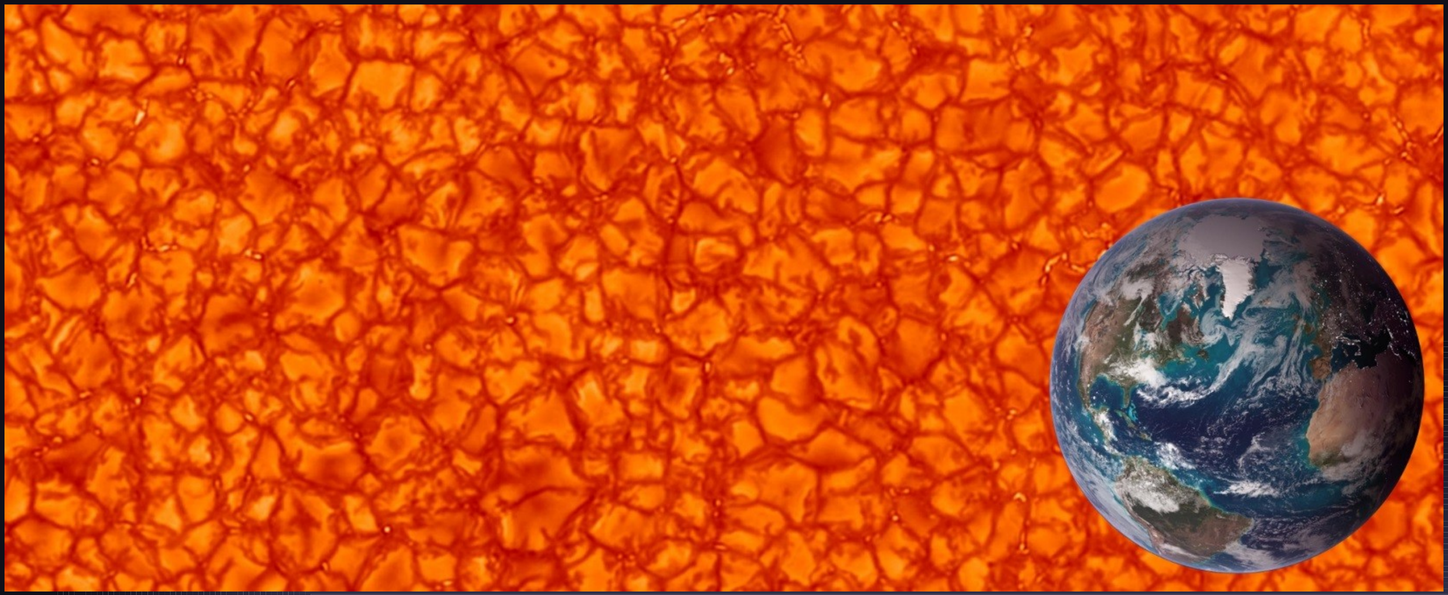


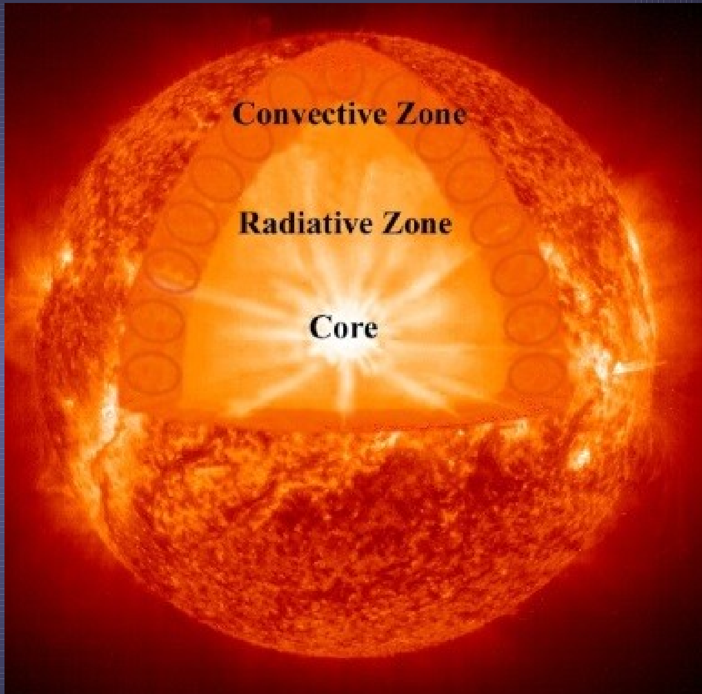
Solar convection

In addition to radiation, convection is the main form of energy transport in solar interior and lower atmosphere. Convection dominates just below the solar surface and produces most structures the lower solar atmosphere

The convection zone

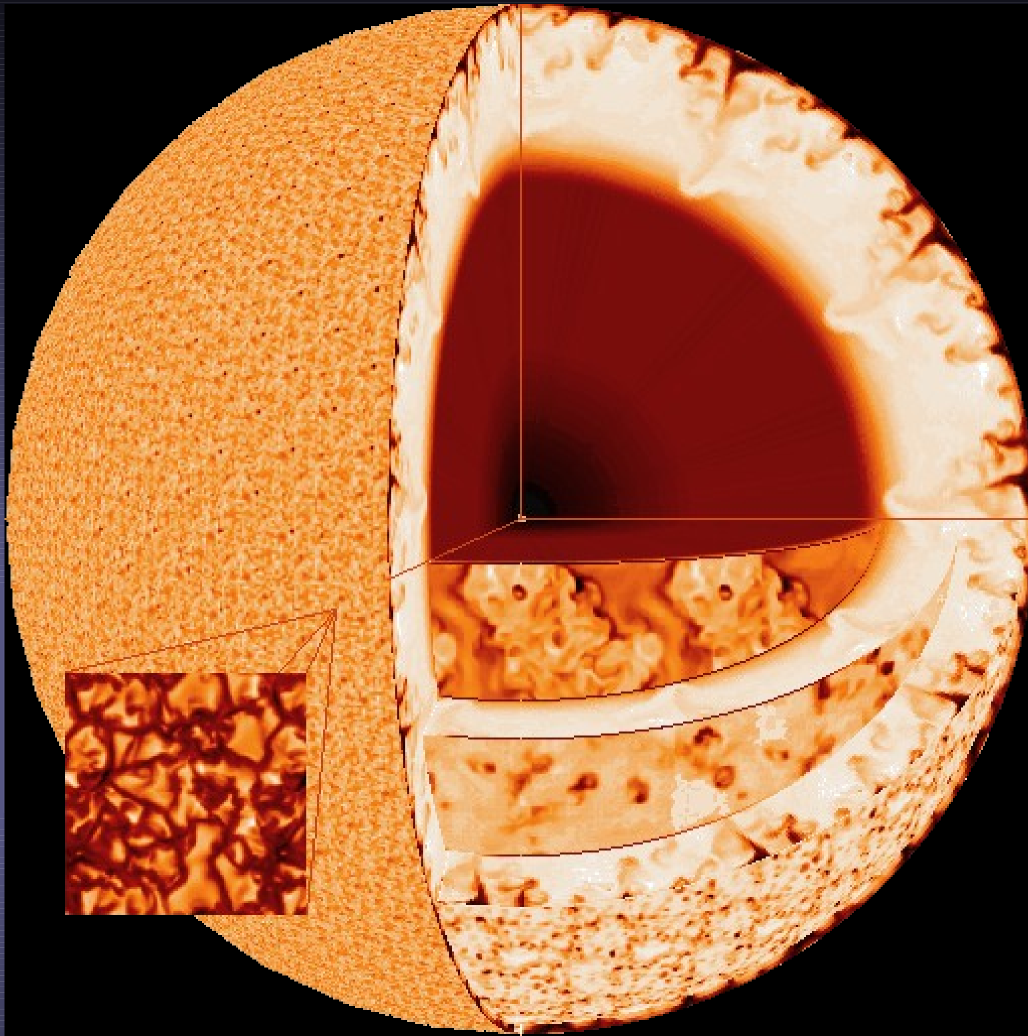
- Through the outermost 30% of solar interior, energy is transported by convection instead of by radiation
- In this layer the gas is convectively unstable.
- I.e. the process changes from a random walk of the photons through the radiative zone (due to high density, the mean free path in the core is well below a millimeter) to convective energy transport
- The unstable region ends just below the solar surface. I.e. the visible signs of convection are actually due to overshooting (see following slides)
- *t* radiative ~ 105 years

  \>> *t*convective ~ weeks



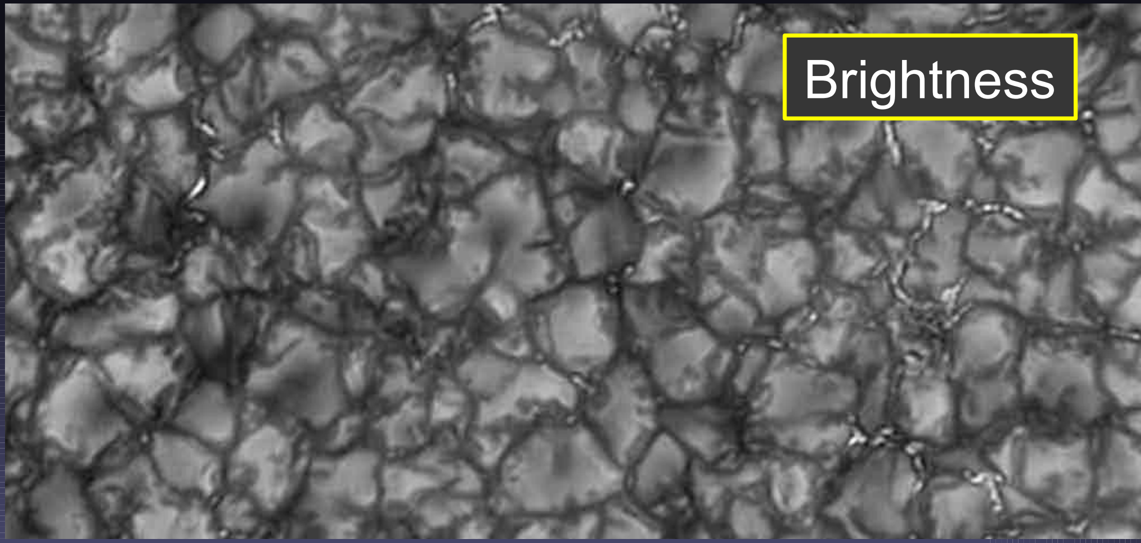
Scales of solar convection

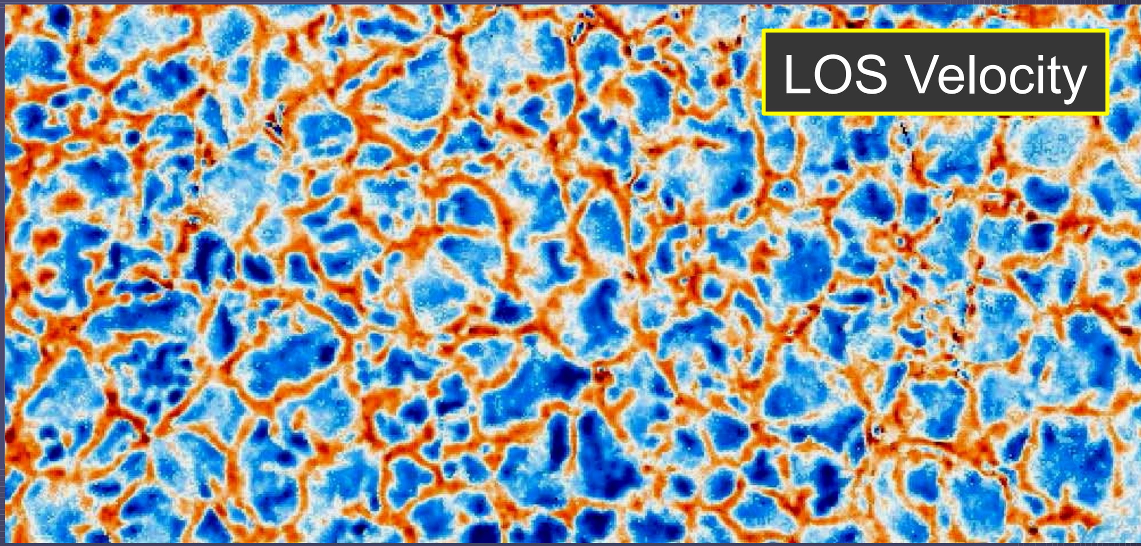
- Observations: 4 main scales
  - **granulation**
  - mesogranulation
  - **supergranulation**
  - giant cells
- **Colour:**
  - **well observed**
  - less strong evidence
- Theory: larger scales at greater depths. Many details are still unclear



Surface manifestation of convection: Granulation

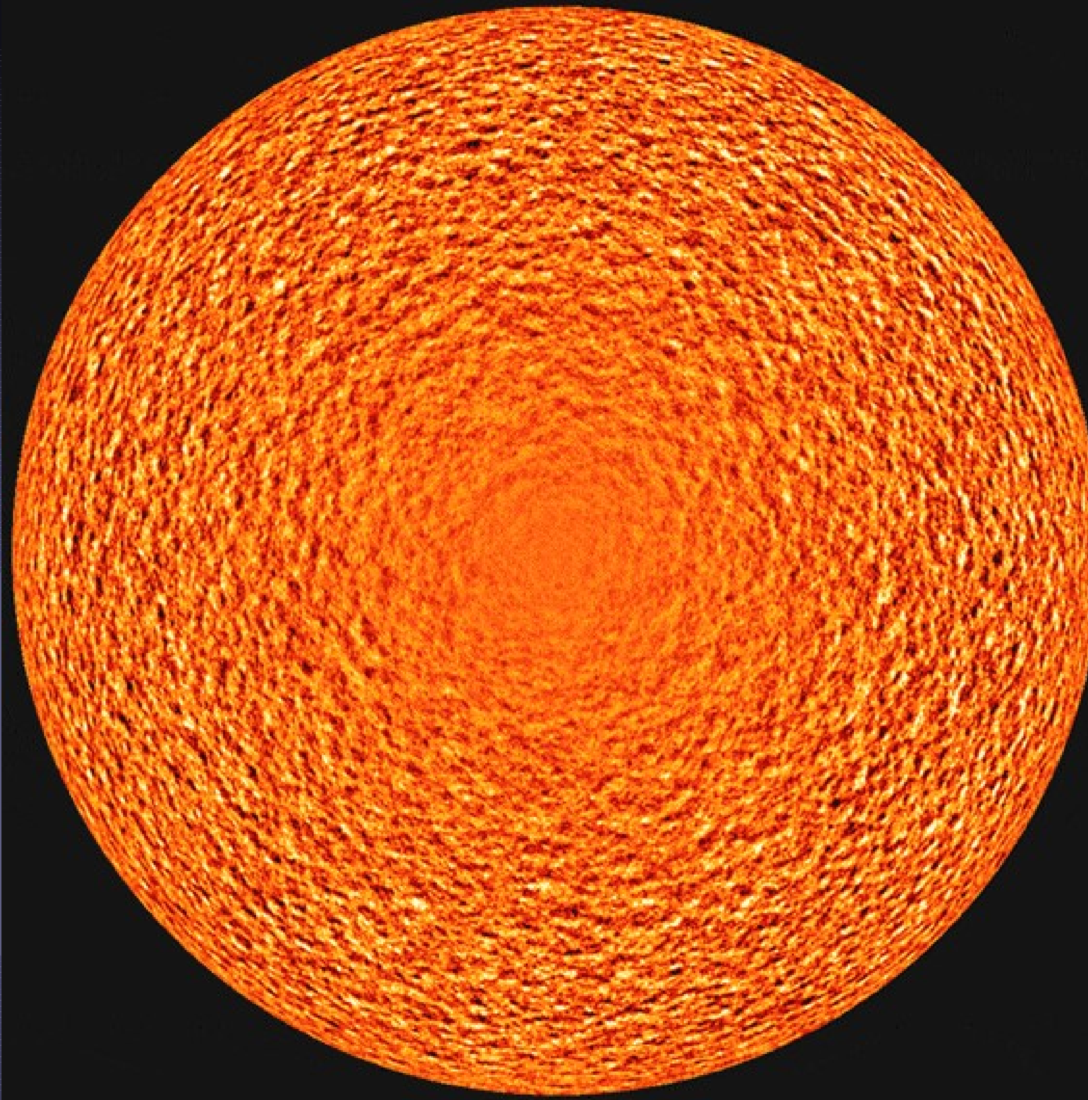
- Typical size: 1-2 Mm
- Lifetime: 5-8 min
- Velocities: 1 km/s (but peak velocities > 10 km/s, i.e. supersonic)
- Brightness contrast: ~15% in visible (green) continuum
- All quantities show a continuous distribution of values
- At any one time 106 granules on Sun





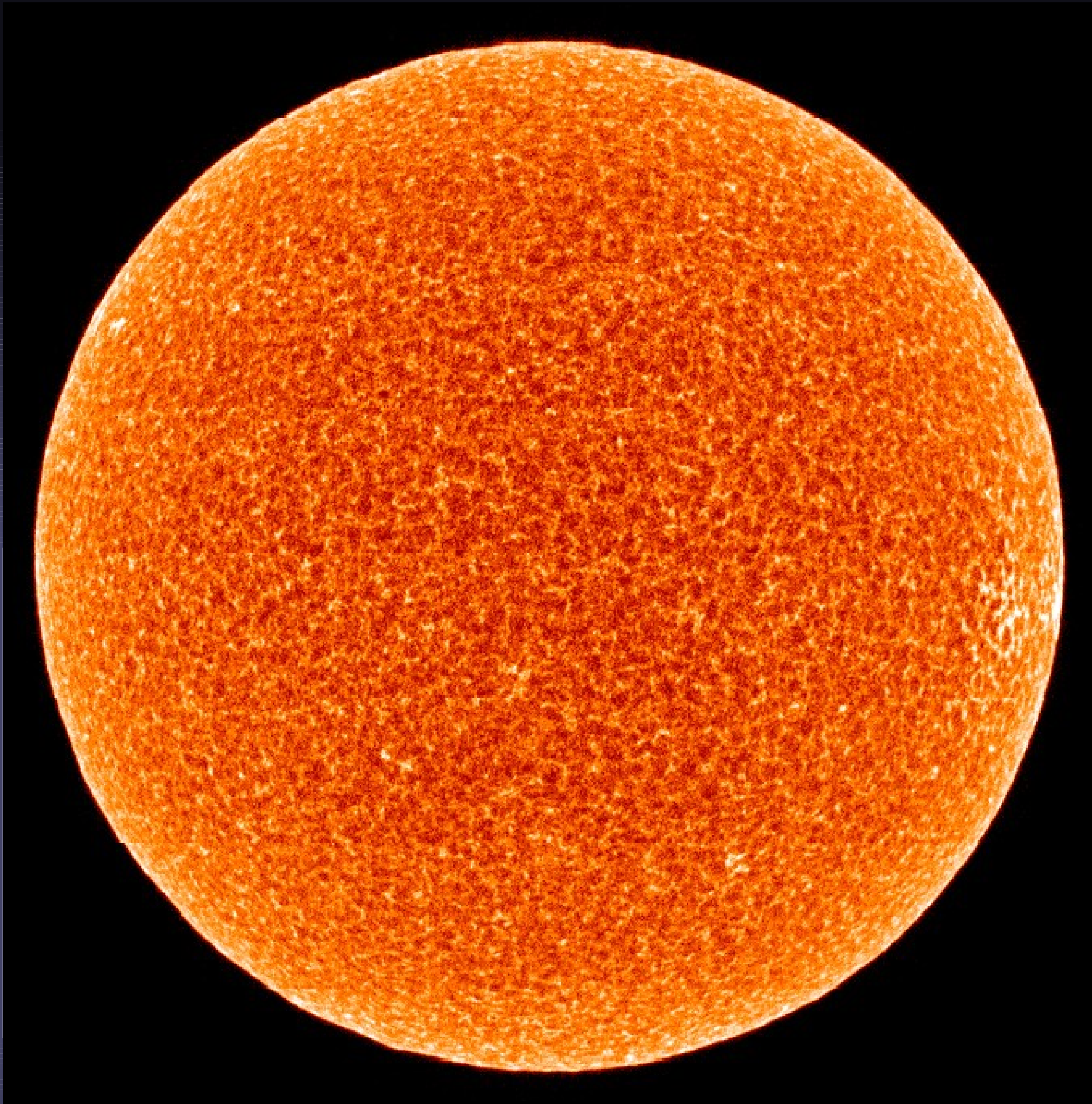
Surface manifesttion of convection: Supergranulation

