Lecture 24

AST5770

Solar and stellar physics

Sven Wedemeyer, University of Oslo, 2023

Assignments

Data Analysis and results

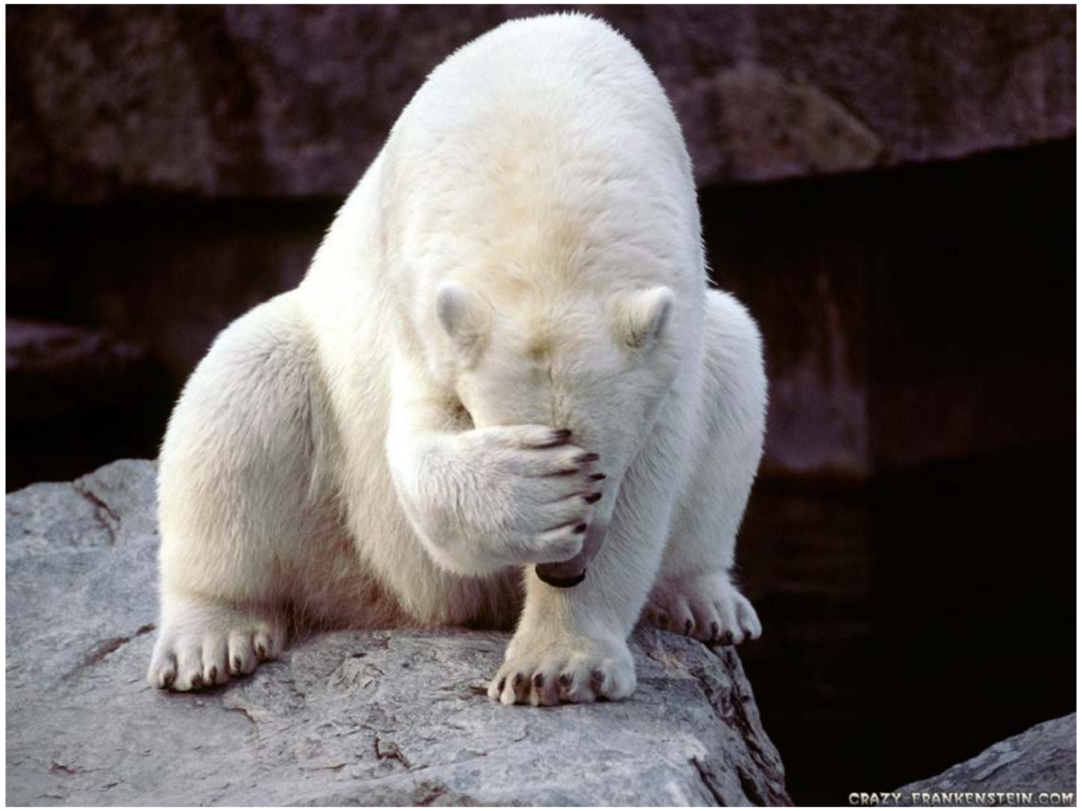


Figure 1. This an animal picture. This colour image shows some animal. Dark colours are background. Some bonus words to reach the word limit.

