Radiation and its Use in Astrophysics

Learners' Space Astronomy

THE ASTRONOMY CLUB
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Introduction

1.1 EM Spectrum

Radiation is one of the primary messengers of information in the Universe, that helps us learn more about the sources that emitted the radiations, their temperature, shape , composition, etc. It is essentially through radiation that we can see the universe, observe it and learn more about it. Therefore developing a sound understanding of radiation and its uses in astrophysics is essential. Electromagnetic Radiations i.e EM waves are composed of orthogonally oscillating Electric and Magnetic Fields, with direction of propagation of the wave given the cross product between Electric Field Vector and Magnetic Field Vector at any given point in time.The entire array of electromagnetic waves comprises the electromagnetic (EM) spectrum.The EM spectrum has been arbitrarily divided into regions or intervals to which descriptive names have been applied. At the very energetic (high frequency; short wavelength) end are gamma rays and x-rays. Followed by Ultraviolet, Visible and Infrared region in increasing order of wavelengths. In long wavelength regions we have Microwaves followed By Radio Waves.Some types of electromagnetic radiation easily pass through the atmosphere, while other types do not. The ability of the atmosphere to allow radiation to pass through it is referred to as its transmissivity, and varies with the wavelength/type of the radiation. The gases that comprise our atmosphere absorb radiation in certain wavelengths while allowing radiation with differing wavelengths to pass through.The areas of the EM spectrum that are absorbed by atmospheric gases such as water vapor, carbon dioxide, and ozone are known as absorption bands. In the figure, absorption bands are represented by a low transmission value that is associated with a specific range of wavelengths.In contrast to the absorption bands, there are areas of the electromagnetic spectrum where the atmosphere is transparent (little or no absorption of radiation) to specific wavelengths. These wavelength bands are known as atmospheric "windows" since they allow the radiation to easily pass through the atmosphere to Earth's surface.Figure below illustrates the concept.This explains why radio telescopes are generally ground based telescopes while X-ray Telescopes are space-based telescopes.

1.2 Definitions

$LUMINOSITY$: Luminosity (L) or power is the rate at which energy is emitted. Its SI units are J s^{-1}1 or watts ($1W = 1Js^{-1}$) and $ergs^{-1}$ in cgs. One calculates a source's luminosity or power by dividing the amount of energy emitted by the length of time over which the energy was emitted.

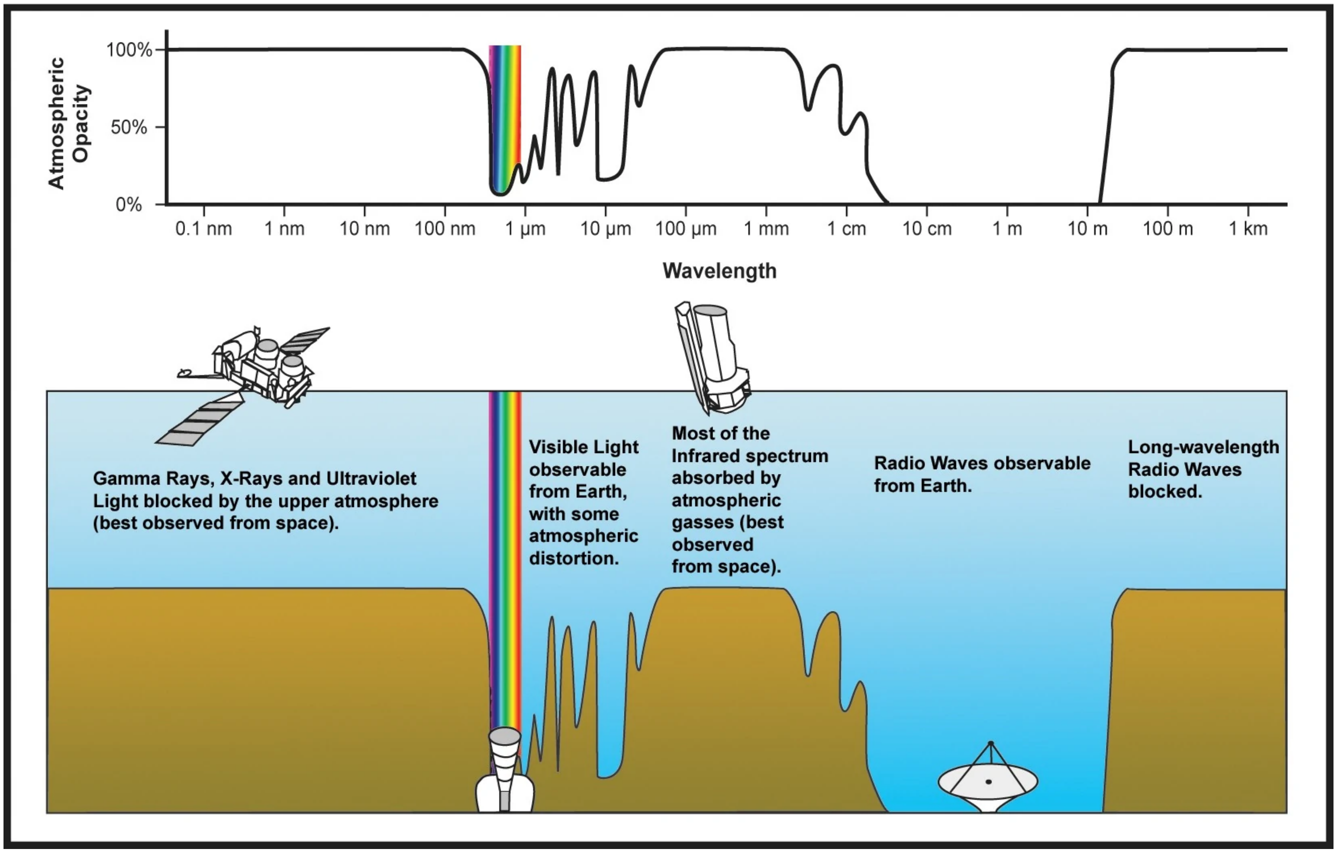

Figure 1.1

This yields the rate of emission.

