo simulation. Using a toy-model MC of a GRB, with 5×10^5 observations starting 2 min after burst trigger, the time-dependent signal $B(t) = \lambda\,t^{-1}$, and the typical IACT background rate (1/min), the following simulation results provided evidence for some additional conclusions (see also Figure 2 of [60]). For a 35-min observation, the mean signi cance over all the simulated observations increased from 4σ to 5σ once time dependence is included, a 25% increase in sensitivity. For a 100-min observation: the simulated observations reveal that the overall signi cance improves from 3.4σ to 5.2σ with

time weighting for that observation time. Comparing 35-min and 100-min simulations, the time-independent method has a reduced sensitivity from 4.6σ to 3.4σ as expected with the increase in background counts, while the time-dependent model increases from 5.1σ to 5.2σ, indicating that the time dependence allows for an improved estimate of the background and signal rates for longer observation times.

The simulation showed that the significance is dependent on the length of the observation. The standard significance [58] peaks and then declines over time as the background rate accumulates counts toward the On region and dilutes the signal counts. However, the time-dependent likelihood ratio grows monotonically if events are weighted to account for the GRB's time dependence, up to as much as 10-25% increase from the time-independent significance (see Fig. 3 of [60].) Some common analysis software packages within the gamma-ray astrophysics field (e.g. CTOOLs) include a temporal model in the maximum likelihood fitting in addition to the spatial and spectral models [62].

2.3.4 Upper limit

Many of the time-domain astrophysical sources within the VHE gamma-ray range are weaker sources than measurable by current IACTs, if they emit gamma rays at all. In this case, rather than reporting the statistical flux of the source with a small statistical significance and very large relative error, an upper limit is reported. The upper limit refers to the largest possible flux attainable by the instrument given the observed background counts, calculated by finding the maximum signal counts within a specified confidence level and known measurement uncertainty. If the instrument is sensitive enough when compared to an expected model, this upper limit may even constrain the model.

Physical Reasoning

The ux upper limit is generally reported as

$$i_{thr} = \int_{thr}^{1} \frac{3\#}{3} 3$$

The observations determine the threshold energy and differential ux, which de ne the instrument in this instance, so these must be solved for in this equation. The differential ux can take the form of a powerlaw or any other common model. For a powerlaw, the normalization needs to be computed using the effective area of the instrument.

$$\frac{3\#}{3} = :_0 \frac{\quad}{_0}^{w}$$

$$:_0 = \frac{hcountsi}{htimei\ harea\ henergyi}$$

$$hcountsi = \_source$$

The upper limit calculations (Rolke, López, and Conrad [63], Helene [64–66], and Cousins, Feldman, and Cousins [67], etc.) use the number of counts in the on and off regions, and a con - dence limit (e.g. 95%) to compute the statistical number of counts possible, $\_thr$. To estimate the largest possible upper limit, these values should use the whole observed On/Off counts and exposure of the instrument, assuming energy independence. The assumption is that these counts are from the source (not background), which follows the given physical spectral model. To calculate the number of counts from a model:

$$\_{source} = \frac{3\#}{3}\,{}^{1}\,{}^{o}3\;3C$$

$$= C\,{}^{\dot{}}{}_{0}\;\frac{\_}{{}_{0}}\;{}^{W}\;{}^{1}\;{}^{o}3$$

Setting $\_{source} = \_{lim}$, we can then nd the upper limit normalization to place into the integral ux i_{thr} at the top.

$$:_{0} = \frac{\_\;\_{lim}}{C\;\frac{\_}{{}_{0}}\;{}^{W}\;{}^{1}\;{}^{o}3}$$

Putting it all together,

$$i_{thr} = \frac{3\#}{3}\,{}^{1}\,{}^{1}\,3$$

$$= :_{0}\;\frac{\_}{{}_{0}}\;{}^{W}\;3\;{}^{1}\,{}^{1}\,{}_{thr}$$

$$= \left[\frac{\_\;\_{lim}}{C\;\frac{\_}{{}_{0}}\;{}^{W}\;{}^{1}\;{}^{o}3}\right]^{3}\;\frac{\_}{{}_{0}}\;{}^{W}\;3\;{}^{1}\,{}^{1}\,{}_{thr} \qquad (2.17)$$

Since this upper limit is model dependent, the nal value could be anything within the model parameter space $f_{0} - W g$. $:_{0}$ is dependent on the count upper limit and so it is possible to vary W to get a distribution of the range of ux upper limits. This process is approximated by trying W to see the spread of ux upper limits, then reporting the energy at which pairs of these uxes are close to each other (the so-called decorrelation energy). If the energies are exceptionally different ($\dot{}$ 30%), it indicates a bias due to the W dependence. The decorrelation energy is then used in the upper limit integral, which approximates the sum of all the possible uxes within the model parameter space $f_{0} - W g$. This is a good estimate for a parameter-independent ux upper limit.