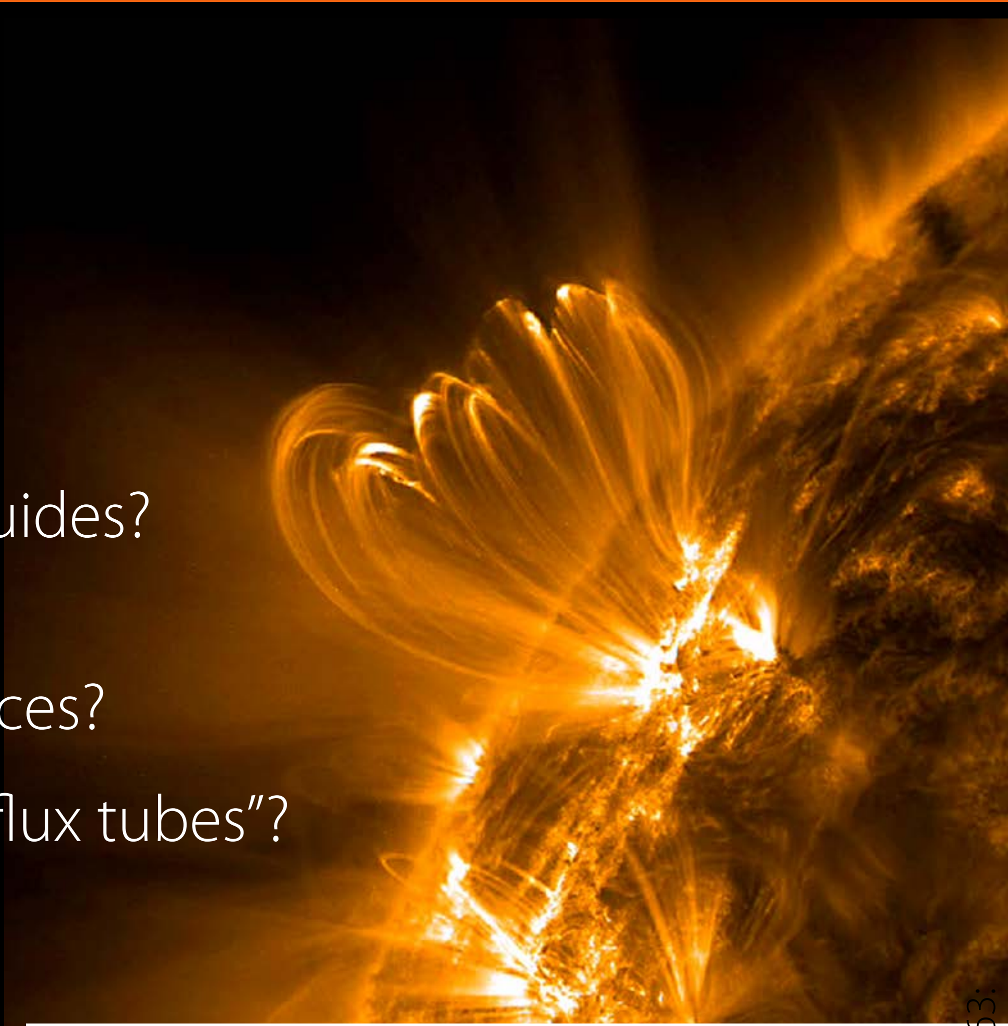
Figure caption:
Describe what is presented.
Describe **all** graphical elements.

Description of figure in main text:
Do not repeat all graphical elements but describe what the figure shows.
What is interesting here?
Do address **details** and provide **numbers**.
Connect it to the analysis before and after to create a “red thread”.

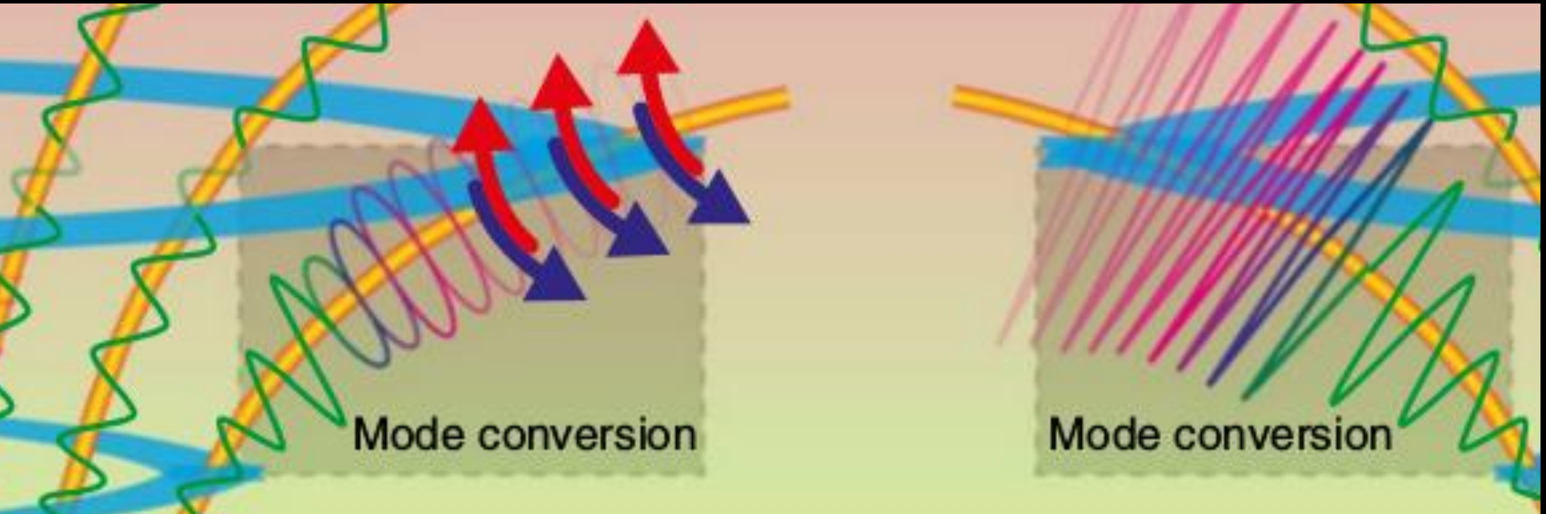
Analysis
Try out different things and go deeper.
Ask questions about your results/figures.
What more could be interesting?