FLUX : The amount of light energy per unit time per unit area. The units of flux are $Js^{-1}m^{-2}$ or Wm^{-2} (SI) and $ergs\ s^{-1}cm^{-2}$ in (cgs). The Flux received by our Telescope is given by the equation :

$$F_{received} = \frac{L}{4\pi d^2}$$

, And the total power of radiation collected by telescope is given by

$$P = L\frac{A_{eff}}{4\pi d^2}$$

FLUX DENSITY : Flux density (F_ν or F_λ) is the flux per unit frequency in the observed spectral range, and it equals the detected flux divided by the width in frequency of the observation.Therefore,

$$F_\nu = \frac{F}{\Delta\nu}$$

where: $\Delta\nu$ is the bandwidth, or the range in frequency of the detected EM waves.The symbol S_ν (rather than F_ν) is often used by radio astronomers to represent flux density per unit bandwidth. When working at visible wavelengths, astronomers tend to measure flux density in terms of wavelength rather than frequency, so flux density is also often defined as flux per unit wavelength.

$$F_\lambda = \frac{F}{\Delta\lambda}$$

Although these two quantities are the same conceptually, they are not the same quantitatively (Their dimensions are different). The Flux Density is the characteristic of the source that we want to infer from the data. The amount of power a telescope collects from a source of given flux density is given by

$$P = F_\nu A_{eff} \Delta\nu$$

We get awkwardly small numbers for Flux Density measurements in SI or CGS Units. Ergo, Radio Astronomers have defined a unit of Flux Density named after Father of Radio Astronomy Karl Jansky. In terms of SI Units, 1 Jansky is defined as :
1 Jansky = 10^{-26} W m^{-2} Hz^{-1}

INTENSITY : Intensity is the Flux Density per unit solid angle. It is also a direct measure of surface brightness of the source. If you know the solid angle of the source (Ω), you can calculate the source's average intensity by dividing the source's Flux density with the solid angle. The units of Flux Density can be given by W m^{-2} Hz^{-1} sr^{-1}. Again I_ν and I_λ, though conceptually same are quantitatively very different. The expression for I_ν :

$$I_\nu = \frac{L}{4\pi d^2 \Delta\nu \Omega}$$

One of the most important aspects of Intensity is that it is independent of the distance from the source and is a direct measurement of the Surface Brightness of the Source.

Blackbody Radiation

2.1 History

Blackbody Radiation was a major challenge for physicists in late 19th Century as Classical Physics failed to provide a complete explanation of the phenomenon. One of the major advances in classical Physics came in the form of Rayleigh-Jeans Law and led to one of the most famous problems in Physics, *The Ultraviolet Catastrophe*.

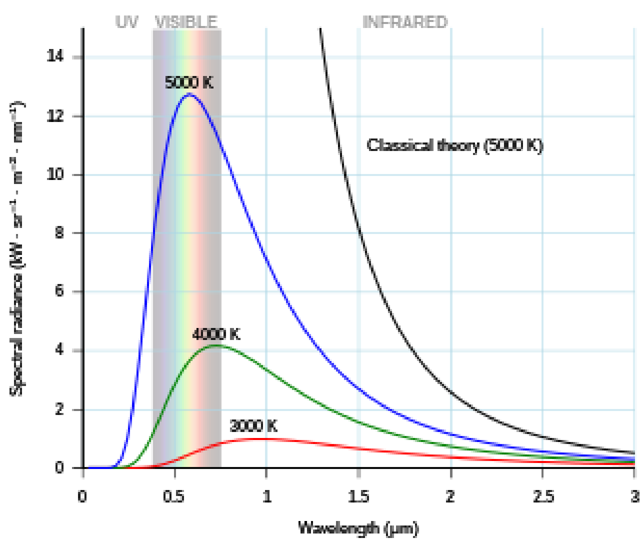

Figure 2.1: Prediction Of Classical Theory : The Ultraviolet Catastrophe

2.1.1 Rayleigh-Jeans Law and The Ultraviolet Catastrophe

The Rayleigh-Jeans law is an approximation to blackbody radiation through classical arguments. For wavelength λ it is given by:

$$B_\lambda = \frac{2ck_BT}{\lambda^4}$$

,Where c is the speed of light,
k_B is the Boltzmann Constant and,

T is the temperature of the Source. For frequency ν it is given by:

$$B_\nu = \frac{2\nu^2 k_B T}{c^2}$$

This formula is obtained from the equipartition theorem of classical statistical mechanics which states that all harmonic oscillator modes (degrees of freedom) of a system at equilibrium have an average energy of $k_B T$. The ultraviolet catastrophe is the expression given to the misbehavior of the Rayleigh Jeans law at higher frequencies, which is the prediction that the intensity of radiation emitted by an ideal black body at thermal goes to infinity as frequency increases (or wavelength decreases).It has been illustrated in the Figure 2.1.Once we learn about the Plank's Function in the succeeding section, we can obtain the Rayleigh-Jeans Law by Taylor expanding the exponent term in the expression for the Plank's Function.

