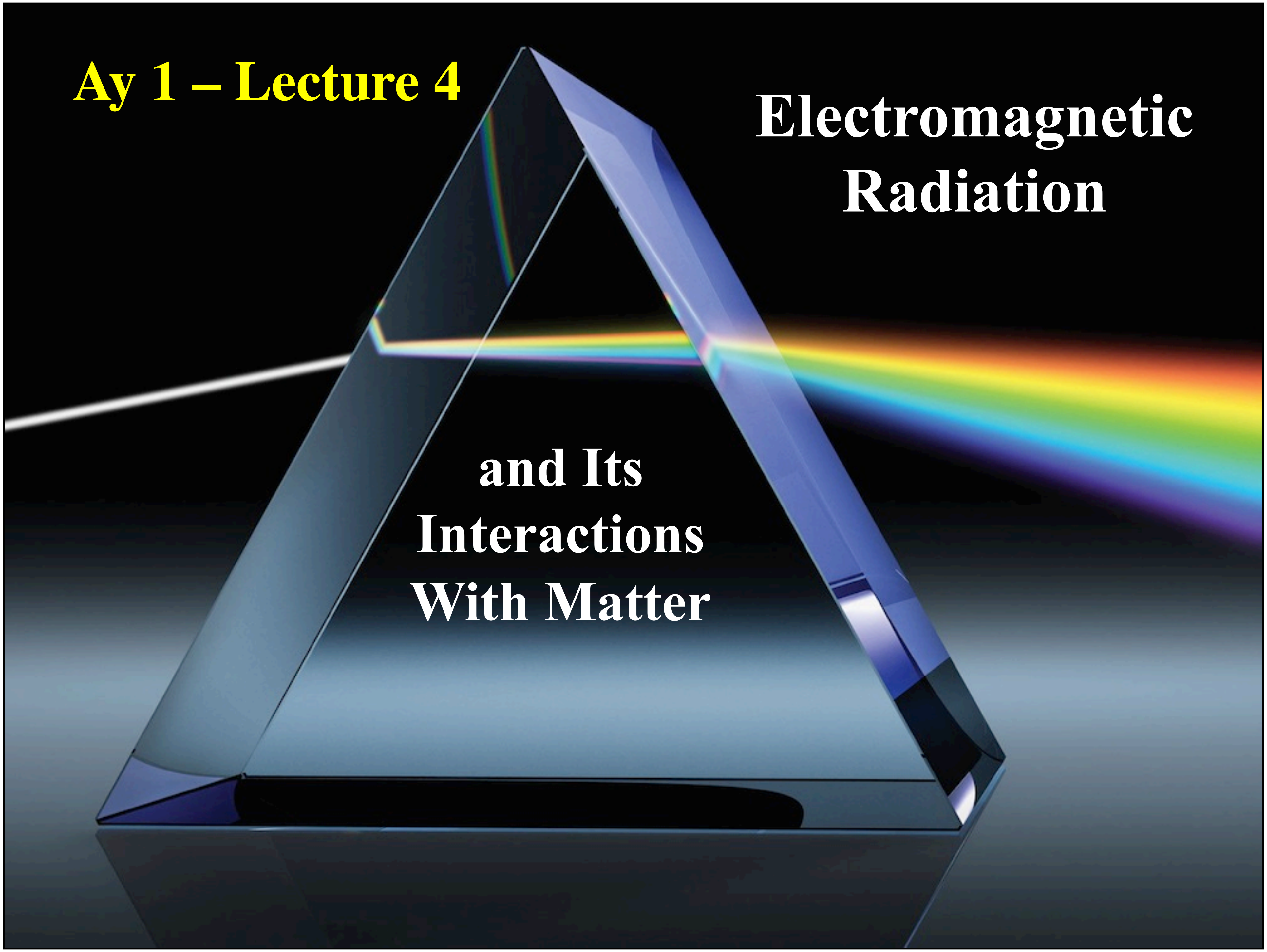
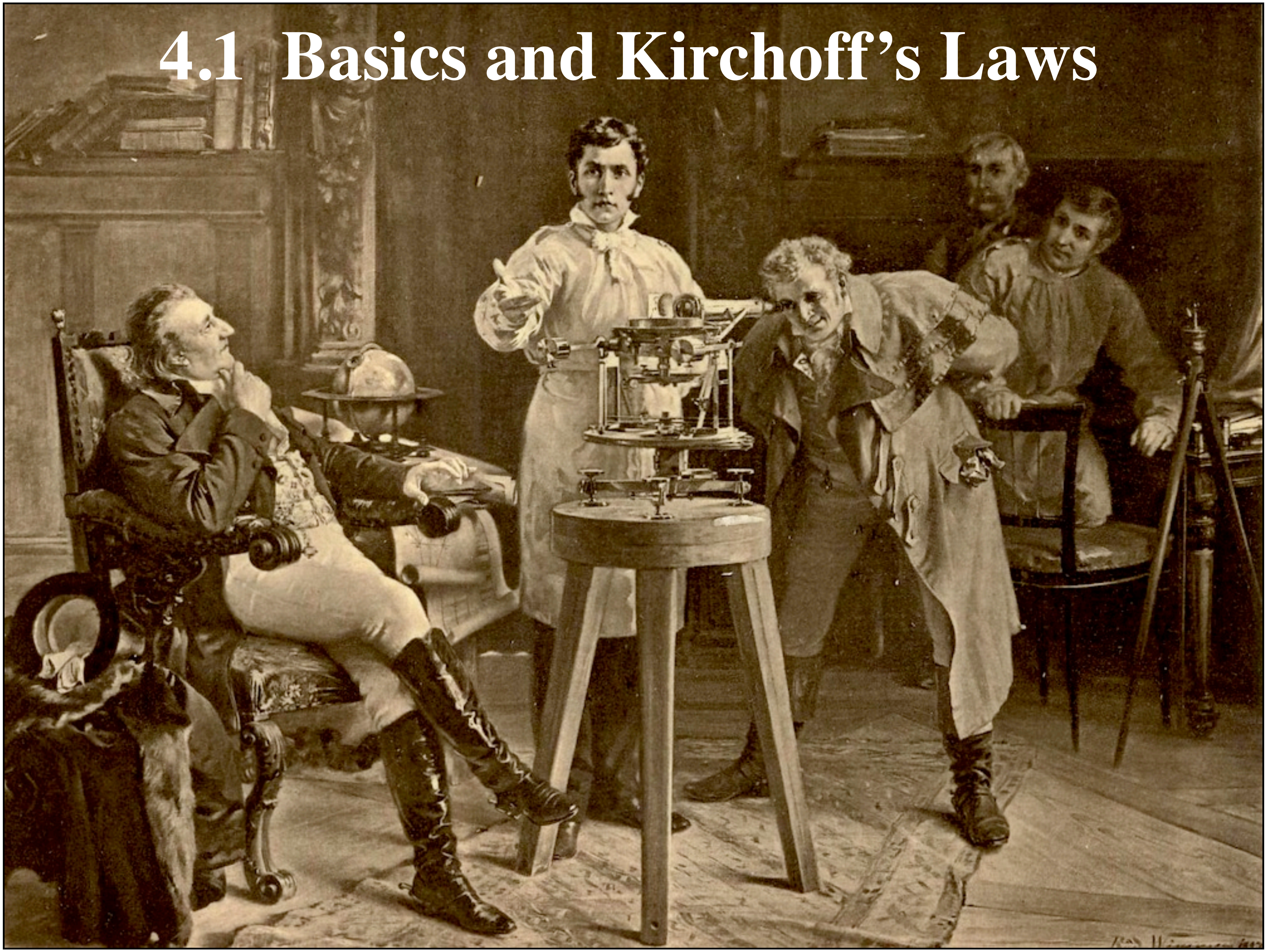

Ay 1 – Lecture 4
Electromagnetic
Radiation
and Its
Interactions
With Matter

4.1 Basics and Kirchoff’s Laws



Photon Energies

Electromagnetic radiation of frequency ν, wavelength λ, in free space obeys:

$$\lambda \upsilon = c$$

Individual photons have energy: $E = h\nu$

h = Planck's constant

c = speed of light

$$h = 6.626 \times 10^{-27} \text{ erg s}$$

$$c = 3.0 \times 10^{10} \text{ cm s}^{-1}$$

Energies are often given in electron volts, where:

$$1 \text{ eV} = 1.6 \times 10^{-12} \text{ erg} = 1.6 \times 10^{-19} \text{ J}$$

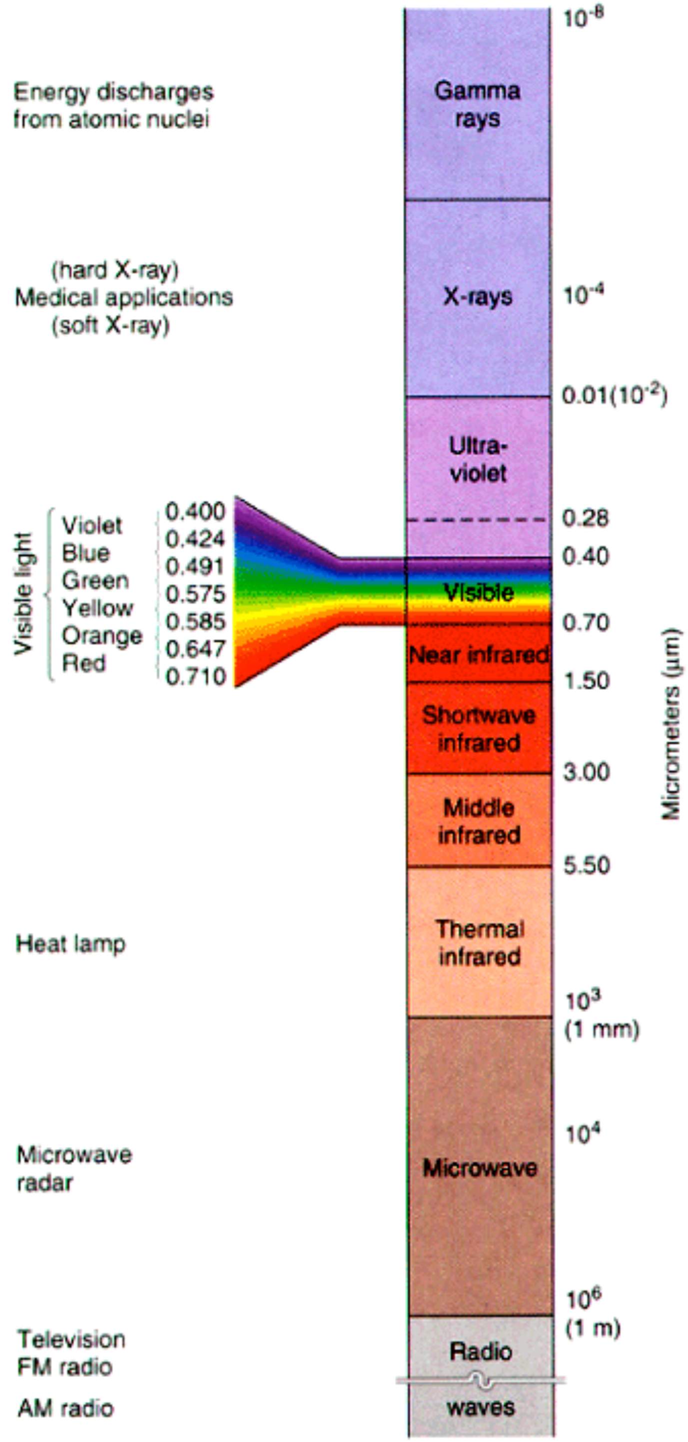
Primary Astrophysical Processes Producing Electromagnetic Radiation

- ***When charged particles change direction (i.e., they are accelerated), they emit radiation***
- ***Quantum systems (e.g., atoms) change their energy state by emitting or absorbing photons***

- Discrete (quantum transitions)
- Continuum
  - Thermal (i.e., blackbody)
  - Nonthermal: Synchrotron
    Free-free
    Cherenkov

Which one(s) will dominate,
depends on the physical conditions of the gas/plasma.
Thus, EMR is a *physical diagnostic*.