- 1 h average of MDI Dopplergrams (averages out oscillations)
- Dark-bright: flows towards/away from observer
- No supergranules visible at disk centre velocity is mainly horizontal
- Size: 20-30 Mm, lifetime: days, horiz. speed: 400 m/s, no contrast in visible



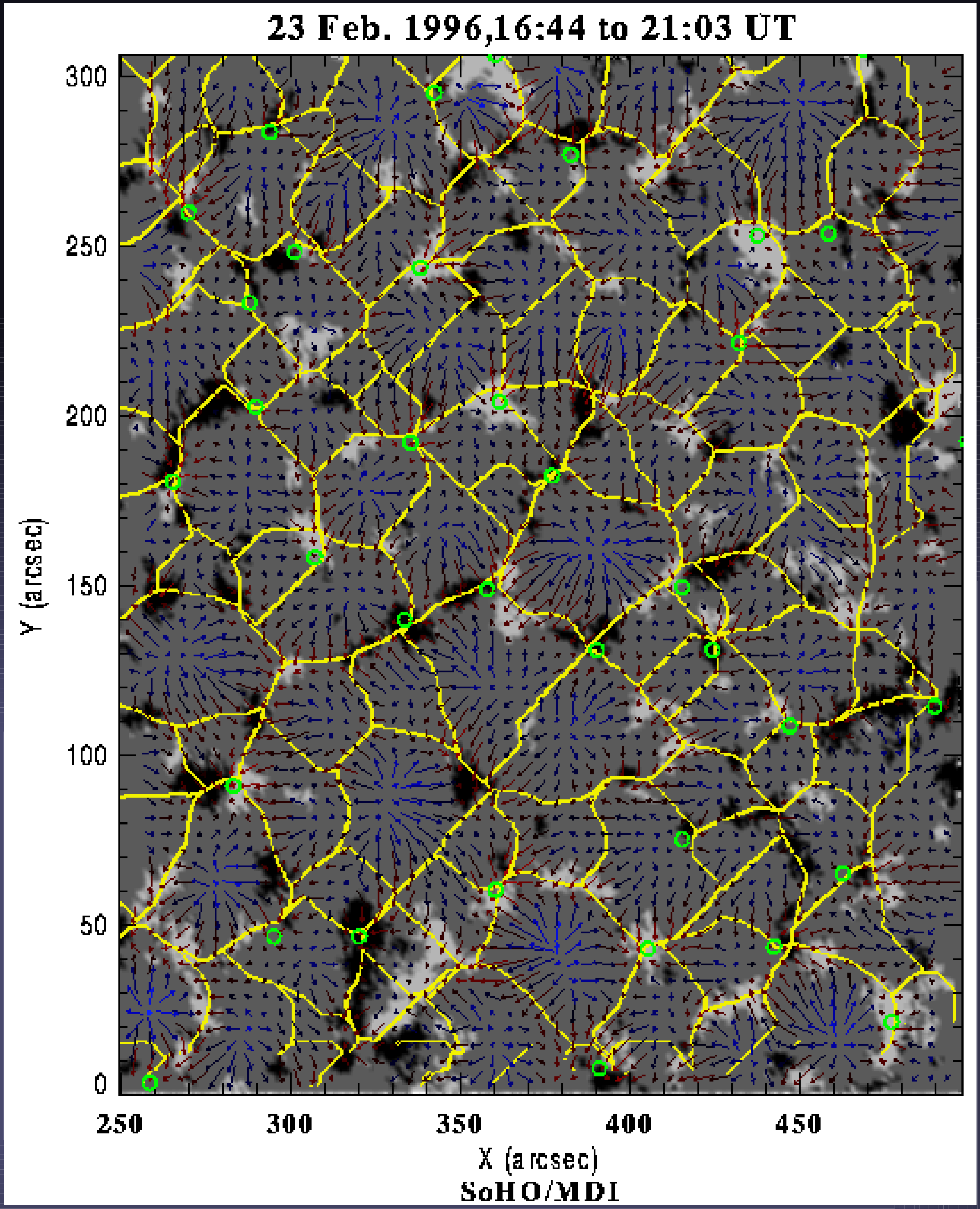
Supergranules seen by SUMER

- Si I 1256 Å full disk scan by SUMER in 1996
- Bright network: found at edges of supergranulation cells
- Darker cells: supergranules



Supergranules & magnetic field

- Why are super-granules seen in chromospheric and transition region lines?
- Network magnetic fields are located at edges of supergranules.
- They appear bright in chromospheric and transition-region radiation (e.g. In UV)



Onset of convection

Schwarzschild's instability criterion

Consider a rising bubble of gas:

bubble	surroundings	depth
ρ^*	$\rho_0 - \Delta\rho$	z-Δz
↑		
ρ	$\rho_0 \ (=\rho)$	z

Condition for convective instability:

$$\boxed{\rho^* < \rho_0 - \Delta\rho}$$

If instability condition is fulfilled ➔ displaced bubble keeps moving ever faster in same direction

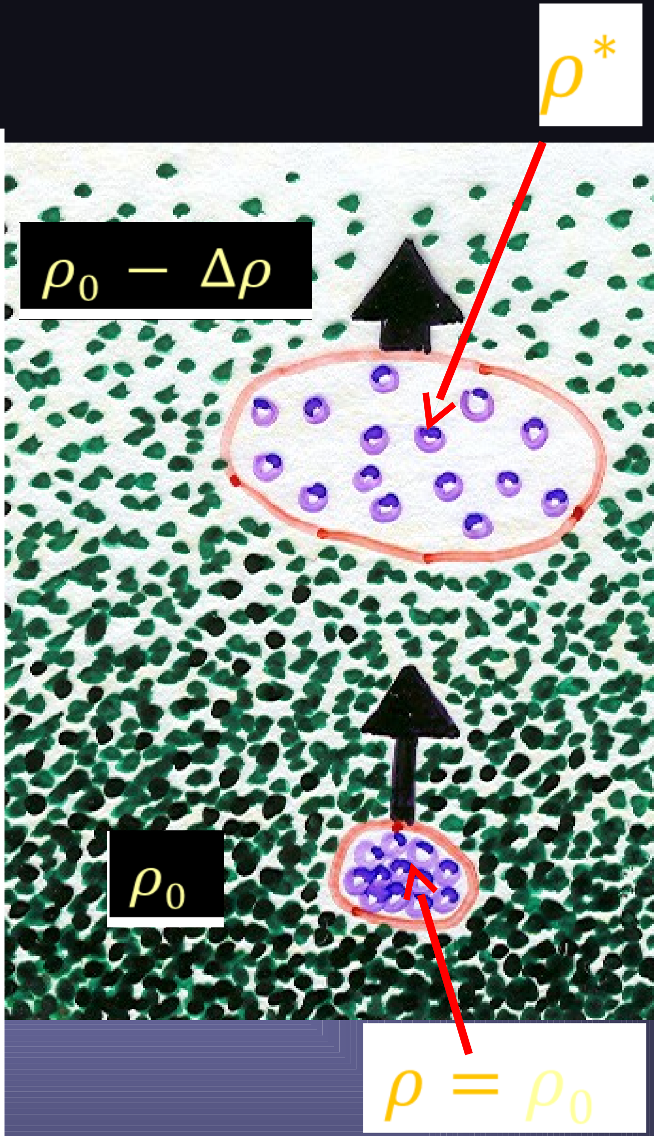
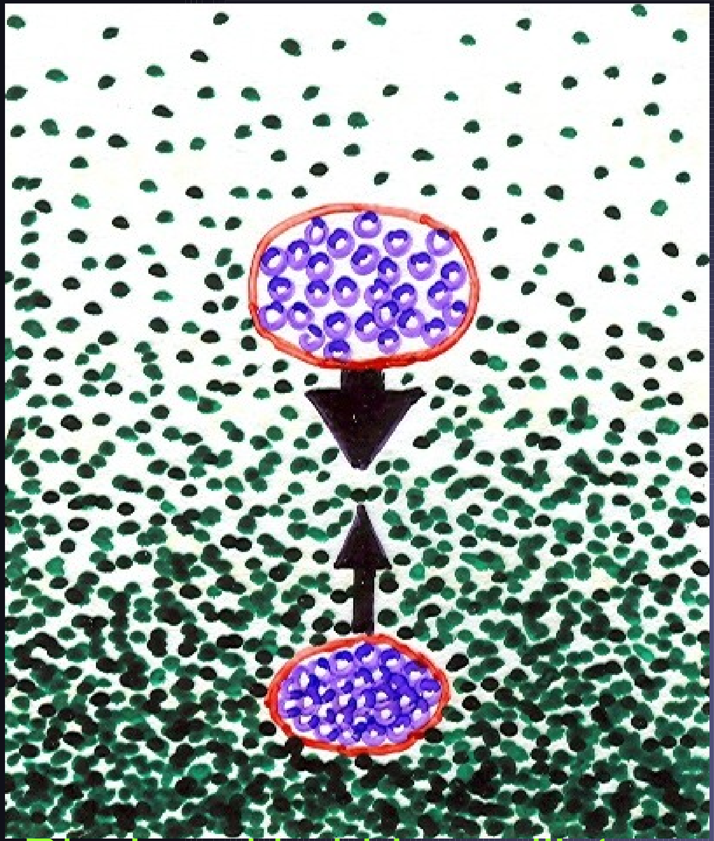


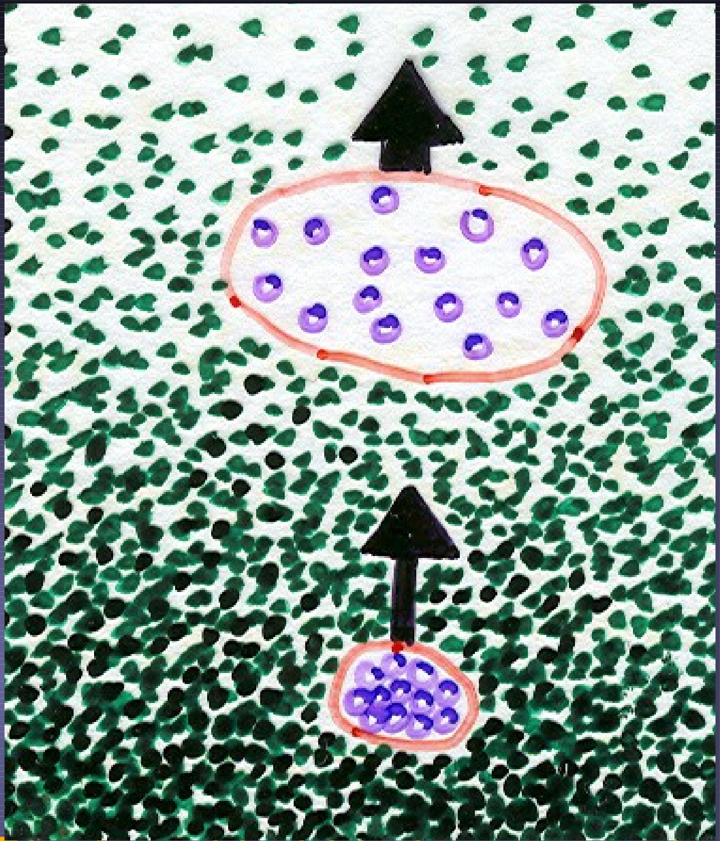
Illustration of convectively stable and unstable situations

Convectively stable



Displaced bubble oscillates: buoyancy or gravity waves

Convectively unstable



Displaced bubble keeps moving: instability □ convection

Onset of convection II

- For small Δz, bubble will not have time to exchange heat with surroundings: adiabatic behaviour. Convectively unstable if:

$$\left[(d\rho/dz)_{rad} - (d\rho/dz)_{adiab}\right] \Delta z < 0$$

$(d\rho/dz)_{rad}$: stellar density gradient in radiative equilibrium

$(d\rho/dz)_{adiab}$: adiabatic density gradient

Often instead of $d\rho/dz$ another gradient is considered: $d\ln T/d\ln P$. A larger $d\rho/dz$ implies a smaller $d\ln T/d\ln P$

Onset of convection III

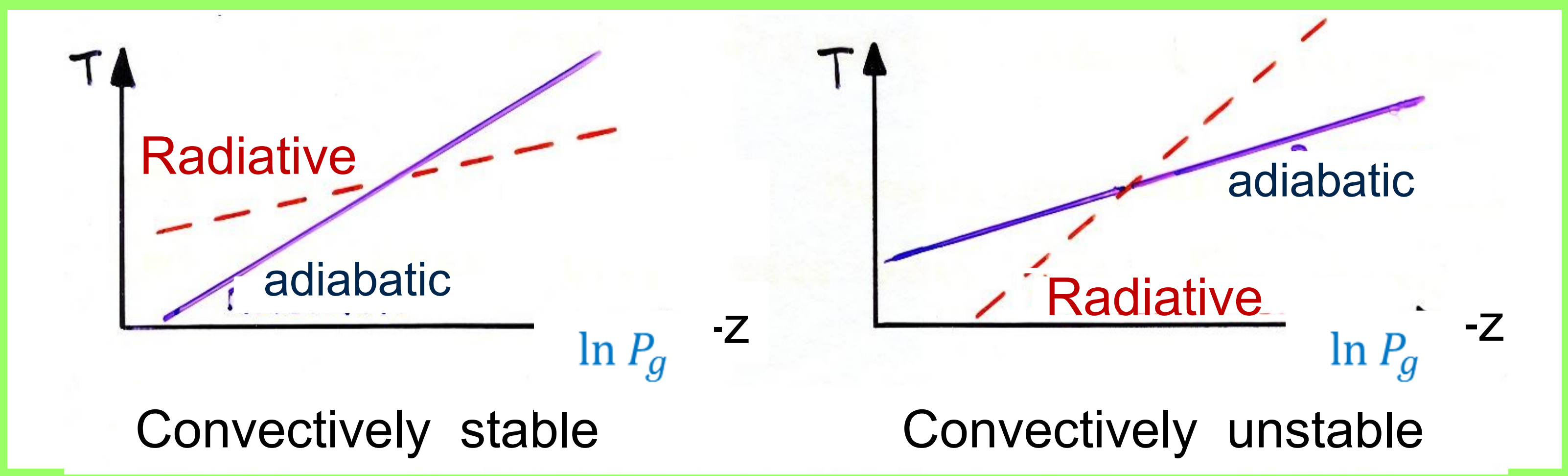
- Rewriting in terms of temperature and pressure:

$$\nabla_{ad} = \left(\frac{d\ln T}{d\ln P_g}\right)_{\mathbf{ad}} = \text{adiabatic temperature gradient}$$

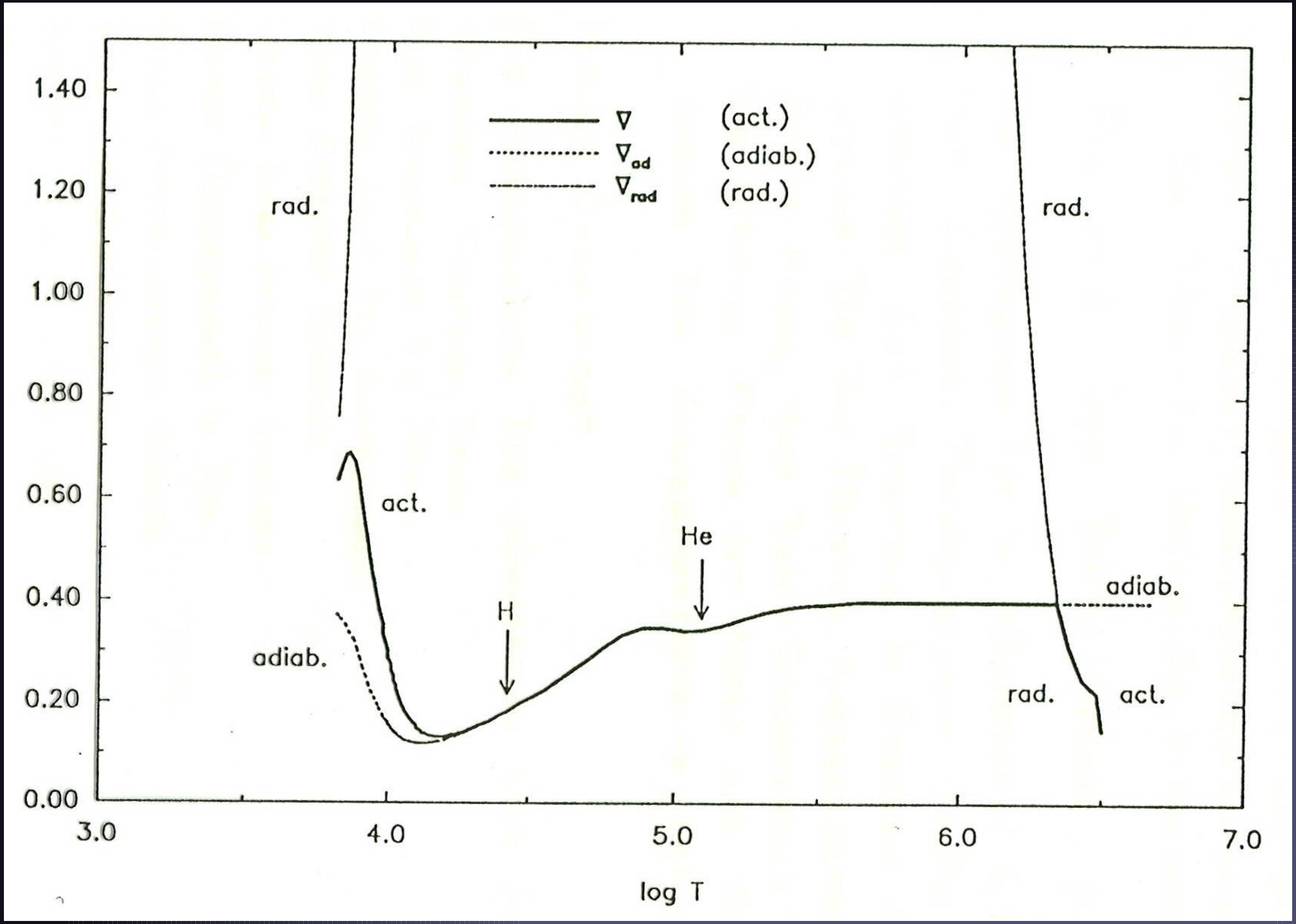
$$\nabla_{rad} = \left(\frac{d\ln T}{d\ln P_g}\right)_{\mathbf{rad}} = T \text{ gradient for radiative equilibrium}$$

Schwarzschild's convective instability criterion:

$$\nabla_{\mathbf{ad}} < \nabla_{\mathbf{rad}}$$



Radiative, adiabatic & actual gradients



Why is radiative gradient so large in convection zone?

Ionisation of H and He

- Radiative gradient is large where opacity increases rapidly with depth. This happens where common elements get ionized (i.e. many electrons are released, increasing f-f opacity)
- Degree of ionisation depends on T and ne :
  - H ionisation happens just below solar surface
  - He → He+ + e- happens 7000 km below surface
  - He+ → He++ + e- happens 30’000 km below surface
- Since H is most abundant, it provides most electrons (largest opacity) and drives convection most strongly
- At still greater depth, ionization of other elements also provides a minor contribution.