2.1.2 Wein's Approximation

Wein's Law is approximation to explain blackbody radiation and accurately describes it for higher frequencies or short wavelengths. Wein derived his law from *thermodynamic* arguments and using *Maxwell-Boltzmann* Distribution for atoms. The approximation is written as :

$$B_\nu = \frac{2h\nu^3}{c^2} e^{-\frac{h\nu}{k_B T}}$$

In terms of wavelength, the law is given by:

$$B_\lambda = \frac{2hc^2}{\lambda^5} e^{-\frac{hc}{\lambda K_B T}}$$

2.2 Planks's Breakthrough

Max Planck's discovery that the spectrum of the radiation emitted by hot, opaque objects can be fit by a model that assumes quantized units of energy led him to win a Nobel Prize in 1918. A body which absorbs all radiation incident on it, reflects none and prevents transmission of light is called a *blackbody*, and is called opaque. The radiation emitted by such a body is called blackbody radiation. Thus, Blackbody Radiation is maximum amount of radiation that an ordinary body of a particular temperature will emit solely as an attempt to cool. The resultant radiation emitted from a body can be modelled as a cloud of photons at the same temperature as the body itself. When the photons have been thermalized, statistics can describe the distribution of energies of the photons, that is, the relative number of photons carrying each small range of energy. In other words, with blackbody radiation, the photons have achieved a thermal distribution of energies. The energy distribution function of photons, (Number of photons as a function of their energy) is directly related to spectral distribution, i.e. Intensity v/s Frequency. The frequency is proportional to the energy (since for photons $E = h\nu$) and the intensity at a given frequency is proportional to the flux of photons at that frequency.

The *Plank's Function* provides a mathematical description of the spectrum of the light emitted by blackbodies. In terms of the emitted flux per unit frequency interval per unit steradian, the Planck function is given by:

$$B_\nu = \frac{2h\nu^3}{c^2}\frac{1}{e^{\frac{h\nu}{kT}}-1}$$

Where,
h is the Plank's Constant,
k is the Boltzmann Constant,
c is the Speed of Light,
ν is the frequency of observation, and
T is the temperature of radiating body in Kelvins

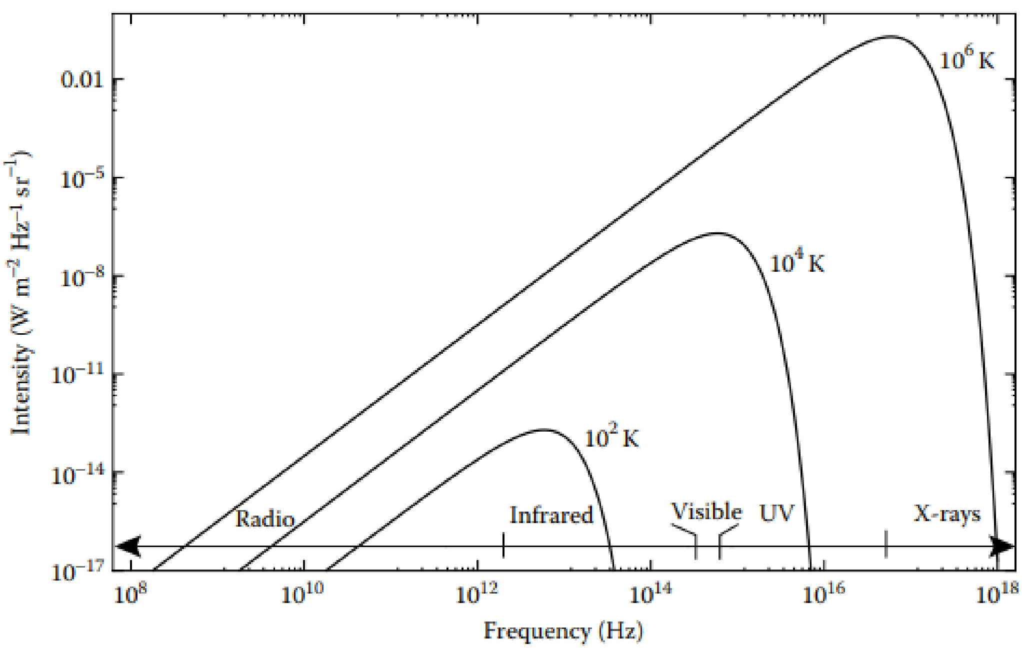

Figure 2.2: Log-log plot of $B_\nu(T)$ versus ν for three different temperatures.

Plank's Function can also be expressed as Flux per unit *wavelength* per steradian, which is B_λ, and is given by :

$$B_\lambda = \frac{2hc^2}{\lambda^5}\frac{1}{e^{\frac{hc}{\lambda kT}}-1}$$

Again, B_λ and B_ν being conceptually the same, are not the same quantities. They are not even the same function. The total flux of radiation (W m^2) emitted by the body can be obtained by integration of the Planck function over frequency and solid angle. The result shows that the total flux is proportional to the fourth power of the body's temperature and is called *Stefan-Boltzmann Law* :

$$F = \sigma T^4$$

where, σ is known as the *Stefan-Boltzmann Constant* = 5.67 x 10^{-8} W m^{-2} K^{-4}. A key feature