Different Physical Processes Dominate at Different Wavelengths



source: Christopherson (2000) *Geosystems*

Nuclear energy levels

Inner shells of heavier elements

Atomic energy levels (outer shells)

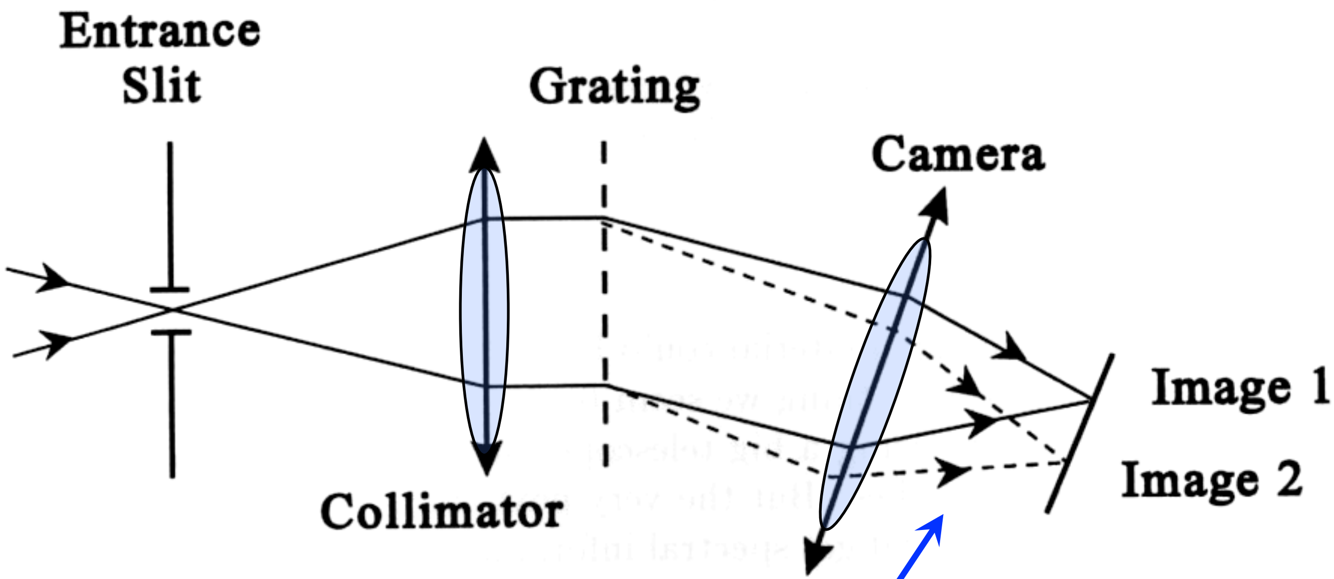
Molecular transitions

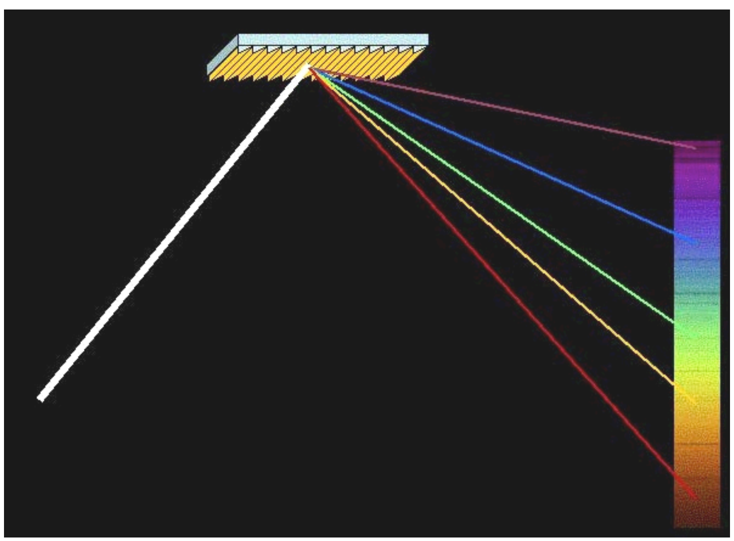
Hyperfine transitions

Plasma in typical magnetic fields

Diffraction Grating Spectrographs

A schematic view of a spectrograph:

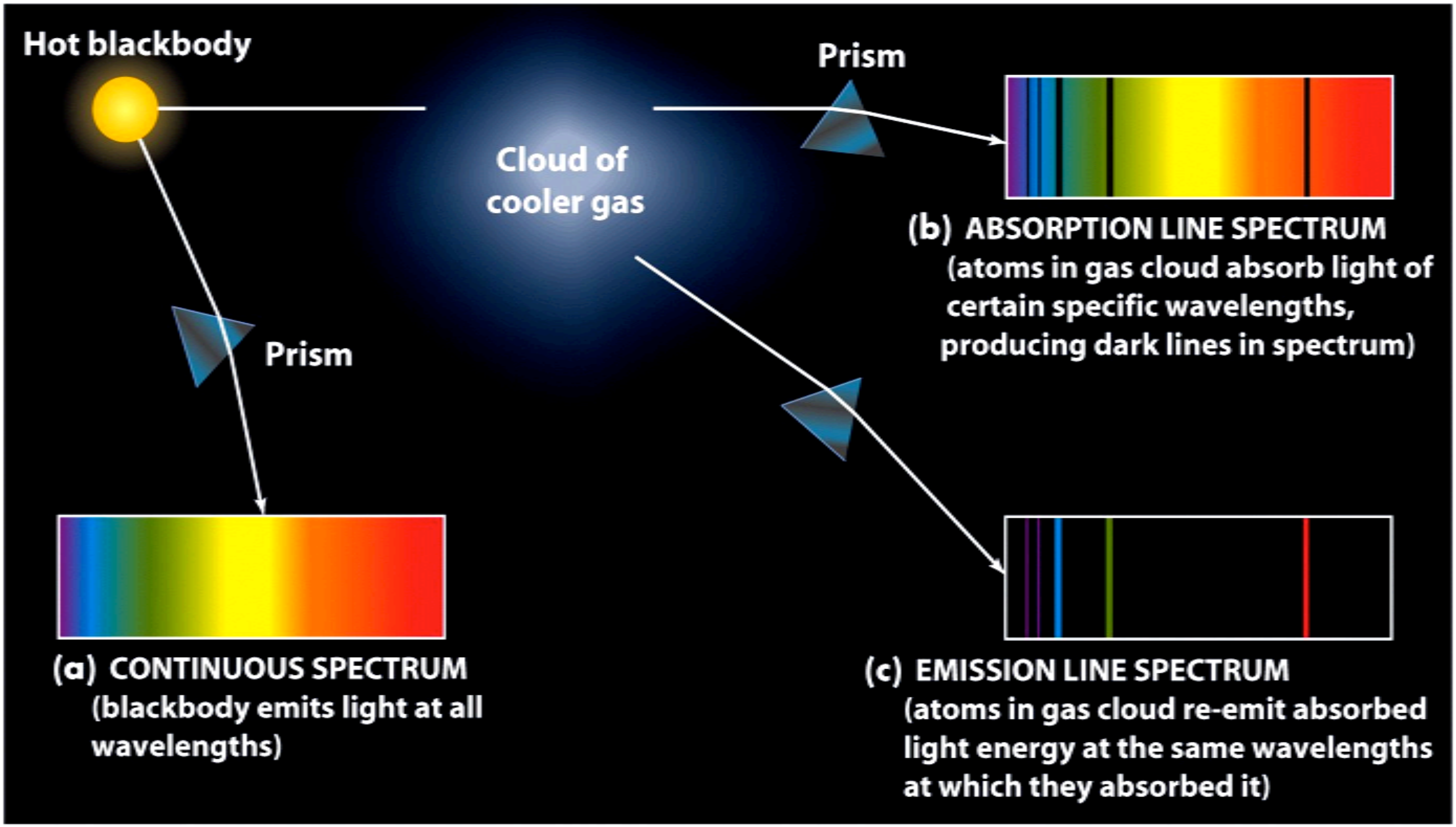




Light of different wavelengths is in phase at different reflection angles from the grating

Detector captures images of the entrance aperture (slit) at different wavelengths

Kirchoff's Laws

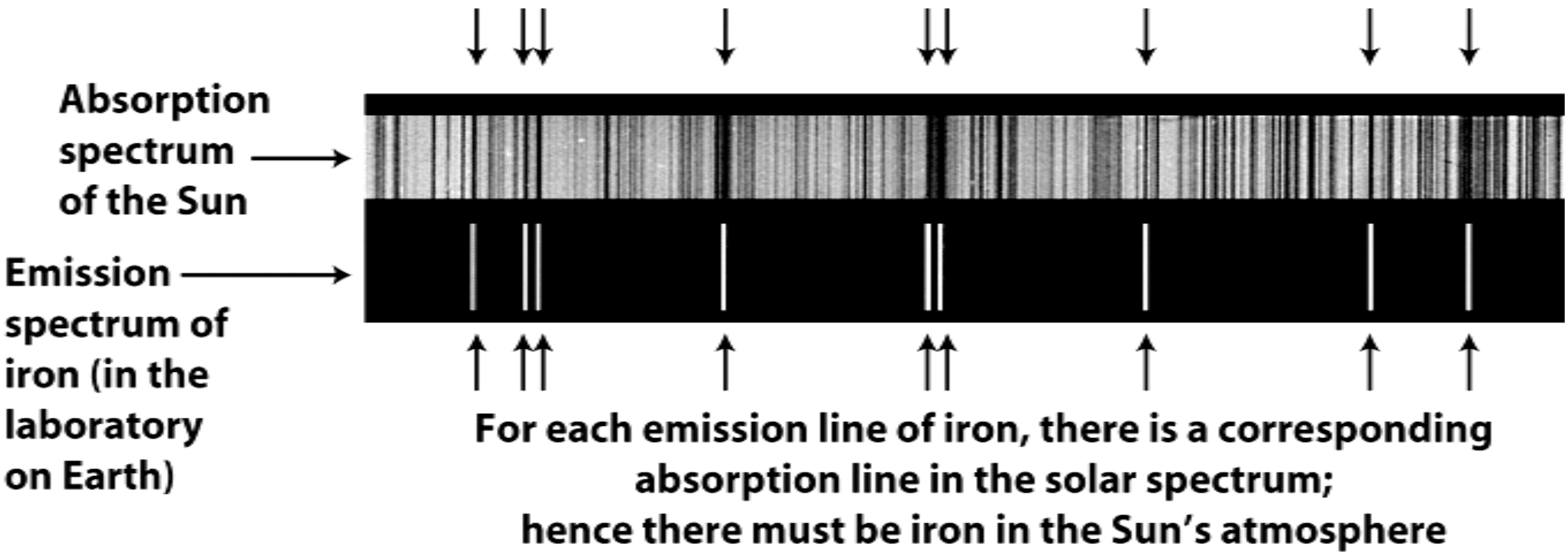


Kirchoff's Laws

1. **Continuous spectrum**: Any hot opaque body (e.g., hot gas/plasma) produces a continuous spectrum or complete rainbow
2. **Emission line spectrum**: A hot transparent gas will produce an emission line spectrum
3. **Absorption line spectrum**: A (relatively) cool transparent gas in front of a source of a continuous spectrum will produce an absorption line spectrum

Modern atomic/quantum physics provides a ready explanation for these empirical rules