Convective overshoot

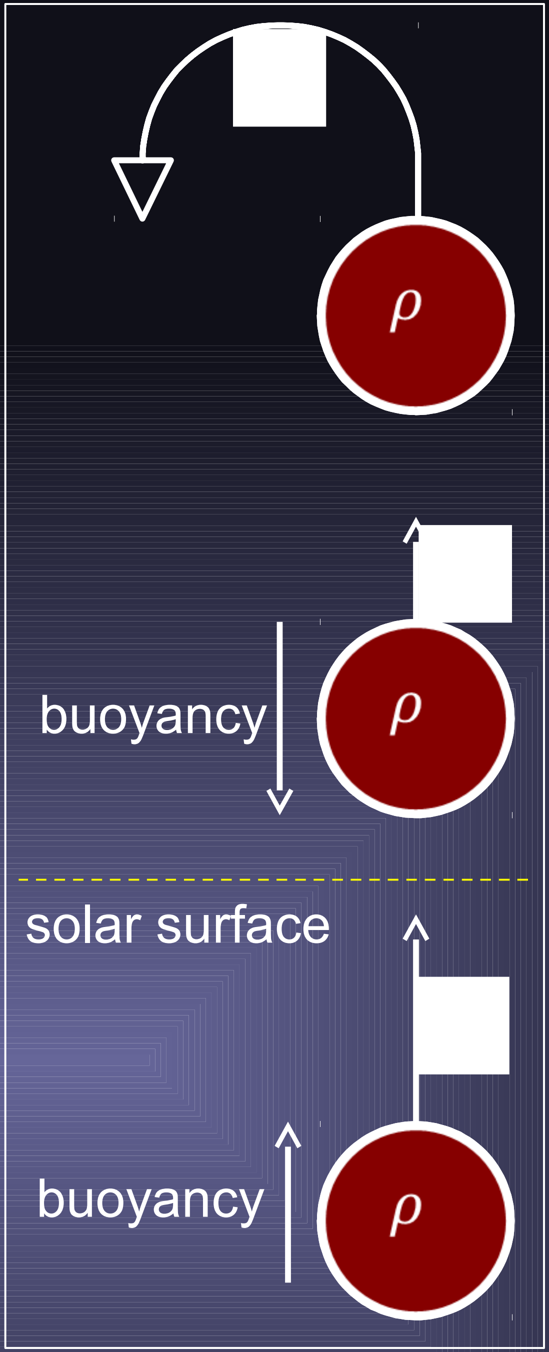
- Due to their inertia, the packets of gas reaching the boundary of convection zone pass into the convectively stable layers, where they are braked & finally stopped

overshooting convection

  - Typical width of overshoot layer: order of H_p
  - This happens at both the bottom and top boundaries of the CZ and is important:

    top boundary: Granulation is overshooting material. $H_p \approx$ 150 km in photosphere

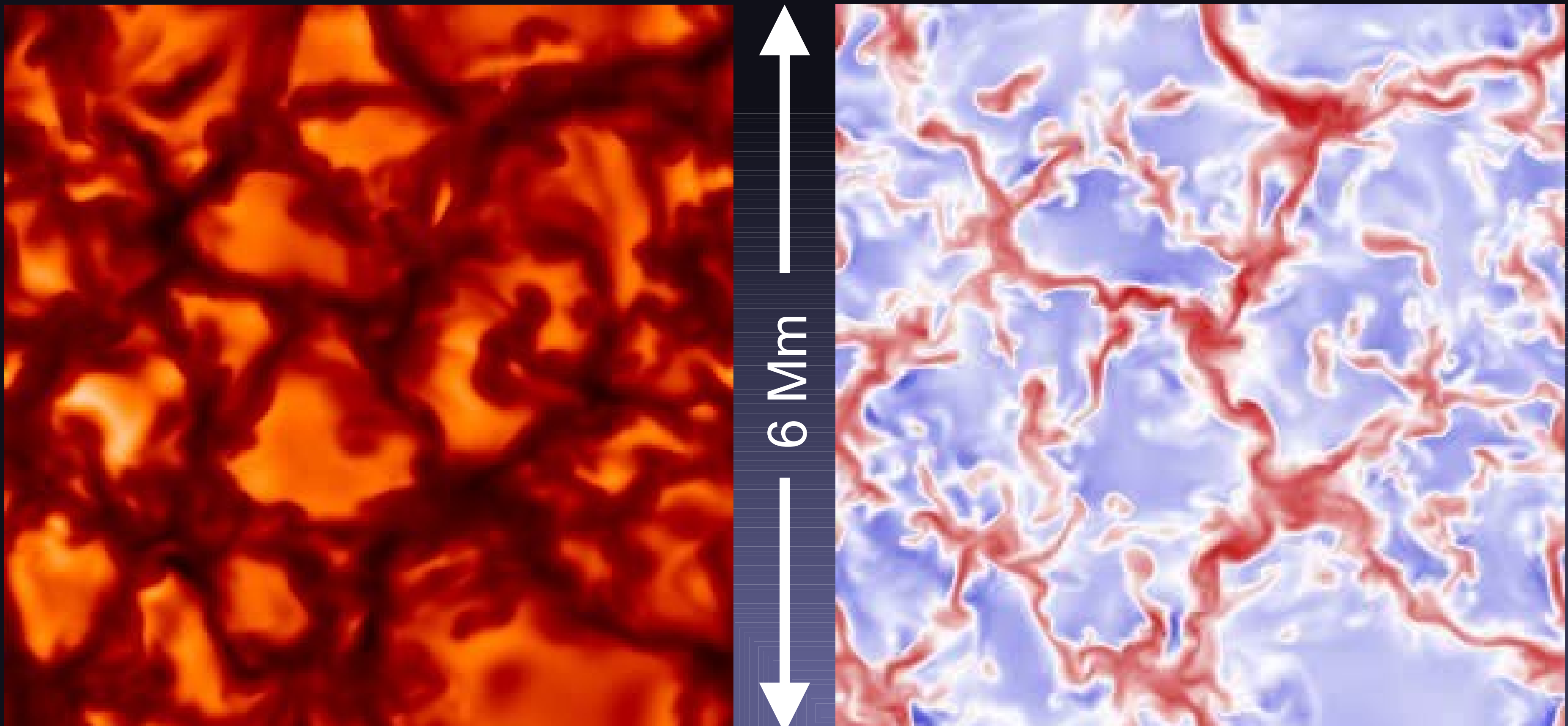
    bottom boundary: the overshoot layer allows B-field to be stored → seat of the dynamo?



Convection Simulations

- 3-D hydrodynamic simulations reproduce a number of observations and provide new insights into solar convection
- These codes solve for mass conservation, momentum conservation (force balance, Navier-Stokes equation), and energy conservation including as many terms as feasible
- Problem: Simulations can only cover 2-3 orders of magnitude in length scale (due to limitations in computing power), while the physical processes on the Sun act over at least 6 orders of magnitude
- Also, simulations can only cover a part of the size range of solar convection, either granulation, supergranulation, or giant cells, but not all

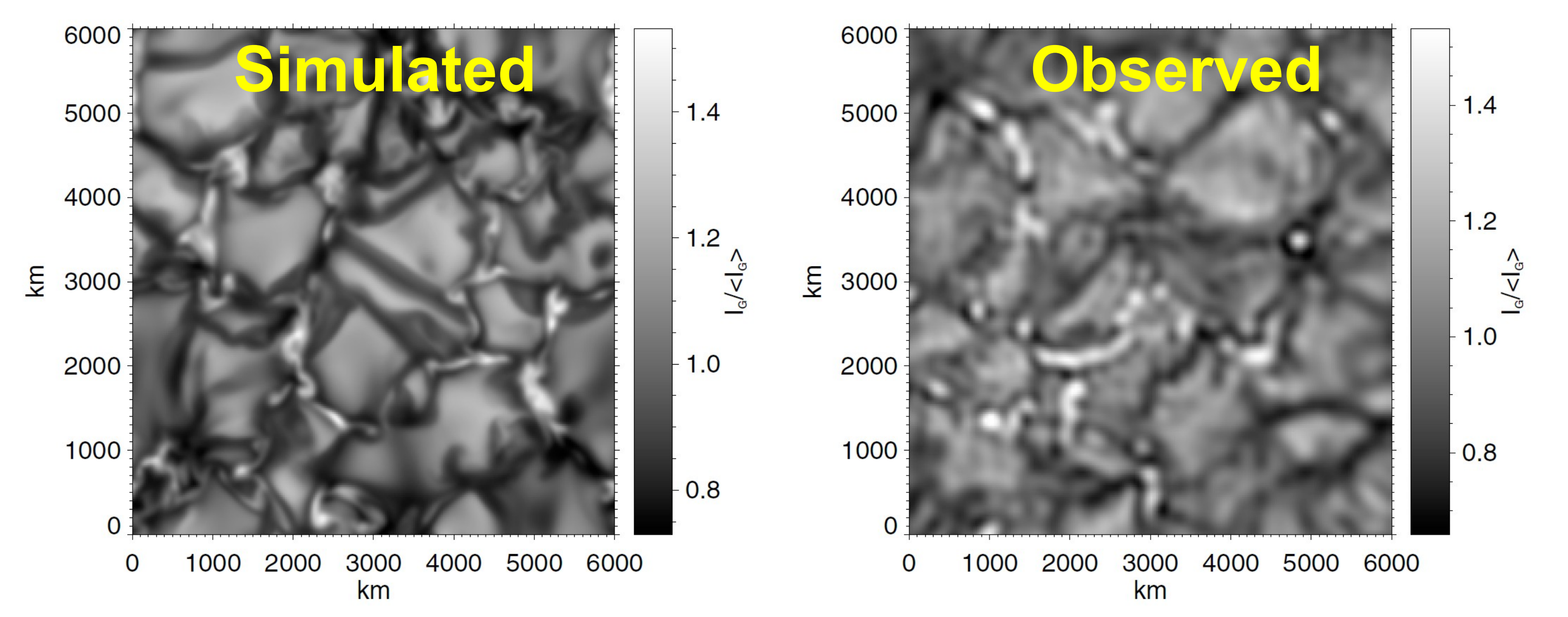
Simulations of solar granulation



Solution of Navier-Stokes equation etc. describing fluid dynamics in a box (6000 km x 6000 km x 1400 km) containing the solar surface. Realistic looking granulation is formed.

Testing the simulations

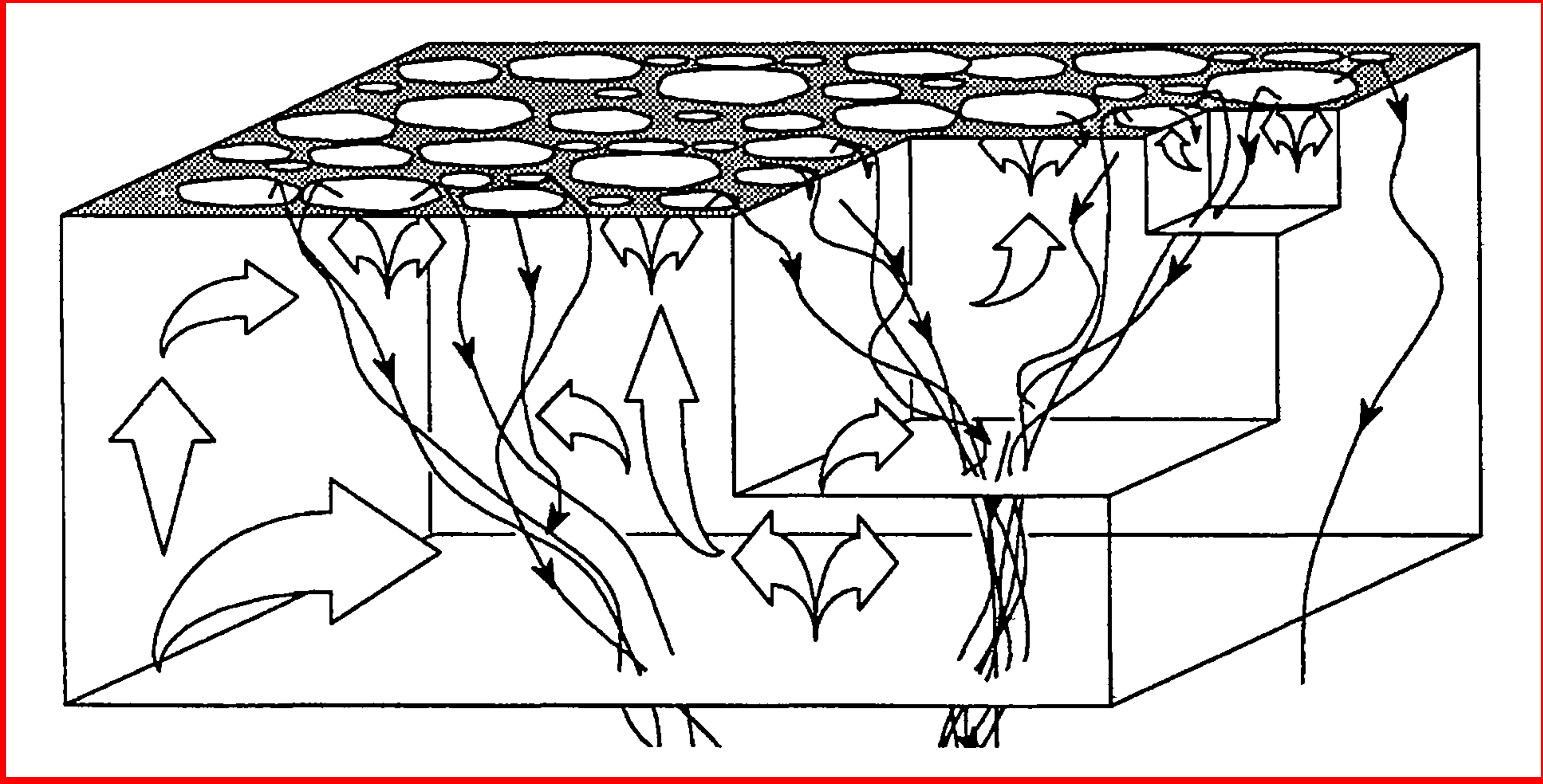
Comparison between observed and computed scenes for spectral lines in yhr Fraunhofer g-band (i.e. lines of an absorption band of the CH molecule): Molecules are dissociated at relatively low temperatures, so that even slightly hotter than average features appear rather bright



Relation between granules and supergranules

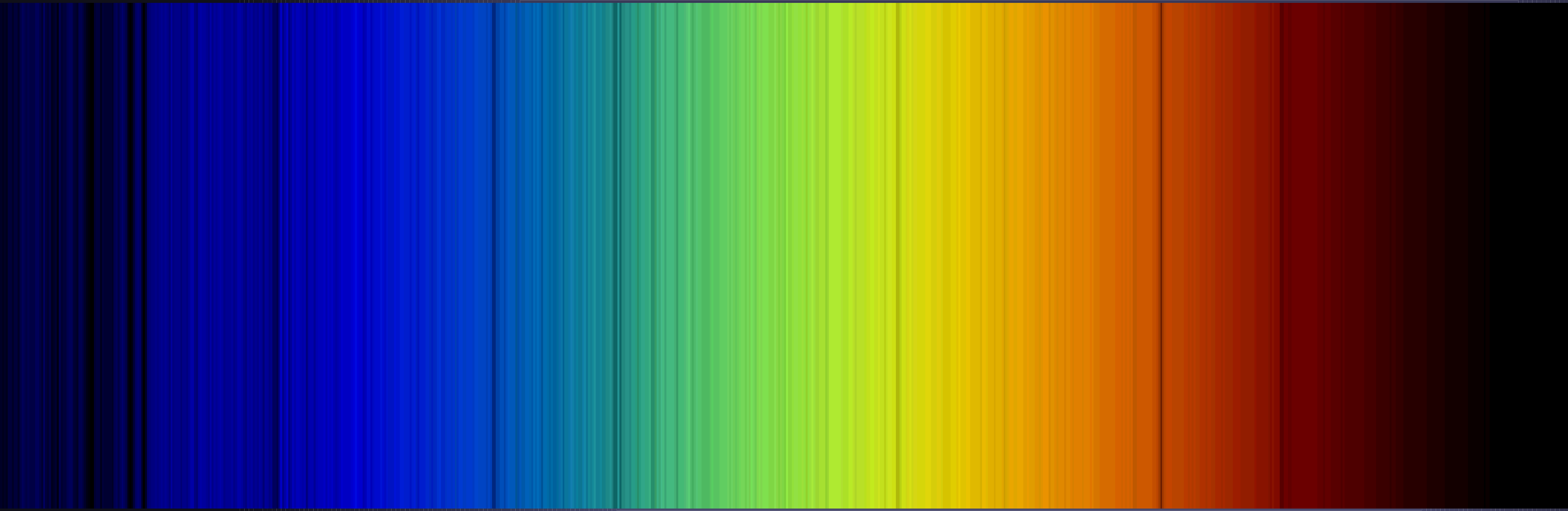
- Downflows of granules continue to bottom of simulations, but the intergranular lanes break up into isolated narrow downflow plumes.
- I.e. topology of flow reverses with depth:
  - At surface: isolated upflows, connected downflows
  - At depth: connected upflows, isolated downflows
- Idea put forward by Spruit et al. 1990: At increasingly greater depth the narrow downflows from different granules merge, forming a larger and less fine-meshed network that outlines the supergranules

Increasing size of convection cells with depth



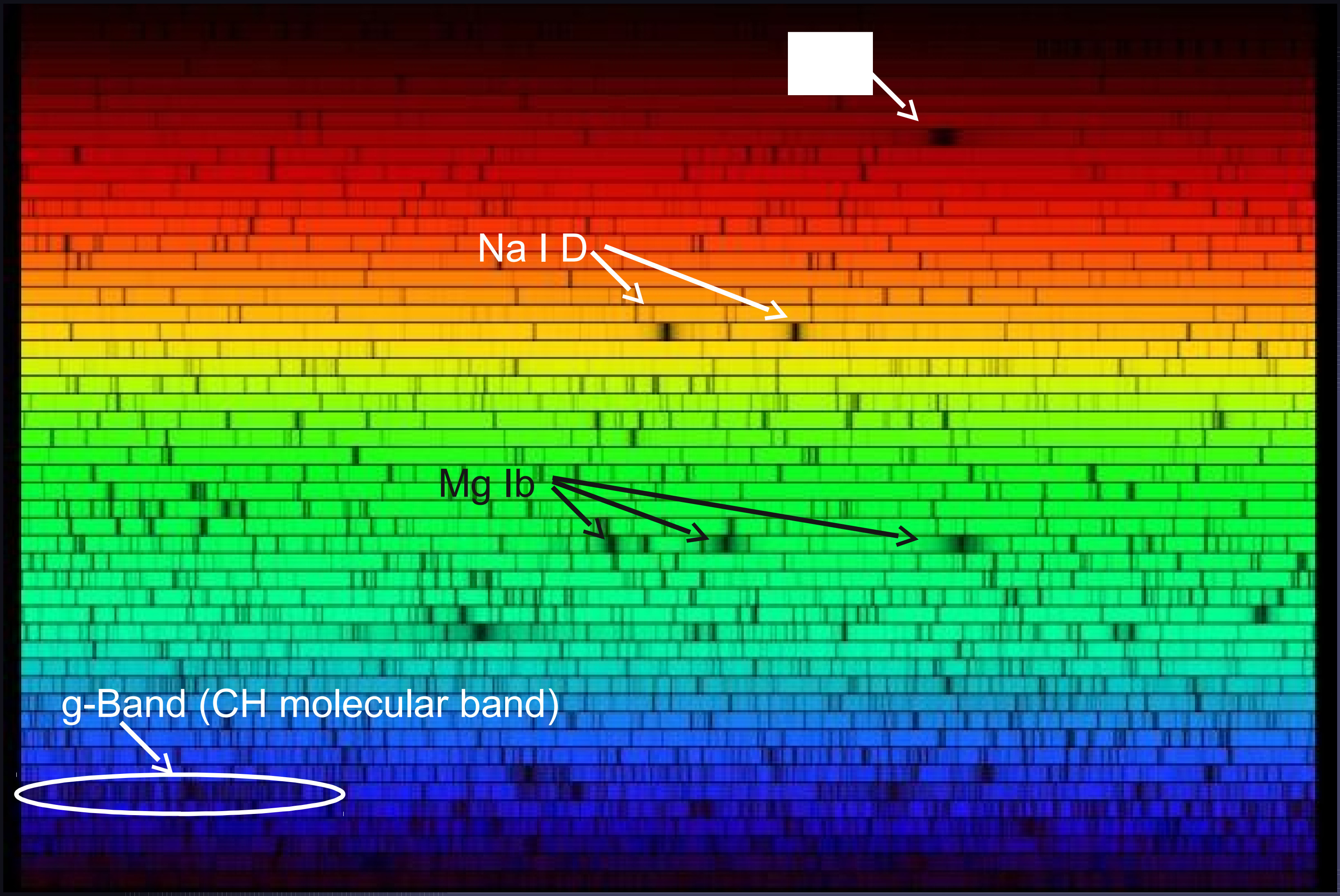
What message does sunlight carry?