The solar atmosphere

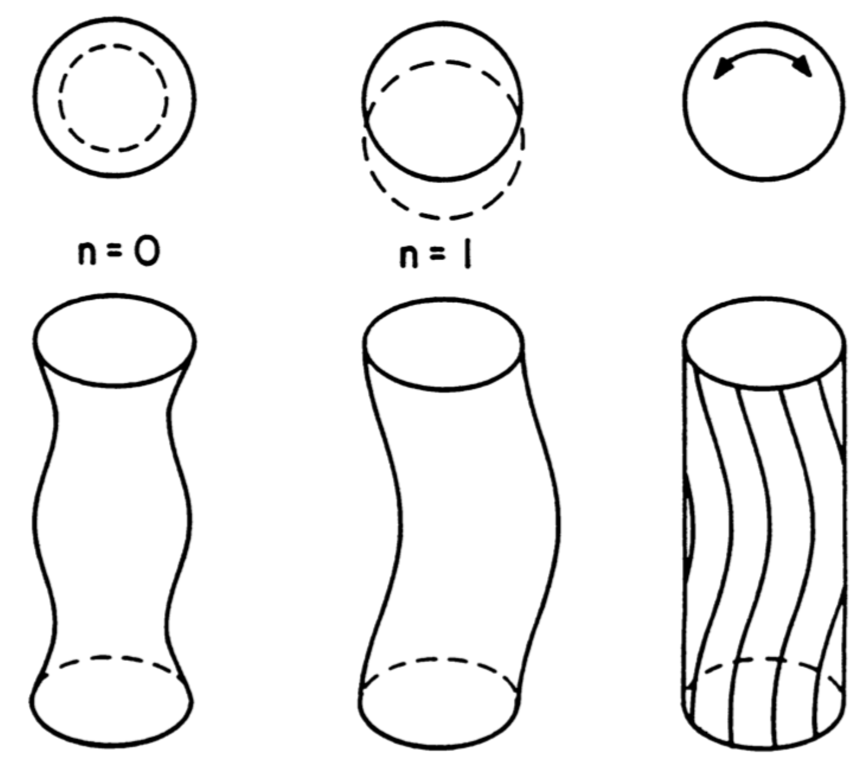
Recap



1. What are spicules? How long do they persist and how often do they occur?
2. Where are they best observed?
3. Which structures in the solar atmosphere serve as wave guides?
4. Which types of wave modes do occur in the solar atmosphere and what are the corresponding restoring forces?
5. What kind of periodic behaviour do we see for magnetic “flux tubes”?
6. What is a typical temperature in the corona?
7. What is coronal rain?