Astronomical Spectroscopy

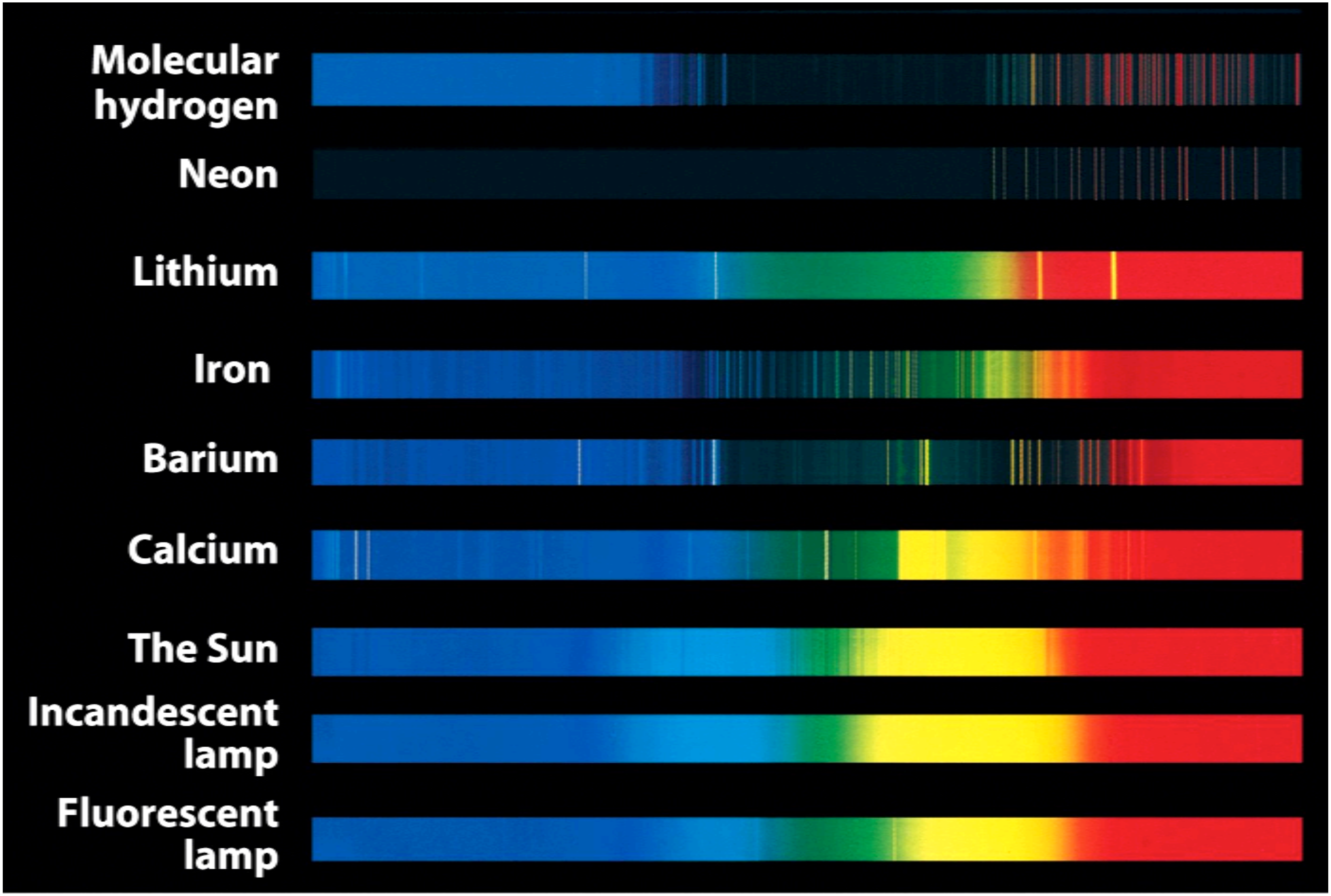


Laboratory spectra ➙ Line identifications in astro.sources

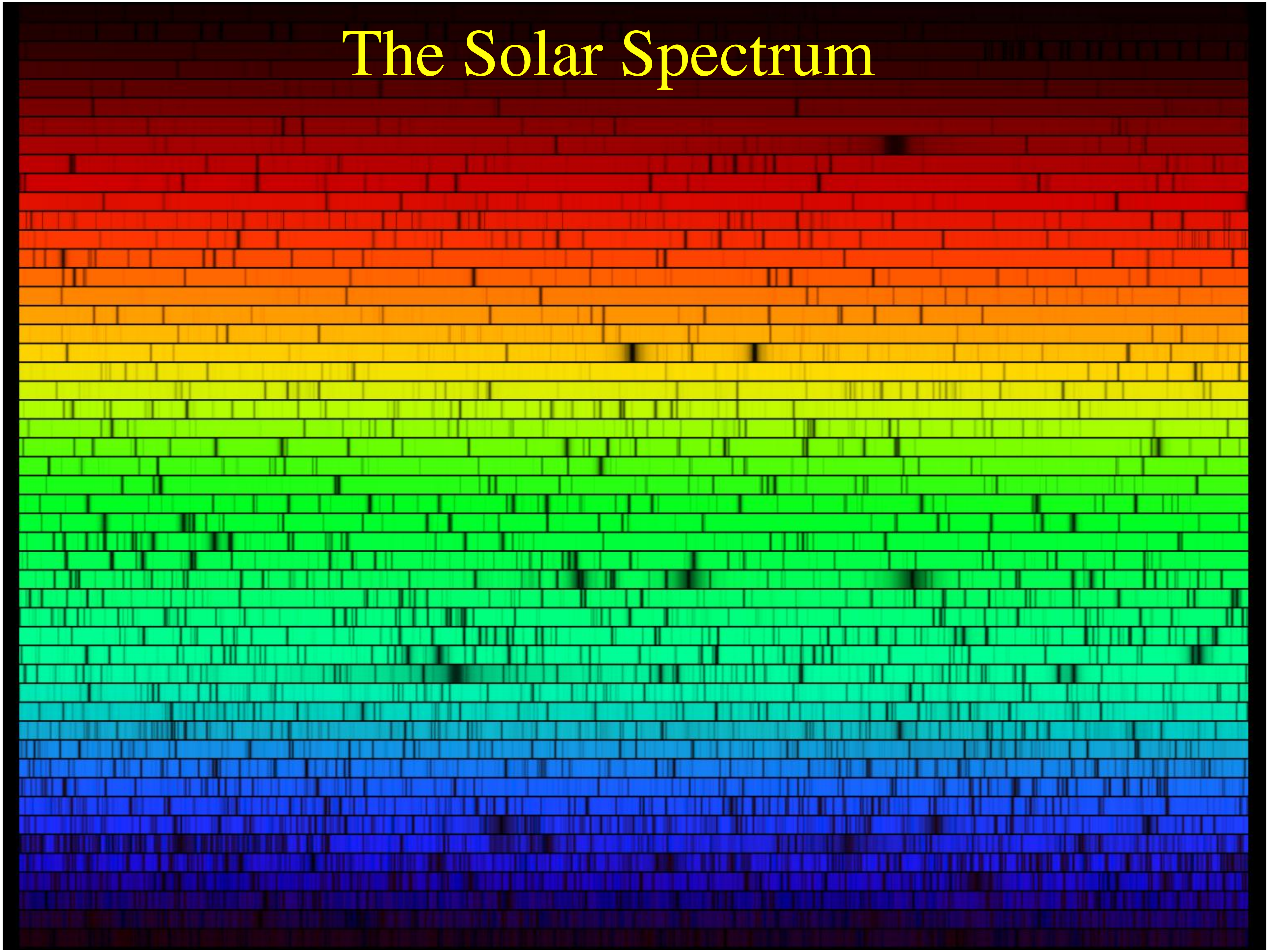
Analysis of spectra ➙ Chemical abundances + physical conditions (temperature, pressure, gravity, ionizing flux, magnetic fields, etc.)

+ Velocities

Examples of Spectra



The Solar Spectrum



Opaque or Transparent?

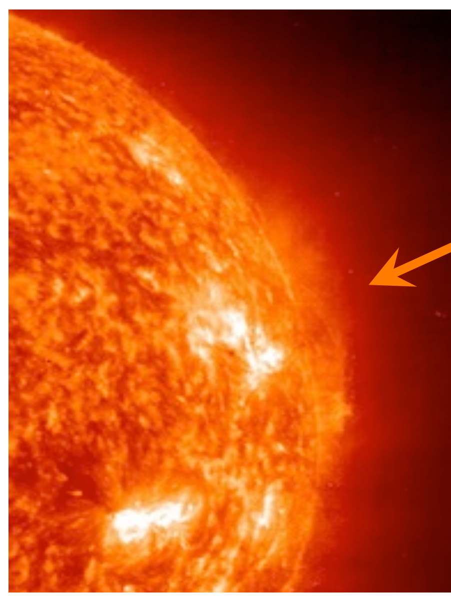
It depends on whether the gas (plasma) is

Optically thick: short mean free path of photons, get absorbed and re-emitted many times, only the radiation near the surface escapes; or

Optically thin: most photons escape without being reabsorbed or scattered

(Optical thickness is generally proportional to density)

Hot plasma inside a star (optically thick) generates a thermal continuum



Cooler, optically thin gas near the surface imprints an absorption spectrum

4.2 The Origin of Spectroscopic Lines

Atomic Radiative Processes

Radiation can be emitted or absorbed when electrons make transitions between different states:

Bound-bound: electron moves between two bound states (orbitals) in an atom or ion. Photon is emitted or absorbed.

Bound-free:

- Bound → unbound: **ionization**
- Unbound → bound: **recombination**

Free-free: free electron gains energy by absorbing a photon as it passes near an ion, or loses energy by emitting a photon. Also called **bremsstrahlung**.

Which transitions happen depends on the temperature and density of the gas → spectroscopy as a physical diagnostic

Energy Levels in a Hydrogen Atom

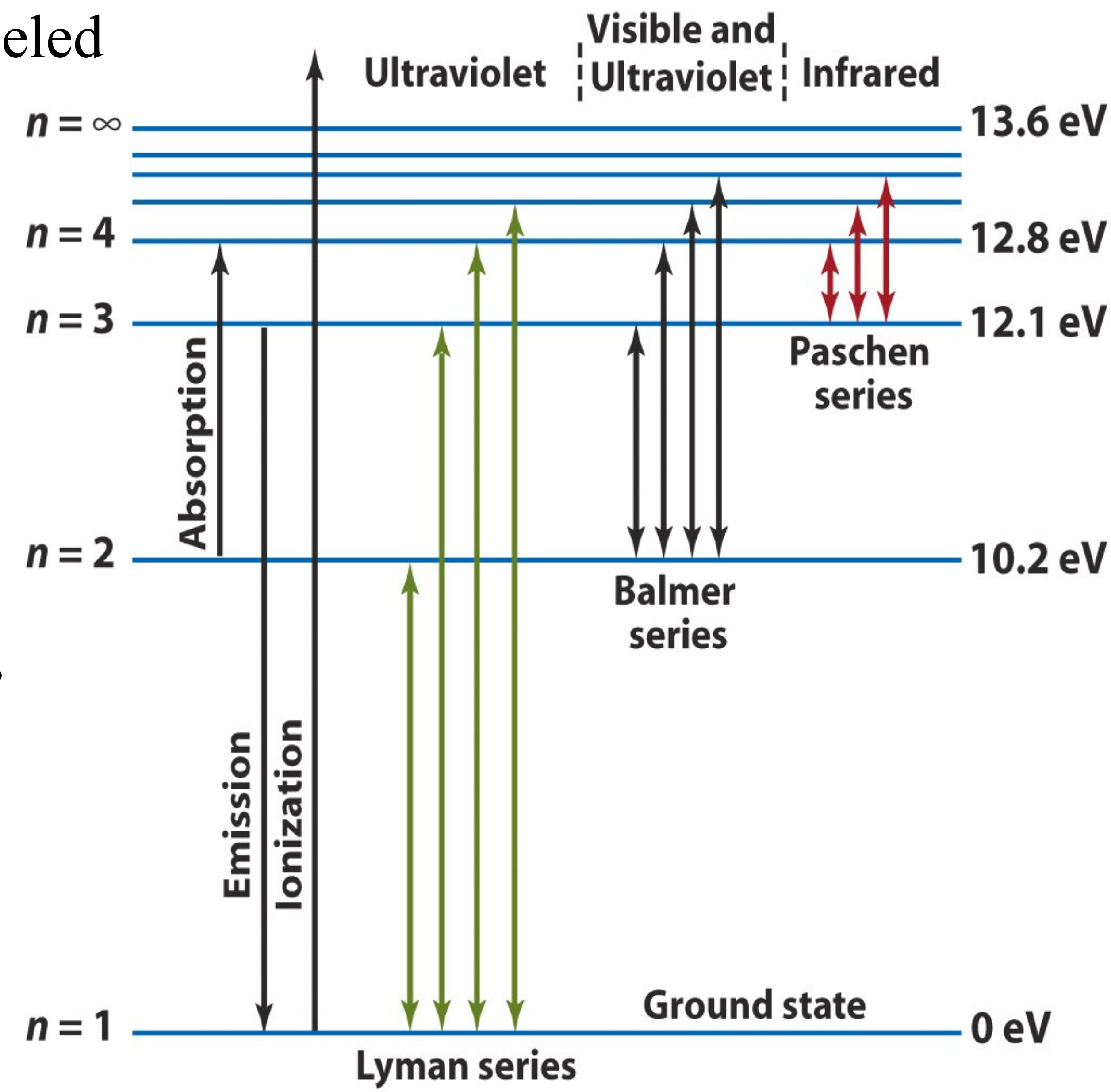
Energy levels are labeled by ***n*** - the *principal quantum number*

Energy of a given level is:

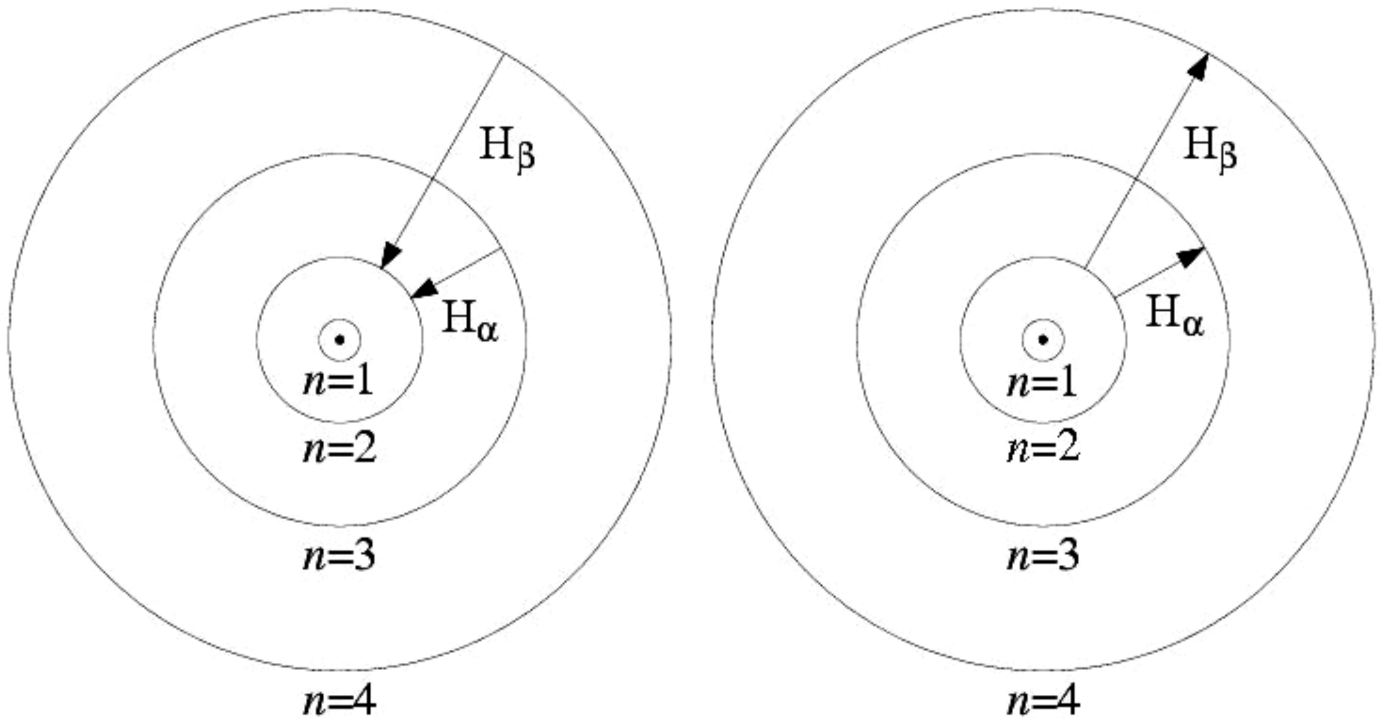
$$E_n = -\frac{R}{n^2}$$

where R = 13.6 eV is a Rydberg's constant

Lowest level, n=1, is the *ground state*



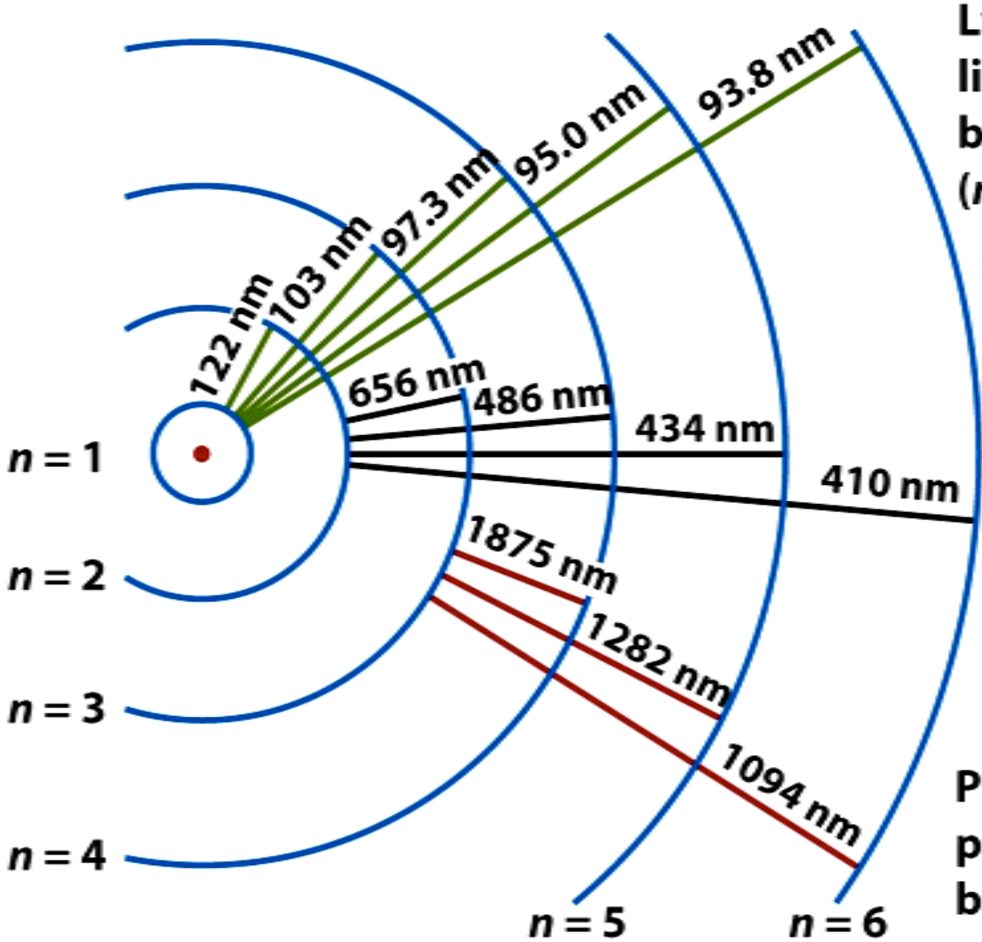
Energy Transitions: The Bohr Atom



Atoms transition from lower to higher energy levels (**excitation / de-excitation**) in discrete quantum jumps. The energy exchange can be **radiative** (involving a photon) or **collisional** (2 atoms)

Families of Energy Level Transitions Correspond to Spectroscopic Line Series

Photon energy: $h\nu = \left| E_i - E_j \right|$

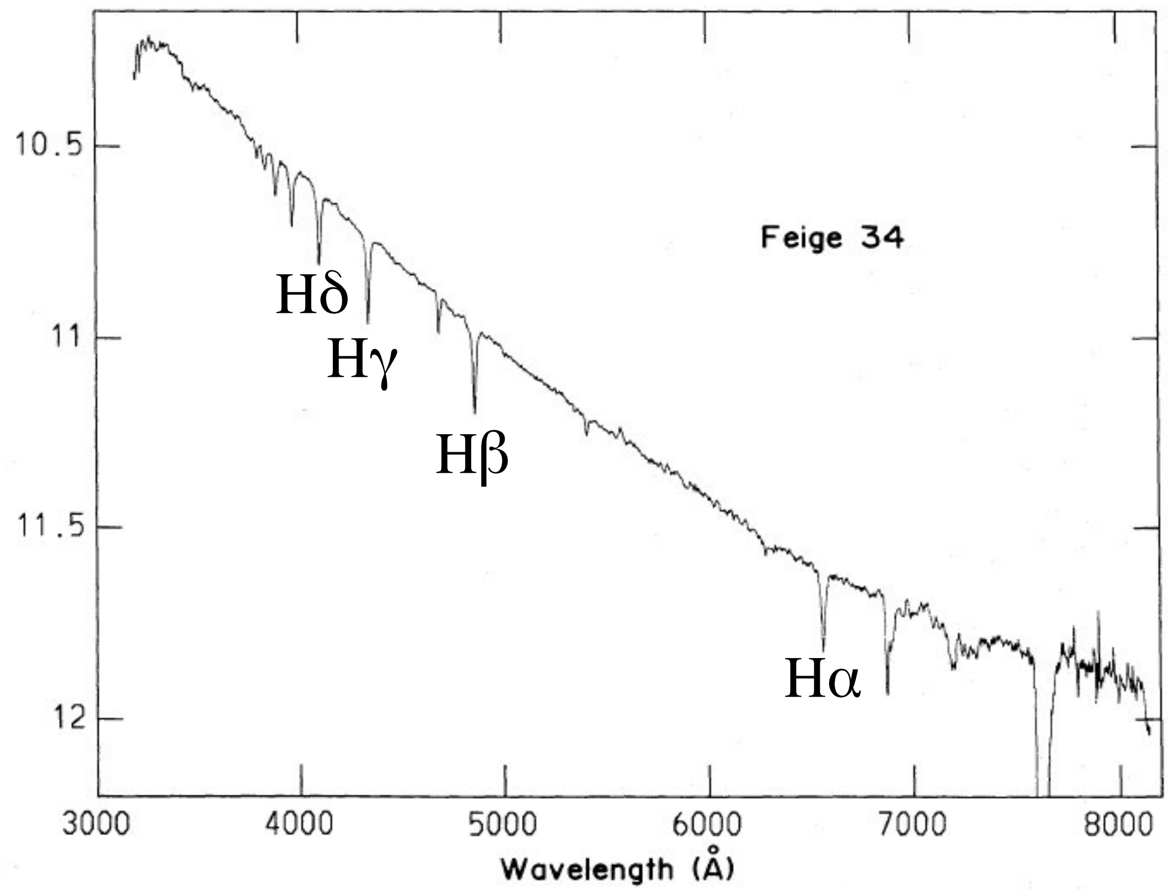


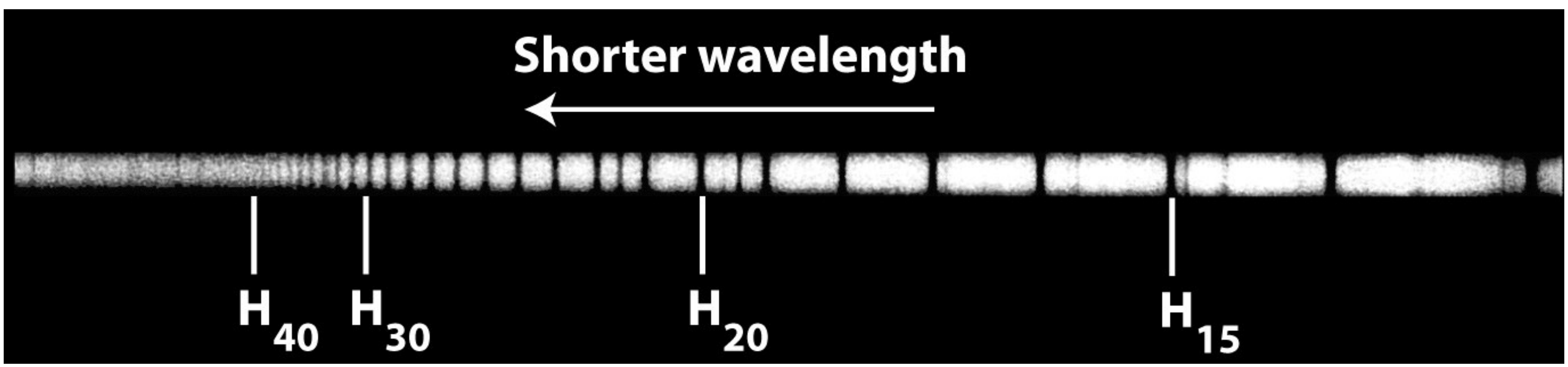
Lyman series (ultraviolet) of spectral lines: produced by electron transitions between the *n* = 1 orbit and higher orbits (*n* = 2, 3, 4, ...)

Balmer series (visible and ultraviolet) of spectral lines: produced by electron transitions between the *n* = 2 orbit and higher orbits (*n* = 3, 4, 5, ...)

Paschen series (infared) of spectral lines: produced by electron transitions between the *n* = 3 orbit and higher orbits (*n* = 4, 5, 6, ...)