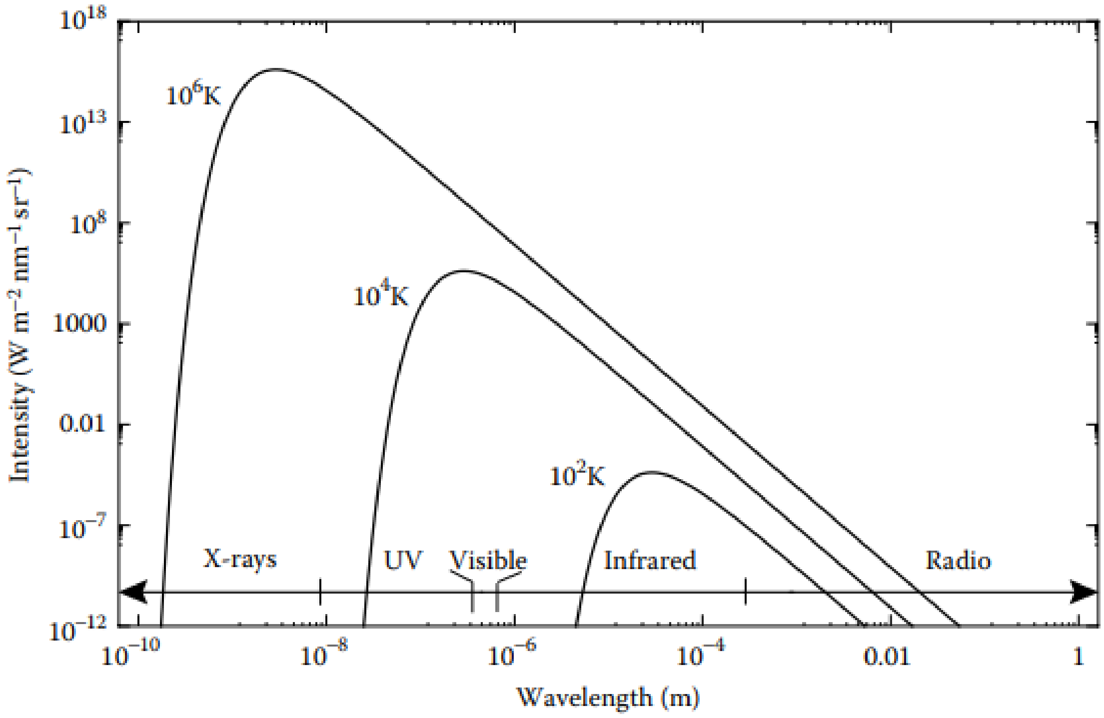

Figure 2.3: Log-log plot of B_λ(T) versus λ for three different temperatures.

of Blackbody Radiation, as apparent in figures 2.1 and 2.2 is that value of B_ν increases with temperature of every frequency, which means that as the temperature of blackbody increases, it emits more energy at every frequency. In practical terms,this has an important implication that at any given frequency, any particular value of intensity corresponds to exactly one temperature. Note that the blackbody curves for different temperatures never cross. Therefore, even if you know the intensity of an opaque object at only a single frequency, and if the radiating source is a blackbody, then you can infer the temperature of the source by this single intensity measurement. The Intensity can be inferred from measurements of both flux density and angular size of the source. The universe's cosmic microwave background (CMB), which fills the entire sky, is a great example of a situation where temperature can be inferred from a single-frequency observation.The value of temperature we get is about 2.7 Kelvins. An important equation to remember, which we do not go into derivation of, is that average Kinetic Energy of Photons at a given temperature of Photon Gas is given by the equation $< E >= 2.7kT$, where k is the *Boltzmann's Constant*.

Figures 2.1 and 2.2 show that peaks of the Plank Spectrum depends only the temperature of the body. As the temperature increases, the peak of the Blackbody spectrum occurs at a higher frequency. The peak can be obtained by differentiating B_ν and B_λ with respect to temperature and setting them to be zero. This gives us Wien's Displacement Law. For B_ν,

$$\nu_{peak} = (5.879 \times 10^{-10})T$$

For B_λ,

$$\lambda_{peak} = (2.898 \times 10^{-3})T$$

An important thing to note here is that the above two equations do not give the same information, which is inherent due to the fact that I_ν and I_λ are two different functions. You can check this by substituting $c = \nu\lambda$ in Planks Function for frequency, you will not get the correct expression for B_λ. The correct conversion factor for converting between I_ν and I_λ is given by :

$$I_\lambda = \frac{c}{\lambda^2} I_\nu$$

If you use this conversion in either of the Planck functions, you will find that one does convert to the other.

2.3 Applications

2.3.1 Applications in Astronomy

In astronomy, stars are often modelled as blackbodies, although it is not always a good approximation. Therefore,the temperature of a star can be deduced from the wavelength of the peak of its radiation curve.
In 1965, the cosmic microwave background radiation (CMBR) was discovered by Penzias and Wilson, who later won the Nobel Prize for their work. It took about 300 000 years for the Universe to cool down to a temperature at which atoms can form (about 3000°C). Matter then became neutral, and allowed the light to travel freely: the Universe became transparent. The relic of that 'first light' is the CMB.Since the time when that radiation was released, the Universe has expanded, becoming at the same time cooler and cooler. The cosmic background has been affected by the same process: it has expanded and cooled down, and today we observe it in the Microwave Spectrum. The radiation spectrum was measured by the COBE satellite and found to be a remarkable fit to a blackbody curve with a temperature of 2.725 K and is interpreted as evidence that the universe has been expanding and cooling for about 13.7 billion years. A more recent mission, WMAP, has measured the spectral details to much higher resolution, finding tiny temperature fluctuations in the early Universe which ultimately led to the large-scale structures we see today. The CMB is the farthest and oldest light any telescope can detect. It is impossible to see further beyond the time of its release because then the Universe was completely 'opaque'. The CMB takes astronomers as close as possible to the Big Bang, and is currently one of the most promising ways we have of understanding the birth and evolution of the Universe in which we live.