Solar radiation and spectrum



Solar spectrum

Approximately 50x better spectrally resolved than previous spectrum

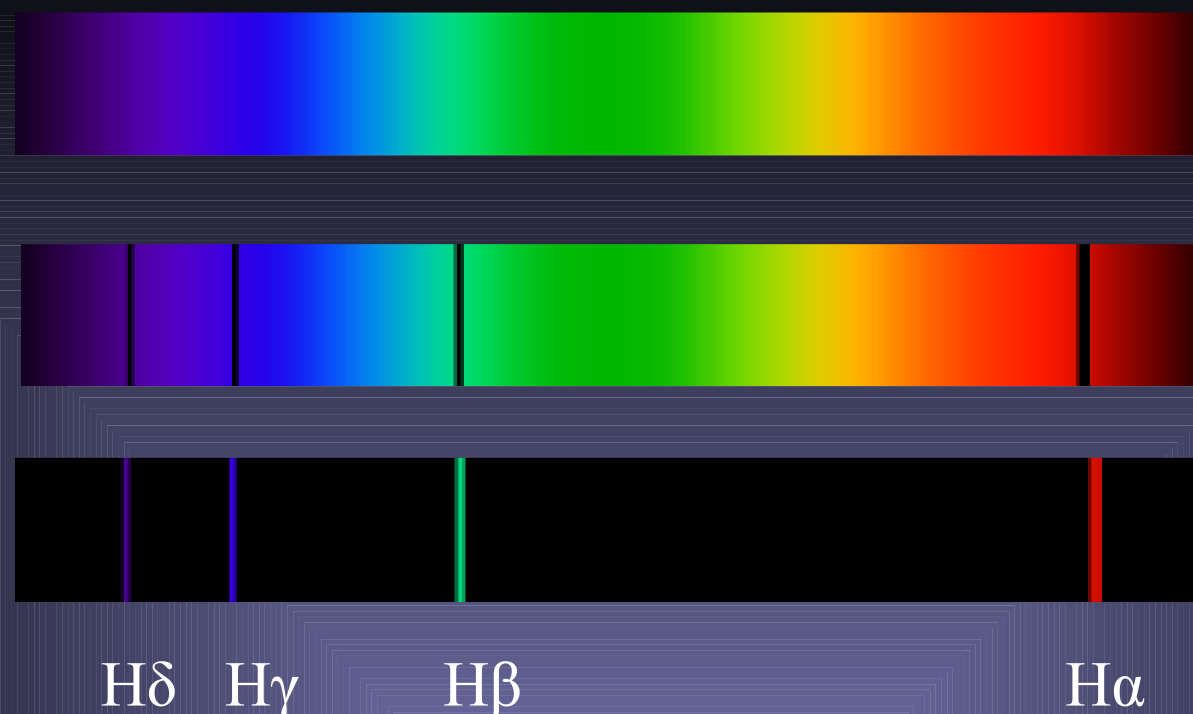


Absorption and emission lines

Continuous spectrum

continuum + absorption lines

Emission lines



Radiation: definitions

- **Intensity** I = energy radiated per wavelength interval, surface area, unit angle and time interval
- **Flux** $F = I$ in direction of observer integrated over the whole Sun:

$$F = 2\pi \int I \mu d\mu, \quad \text{where} \quad \mu = \cos\theta$$

- $F\, r_{\odot}^{\,2}$ = radiation emitted by Sun (or star) in direction of observer
- **Irradiance** S_λ = solar flux at 1AU = $F\, r^2{}_{\odot}/R^2{}_{orbit}$
  - Spectral irradiance S_λ = irradiance per unit wavelength
  - Total irradiance S_{tot} = irradiance integrated over all λ

Hydrostatic equilibrium

- Sun is (nearly) hydrostatically stratified (this is the case even in the turbulent convection zone). I.e. gas satisfies:

$$dP/dz = -g\rho = -gP/\mu RT$$

(P pressure, g gravitational accelaration, ρ density, μ mean molecular weight, R gas constant)

- Solution for constant temperature:

$$P = P_0 \exp(-z/H)$$ → pressure drops exponentially with height

- Here H is the pressure scale height. $H \sim T^{1/2}$ and varies between 100 km in photosphere (solar surface) and ~10^4 km at base of convection zone and in corona

Optical depth

- Let axis z point in the direction of light propagation
- Optical depth: $\Delta\tau_\nu = -\kappa_\nu \Delta z$,

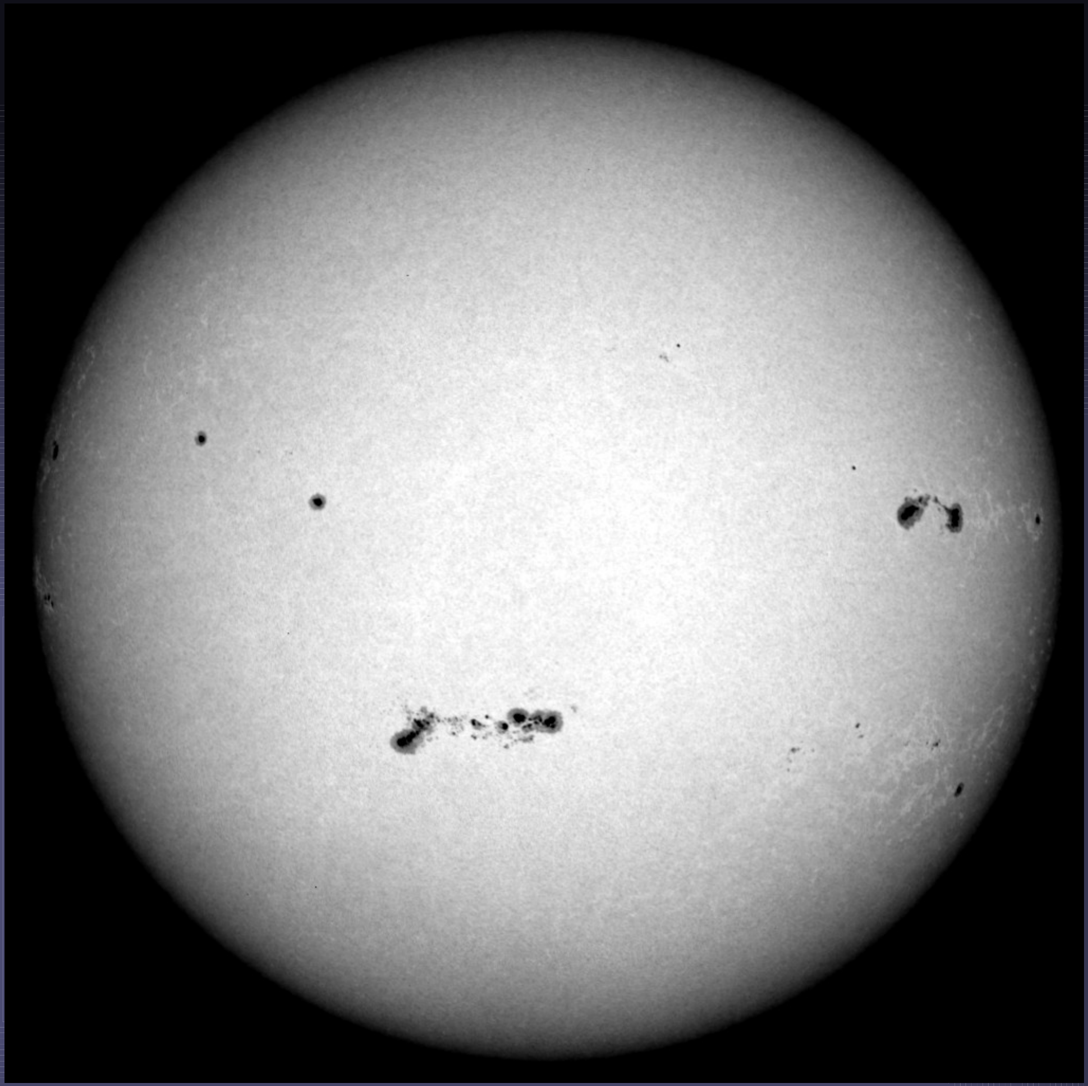
  where κ_ν is the absorption coefficient [cm^{-1}] and ν is the frequency of the radiation. Light only knows about the τ_ν scale and is unaware of z

- Integration over z: $\tau_\nu = -\int \kappa_\nu(z)dz$

  Every ν has its own τ_ν scale (note that the scales are floating, no constant of integration is fixed)

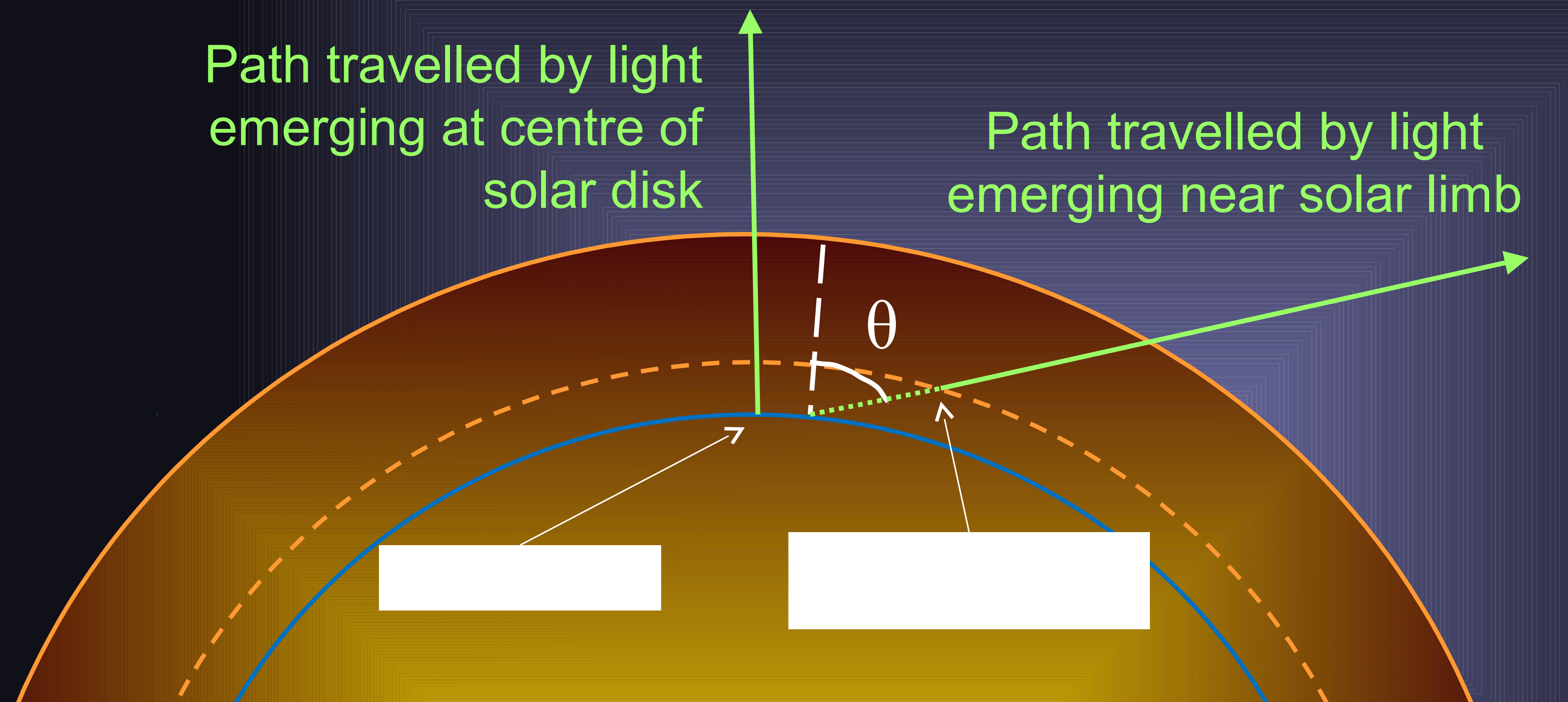
The Sun in white light: Limb darkening

- In the visible, Sun's limb is darker than solar disc centre (□ limb darkening)
- Since intensity ~ Planck function, $B\nu(T)$, T is lower near limb
- Due to grazing incidence we see higher near limb: T decreases outward



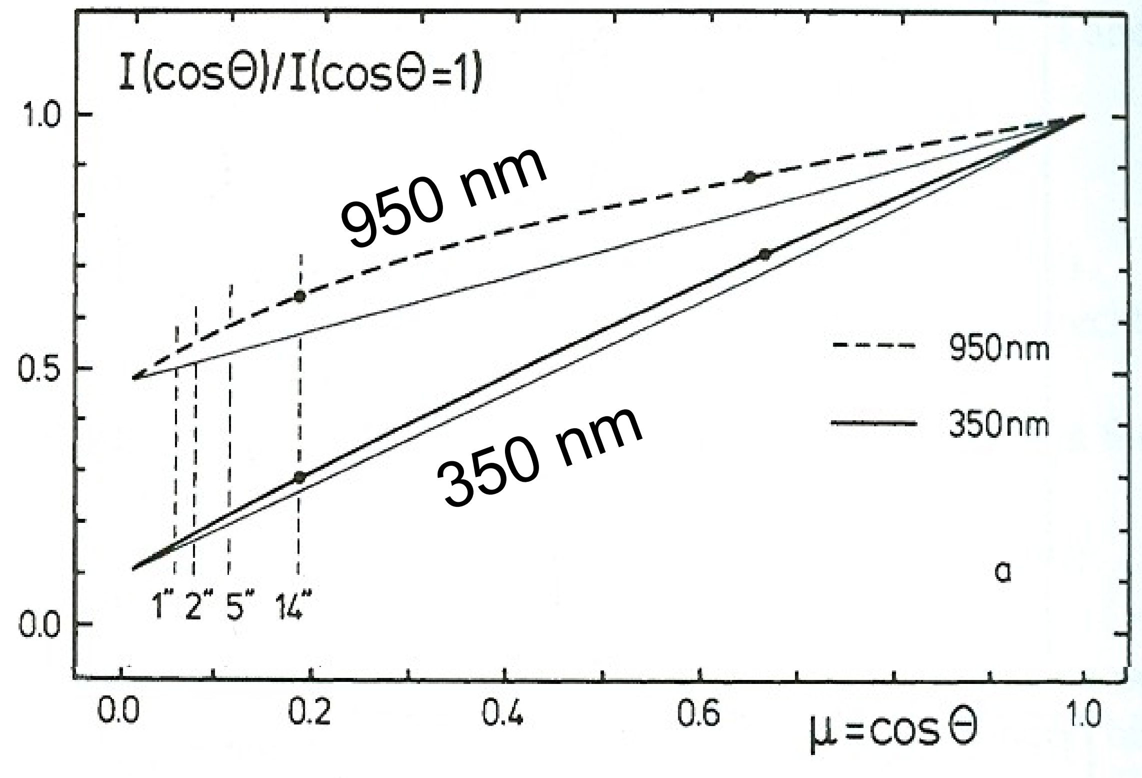
Rays emerging from disk-centre and limb

Rays near solar limb originate higher in atmosphere since they travel 1/cosθ longer path in atmosphere same number of absorbing atoms along path is reached at a greater height



Limb darkening vs. wavelength λ

- Short λ: large limb darkening;
- Long λ: small limb darkening
- Departure from straight line: limb darkening is more complex than $I(\theta) \sim \cos(\theta)$

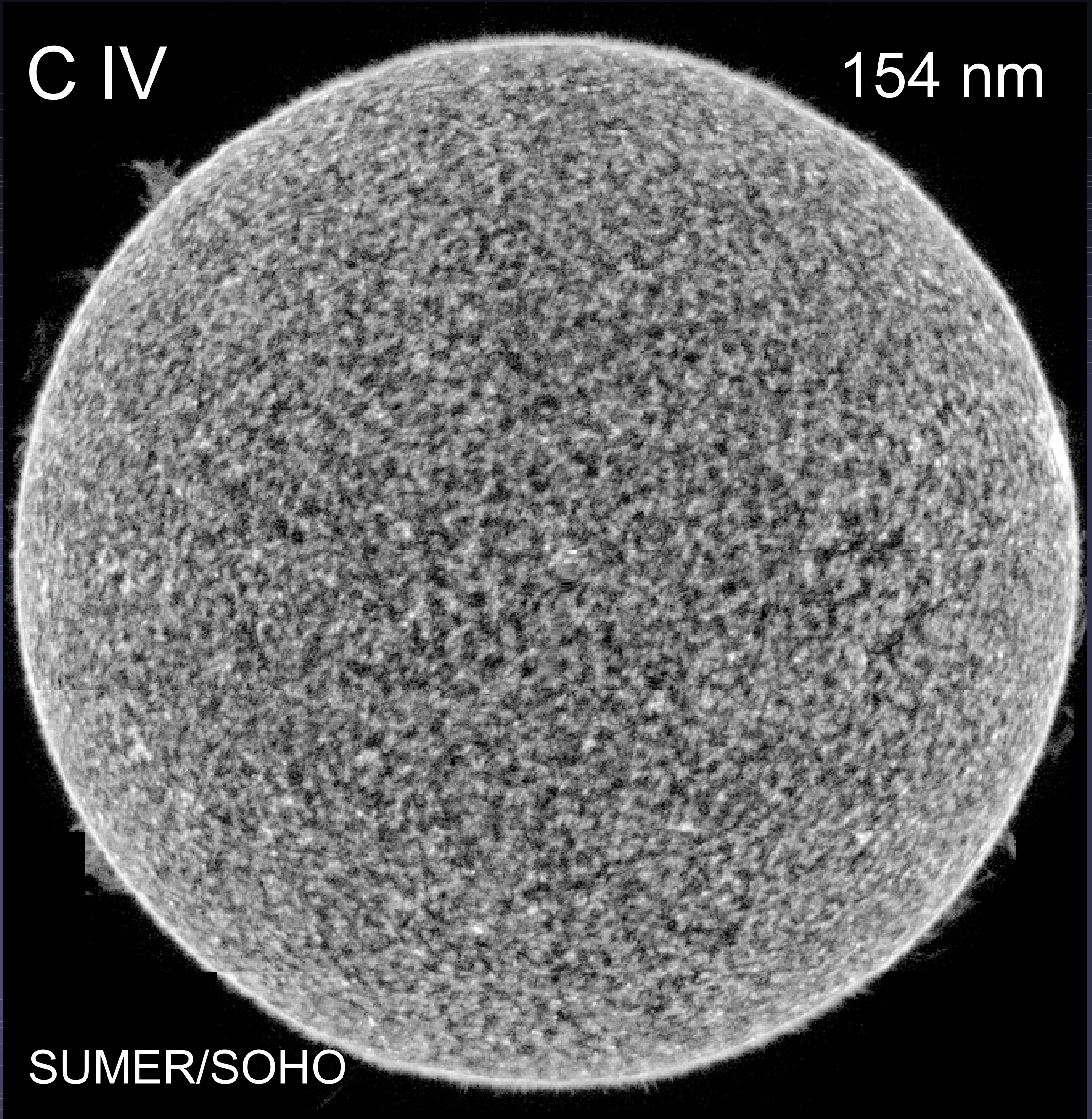


- For most purposes it is sufficient to add a quadratic term $I(\theta) = a_1 \cos\theta + a_2 \cos^2\theta$