Grant et al. 2018

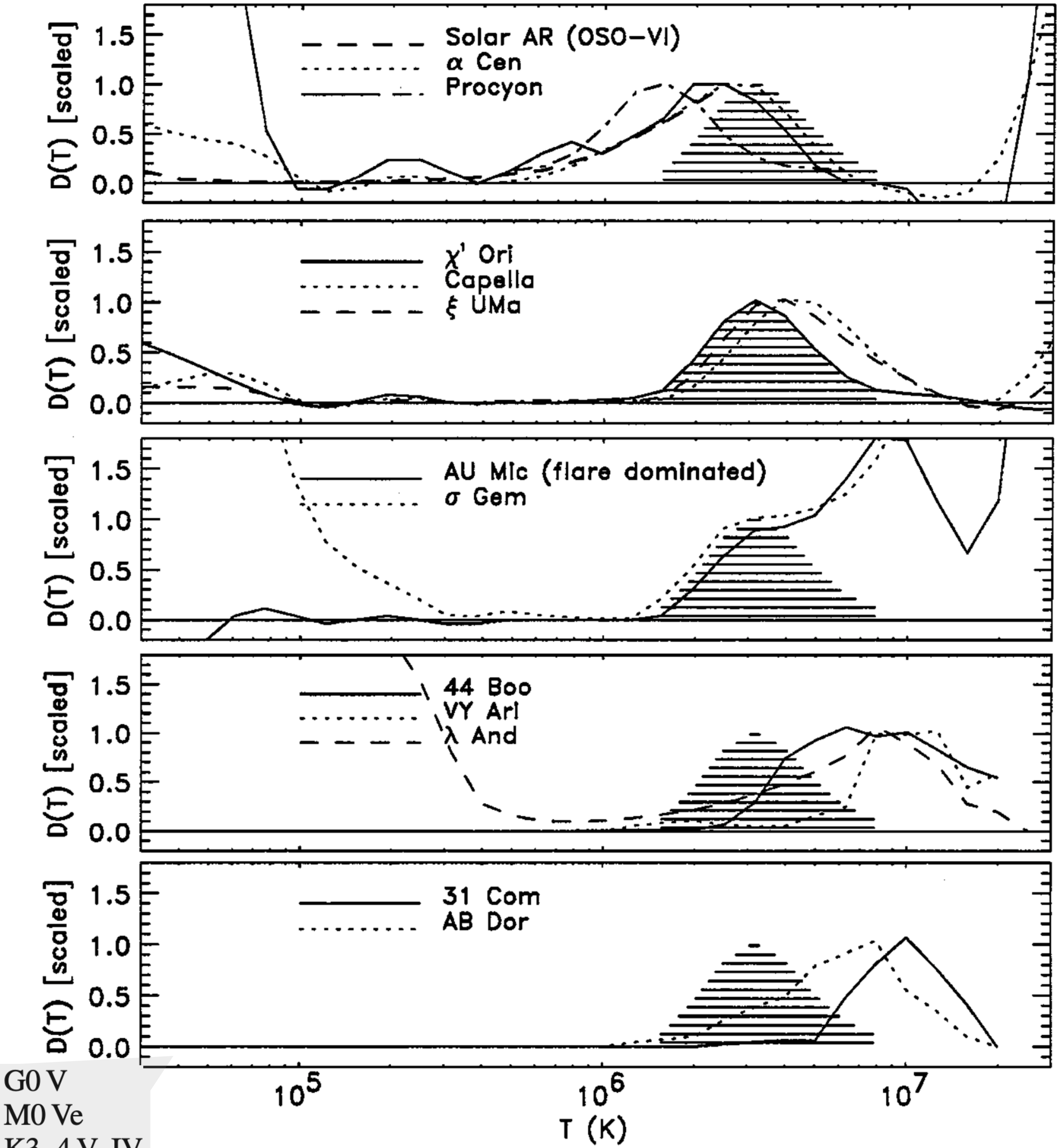


Solanki, S. K., 1993, Space Science Reviews 63:

Stellar coronae

Observations

- Spectral line observations of other stars (incl. differential emission measure) compared to AR/Sun
- Clear indications for existence of coronae for late-type stars (spectral type F and cooler)
- Also observed emission at shorter wavelengths (EUV, X-rays) and radio wavelengths consistent with coronal emission expected for this type of stars



Dashed region = χ^1 Ori, repeated for reference

α Cen	G2 V + K2 V	Procyon	F5 IV–V	χ^1 Ori	G0 V
Capella	G5 III+G0 III	ξ Uma B	G0 V	AU Mic	M0 Ve
σ Gem	K1 III	44 Boo	G0 V	VY Ari	K3–4 V–IV
λ And	G8 IV–III	31 Com	G0 III	AB Dor	K0 V*

Schrijver et al (1996), S; Data from Mewe et al (1995), Schrijver et al (1995), Schmitt et al (1996), Dere(1982)

Stellar coronae

Observations

- Plasma properties can be derived, e.g. electron densities from EUV observations
- Note: Stellar observations are spatially unresolved

  ➡ Average properties only
- Caution: How are these properties affected by unresolved variations?!?