2.3.2 Calculating Temperature of the Earth

Here, we will apply the Stefan's Law mentioned above to theoretically calculate the Temperature of the Earth. We assume the Earth to be a perfect blackbody. The radiation from the Sun accounts for almost the entire incoming radiation received by the Earth. By Stefan's Law, we can calculate the Luminosity the Sun:

$$L_{sun} = 4\pi R_{sun}^2 \sigma T_{sun}^4 = (5.67 \times 10^{-8}) \times (5772)^4 \approx 3.83 \times 10^{26} W$$

, Where $T_{sun} = 5772K$, is taken to be the *surface temperature* of the Sun, and
R_{sun} is the radius of the Sun = 695700 km.
The total Power received the Earth can then be calculated as :

$$P = L_{sun} \frac{A_{eff}}{4\pi d^2} = L_{sun} (\frac{R_{earth}}{2d})^2$$

, where, A_{eff} of earth is taken to be πR_{earth}^2,
R_{earth} = 6378km,
d is the distance between Earth and the Sun = 149,597,871 Km.
Now, as we have assumed Earth to be a blackbody, the power of radiation emitted by Earth will be equal to the power of Radiation absorbed the Earth. By Stefan's Law, we can write the power radiated by Earth as :

$$P = 4\pi R_{earth}^2 \sigma T_{earth}^4 = 4 \times A_{eff} \sigma T^4$$

Equating, the above two equations, we get

$$T_{earth} = \sqrt{\frac{R_{sun}}{2d}} \times T_{sun} \approx 279K.$$

It is no surprise that the surface temperature of Earth we get assuming it to be a perfect blackbody is much lower than we expect, because in reality, the atmosphere of the Earth absorbs heat and keeps Earth Warm for life to thrive.

Photometry

Photometry is the science of measuring the amount of light received from a star. The word comes from the Greek 'photo' meaning light and 'metron' to measure. Photometry measures the flux of the star. This is the amount of energy from the star that has reached us. The units of flux are Watts per square metre (Wm^2).Photometry is often used together with called spectroscopy. In photometry, astronomers use filters to measure the amount of light received within specific wavelength ranges, allowing them to quantify the brightness of objects in different passbands.we will explore concepts of apparent magnitude and absolute magnitude. We will also discuss passbands, specifically the uvriz and JHK filters, and their role in inferring the temperature of stars.

3.1 Apparent Magnitude

Apparent magnitude m of a star is a number that tells how bright that star appears at its great distance from Earth. The scale is "backwards" and logarithmic. Larger magnitudes correspond to fainter stars. Note that brightness is another way to say the flux of light, in Watts per square meter, coming towards us. On this magnitude scale, a brightness ratio of 100 is set to correspond exactly to a magnitude difference of 5. As magnitude is a logarithmic scale, one can always transform a brightness ratio $\frac{B2}{B1}$ into the equivalent magnitude difference $m2 - m1$ by the formula:

$$m2 - m1 = -2.50log(\frac{B2}{B1})$$

You can check that for brightness ratio $\frac{B2}{B1} = 100$, we have $log(\frac{B2}{B1}) = log(100) = log(10^2) = 2$, and then $m2 - m1 = -5$, the basic definition of this scale (brighter is more negative m).

3.2 Absolute Magnitude

Absolute magnitude M_v is the apparent magnitude the star would have if it were placed at a distance of 10 parsecs from the Earth. Doing this to a star, will either make it appear brighter or fainter. From the inverse square law for light, the ratio of its brightness at 10 parsecs to its brightness at its known distance d (in parsecs) is :

$$\frac{B_{10}}{B_d} = (\frac{d}{10})^2$$

Then we say, the formula for absolute magnitude is :

$$M_v = m - 5log(\frac{d}{10})$$

3.3 Passbands and Photometric Filters

Passbands are specific wavelength ranges in which astronomers observe celestial objects. Photometric filters are used to select and isolate light within these passbands for measurement. Different filters are denoted by letters representing the colors they capture. Some common filters include:

1. UVRIZ : These filters are used in the ultraviolet, violet, red, infrared, and near-infrared regions of the electromagnetic spectrum, respectively.

2. JHK : These filters are used in the near-infrared to study cooler objects like stars with lower temperatures.

3.3.1 Inferring Temperatures of Stars

Photometric filters, especially the UVRIZ and JHK passbands, are valuable tools for inferring the temperature of stars. The concept behind this is related to the blackbody radiation curve, which states that hotter objects emit more energy at shorter (bluer) wavelengths, while cooler objects emit more energy at longer (redder) wavelengths.By observing stars through various passbands, astronomers can measure their fluxes in different wavelength ranges. For example, a star's flux measured through the u filter will primarily depend on its ultraviolet emission, while the flux through the J filter will predominantly depend on its near-infrared emission. Comparing the fluxes measured in different passbands, astronomers can construct a color-color diagram or a color-magnitude diagram. The positions of stars in these diagrams can then be compared to theoretical models to estimate their temperatures. You will learn more about this in-depth in Stellar Physics Module. For instance, if a star appears brighter in the ultraviolet (u) passband compared to the infrared (K) passband, it is likely a hot star emitting more ultraviolet radiation, suggesting a higher temperature. Conversely, if a star appears brighter in the infrared (K) passband compared to the ultraviolet (u) passband, it is likely a cooler star emitting more infrared radiation, suggesting a lower temperature.

Spectroscopy

A powerful tool for analyzing the detected radiation is to examine its spectrum (a plot or display of the amount of radiation vs. frequency or wavelength), as the details of a source's spectrum contain much valuable information about the physics of the source. The three basic types of spectra are as follows :

Continuous Spectra : When a radiation source emits at all frequencies over a range without breaks, the spectrum is called a continuous spectrum and the emitting object is called a continuum source. A classic example of a continuum source is an incandescent lamp. A plot of a continuous spectrum, for example, will show intensities over a broad range of frequencies, although the intensity of the emission can vary significantly. A qualitative example of a continuous spectrum is shown in Figure 3.