This equation forms the strategy implemented in VEGAS in the next subsection, i_{thr} is re-

named to $\frac{3\#}{3}$ and $\frac{3\#}{3}$ to q¹ º.

Extragalactic Background Light

The universe is lled with extragalactic background light (EBL) that interacts with high energy photons. Unlike lower energy photons and neutrinos that pass directly through the universe unimpeded and in a straight line, VHE photons above 1 TeV interact with optical (a few eV energy) photons via photon-photon pair production. Given a probability of interacting with the EBL, VHE gamma rays traveling a larger distance increase the overall absorption factor as they are more likely to interact with light along the way [68]. The standard energy and redshift dependent EBL absorption factor g¹I– º decreases the detected ux by g¹I– º.

Applying the absorption formula to equation 2.17 above:

$$q_{?-8=\bar{c}} = \frac{1}{\text{thr}} \int_{\#_8}^{\#_{*!}} \frac{h\, i_8}{\int_{;-8}^{D-8} \frac{1}{0} o\, w_4\, g^1 I - {}^{o\rho} 3\, {}_0} \frac{i}{0} \frac{w}{0} 4\, g^1 I - {}^{o} 3$$

In the denominator, this changes the spectrally-averaged effective area $\ddot{\bar{e}}_{\bar{c}}$ of the instrument at the high energies by a small amount. EBL deabsorption hardens the spectrum as the energy increases, so each ne bin will increase by a small amount. Overall, it is dif cult to predict the standard total change (whether greater or less than original ux) because the actual change in the spectrally-averaged effective area has to compete with the change in ux integral in the numerator.

Time Dependence

Assume there is a time dependence on the spectral model, $5^1 C^{\rho}$ generally a power law with V ' 1•3

$$\frac{3\#}{3} = :_0 \frac{w}{0} 5^1 C^{\rho}$$

$$= :_0 \frac{w}{0} C^{\nu}$$

The estimate source counts is then

$$\_{source} = \frac{3\#}{3} \; {}^{1} \; {}^{o}3 \; 3C$$

$$= : \; {}^{0} \; \frac{}{} \; {}^{0} \; \frac{}{} \; C^{V} \; {}^{1} \; {}^{o}3 \; 3C$$

$$= \begin{cases} : \; {}^{0} \frac{C^{1, V}}{1, V} \; \frac{}{} \; {}^{0} \; \frac{}{} \; {}^{1} \; {}^{o}3 & \text{for } V < 1 \\ : \; {}^{0} \ln{}^{1}C^{o} \; \frac{}{} \; {}^{0} \; \frac{}{} \; {}^{1} \; {}^{o}3 & \text{for } V = 1 \end{cases}$$

Fast forwarding the same analysis above to its nal form (for now $w_0 = 1$) 4+ and E in units of TeV, so that it simpli es the text):

$$q_{?-8} = \frac{}{} \; {}^{1} \; {}^{1} \; \frac{}{} \; {}^{1} \; \frac{}{} \; \frac{h}{\frac{C^{1, V}}{1, V}} \; \frac{}{} \; \frac{\#*!}{\frac{\#}{8} \; h \; i_8 \; {}^{D-8} \; {}^{0} \; w_3 \; {}^{0}} \; i \; {}^{1} \; {}^{o} \; w_V \; C3 \; 3C$$

$$= \frac{}{} \; {}^{1} \; {}^{1} \; \frac{}{} \; {}^{1} \; \frac{}{} \; \frac{h}{\frac{\#}{8} \; h \; i_8 \; {}^{D-8} \; {}^{0} \; w_3 \; {}^{0}} \; \frac{\#*!}{} \; {}^{1} \; \frac{1, V^{o}}{;-8} \; i \; {}^{1} \; {}^{o} \; w \; \frac{C^{V}}{C^{1, V}} \; 3 \; 3C$$

Assuming $C = C^{o}$, that can be integrated together

$$= \begin{cases} {}^{1}1, V^{o} \ln{}^{1}C^{o} \; \frac{}{} \; {}^{1} \; \frac{}{} \; \frac{h}{\frac{\#}{8} \; h \; i_8 \; {}^{D-8} \; {}^{0} \; w_3 \; {}^{0}} \; \frac{\#*!}{;-8} \; i \; {}^{1} \; {}^{o} \; w_3 & \text{for } V < 1 \\ \frac{}{} \; {}^{1} \; \frac{}{} \; \frac{h}{\frac{\#}{8} \; h \; i_8 \; {}^{D-8} \; {}^{0} \; w_3 \; {}^{0}} \; \frac{\#*!}{;-8} \; i \; {}^{1} \; {}^{o} \; w_3 & \text{for } V = 1 \end{cases}$$