Just as an example…

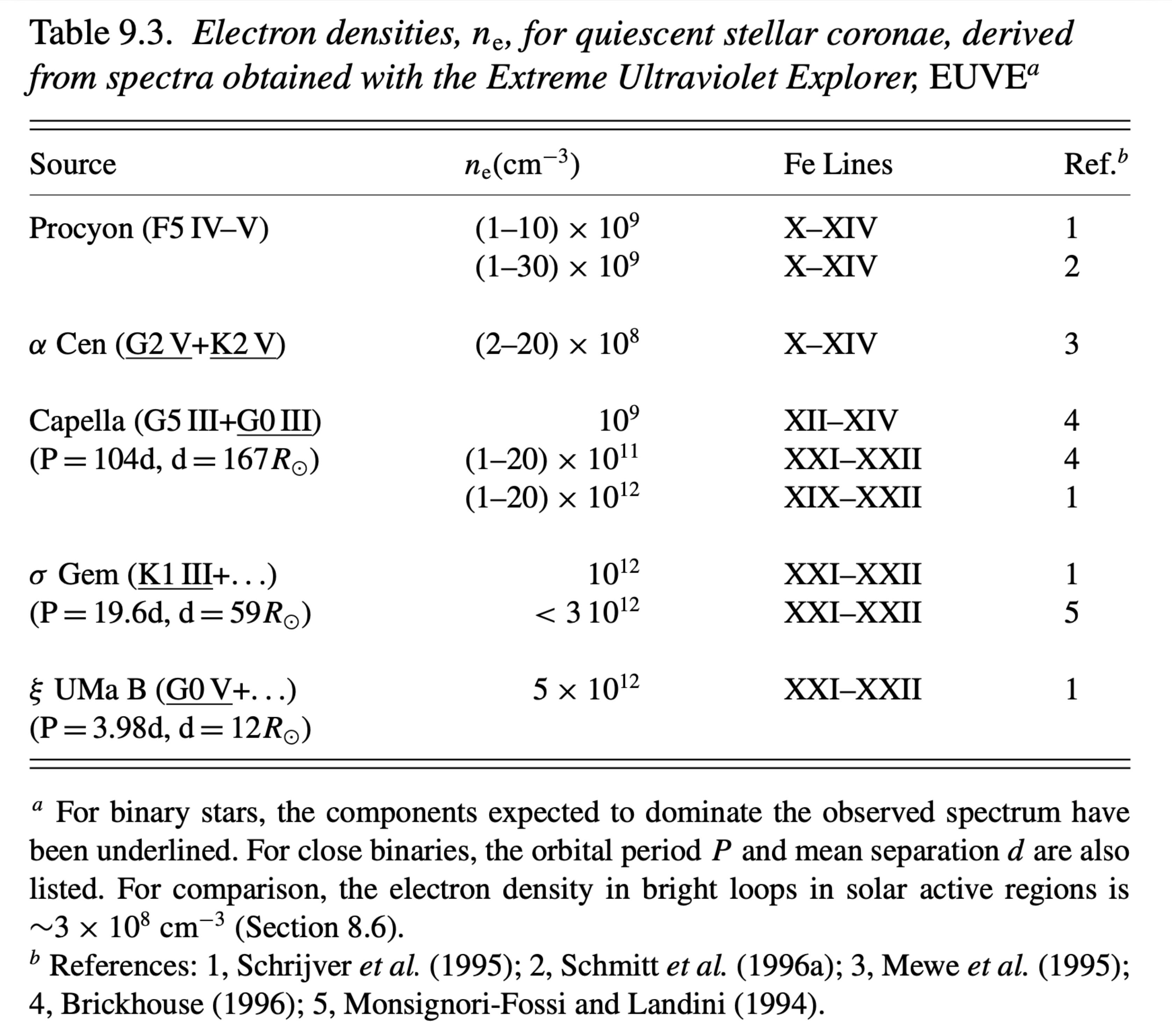


Table 9.3. *Electron densities, n_e, for quiescent stellar coronae, derived from spectra obtained with the Extreme Ultraviolet Explorer,* EUVE[a]

Source	$n_e(\mathrm{cm}^{-3})$	Fe Lines	Ref.[b]
Procyon (F5 IV–V)	$(1\text{–}10) \times 10^9$	X–XIV	1
	$(1\text{–}30) \times 10^9$	X–XIV	2
α Cen (G2 V+K2 V)	$(2\text{–}20) \times 10^8$	X–XIV	3
Capella (G5 III+G0 III)	10^9	XII–XIV	4
(P = 104d, d = 167 $R_\odot$)	$(1\text{–}20) \times 10^{11}$	XXI–XXII	4
	$(1\text{–}20) \times 10^{12}$	XIX–XXII	1
σ Gem (K1 III+…)	10^{12}	XXI–XXII	1
(P = 19.6d, d = 59 $R_\odot$)	$< 3\,10^{12}$	XXI–XXII	5
ξ UMa B (G0 V+…)	5×10^{12}	XXI–XXII	1
(P = 3.98d, d = 12 $R_\odot$)			