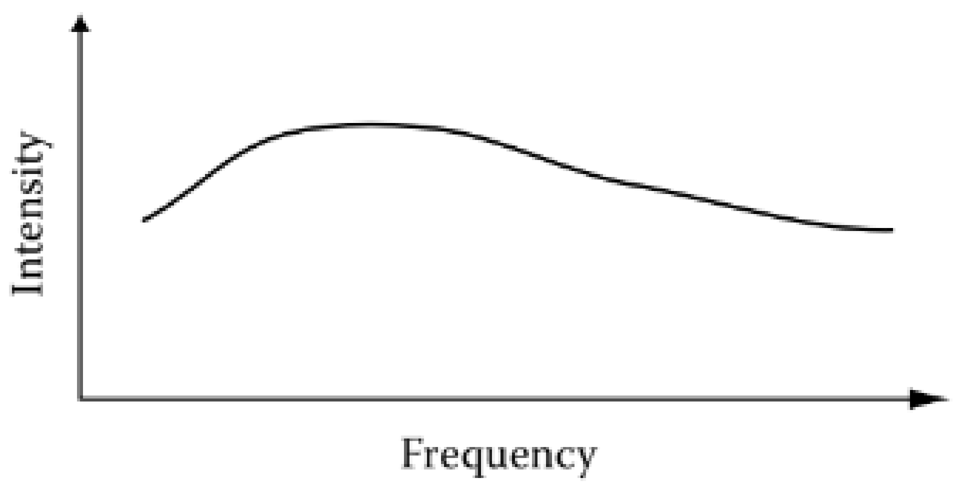

Figure 4.1: Graphical representation of Continuous Spectrum

Emission Spectra : When a radiating object emits radiation only at some very specific frequencies, or wavelengths, the spectrum contains a set of discrete bright lines. These lines of light are called emission lines. The reason that a light source would emit only at some very specific frequencies is due to the quantum physics of atomic and molecular structures. An atom or molecule can be in an excited energy state and then spontaneously drop to a lower energy state. To conserve energy, it gives off a photon that carries the exact amount of energy that the atom or molecule loses. However, due to the laws of quantum mechanics, the internal energy levels of atoms and molecules are restricted to a set of discrete values. Thus, changes in energy (and hence the emitted photon energy or frequency) can only have certain specific values.

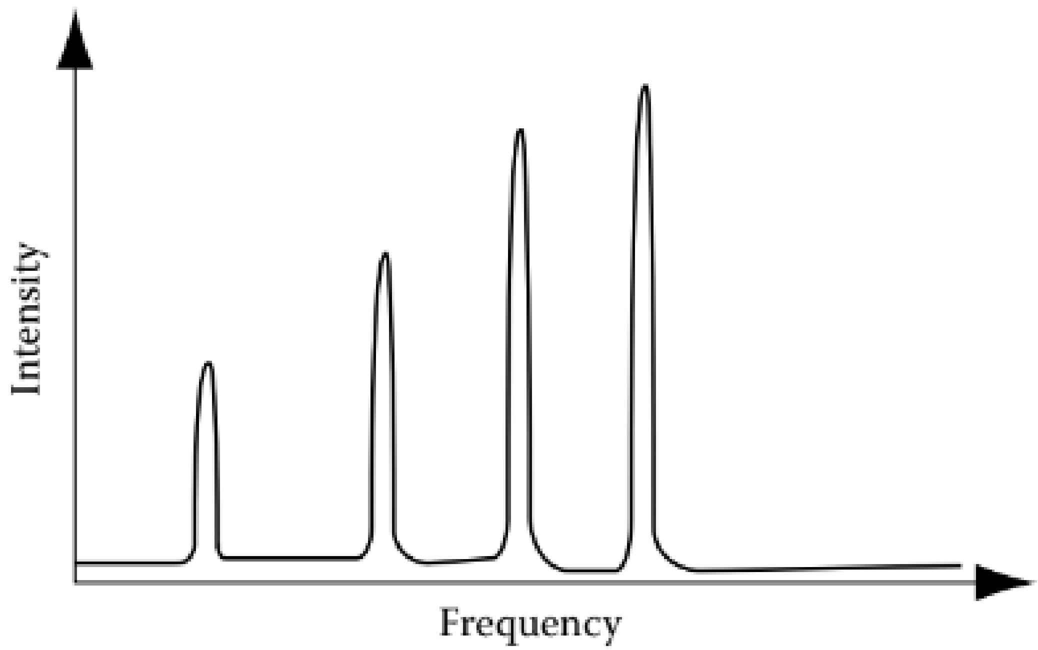

Figure 4.2: Graphical representation of Emission Spectrum

Absorption Spectra : An interesting case arises when the radiation from an intense continuum source passes through a cool, transparent gas of atoms or molecules. Some of the atoms or molecules in the gas can absorb photons from the continuum to raise them into a higher allowed energy level. The photons that they absorb must have exactly the same amount of energy that the atom or molecule gains, and so only the photons of certain, specific frequencies will be removed from the continuum. To an observer stationed beyond the cool gas, the spectrum will show the continuous spectrum of the background source with the radiation at these specific frequencies appearing darkened, thus the dark line nomenclature. In a plot of intensity versus frequency, they will appear as dips or decreased intensity, as shown in Figure 5. These dips are called absorption lines. The atomic or molecular absorption frequencies are the same as the corresponding emission frequencies, and so again, the chemical composition of a gas cloud can be identified by the frequencies of the absorption lines that it produces.