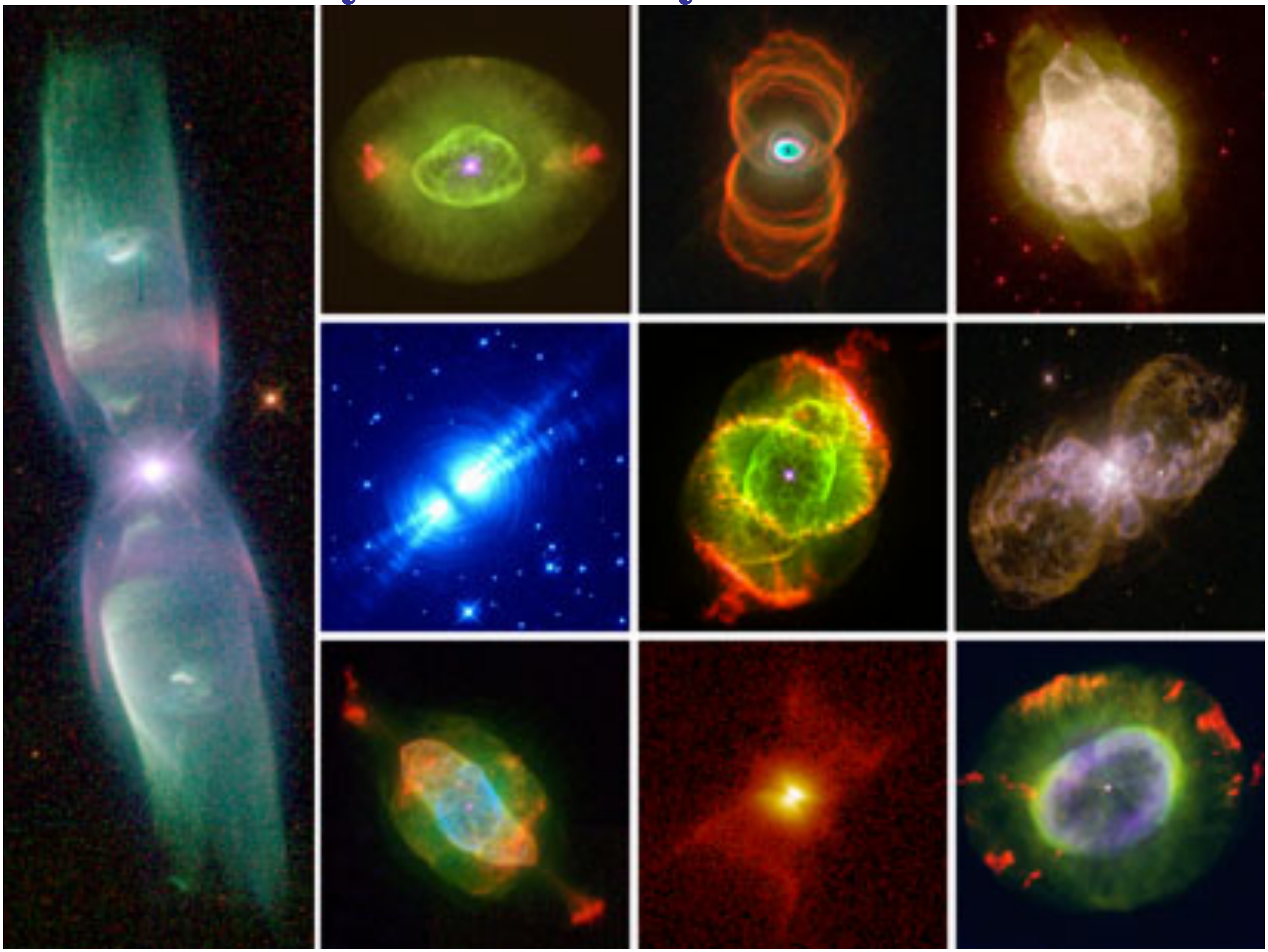
Balmer Series Lines in Stellar Spectra



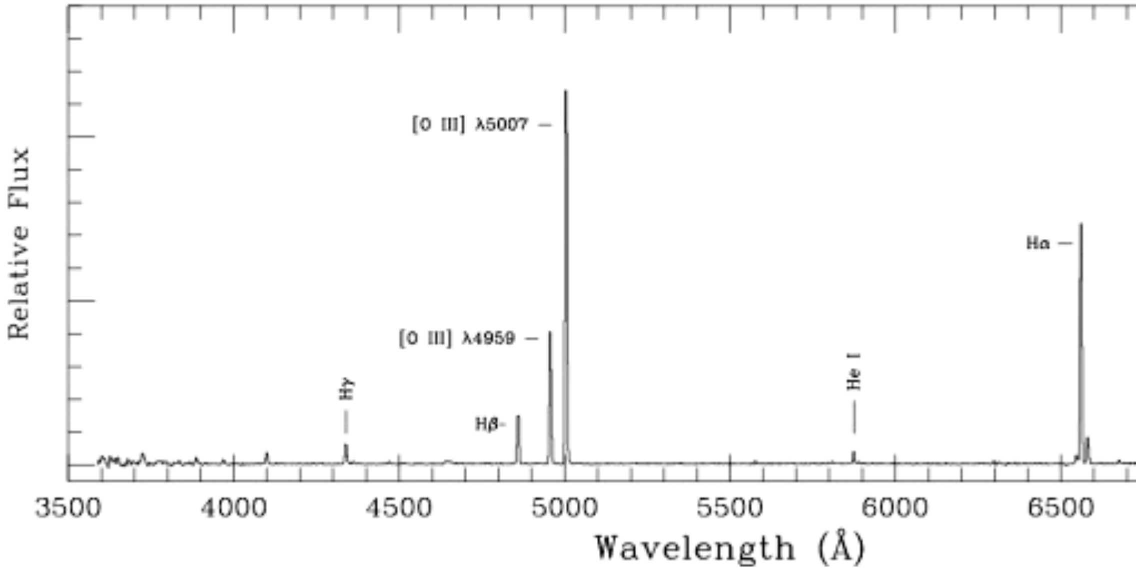


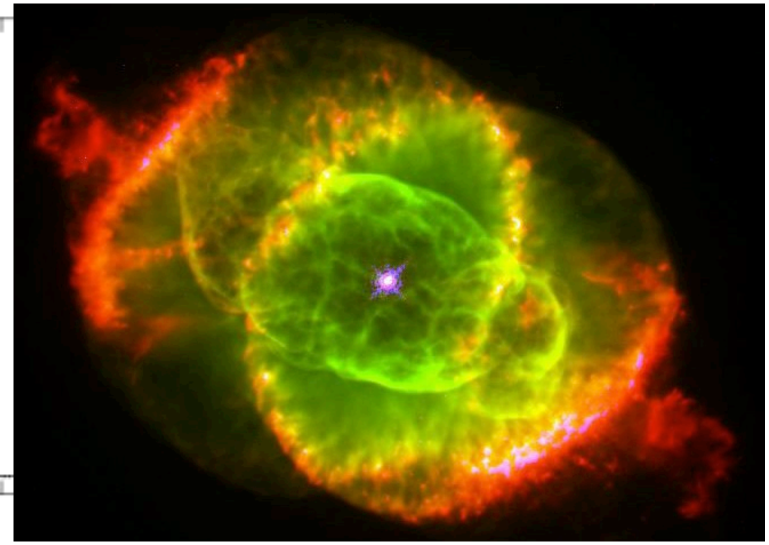
An Astrophysical Example: Photoionization of Hydrogen by Hot, Young Stars

An Astrophysical Example: Photoionization of Planetary Nebulae by Hot Central Stars



"Forbidden" Lines and Nebulium





Early spectra of astronomical nebulae have shown strong emission lines of an unknown origin. They were ascribed to a hypothetical new element, "nebulium".

It turns out that they are due to excited energy levels that are hard to reproduce in the lab, but are easily achieved in space, e.g., doubly ionized oxygen. Notation: [O III] 5007 ← Wavelength in Å

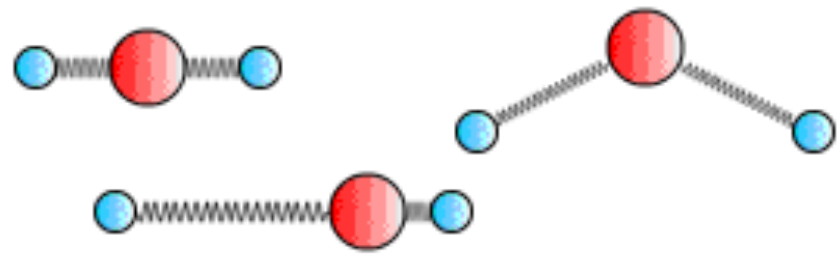
Brackets indicate "forbidden"

Element

Ionization state: III means lost 2 e's

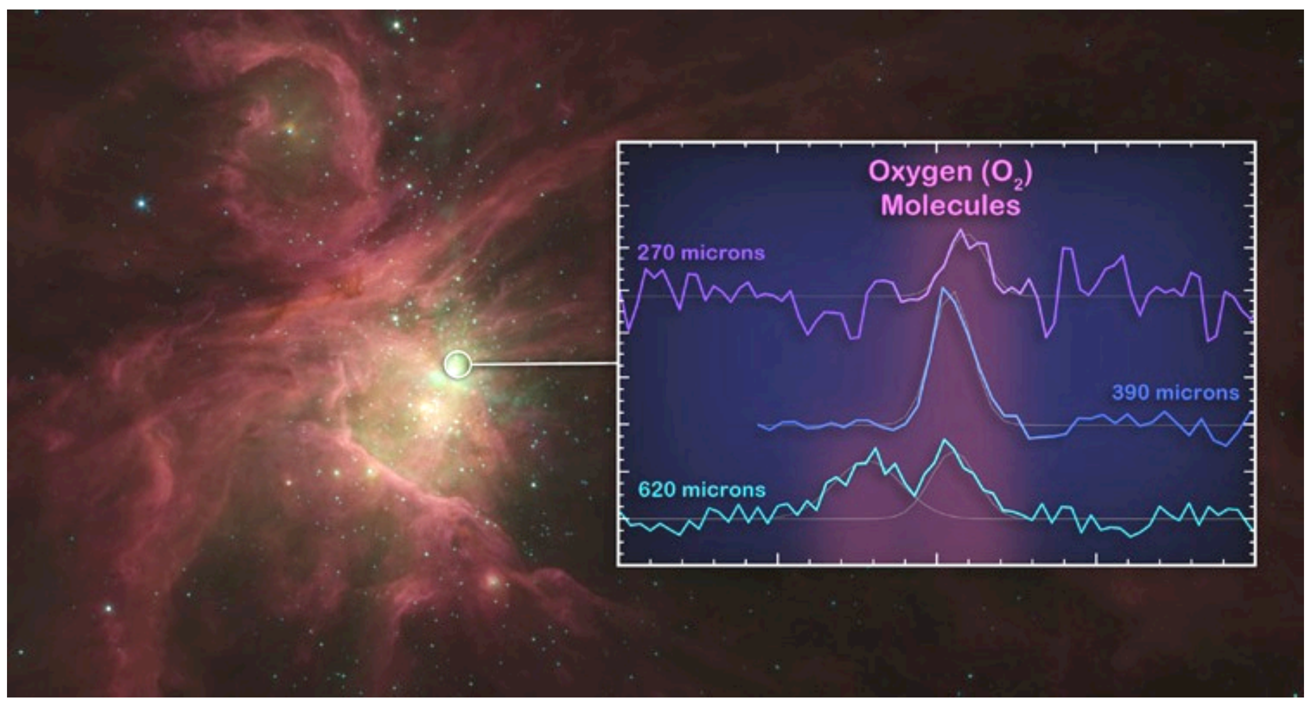
Spectra of Molecules

They have additional energy levels due to vibration or rotation



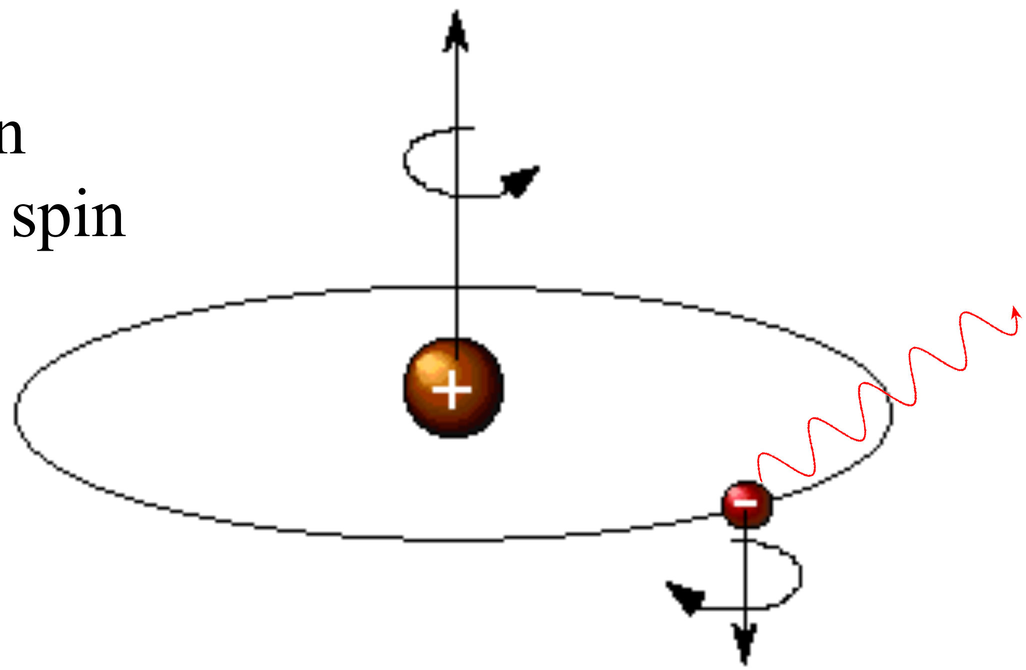
These tend to have a lower energy than the atomic level transitions, and are thus mostly on IR and radio wavelengths