The Sun in the FUV: Limb brightening

- At $\lambda < 150$ nm, Sun's limb is brighter than disc centre (→ limb brightening)
- Most FUV spectral lines are optically thin
- Optically thin radiation (i.e. $\tau \ll 1$ throughout atmosphere) comes from the same height at all θ

→ Intensity ~ thickness of layer contributing to it. Near limb this layer is thicker → $I \sim 1/\cos\theta$



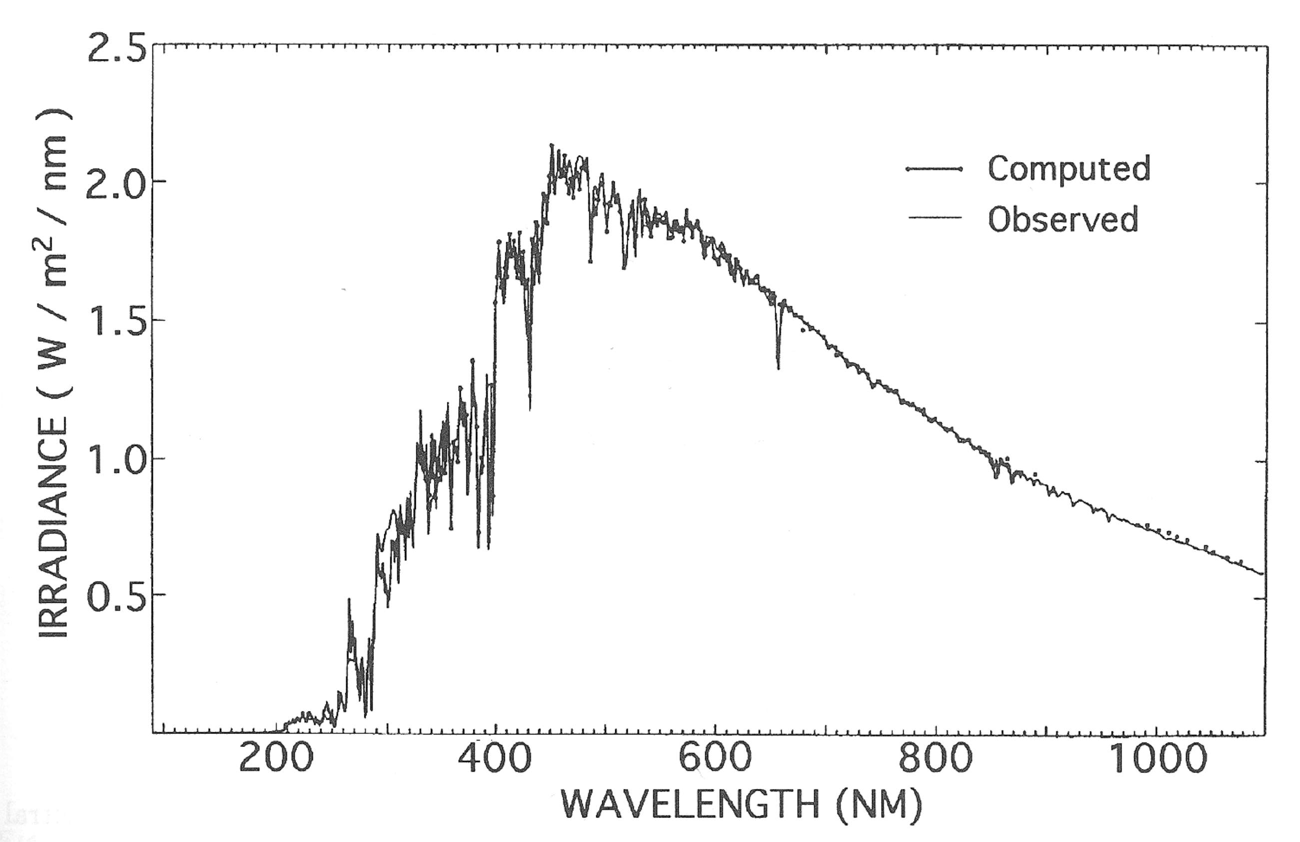
The Sun in the FUV: Limb brightening

- Limb brightening in optically thin lines does NOT imply that temperature increases outwards (although by chance it does in these layers....)

Solar irradiance spectrum

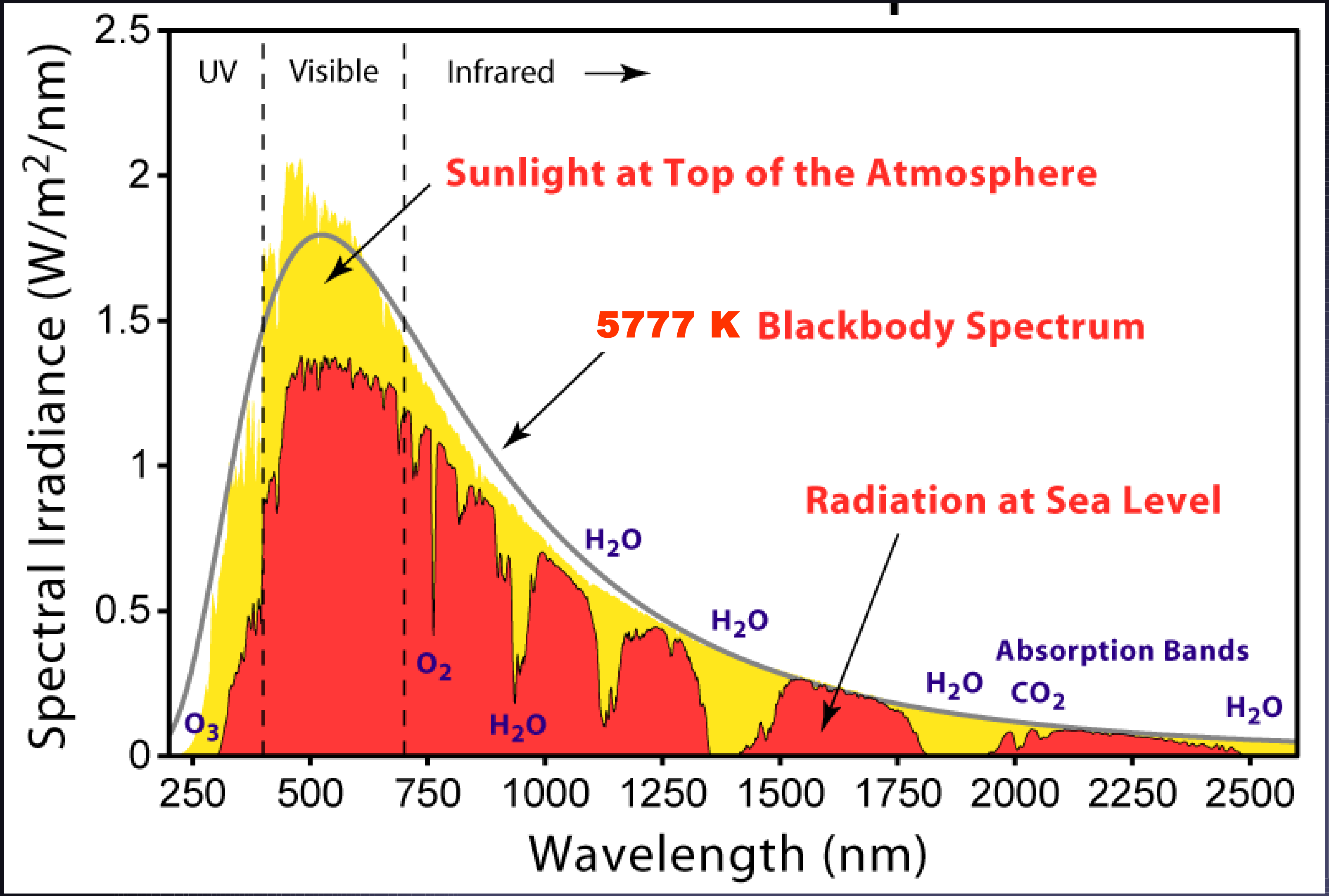
Irradiance = solar flux at 1AU

Spectrum is similar to, but not equal to Planck function
 Radiation comes from layers with diff. temperatures.



Often used temperature measure for stars:
Effective temp: $\sigma T4\mathrm{eff}$ = Area under flux curve

Solar spectrum above and under the Earth's atmosphere



Absorption in the Earth's atmosphere

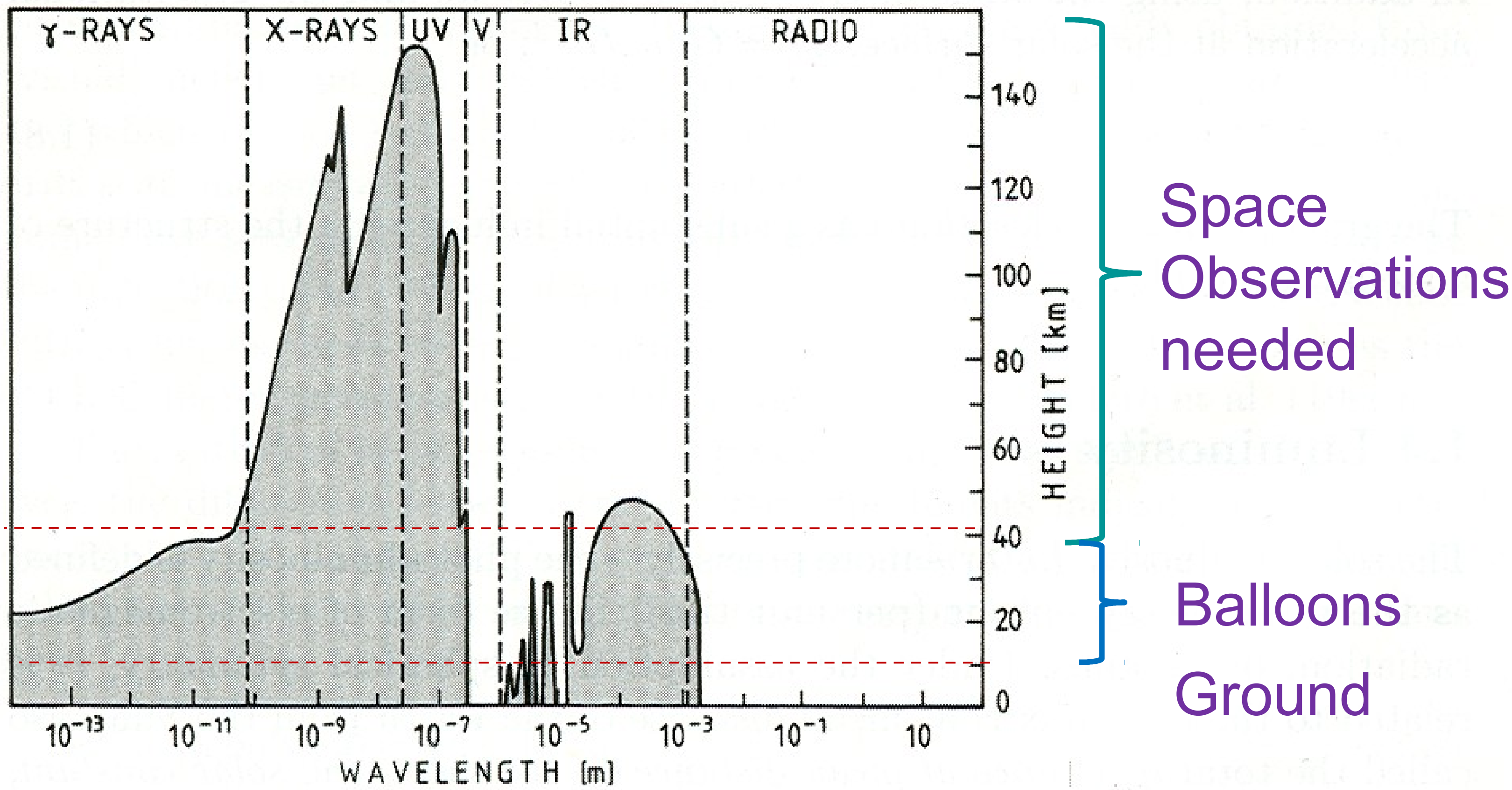
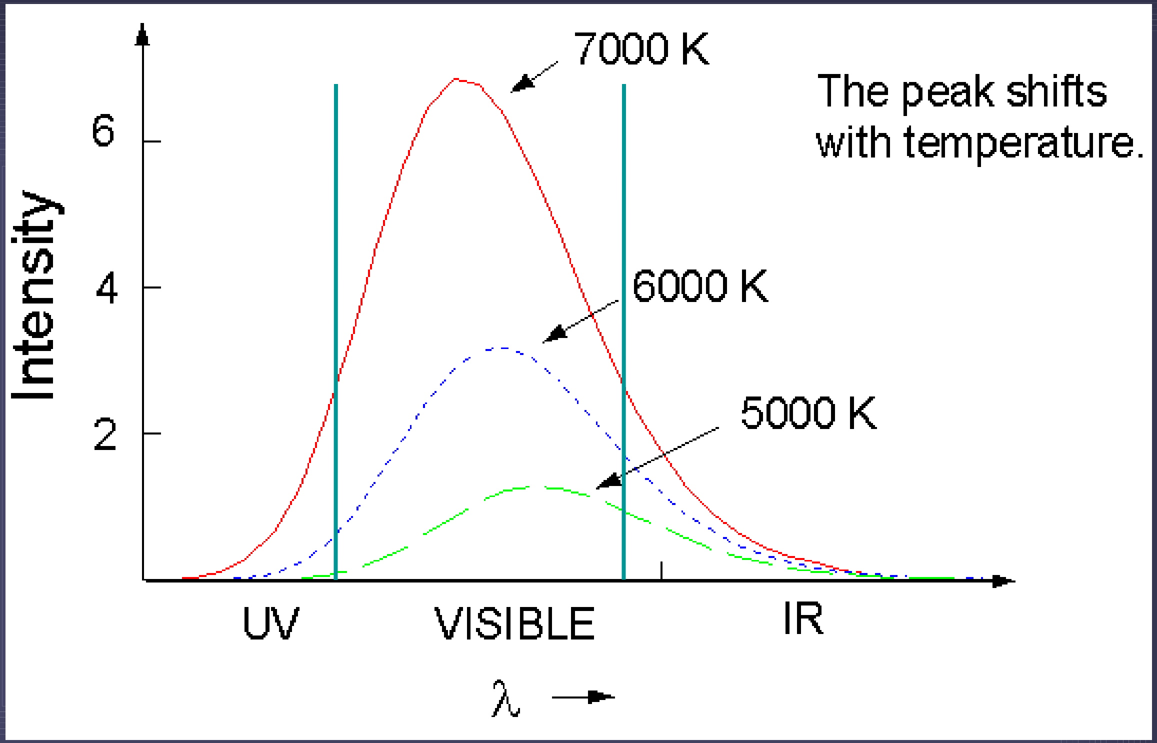


Fig. 1.2. Absorption in the Earth's atmosphere. The edge of the shaded area marks the height where the radiation is reduced to 1/2 of its original strength. UV ultraviolet; V visible; IR infrared

Planck's function

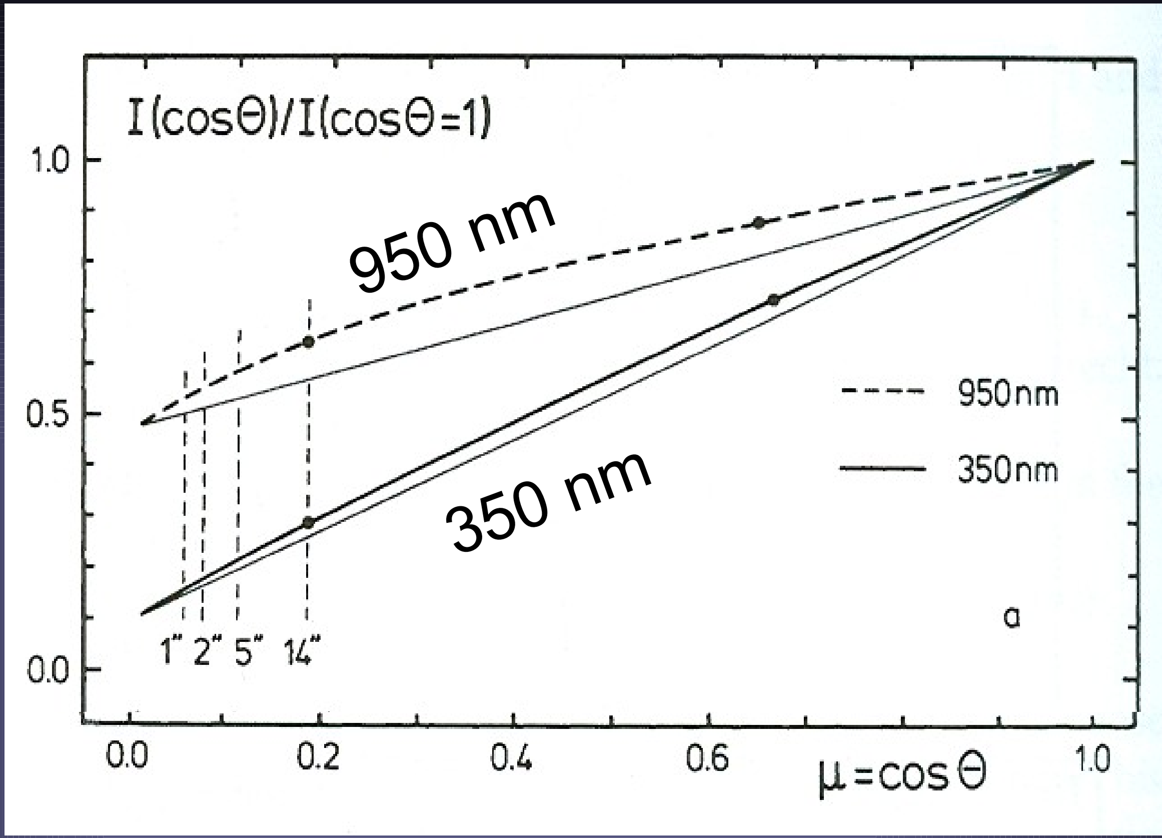
- Amplitude increases rapidly with temperature, area increases $\sim T^4$ (Stefan-Boltzmann law) → from λ-integrated intensity we get (effective) temperature
- Wavelength of maximum changes linearly with temperature (Wien's law)