4.1 Doppler Shift in Light

When we observe the emission or absorption spectra of stars in the universe, we observe their spectral lines to be shifted. This is a consequence of *relativistic* doppler effect in light coming

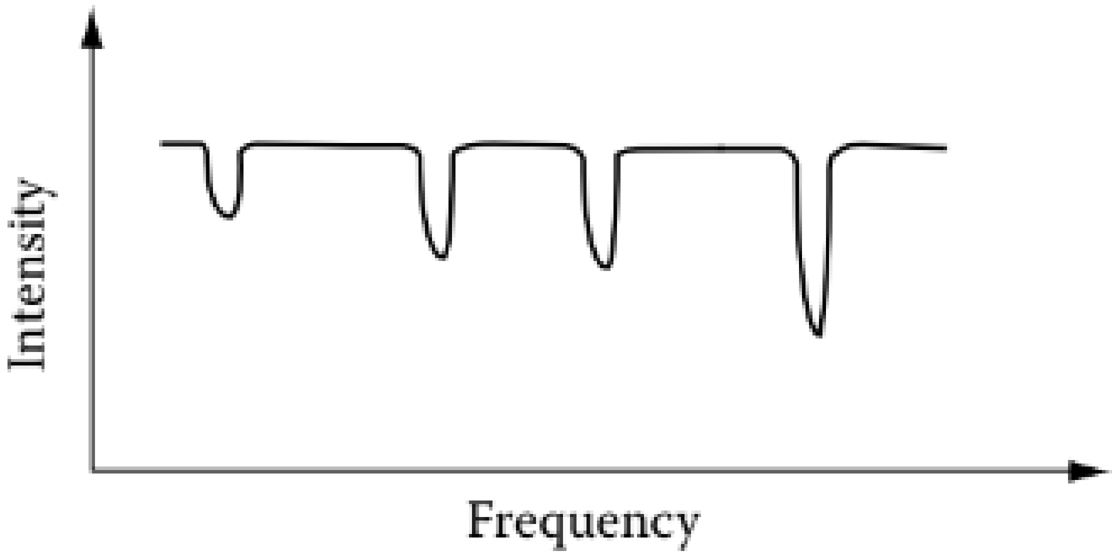

Figure 4.3: Graphical representation of Absorption Spectrum

from the stars. If the source is moving away from us, the wavelength of the radiated radiation is observed to longer than at the time it was radiated, and the radiation is said to be *redshifted*. Whereas, if the source is moving towards us, the wavelength of the radiation is observed to be shorter than at the time it was radiated, and the radiation is said to be *blueshifted*. Astronomers encapsulate the redshift of incoming radiation by a single number z.

$$z = \sqrt{\frac{c+v}{c-v}} - 1$$

,where c is the velocity of light and
v is the recession velocity of the source.
For low recessional velocities, the expression of redshift can be approximated as :

$$z = \frac{v}{c}$$

The recession of velocity of a source is the rate at which the source recedes from the observer as a result of the expansion of the universe. Apart from playing an important role in broadening of spectral lines, doppler shift in light is used to measure distance to stars. This cam be done using the Hubble's Law. In 1929, Edwin Hubble gave the first inconvertible proof that their galaxies outside the Milky Way. Hubble discovered, that the galaxies are not only moving away from us, but their velocity of *recession* is directly proportional to their distance from us, and this proportionality is encapsulated by the Hubble's Constant H_0.

$$v = H_0 d$$

where, v is the recessional velocity of the star

H_0 is the Hubble's Constant = 70 (km/s)/Mpc, and
d is the distance between the source and the observer. Therefore, by observing the redshift of spectral lines in emission and absorption spectra, we can find distance to the stars. Some important spectral lines like the *21 cm line* and the H_α line are often used in such calculations. These lines are discussed in a separate section.

4.2 Broadening of Spectral Lines

4.2.1 Intrinsic Broadening

Natural broadening is directly related to the principle of Heisenberg. This is one of the basic principles of quantum mechanics, and is characteristic of particle waves behavior. The principle, in its energy-time version, states that in a physical phenomenon the product of the uncertainties ΔE and Δt assumes the minimum value:

$$\Delta E \Delta t >= \frac{h}{4\pi}$$

Now we know that the lifetime of excited atomic states is finite (of the order of microseconds-nanoseconds), so also the uncertainty Δt has this value. From Heisenberg's relation it can be deduced that $\Delta E \neq 0$ (would be null in the case of infinite lifetime) and therefore also the frequency of the photon emitted has an intrinsic uncertainty, resulting in thickening of spectral lines. The value of uncertainity in most cases is very low and almost negligible.

4.2.2 Thermal Doppler Broadening

Thermal broadening is due to the doppler effect that occurs when the radiating atoms have a movement relative to the observer. The random atomic movement of the atoms is directly related to the temperature, which is why this broadening mechanism is called thermal. As we know from the Maxwell-Boltzmann statistics the speed of gas atoms follow a gaussian profile and the mean energy is on the order of kT, where k is the Boltzmann constant and T is the temperature. Mathematically it can be shown that the spectral line broadening profile is also Gaussian, with a value given by:

$$\Delta \nu = \frac{\nu_0}{c}\sqrt{\frac{2kT}{m}}$$

This equation shows that broadening increases with temperature and is higher for lighter atoms.

4.2.3 Rotational-Doppler Broadening

For a star rotating with time period of rotation T, the doppler velocity can be calculated as :

$$v = \frac{2\pi r}{T}$$

,where r is the radius of the star. Using the doppler veocity, we can calculate $\Delta \nu$ as :

$$\Delta \nu = \nu \frac{v}{c}$$

4.2.4 Collision Bradening

The last phenomenon that causes broadening of spectral lines is the collisions between excited atoms which can become predominant in high-density gas plasma emissions (high pressures). In this case, the spectral line profile is not Gaussian but Lorentzian, characterized by a narrower spike and longer wings. In the context of high pressure and high temperatures, the phenomenon of resonant self-absorption can also occur. This phenomenon consists in the auto-absorption of the emitted radiation by colder external gas layers. This phenomenon can also be easily observed with our instrumentation using a high pressure sodium lamp.