They can thus probe cooler gas, e.g., interstellar or protostellar clouds



Hydrogen 21cm Line

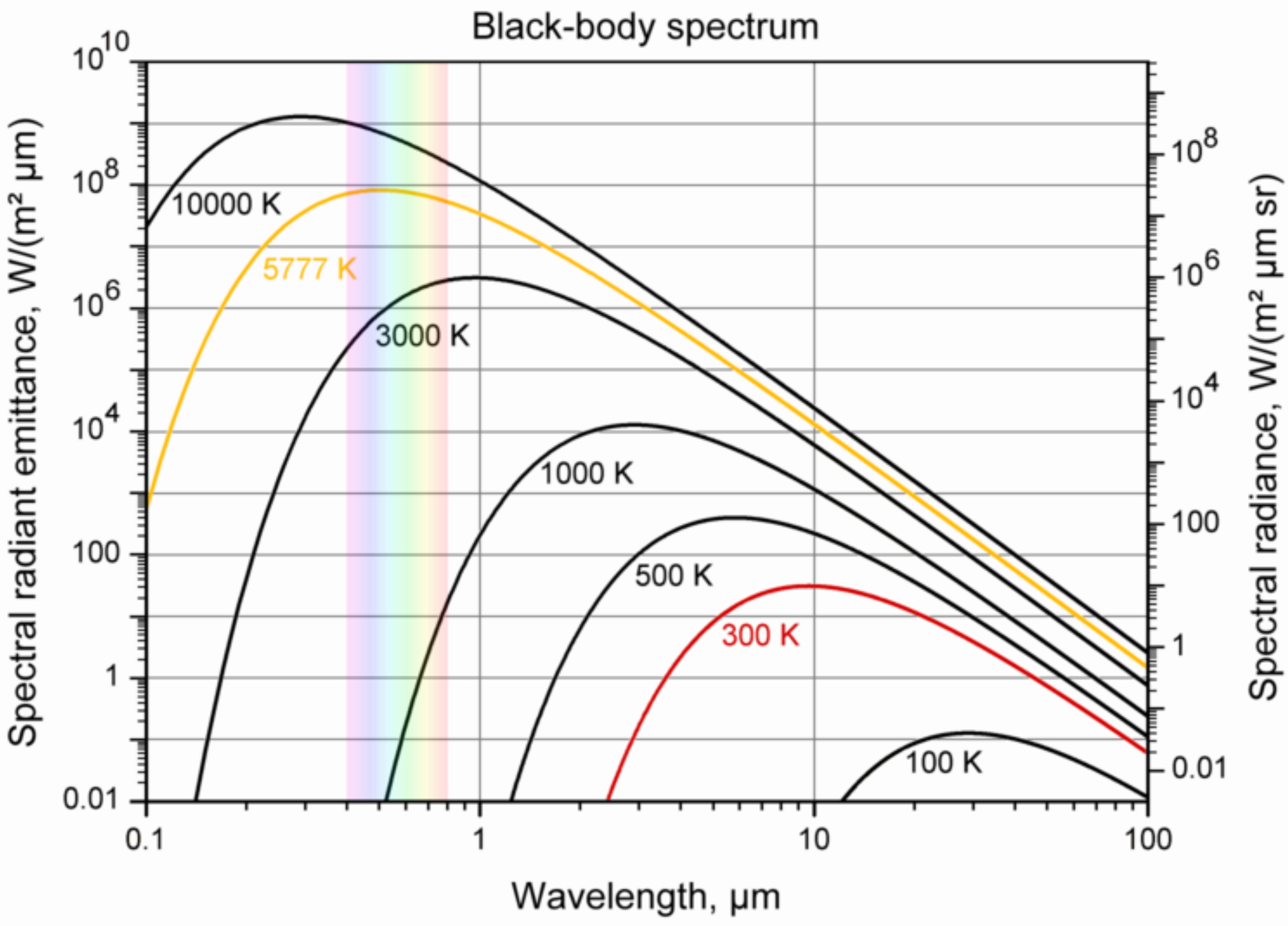
Corresponds to different orientations of the electron spin relative to the proton spin



Transition probability $= 3\times10^{-15}\ s^{-1}$ = once in 11 Myr per atom

Very important, because neutral hydrogen is so abundant in the universe. This is the principal wavelength for studies of interstellar matter in galaxies, and their disk structure and rotation curves

4.3 Blackbody Radiation and Other Continuum Emission Mechanisms

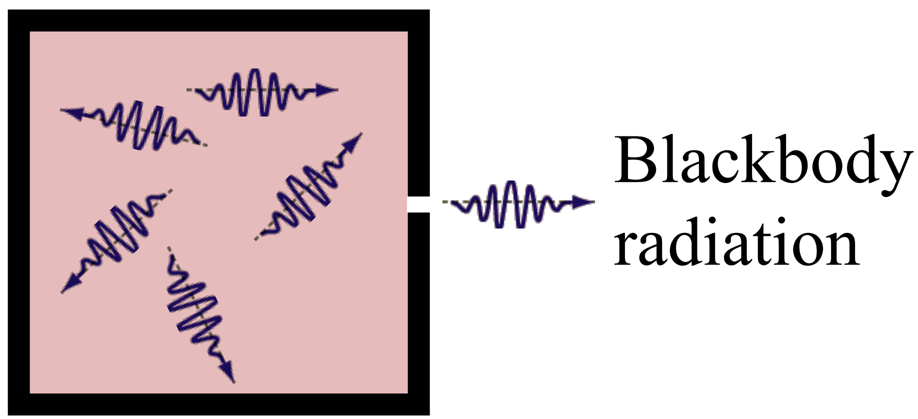


Blackbody Radiation

This is radiation that is in ***thermal equilibrium*** with matter at some temperature T.

Blackbody is a hypothetical object that is a perfect absorber of electromagnetic radiation at all wavelengths

Lab source of blackbody radiation: hot oven with a small hole which does not disturb thermal equilibrium inside:



Important because:

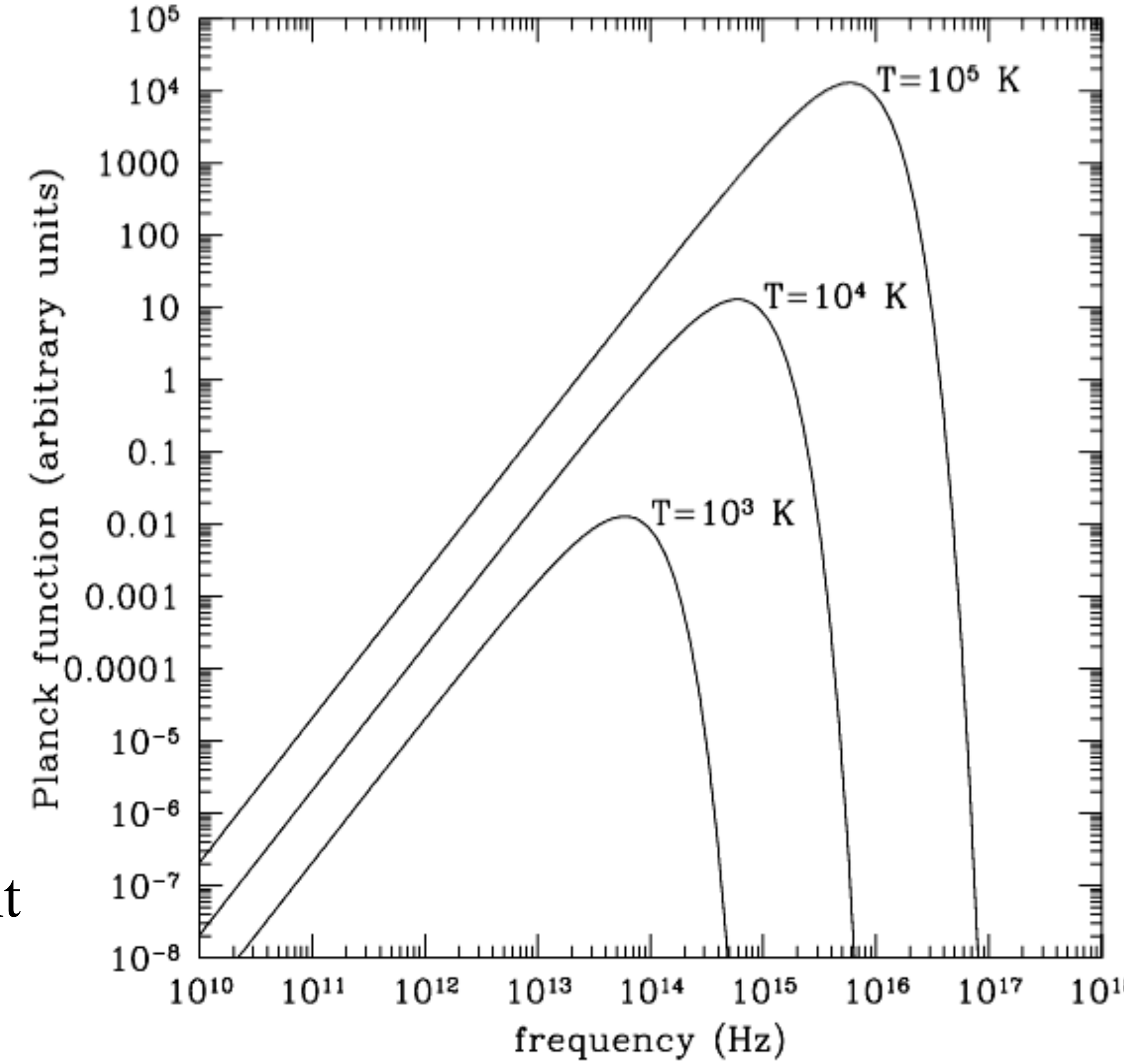
- Interiors of stars (for example) are like this
- Emission from many objects is roughly of this form.

Blackbody Spectrum

The frequency dependence is given by the **Planck function**:

$$B_\nu(T) = \frac{2h\nu^3/c^2}{\exp(h\nu/kT) - 1}$$

h = Planck's constant
k = Boltzmann's constant



Same units as specific intensity: erg s^{-1} cm^{-2} $sterad^{-1}$ Hz^{-1}

Blackbody Spectrum

The Planck function peaks when $dB_n(T)/d\nu = 0$:

$$h\nu_{\max} = 2.82kT$$

$$\nu_{\max} = 5.88 \times 10^{10} T \text{ Hz K}^{-1}$$

This is ***Wien displacement law*** - peak shifts linearly with Increasing temperature to higher frequency.

Asymptotically, for low frequencies $h\,\nu \ll kT$, the ***Rayleigh-Jeans law*** applies:

$$B_\nu^{RJ}(T) = \frac{2\nu^2}{c^2} kT$$

Often valid in the radio part of the spectrum, at freq's far below the peak of the Planck function.

Blackbody Luminosity

The **energy density** of blackbody radiation:

$$u(T) = aT^4$$

$a = 7.56 \times 10^{-15}$ erg cm^{-3} K^{-4} is the radiation constant.

The **emergent flux** from a surface emitting blackbody radiation is:

$$F = \sigma T^4$$

$\sigma = 5.67 \times 10^{-5}$ erg cm^{-2} K^{-4} s^{-1} = Stefan-Boltzmann const.

A sphere (e.g., a star), with a radius R, temperature T, emitting as a blackbody, has a **luminosity**:

$$L = 4\pi R^2 \sigma T^4$$

Effective Temperature

Emission from most astronomical sources is only roughly described by the Planck function (if at all).