- Planck function is more sensitive to T at short λ than at long λ

Limb darkening vs. wavelength λ

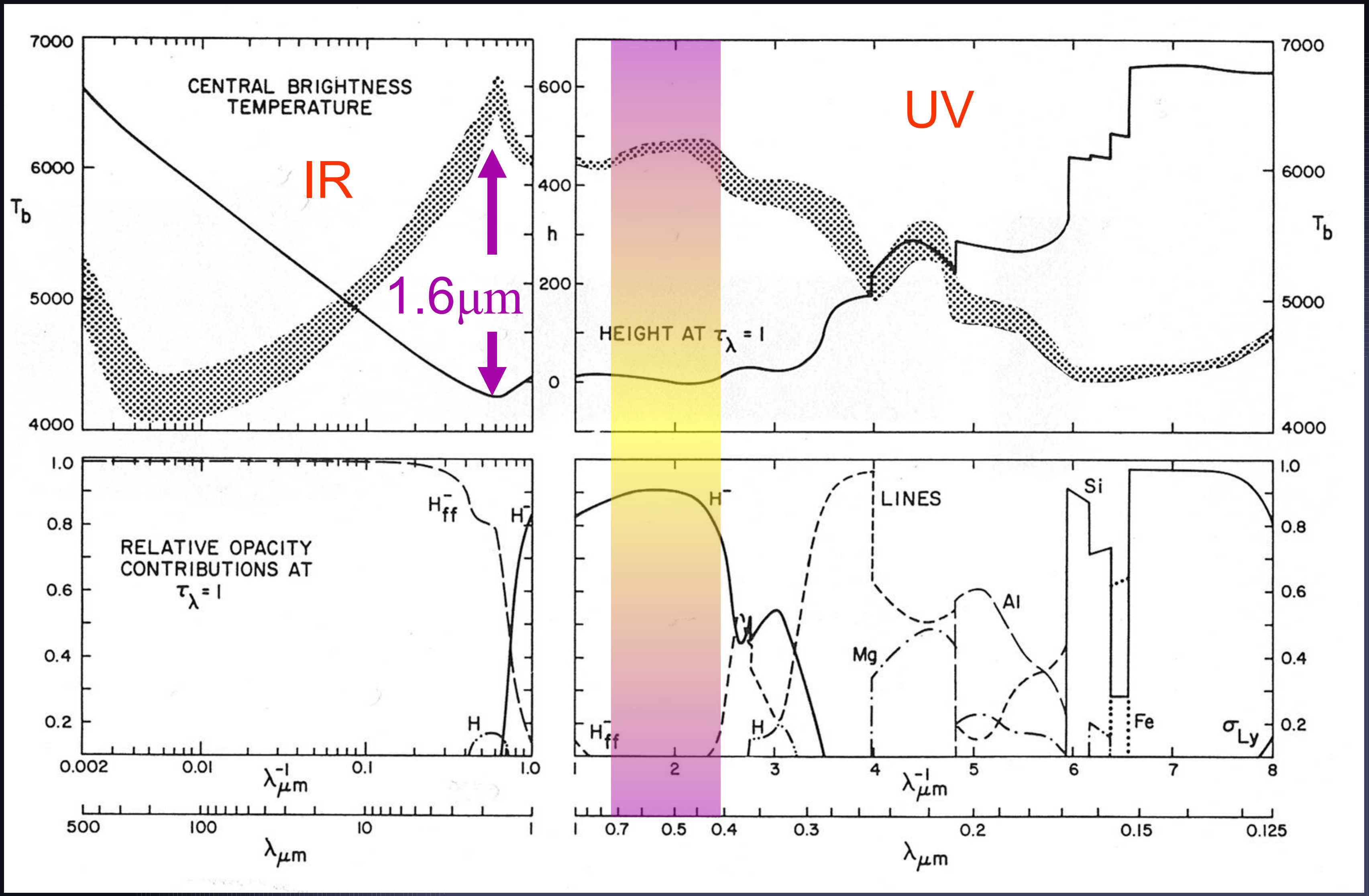
- Short λ: large limb darkening;
- Long λ: small limb darkening
- Due to T-dependence of Planck function, both are consistent with a single $T(z)$ profile



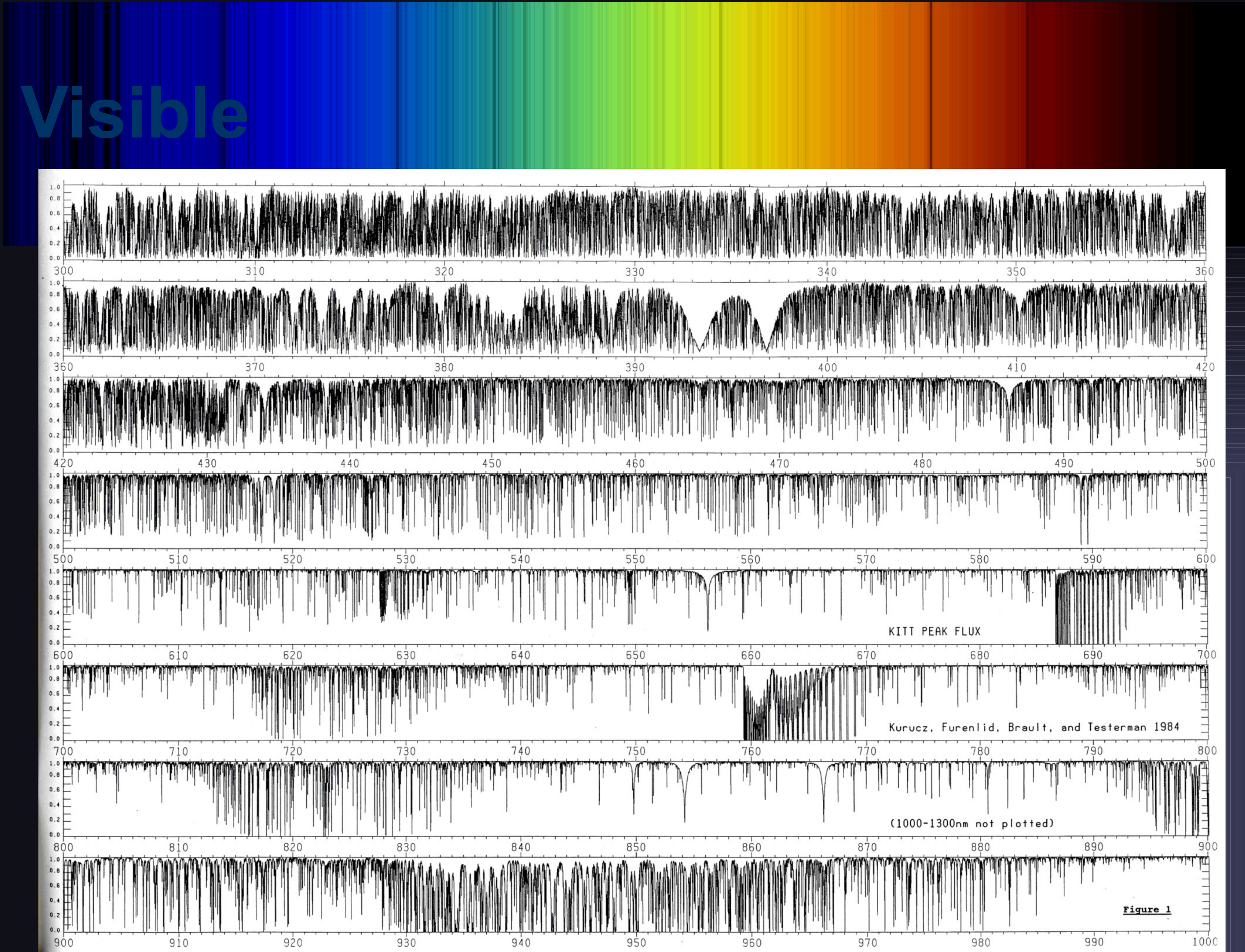
Optical depth and solar surface

- Radiation at frequency ν escaping from the Sun is emitted mainly at heights around $\tau\nu \approx 1$.
- At wavelengths at which $\kappa\nu$ is larger, the radiation comes from higher layers in the atmosphere.
- In solar atmosphere $\kappa\nu$ is small in visible and near IR, but large in UV and Far-IR □We see deepest in visible and Near-IR, but sample higher layers at shorter and longer wavelengths.

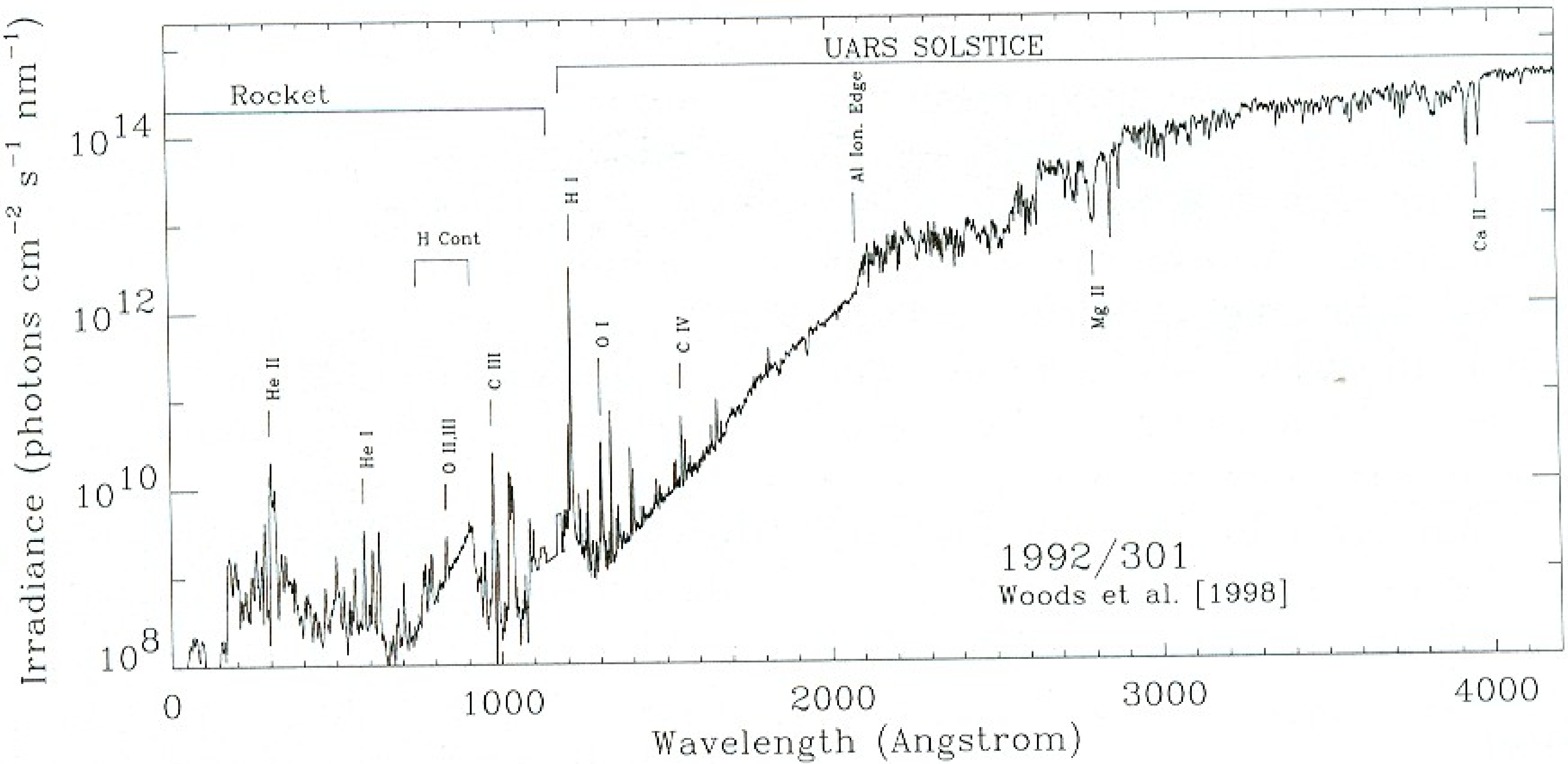
Height of τ = 1, brightness temperature and continuum opacity vs. λ



Visible


KITT PEAK FLUX
Kurucz, Furenlid, Brault, and Testerman 1984
(1000-1300nm not plotted)
Figure 1

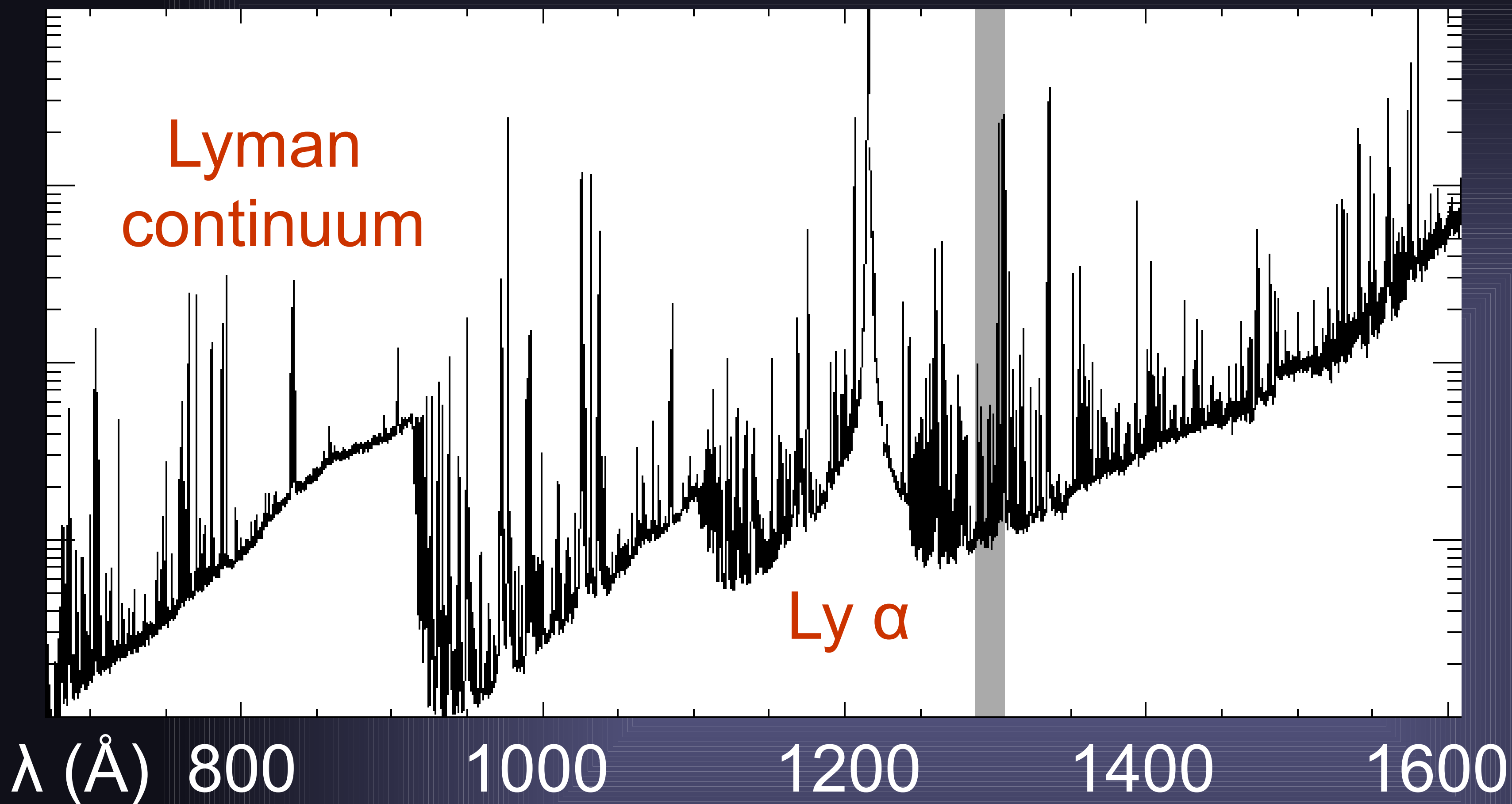
Solar UV spectrum



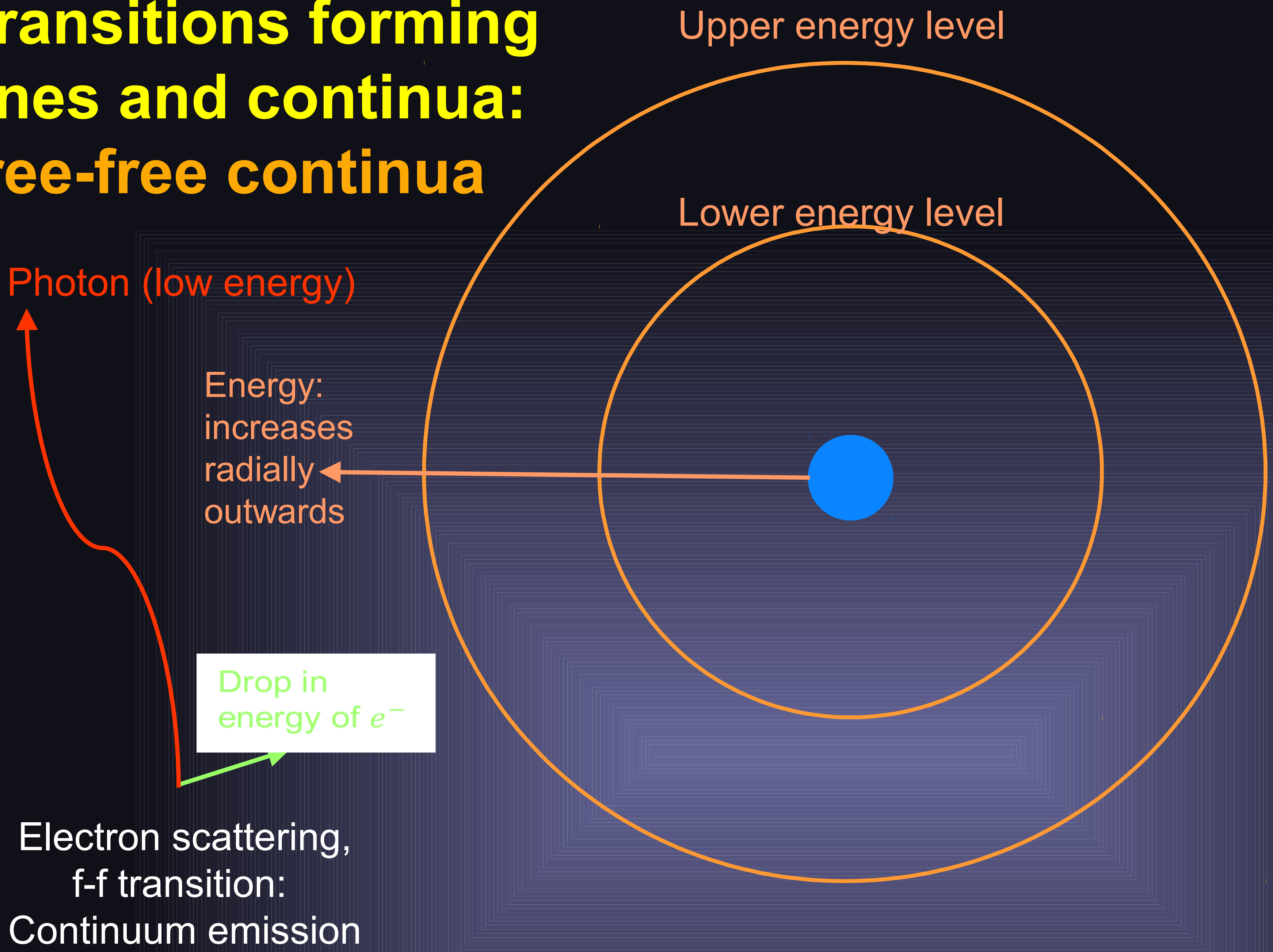
Note the transition from absorption lines (for λ > 2000Å) to emission lines (for λ < 2000Å)