[a] For binary stars, the components expected to dominate the observed spectrum have been underlined. For close binaries, the orbital period *P* and mean separation *d* are also listed. For comparison, the electron density in bright loops in solar active regions is $\sim 3 \times 10^8\ \mathrm{cm}^{-3}$ (Section 8.6).

[b] References: 1, Schrijver *et al.* (1995); 2, Schmitt *et al.* (1996a); 3, Mewe *et al.* (1995); 4, Brickhouse (1996); 5, Monsignori-Fossi and Landini (1994).

Chromospheric / coronal heating

Chromospheric / coronal heating

- Observations of spectral lines for high ionisation stages (Grotrian 1939, Edlén 1940)

➡ Implies gas temperature T > 1 000 000 K in the outer layers of the Sun.

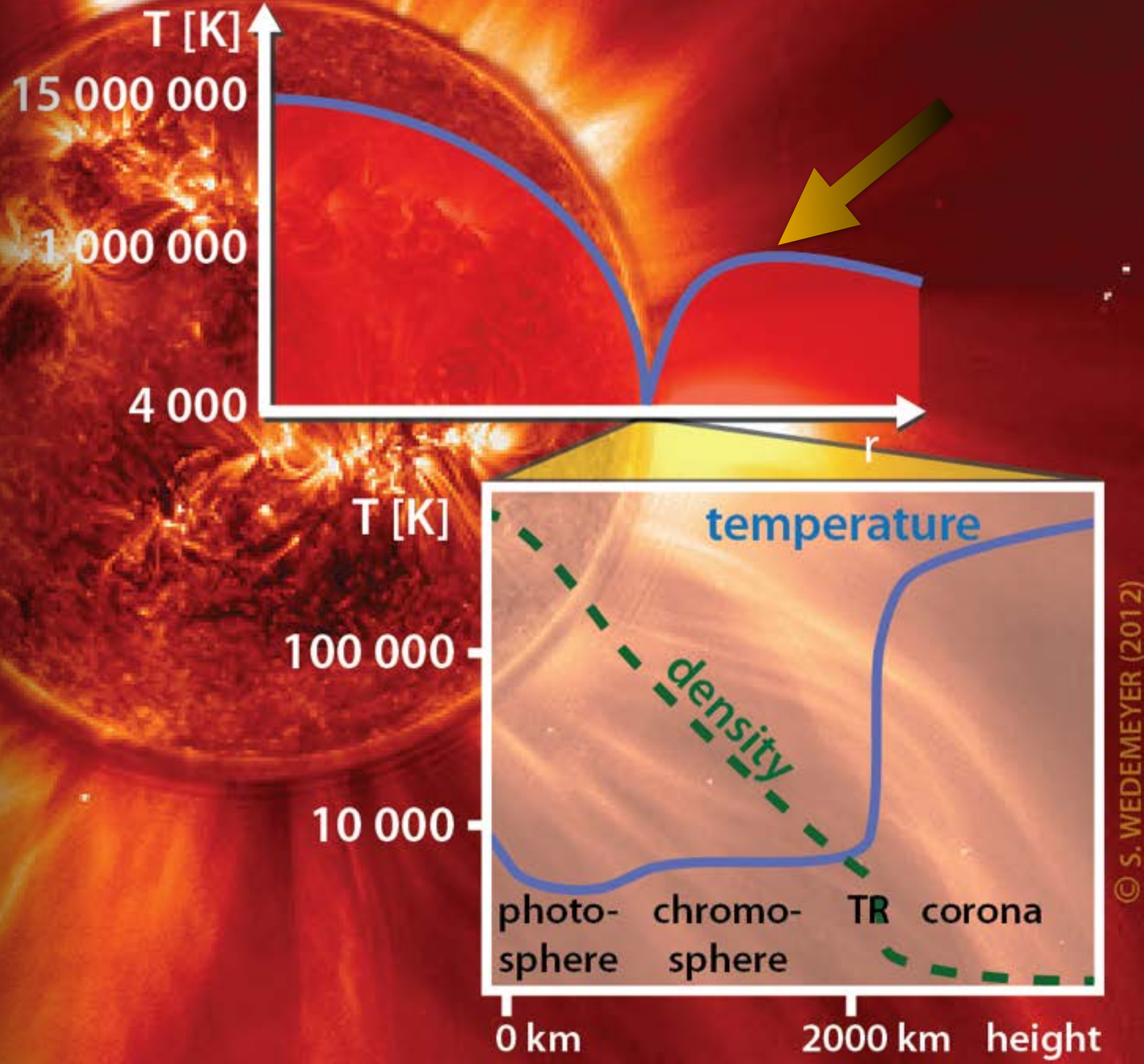
- A long-standing central problem in solar / stellar astrophysics (known and unsolved for > 80 yr)
- Temperatures far above radiative equilibrium values