Statistics

The upper limit of counts is found by calculating the con dence region of a distribution of signal counts. The counts are Poissonian, with some rate given by background and signal physics (all kinds of physical and instrumental reasons will de ne the rate itself). The signal counts are found by the difference in counts for different observing modes, on and off. With a Poissonian

distribution, likelihood functions for null and hypothesis are created, and then the con dence range is developed with these functions, whether by likelihood ratio or numerically. The three main upper limit calculations are Rolke [63], Feldman & Cousins [67], and Helene [64], described in the following subsections.

Rolke The Rolke method adds nuisance parameters to the basic set up of the pro le likelihood to allow for statistical uncertainty. Speci cally, say x events are run through a Monte Carlo (without background) and nds s events surviving. Then, the ef ciency ϵ is modeled as a binomial random variable based on the original number of true events x and ef ciency parameter τ and nds the complete model to be

$$\#_{on} = \mathrm{Poiss}(\tau B + b)$$

$$\#_{o} = \mathrm{Poiss}(\tau \cdot U)$$

$$\epsilon = \mathrm{Binomial}(x, \tau)$$

The modi ed log-likelihoods $\ln(B, \tau, \#_{on}, \#_{o}, \epsilon)$ are maximized over b and τ, the background counts and ef ciency, respectively, eliminating the nuisance parameter ϵ. The log-likelihoods are used to calculate the pro le likelihood function, parameterized by the signal counts. If the B counts are not small, the con dence range is found by varying B until the pro- le likelihood value has changed by the con dence limit (C.L). For example, if C.L. = 90%, the pro le likelihood will change by 2.706 from the minimum point (at the estimator $\hat{B}$). This technique is unbounded by any constraints to the pro le likelihood, so it is the default method used when computing the Rolke upper limit.

There are a few cases when the resulting signal counts upper limit is negative – that is nonphysical – which contain a few available strategies to respond. If the B counts is negative on the lower end of the uncertainty range, the B counts lower limit is set to 0. In another strategy, the pro le likelihood function has a negative estimator $\hat{B}$, and the upper limit (the estimator plus the C.L. increase in pro le likelihood) is also negative. One count is added to

Figure 2.10: A direct reproduction of a pro le likelihood example from Rolke, López, and Conrad [63]. In both cases, the optimal signal rate B is likely below 0 and non-physical, requiring special care.

$\#_{on}$ and the pro le likelihood function to nd the next upper limit recomputed. This addition of one count to $\#_{on}$ is repeated until the upper limit is positive. Special care is also needed when $\#_{on} = 0$ and/or $\#_{o} = 0$. The simplest method adds one count to either or both $\#_{on}$ and $\#_{o}$, and a linear extrapolation is performed on the pro le likelihood to nd the upper limit. See gure 2.10 for the pro le likelihood example with a negative B. In the computation of the “bounded" case of the Rolke method within VEGAS, the chosen strategy adds the C.L. to the pro le likelihood when B is set to zero.

Feldman & Cousins Instead of using the likelihood ratio to nd the con dence region, this method

uses the probability distribution of the true parameter θ to find the confidence region with appropriate coverage. It uses an ordering method based on the ratio

$$R = \frac{P(n|\theta)}{P(n|\theta_{best})}$$

$$P(n|\theta) = \frac{(\theta + b)^n e^{-(\theta + b)}}{n!}$$

Cousins, Feldman, and Cousins [67] makes a lookup table for all possible θ and n values, filling in $P(n|\theta)$ and $P(n|\theta_{best})$ (found by maximizing P by varying θ). With the table completed for many θ, the procedure then ranks the table by R. The confidence limit is found by summing all $P(n|\theta)$ in order of R until the confidence limit U (e.g. 90%) is reached for the measured background b. The largest and smallest θ for given n are the confidence intervals in the table.

Helene The basic assumption of Helene is to pick a probability distribution for the signal counts $P(\theta)$, then solve directly for the upper limit θ:

$$U = \int_{\theta}^{1} P(\theta)\, d\theta$$

$$P(\theta) = \begin{cases} N_1 \dfrac{(\theta + b)^n e^{-(\theta + b)}}{n!} & \text{Poisson s, known b} \\ N_2 \dfrac{e^{-\frac{(\theta + b - s)^2}{2\sigma_n}}}{\sqrt{2\pi\sigma_n}} & \text{Gaussian s, known b} \\ N_3 \dfrac{e^{-\frac{(\theta + b - s)^2}{2\sigma_f}}}{\sqrt{2\pi\sigma_f}} & \text{Poisson s, Gaussian b} \\ \vdots \end{cases}$$

A key nuance: θ is never 0, since the first integral is unity (note the normalization N_1, N_2, N_3, which are set such that $\int_0^1 P(\theta)\, d\theta = 1$.)