FUV spectrum

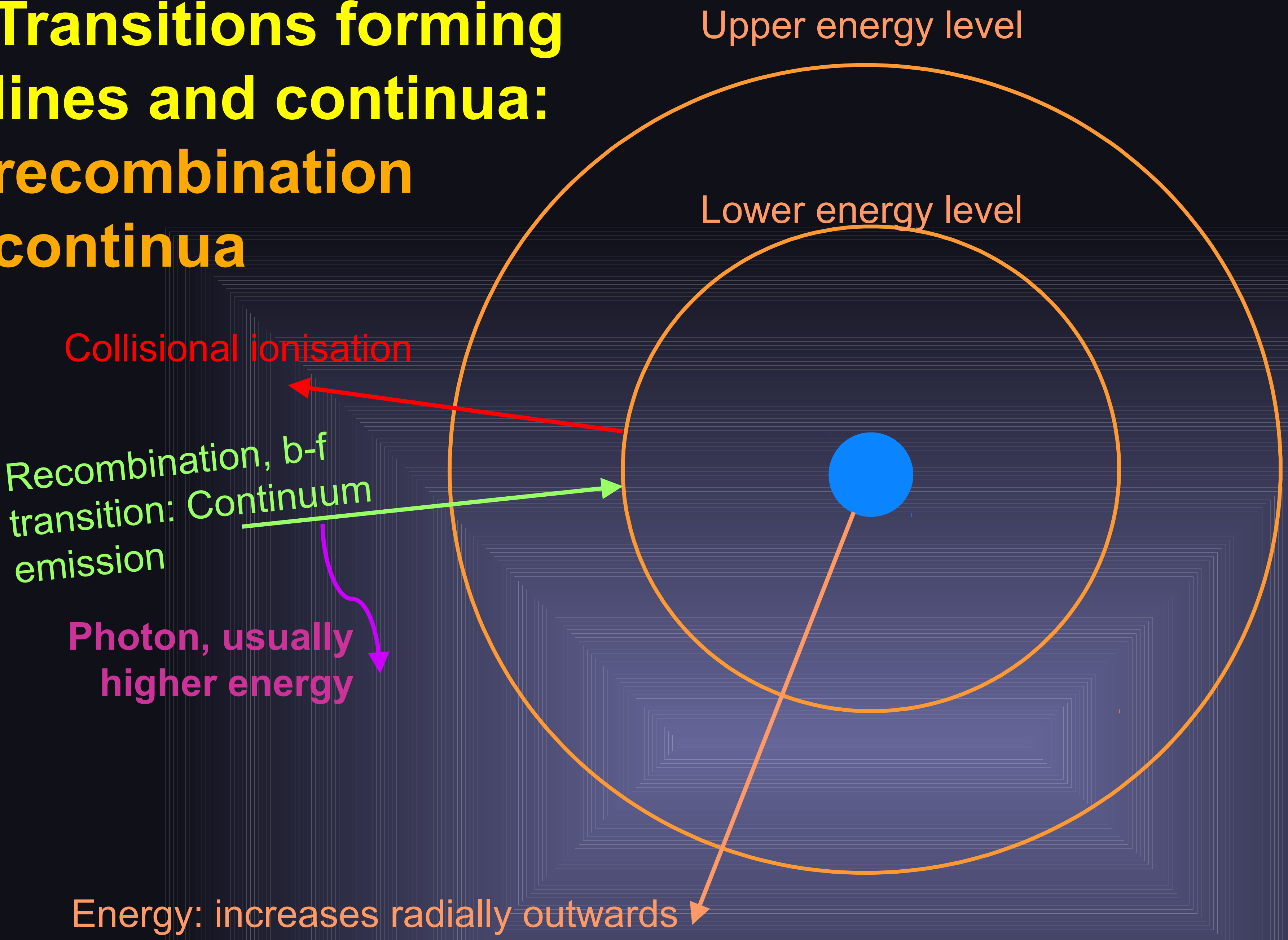
The solar spectrum from 670 Å to 1620 Å measured by SUMER on SOHO (logarithmic intensity scale)



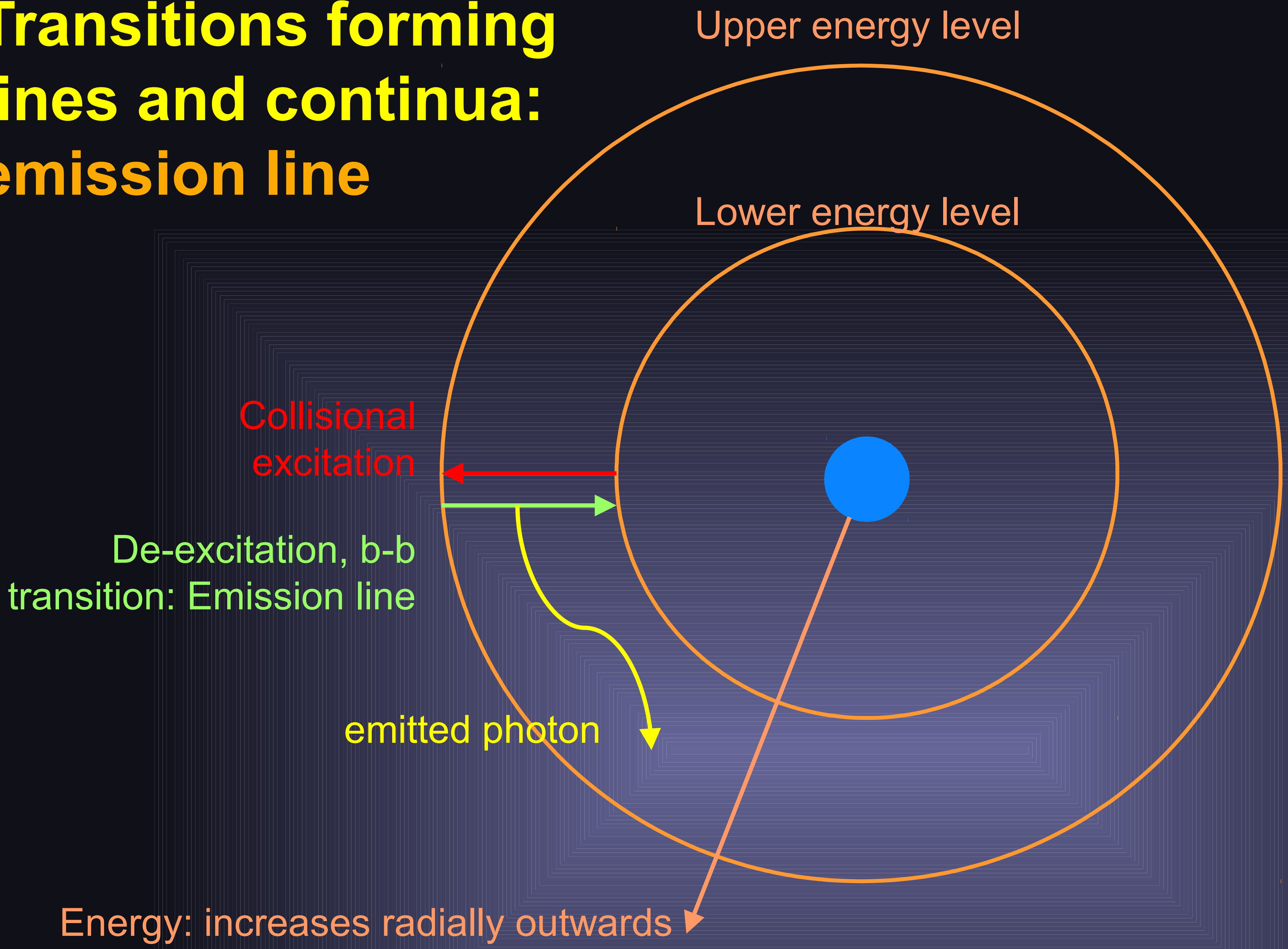
Transitions forming lines and continua: free-free continua



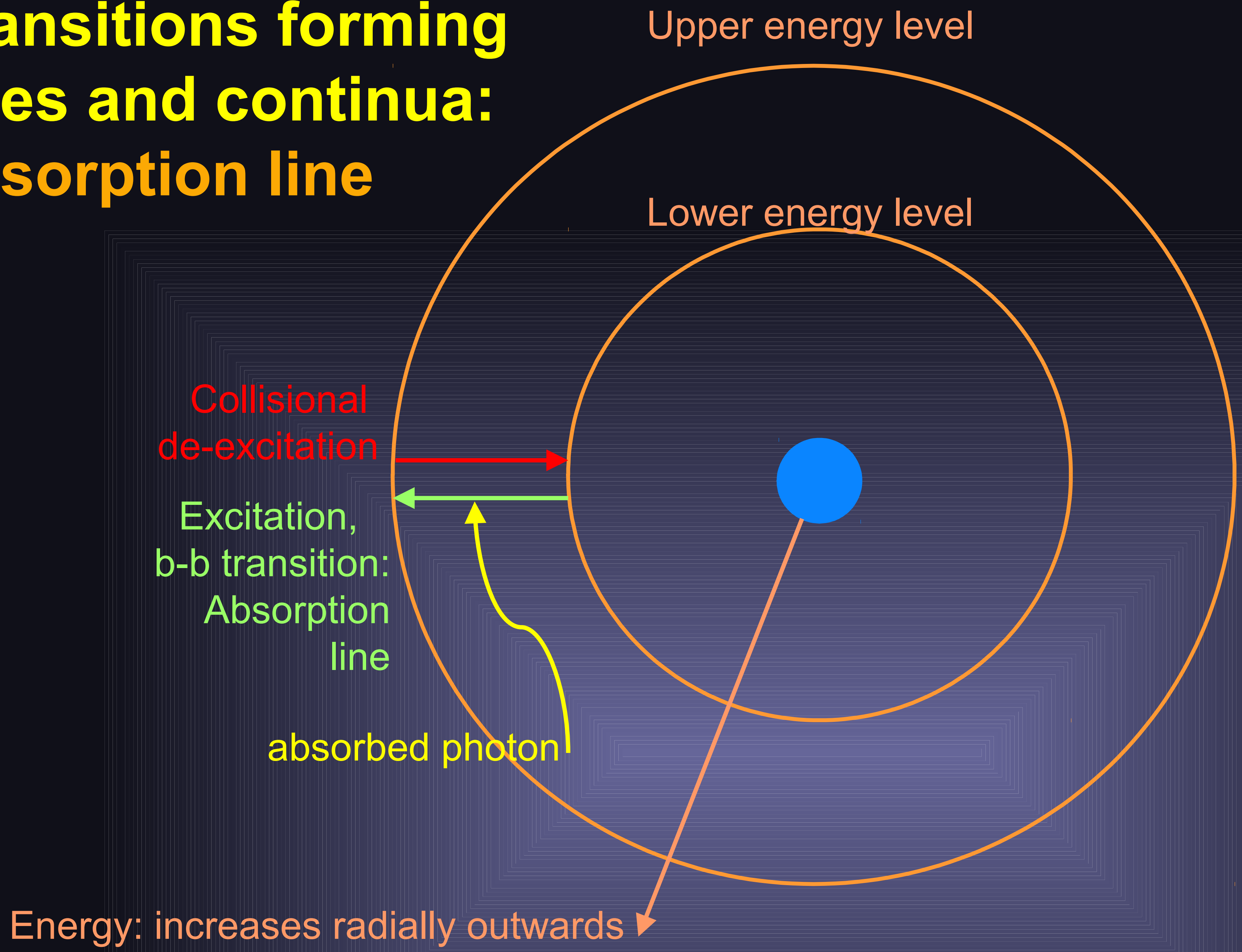
Transitions forming lines and continua: recombination continua



Transitions forming lines and continua: emission line



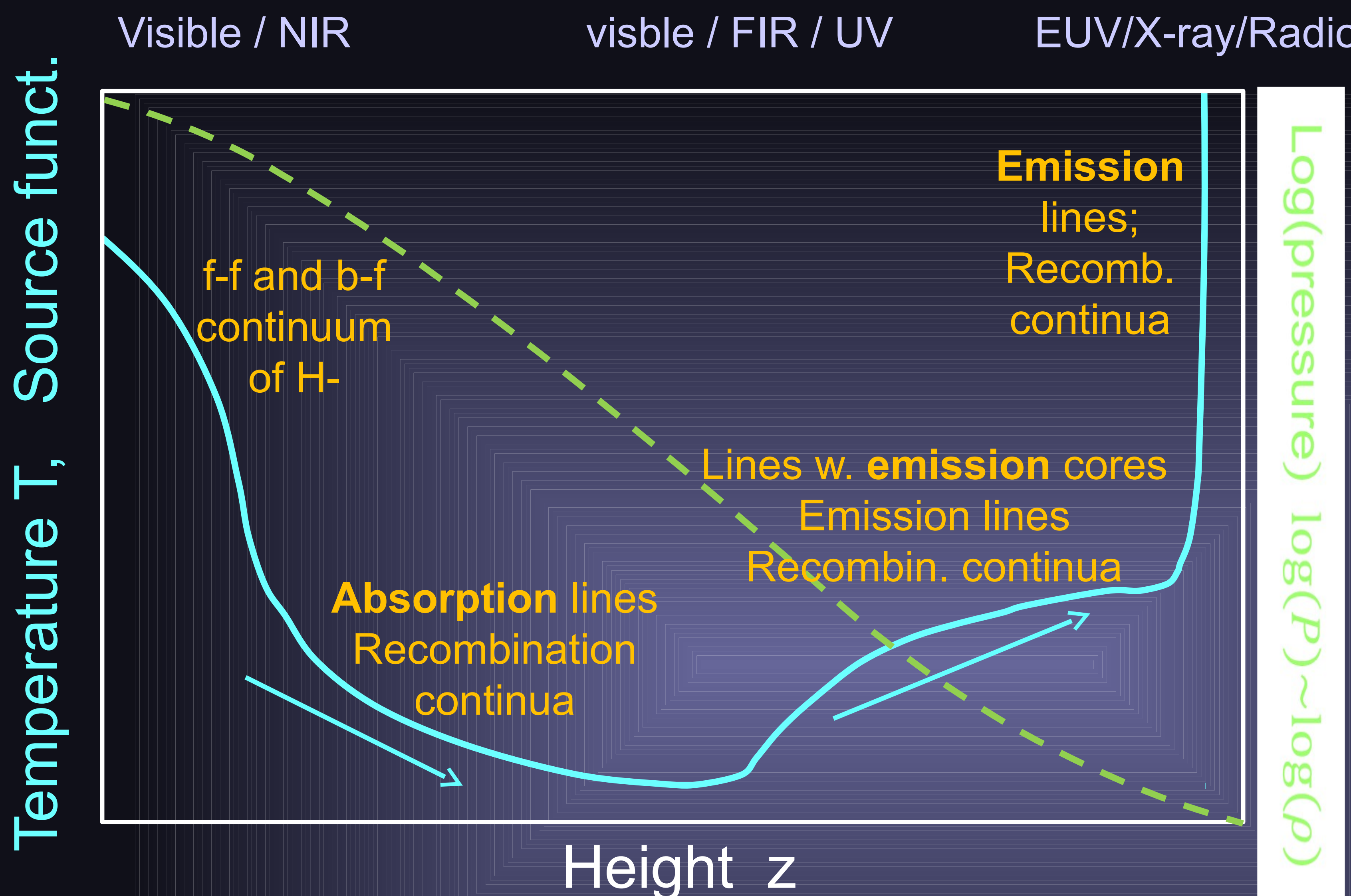
Transitions forming lines and continua: absorption line



The solar spectrum: continua with absorption and emission lines

- Solar spectrum changes in character at different λ
- X-rays: Emission lines of highly ionized species
- EUV - FUV: Emission lines of neutral to multiply ionized species plus recombination continua
- NUV: stronger recombination (bound-free, or b-f) continua and absorption lines
- Visible: H^- b-f continuum with absorption lines
- FIR: H^- f-f (free-free) continuum, increasingly cleaner (i.e. less lines as λ increases, except molecular bands)
- Radio: thermal and, at longer λ, increasingly non-thermal continua

Typical scenario in solar atmosphere
(not to scale)



When are emission, when absorption lines formed?

- Continua are in general formed the deepest in the atmosphere (or at similar heights as the lines)
- Absorption lines are formed when the continuum is strong and the temperature (source function) drops outwards (most effective for high density gas)

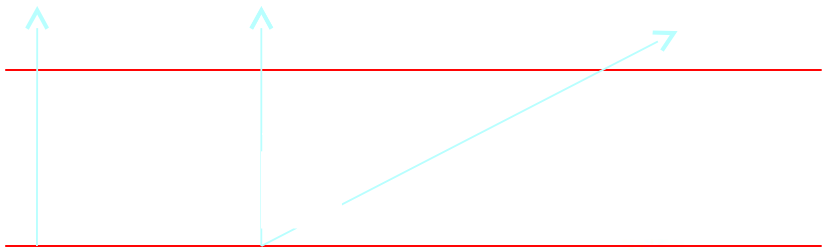
  Photons excite the atom into a higher state. For a high gas density, atoms are de-excited by collisions & the absorbed photon is destroyed

- Emission lines are formed when the temperature (source function) rises outwards and gas density is low:

  If the gas density is low an excited atom decays spontaneously to a lower state, emitting a photon

Radiative Transfer

- If a medium is not optically thin, then radiation interacts with matter and its transfer must be taken into account
- Geometry:



- Equation of radiative transfer: $\mu \frac{dI_\nu}{d\tau_\nu} = I_\nu(\tau_\nu) - S_\nu(\tau_\nu)$

  where I_ν is intensity and S_ν is the source function. $S_\nu = \varepsilon_\nu/\kappa_\nu$. Here emissivity ε_ν, abs. coefficient κ_ν . $\mu = \cos\theta$, where θ = angle of line-of-sight to surface normal
- The physics is hidden in ε_ν and κ_ν, i.e. in τ_ν and S_ν . They are functions of T, p, ν and elemental abundances
- Simple expression for S_ν valid in solar photosphere: $S_\nu = B_\nu(T)$ = Planck function

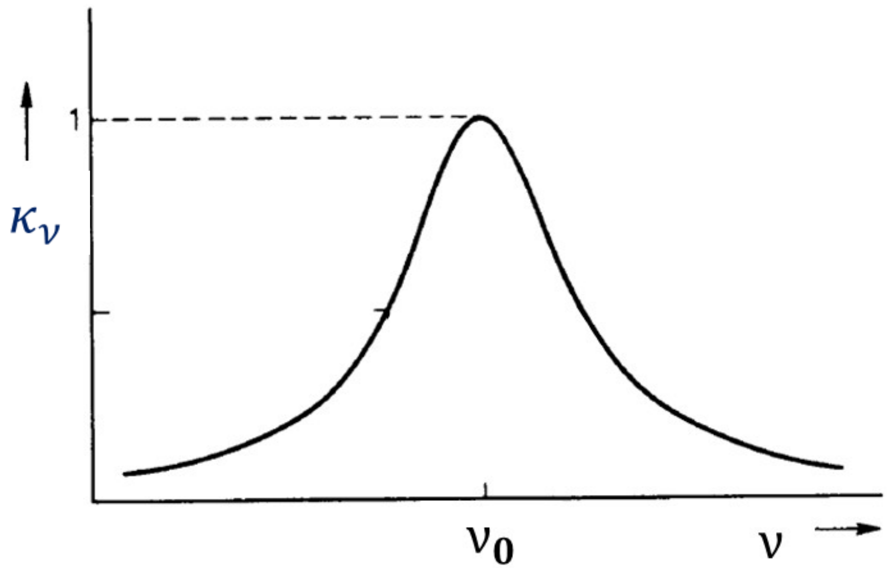
Radiative Transfer in a spectral line

- A spectral line has extra absorption: let κ_ν , κ_C be absorption coefficient in spectral line & in continuum, respectively

- Equation of radiative transfer for the line:

$$\mu \frac{dI_\nu}{d\tau_c} = \frac{\kappa_\nu + \kappa_c}{\kappa_c} I_\nu - S_L = (1 + \eta_\nu) I_\nu - B_\nu$$

where S_L is the line source function, which we assume here to be equal to the Planck function B_ν



Milne-Eddington solution

- A simple analytical solution for a spectral line exists for a Milne-Eddington atmosphere, i.e. $\eta_0 = (\kappa_\nu / \kappa_C)$ independent of τ_C and $\mathbf{S}_L$ depending only linearly on τ_ν: $S_L = B_\nu = \alpha + \beta\tau_C$

$$\frac{I(\mu)}{I_{c(\mu)}} = \frac{1}{\alpha+\beta\mu}\left(\alpha + \frac{\beta\mu}{1+\eta_\nu}\right)$$

$I(\mu)/I_C(\mu)$ = continuum-normalized emergent intensity (residual intensity), where $\eta_\nu = (\kappa_\nu/\kappa_C) = \eta_0\Phi_\nu$ Here η_0 is line center absorption coefficient, Φ_ν is line profile shape

- The term $(1+\eta_\nu)$ takes care of line saturation.