For a source with a bolometric flux ***F***, define the **effective temperature *T*$_e$** via:

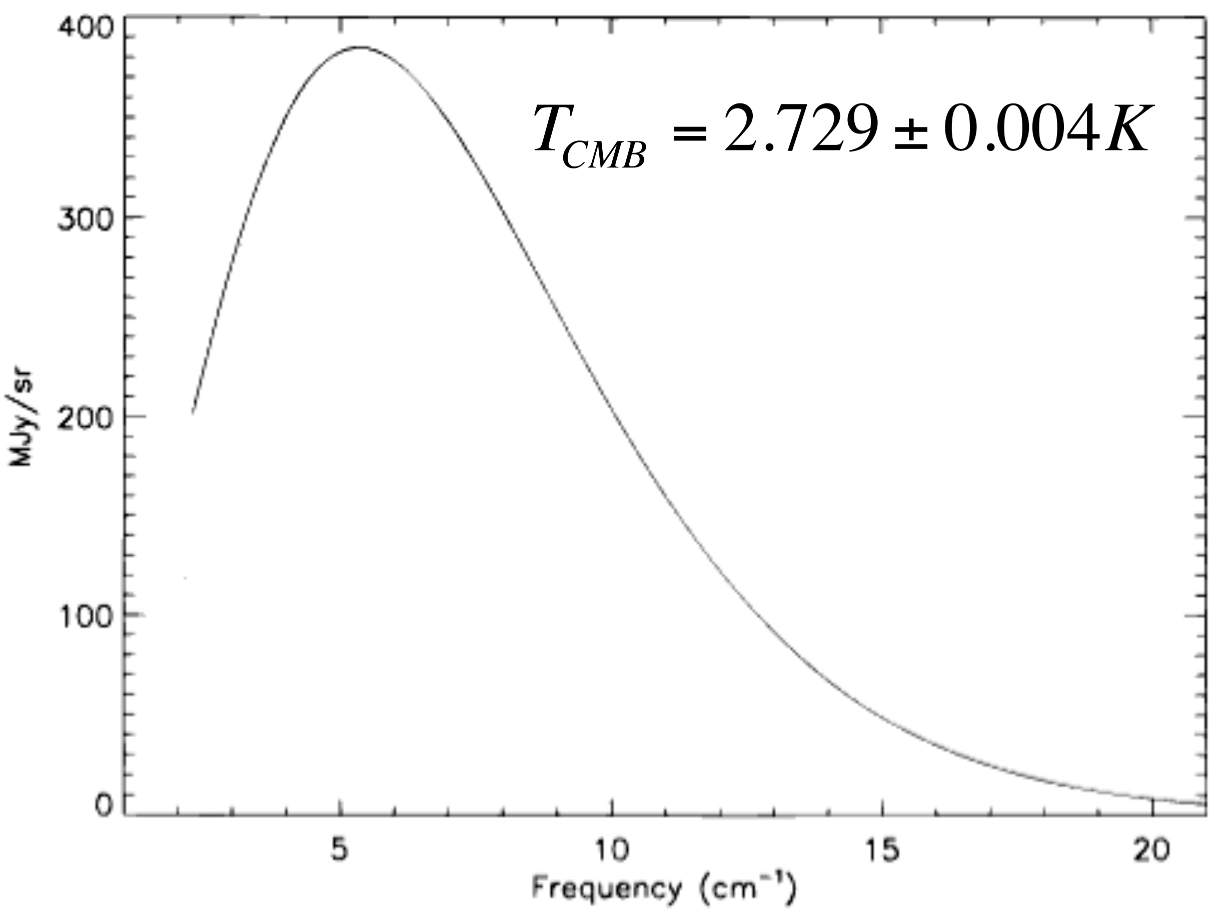
$$F \equiv \sigma T_e^4$$

e.g., for the Sun: $L_{sun} = 4\pi R_{sun}^2 \sigma T_e^4$ …find T_e = 5770 K.

Note: effective temperature is well-defined even if the spectrum is nothing like a blackbody.

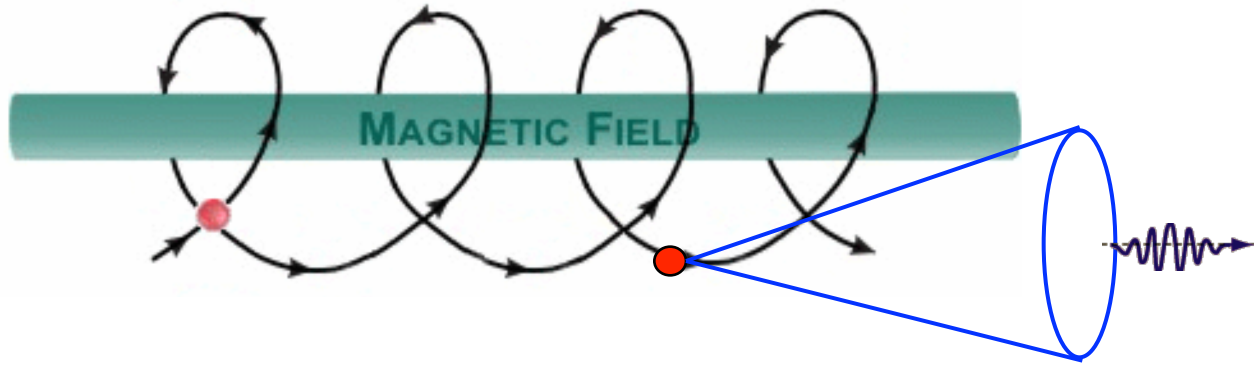
Big bang model - Universe was hot, dense, and in thermal equilibrium between matter and radiation in the past.

Relic radiation from this era is the **cosmic microwave background** radiation. Best known blackbody:



No known distortions of the CMB from a perfect blackbody!

Synchrotron Emission

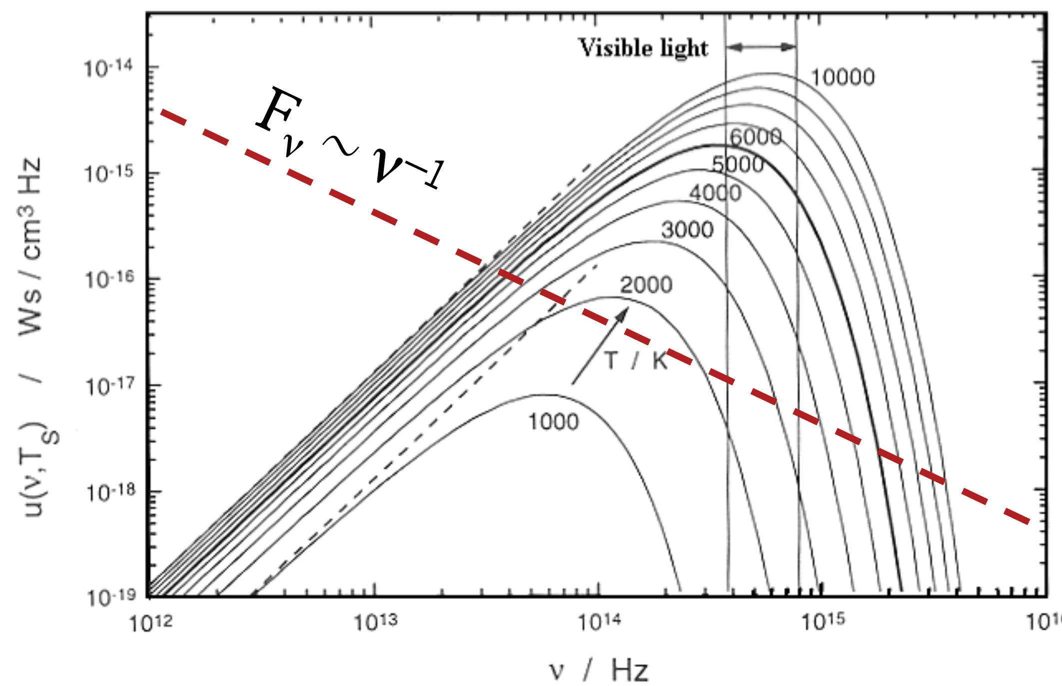


- An electron moving at an angle to the magnetic field feels Lorentz force; therefore it is accelerated, and it radiates in a cone-shaped beam

- The spectrum is for the most part a power law:

$$F_\nu \sim \nu^\alpha \,, \ \ \alpha \sim -1$$

(*very* different from a blackbody!)



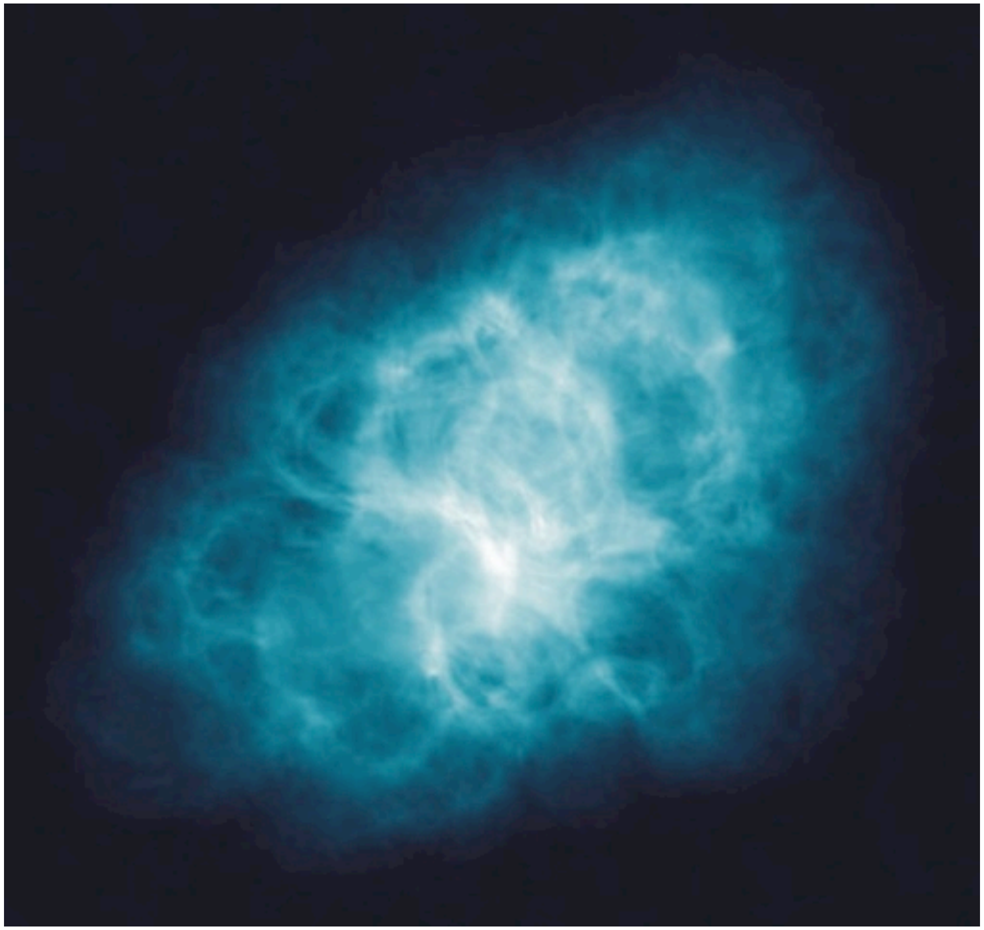
Examples of Synchrotron Radiation:

Radio galaxy Cygnus A at 5 GHz

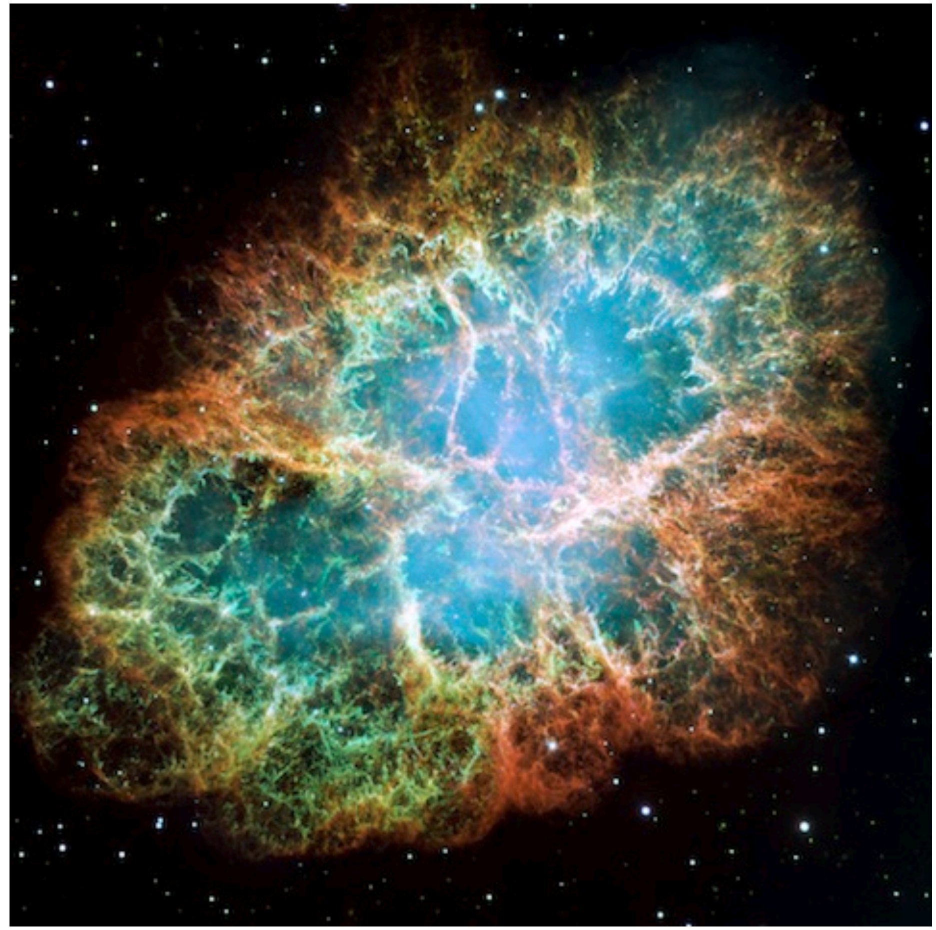
Jet of M87 in the visible light

Examples of Synchrotron Radiation:

Crab nebula in radio

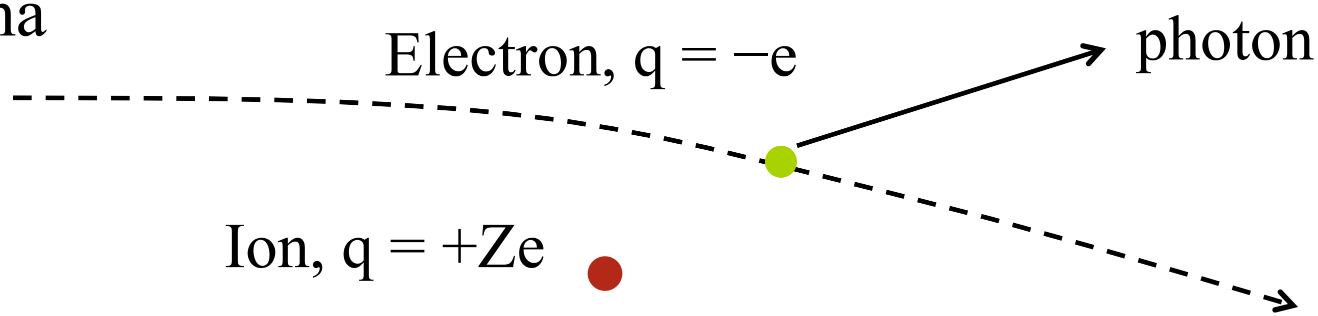


Crab nebula in visible light

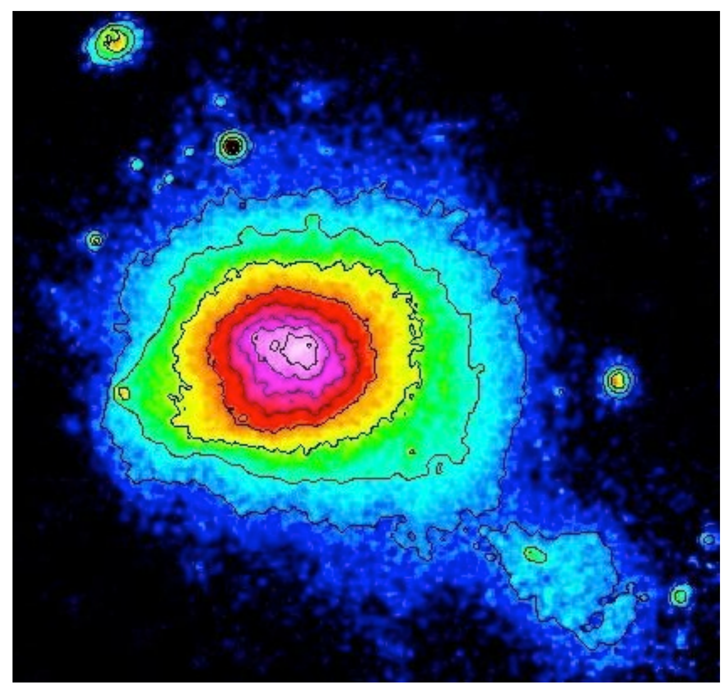


Thermal Bremsstrahlung

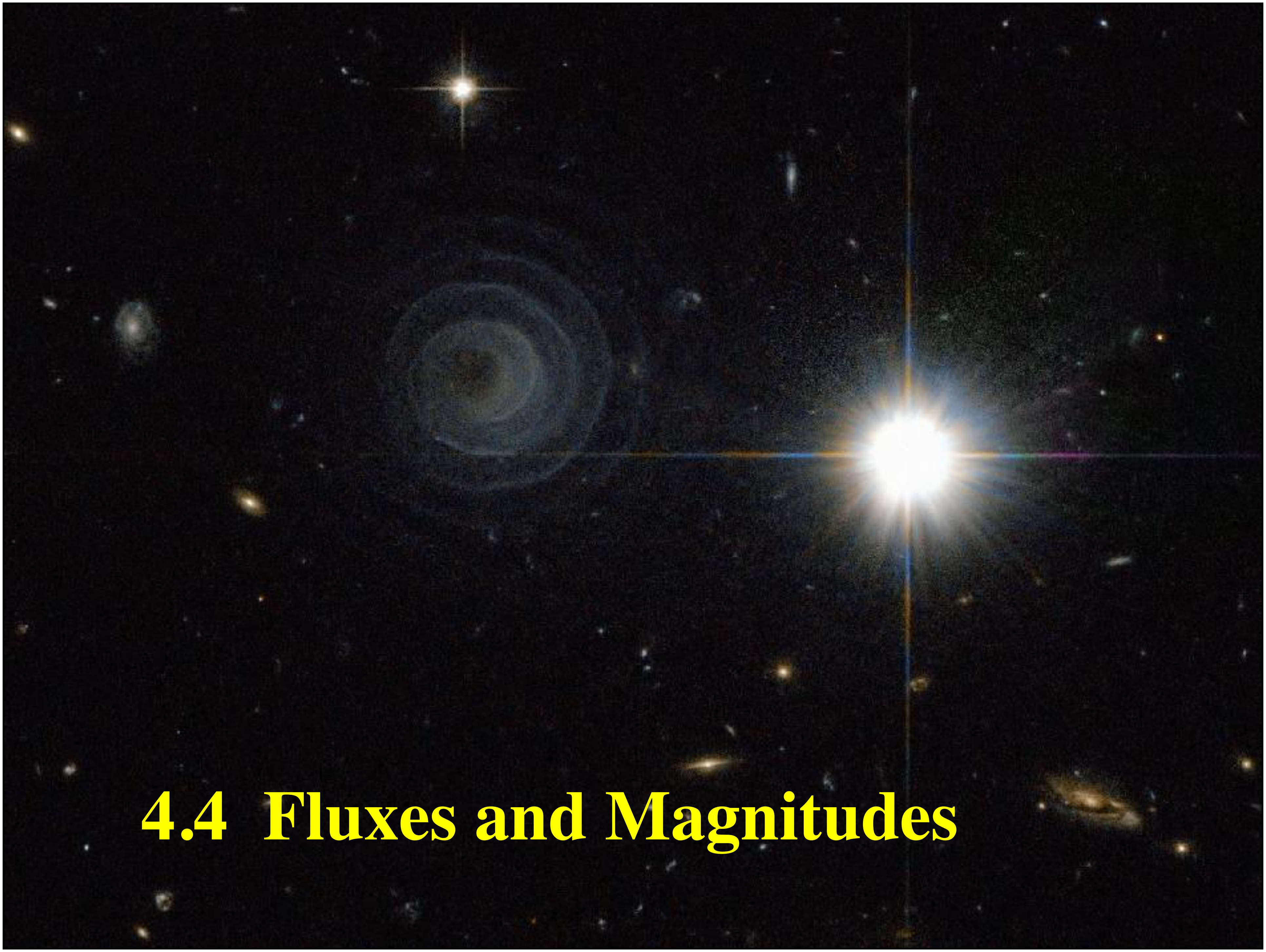
A free-free emission from electrons scattering by ions in a very hot plasma



Example: X-ray gas in clusters of galaxies



4.4 Fluxes and Magnitudes



Measuring Flux = Energy/(unit time)/(unit area)

Real detectors are sensitive over a finite range of λ (or ν). Fluxes are always measured over some finite bandpass.

Total energy flux:

$$F = \int F_\nu(\nu)\,d\nu$$

Integral of f_ν over all frequencies

Units: erg s^{-1} cm^{-2} Hz^{-1}

A standard unit for specific flux (initially in radio, but now more common):

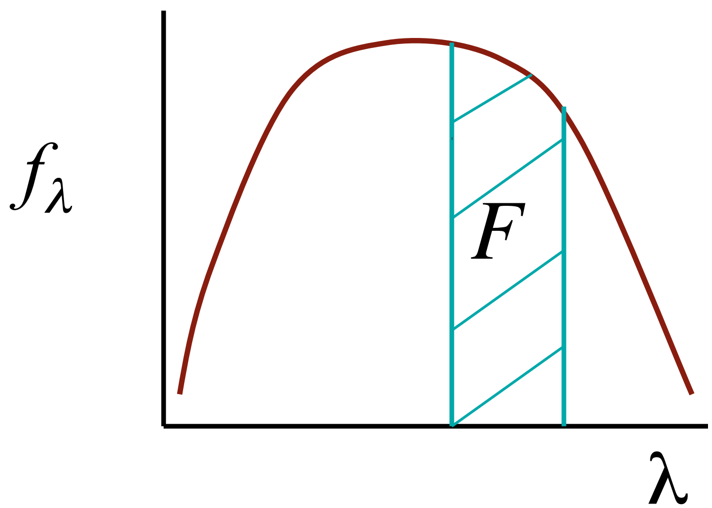
$$1 \text{ Jansky (Jy)} = 10^{-23} \text{ erg s}^{-1} \text{ cm}^{-2} \text{ Hz}^{-1}$$

f_ν is often called the ***flux density*** - to get the ***power,*** one integrates it over the bandwith, and multiplies by the area

Fluxes and Magnitudes

For historical reasons, fluxes in the optical and IR are measured in **magnitudes**:

$$m = -2.5 \log_{10} F + \text{constant}$$



If F is the total flux, then m is the bolometric magnitude. Usually instead consider a finite bandpass, e.g., V band ($\lambda_c \sim$ 550 nm, $\Delta\lambda \sim$ 50 nm)

Magnitude zero-points (constant in the eq. above) differ for different standard bandpasses, but are usually set so that Vega has $m = 0$ in every bandpass.

Vega calibration ($m = 0$): at $\lambda = 5556$:

$f_\lambda = 3.39 \times 10^{-9}$ erg/cm^2/s/Å
$f_\nu = 3.50 \times 10^{-20}$ erg/cm^2/s/Hz
$N_\lambda = 948$ photons/cm^2/s/Å

Using Magnitudes

Consider two stars, one of which is a hundred times fainter than the other in some waveband (say V).

$$m_1 = -2.5\log F_1 + \text{constant}$$

$$\begin{aligned} m_2 &= -2.5\log(0.01F_1) + \text{constant} \\ &= -2.5\log(0.01) - 2.5\log F_1 + \text{constant} \\ &= 5 - 2.5\log F_1 + \text{constant} \\ &= 5 + m_1 \end{aligned}$$

Source that is 100 times **fainter** in flux is five magnitudes fainter (**larger** number).

Faintest objects detectable with *HST* have magnitudes of ~ 28 in R/I bands. The sun has m_V = -26.75 mag

Apparent vs. Absolute Magnitudes

The absolute magnitude is defined as the apparent mag. a source would have if it were at a distance of 10 pc:

$$M = m + 5 - 5 \log d/\text{pc}$$

It is a measure of the **luminosity** in some waveband.

For Sun: $M_{\odot B} = 5.47$, $M_{\odot V} = 4.82$, $M_{\odot bol} = 4.74$

Difference between the apparent magnitude m and the absolute magnitude M (any band) is a ***measure of the distance*** to the source

$$\underbrace{m - M}_{\text{Distance modulus}} = 5\log_{10}\left(\frac{d}{10\ \text{pc}}\right)$$

Distance modulus