➡ **Outer layers heated!**

- But how?!

➡ **"Coronal heating problem"**

- Applies to (solar-like) stars in general



Chromospheric / coronal heating

CURRENT STATUS OF THE FIELD

- Many processes known that can potentially provide enough heating.
- Many attempts but no comprehensive solution yet!
- Major problem:
  - Available diagnostics only indirect and/or difficult to interpret.

  ➡ Remaining uncertainties when deriving state of atmospheric gas

THE STILL OPEN KEY QUESTIONS

1. Which are the most important physical processes?
2. How much do the individual processes contribute to the heating?

Chromospheric / coronal heating

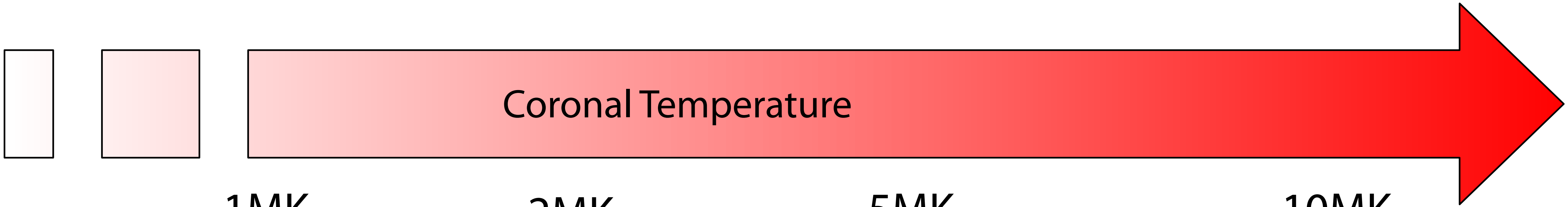
Multi-temperature structure of active regions

- Large difference in temperature between different parts of the corona (Quiet Sun, Active Regions, Coronal Holes — Flares)

➡Required heating input to compensate for radiative losses is one order of magnitude different

  - Hot (T > 2MK) $\sim 10^7$ erg cm-2 s-1
  - Cool (T ~ 1MK) $\sim 10^6$ erg cm-2 s-1

- “One size won’t fit all”? Scalable heating process? Or different processes dominating / contributing differently in different situations/regions (with different temperature)?



“Cool”

Hot

Stationary/continuous heating

Transient heating (localised)

Many candidates

Not understood yet.

Magnetic reconnection
(flares, microflares)
Required due to the large amount of released energy

Chromospheric / coronal heating

Candidates for continuous heating mechanims

- Previously thought: **acoustic heating** accounts for "basal flux" (remember stellar activity!), whereas processes involving magnetic fields would add varying contribution on top.

- Two major categories — AC-DC:
  - **AC — MHD waves**: High frequency slow MHD waves, fast waves, Alfvén(ic) waves; combined with linear mechanisms (e.g., mode conversion, resonant absorption and phase mixing) or nonlinear mechanisms (e.g., dynamic instabilities and turbulence, wave-to-wave interaction, parametric decay)
  - **DC — Small-scale magnetic reconnection**: Frequent nanoflares outside active regions (continuous non-thermal heating contributions); stress-induced, field line braiding

- Other possible contributors:
  - Gravity waves
  - Multi-fluid effects
  - Plasma instabilities
  - …