- As line opacity, ~$\eta_{0,}$ increases, I/I_c initially decreases, but for large η_0 saturates around $\alpha/(\alpha+\beta\mu)$

Illustration of line saturation: weak and strong spectral lines

- 4 spectral lines computed using Milne-Eddington assumption
- Different strengths, i.e. different number of absorbing atoms along LOS. Parameterized by $\eta 0$
- As $\eta 0$ increases the line initially becomes deeper, then wider, finally showing prominent line wings

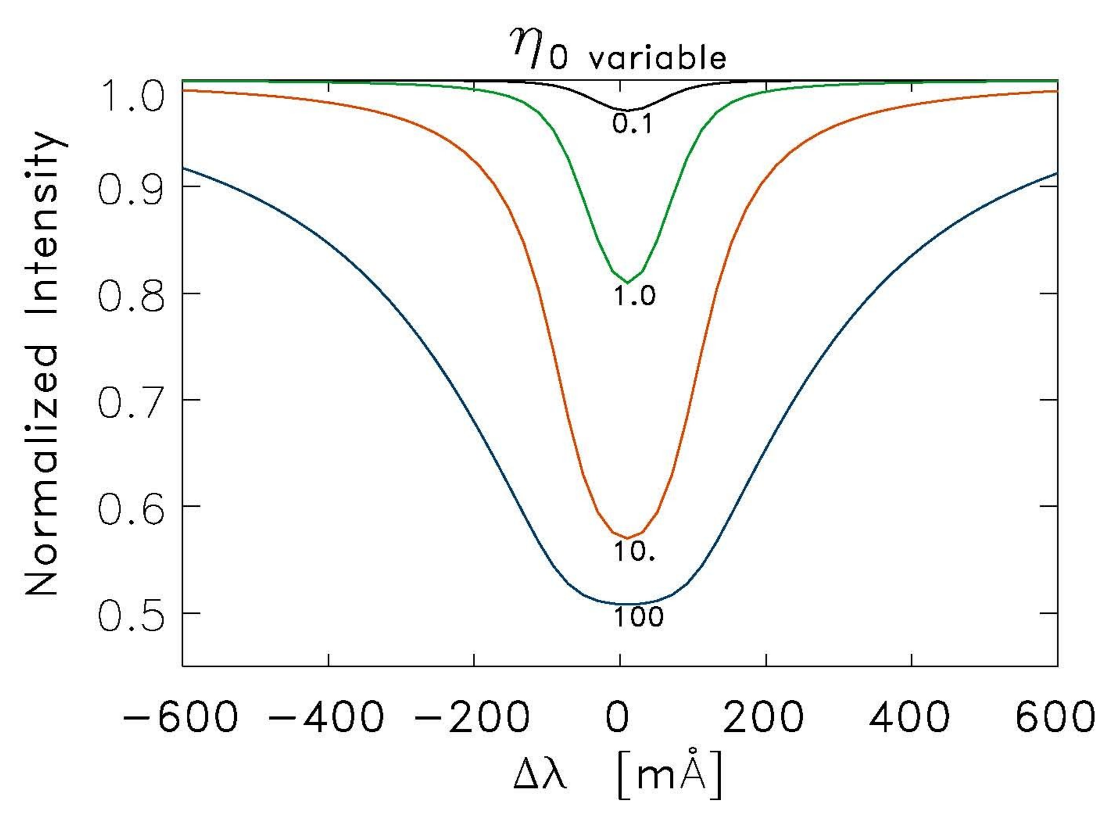
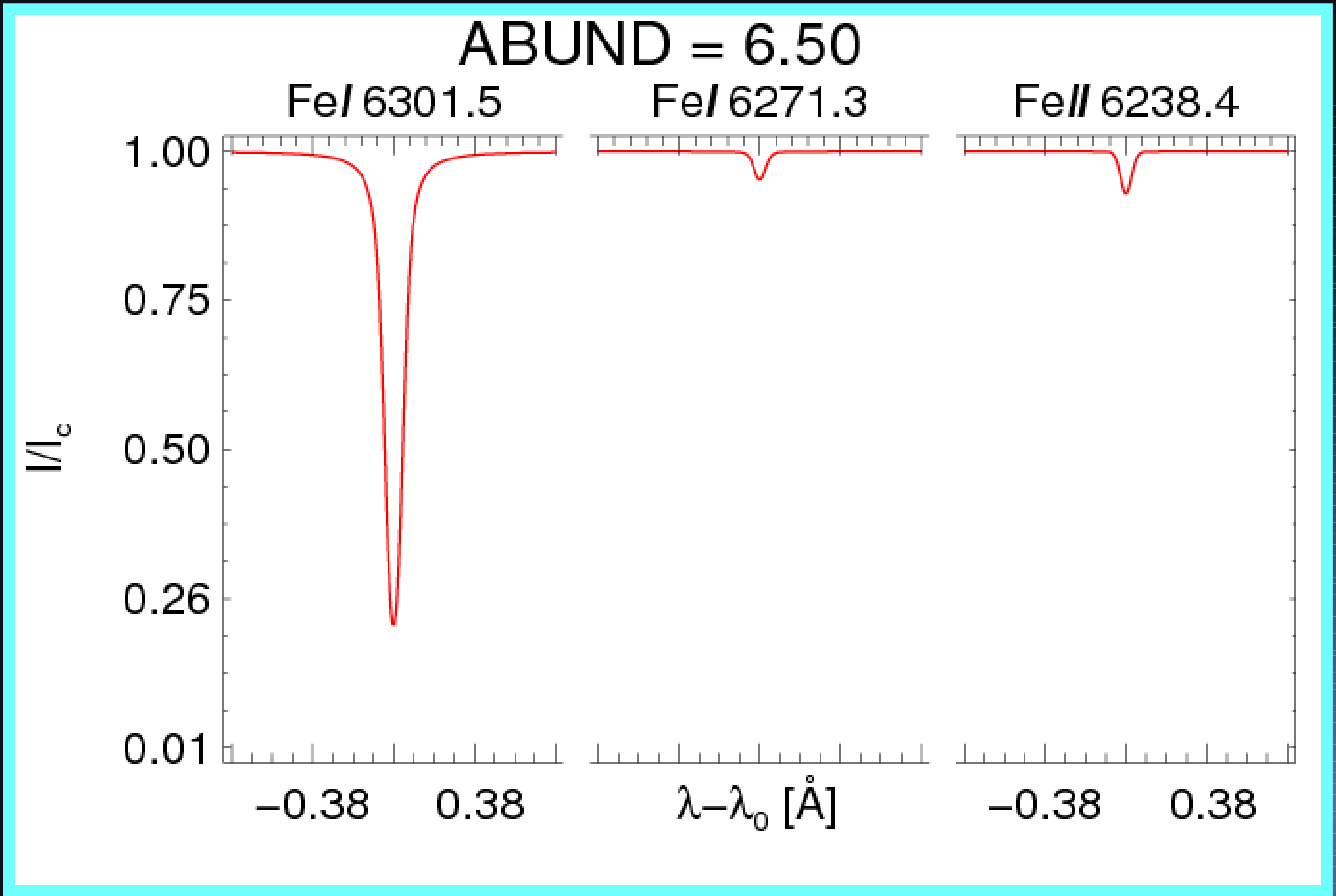


Figure kindly provide by J.M. Borrero

Diagnostic power of spectral lines

- Different parameters describing line strength and line shape contain information on physical parameters of the solar/stellar atmosphere:
  - Doppler shift of line: (net) flows in the LOS direction
  - Line width: temperature and turbulent velocity
  - Area under the line (equivalent width): elemental abundance, temperature (via ionisation and excitation balance)
  - Line depth: temperature (and temperature gradient)
  - Line asymmetry: velocity gradients, v, T inhomogenieties
  - Wings of strong lines: gas pressure
  - Polarisation and splitting: magnetic field

Effect of changing abundance

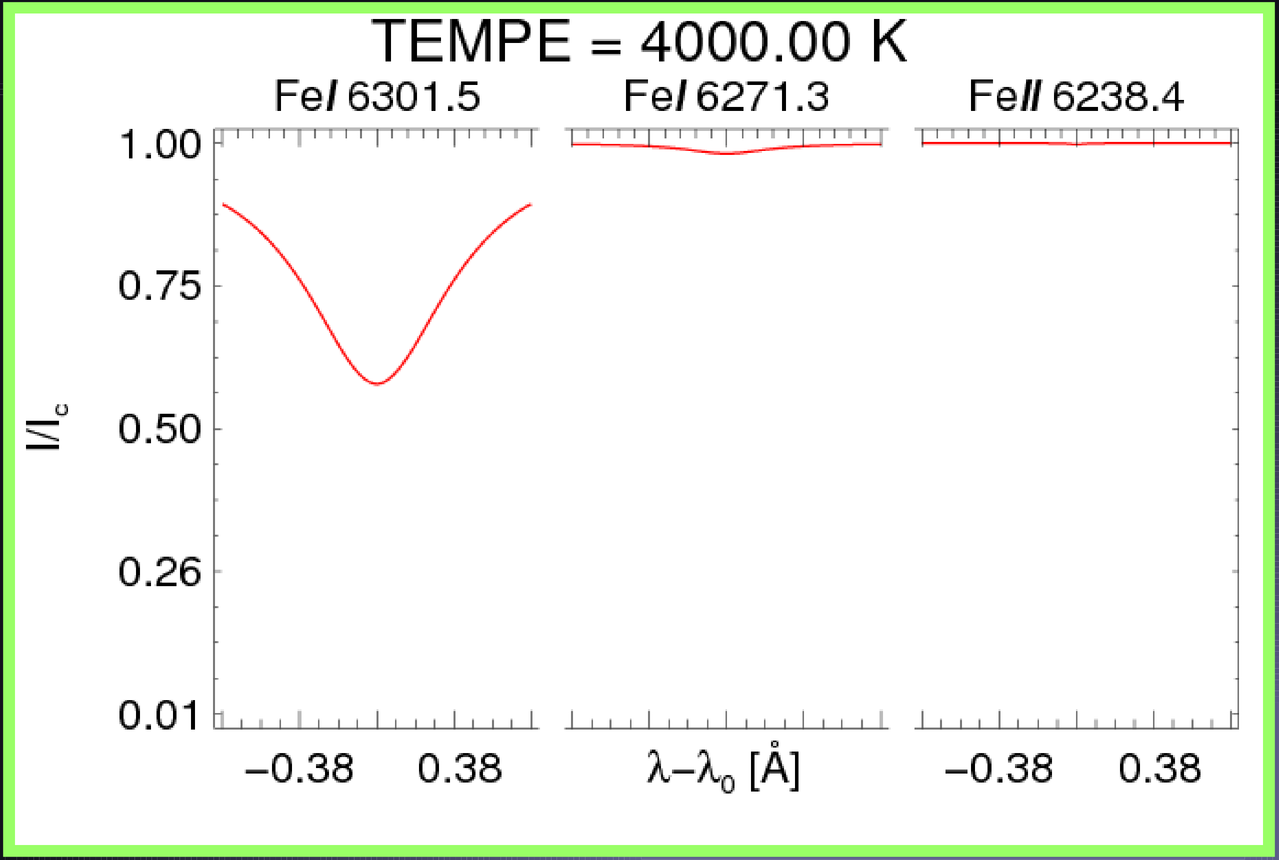


Fe I = spectral line of neutral iron

Fe II = spectral line of singly ionized iron: Fe+

Higher abundance = more absorbing atoms, stronger absorption
Abundance given on logarithmic scale with abundance (H) = 12

Effect of changing temperature on absorption lines

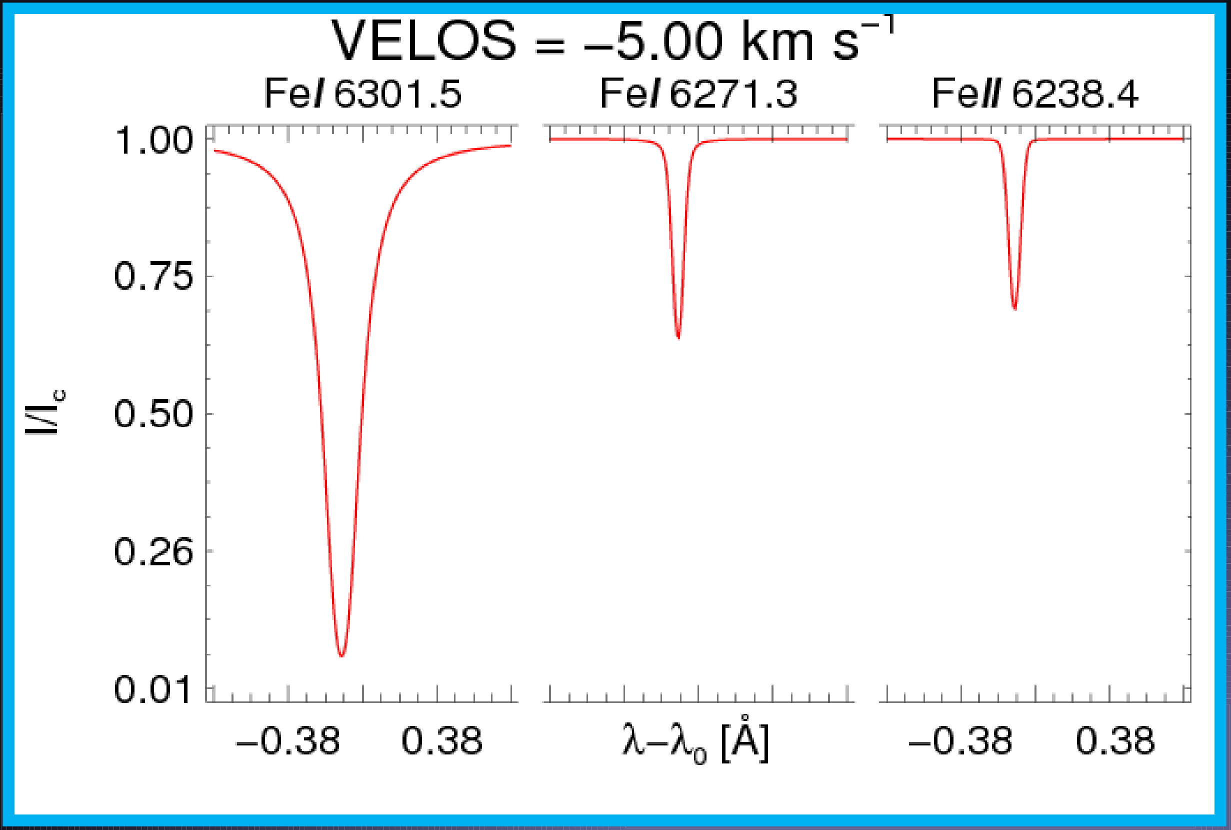


Low temperatures: Fe is mainly neutral strong Fe I, weak Fe II lines

High temperatures: Fe gets ionized Fe II strengthens relative to Fe I

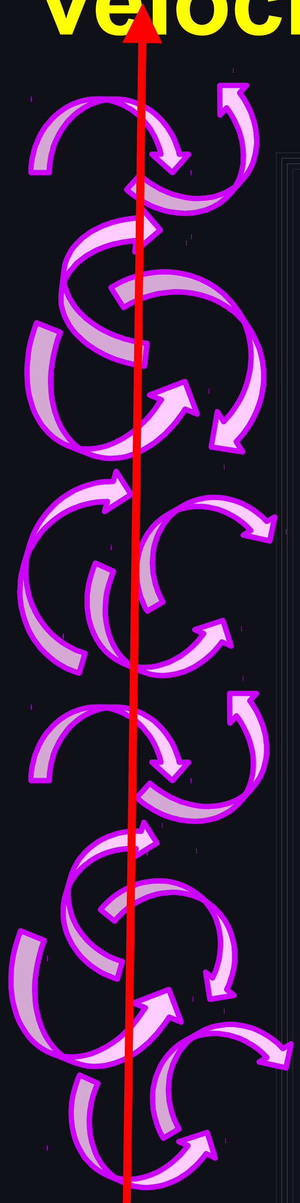
Very high temperatures: Fe gets doubly ionized Fe II also gets weak

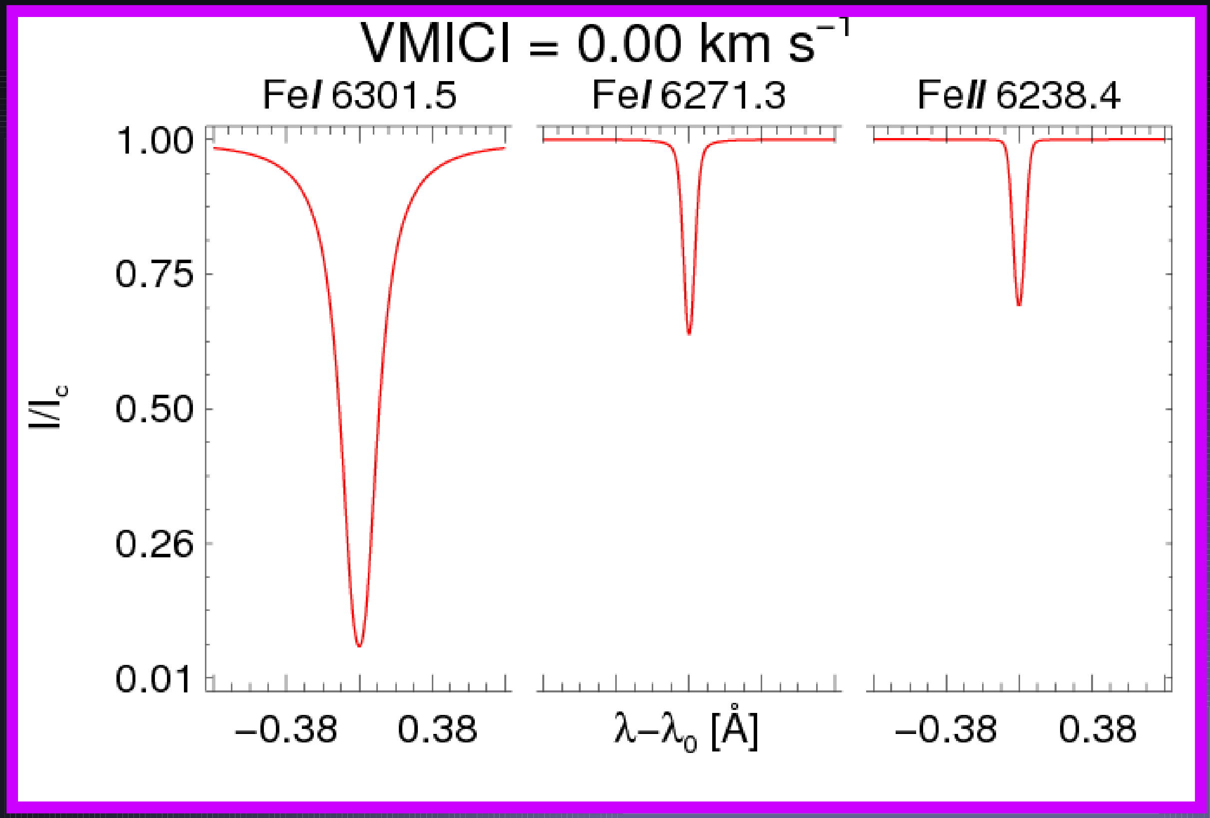
Effect of changing line-of-sight velocity on absorption lines

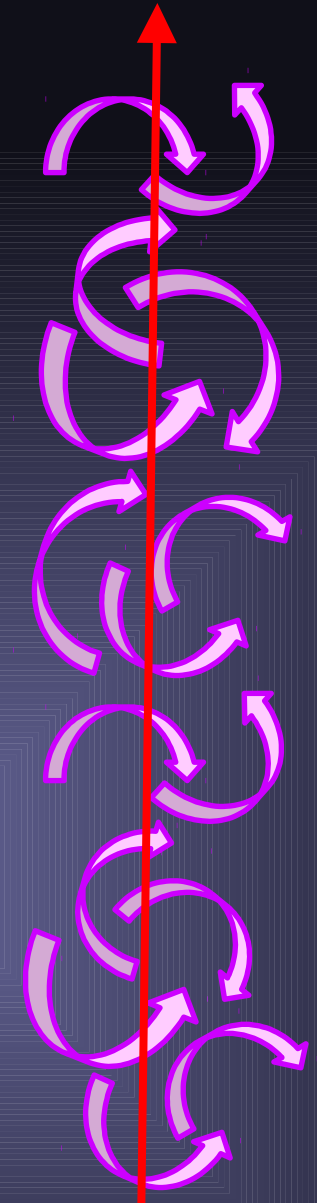


Shift in line profiles due to Doppler effect
Same magnitude of shift for all lines

Effect of changing microturbulence velocity on absorption lines







In a turbulent medium the sum of all Doppler shifts leads to a line broadening