4.2.5 Surface Gravity of Star

Many of the strongest spectral lines (e.g. Balmer absorption lines and resonance lines of metals) are very sensitive to the surface gravity of the star. This enables a distinction between main sequence dwarfs and giants because a giant star's surface gravity is factors of $\sim$100 lower than that of a dwarf star of the same temperature and has narrower absorption lines. Conversely, white dwarfs have much broader lines, because their surface gravities are $\sim 10^4$ times larger than a main sequence star.

Different atoms and molecules emit different sets of frequencies, and so the composition of a radiation source can be identified by determining the frequencies of the emission lines. The science of spectroscopy is quite sophisticated. From spectral lines astronomers can determine not only the element, but the temperature and density of that element in the star. The spectral line also can tell us about any magnetic field of the star. The width of the line can tell us how fast the material is moving. We can learn about winds in stars from this. If the lines shift back and forth we can learn that the star may be orbiting another star. We can estimate the mass and size of the star from this. If the lines grow and fade in strength we can learn about the physical changes in the star.

4.3 Some Important Spectral Lines

In all the Universe, the most common atom of all is hydrogen, with just one proton and one electron. Wherever new stars form, hydrogen atoms become ionized, becoming neutral again if those free electrons can find their way back to a free proton. Although the electrons will typically cascade down the allowed energy levels into the ground state, that normally produces only a specific set of infrared, visible, and ultraviolet light. Among the spectra of lines produced, we will talk to very special lines, the 21 cm line and the H_α line in Hydrogen spectrum.
The H_α line is the first line in Balmer Series of Hydrogen Atom spectra, is one of the most prominent lines in optical region spectral lines of stars. Therefore, it is one of the easiest ways for astronomers to trace out the hydrogen content of gas clouds in Universe. Moreover as discussed above, it is often used to properly measure doppler shift broadening of spectral lines and in calculating recessional velocity of the source.
The 21 cm line is often called the *the magic length* in Astronomy and has potential to unlock deppest mysteries of the Universe. The phenomenon of *Quantum Tunneling* makes it possible for many forbidden transitions in Quantum Mechanics to occur. The invocation of the phenomenon of Quantum Tunneling allows use to go beyond the mere "fine-structure" of the Atomic Spectra to hyperfine structure,where the spin of the atomic nucleus and one of the electrons that orbit

it begin in an “aligned” state and transition to an "anti-aligned" state. In "aligned" state, the spins of the atomic nucleus and the orbiting electron are in the same direction, whereas in the "anti-aligned" state, they are reversed.

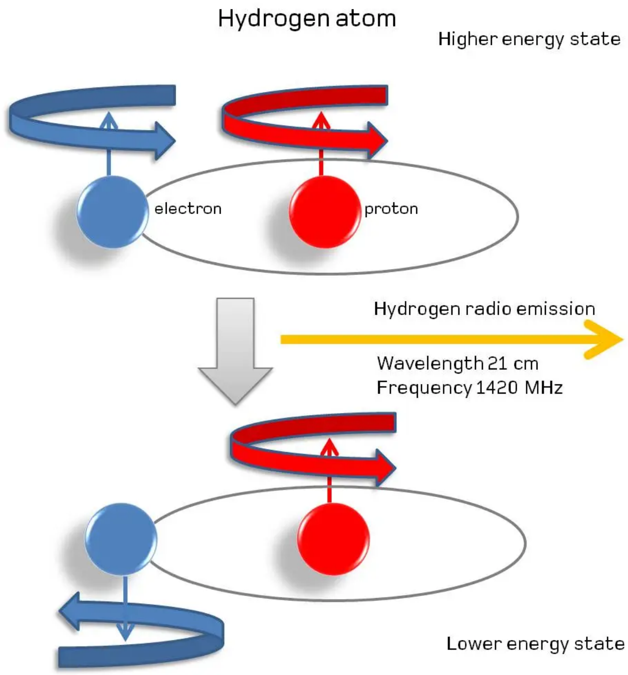

Figure 4.4: Transition from aligned to anti aligned state

The most famous of these transitions occurs in the simplest type of atom of all: hydrogen (Figure 1.6). With just one proton and one electron, every time you form a neutral hydrogen atom and the electron cascades down to the ground (lowest-energy) state, there’s a 50 percent chance that the spins of the central proton and the electron will be aligned, with a 50 percent chance that the spins will be anti-aligned. If the spins are anti-aligned, that’s truly the lowest-energy state; there’s nowhere to go via transition that will result in the emission of energy at all. But if the spins are aligned, it becomes possible to quantum tunnel to the anti-aligned state: even though the direct transition process is forbidden, tunneling allows you to go straight from the starting point to the ending point, emitting a photon in the process.This transition, because of its “forbidden” nature, takes an extremely long time to occur: approximately 10 million years for the average atom. However, this long lifetime of the slightly excited, aligned case for a hydrogen atom has an upside to it: the photon that gets emitted, at 21 centimeters in wavelength and with a frequency of 1420 megahertz, is intrinsically, extremely narrow. In fact, it’s the narrowest, most precise transition line known in all of atomic and nuclear physics!
The large star forming molecular clouds in space are very cold and they do not have any spectral lines in shorter wavelength regions of EM Spectrum such as visible. Therefore, astronomers rely heavily on the 21cm line to identify the gas clouds and also learn more about their shape and extent. Using this line, along with the doppler shift of light, one can measure the rotational velocity of galaxies. This eventually led to discovery of dark matter, where the galaxy rotation curve obtained from the measurements 21cm line revealed that stars at boundaries of galaxies revolve around the galactic centre at much higher velocities than expected.
Trivia: Even though this line is popularly known as 21cm line, its *exact* wavlength is 21.106cm.