For the gaussian approximation,

$$^{1}I^{o} = p\frac{1}{2c}\,{}^{1}\,{}^{1}\,4\,{}^{G^{2}\bullet 2}3G$$

$$U = \frac{1\,\frac{B}{f}\,\frac{B}{o}}{1\,\frac{B}{f}\,\frac{B}{o}}$$

Out of these three upper limit methods, Rolke is most commonly use within the eld due to its reliance on the log-likelihood ratio, as a natural extension of the Li & Ma signi cance technique [58]. It is also the default method in the VEGAS analysis package, and used throughout this thesis with the bounded method.

Chapter 3: Prototype Schwarzschild-Couder Telescope

3.1 Introduction

The recent observations of gravitational waves and detections of very-high-energy (VHE) neutrinos and gamma rays have expanded the eld of multi-messenger astrophysics. To further explore the new sources found by these detectors, which encompass high-energy, transient, point-like or extended sources, improvements are necessary to allow coincident observations by VHE detectors for photons with energies greater than 100 GeV.

The present-day IACTs are based on conventional Davies-Cotton (DC) design of a single parabolic re ector [69]. The Davies-Cotton design is a well-tested design for detecting atmospherical Cherenkov light emissions; however, due to spherical and comatic aberrations the DC design is constrained to the larger pixels using photo-multiplier tubes (PMTs) with limited imaging resolution. These limitations of the DC design constrains the potential of the next-generation world-wide observatory Cherenkov Telescope Array (CTA) [70, 71].

The arrays of imaging atmospheric Cherenkov telescopes (IACTs) indirectly observe VHE photons by stereoscopic imaging of the shower cascades and reconstructing their energy and direction from the ground. Current-generation of small arrays of IACTs use prime-focus optical system (OS) based on the DC or parabolic design, which consists of a single segmented mirror with large aperture (>10 m) and f-stop (#) in the range of 1•2 1•4. As an alternative, the fast Schwarzschild-Couder (SCT) OS (0•578) uses two segmented aplanatic mirrors to achieve the best possible imaging resolution in the CTA core energy range (0.1 TeV - 10 TeV). The focal length of the fast SC OS is 5•586 m. These optical improvements enable a high-resolution SiPM camera (see [72]), but require sub-mm and sub-mrad alignment precision.

The SCT OS is designed to fully correct for spherical and comatic aberrations in the full

degrees field of view (FoV). Astigmatism dominates the PSF at the edge of the FoV and it is minimized by implementing a curved focal plane. The choice of a two-mirror design with a de-magnifying secondary mirror reduces the focal plane (FP) plate scale to provide compatibility with the SiPM detectors [73]. The full camera has 11328 pixels, a factor of 6 more than cameras in DC-based medium-sized telescope designs.

Located at the Fred Lawrence Whipple Observatory in Southern Arizona, the construction of the prototype SCT (pSCT) started in late 2015 at the old site of the VERITAS T1 telescope. The optical support structure (OSS) was built by August 2016. The mirrors and camera were assembled and installed by August 2018; however the current installation contains only the central backplane, consisting of 24 camera modules (rather than the full 11328 pixels). The telescope was inaugurated and entered commissioning in January 2019. Important technical advances were made during this time period, as reported by [74–77], with many of those results represented here.

To achieve the point spread function (PSF) of the OS in the FoV compatible with the SiPM pixel size (6 mm), sub-mm and sub-mrad alignment is required [79]. The pSCT has achieved an on-axis PSF within the design goal of 0.26^{0}, allowing it to detect the Crab Nebula in 2020 [78]. The commissioning work to achieve that milestone, and the more recent activity toward global and off-axis alignment, is described in detail in this chapter.

3.2 Optical System Hardware

The SCT is a novel dual mirror optical design. The prototype SCT is built on top of the MST optical support tower with a modified counterweight to contain the larger optical support structure (OSS) needed for two mirrors. Nearby are two support trailers, a maintenance shed and a carport for protecting the camera chiller system. Within the main support trailer, the power, ethernet and optical fiber cables routed from the telescope via underground trenches meet the main computer stack, containing the alignment computer (cta.alignment), camera computer (cta.camera), database computer (cta.database) and network switches. Also present is the main control room containing 8 monitors and two desktop computers (ctt01 and