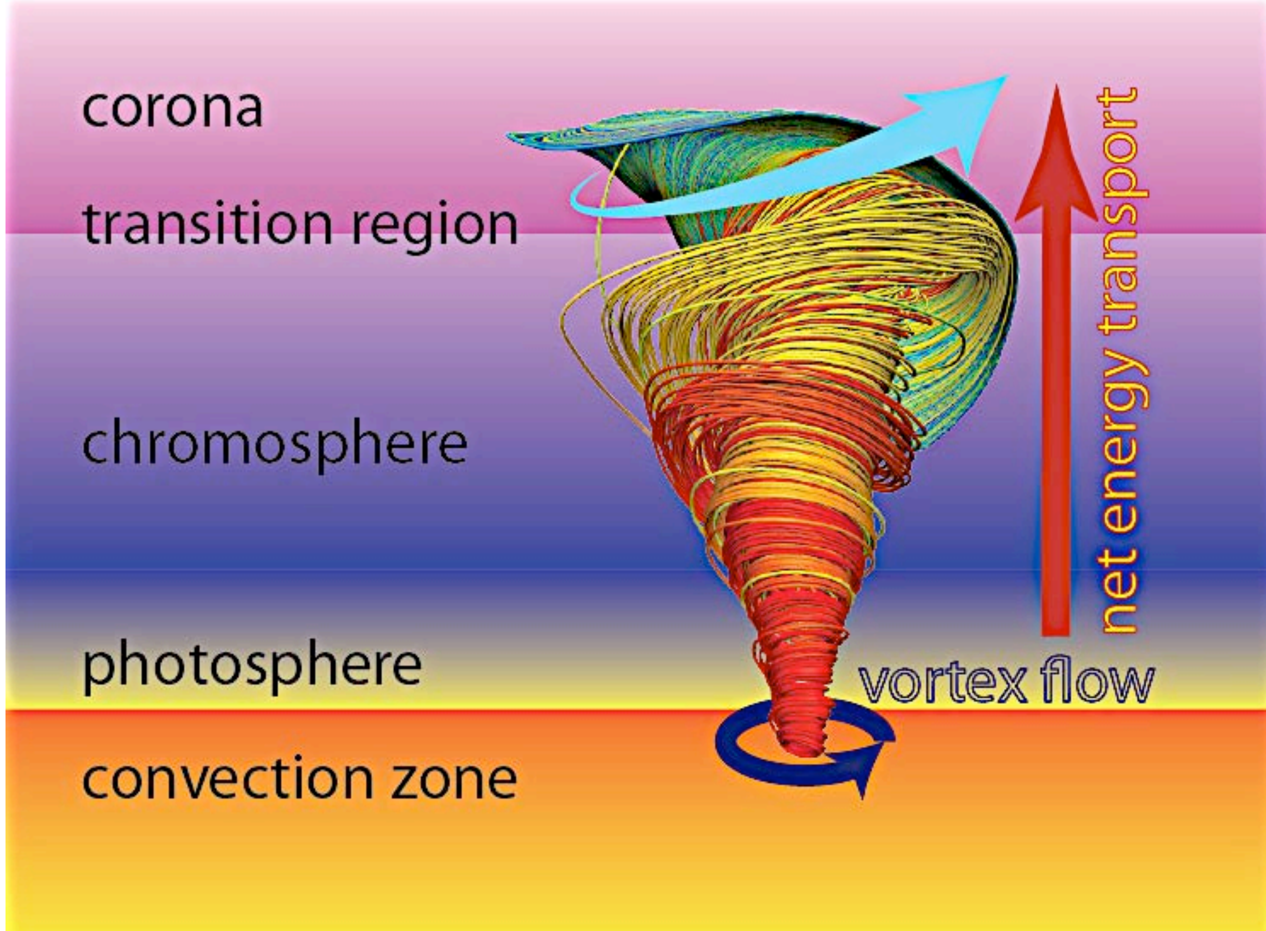
"Updated coronal heating problem"

- Many known candidates that can provide more than enough energy in chromosphere/corona (alone or combined)
- Question(s) now:
  - How is the energy **dissipated**?!
  - Which process is contributing how much in different types of region?

Chromospheric / coronal heating

Candidates for continuous heating mechanims

- **Hot candidate: Alfvén waves**
  - Once generated, Alfvén waves propagate easily along the magnetic field structures into upper atmosphere
  - Just dissipation not that easy and not sufficiently understood yet.
  - Wave guides:
    - Loops
    - Spicules
    - Magnetic tornadoes
    - …

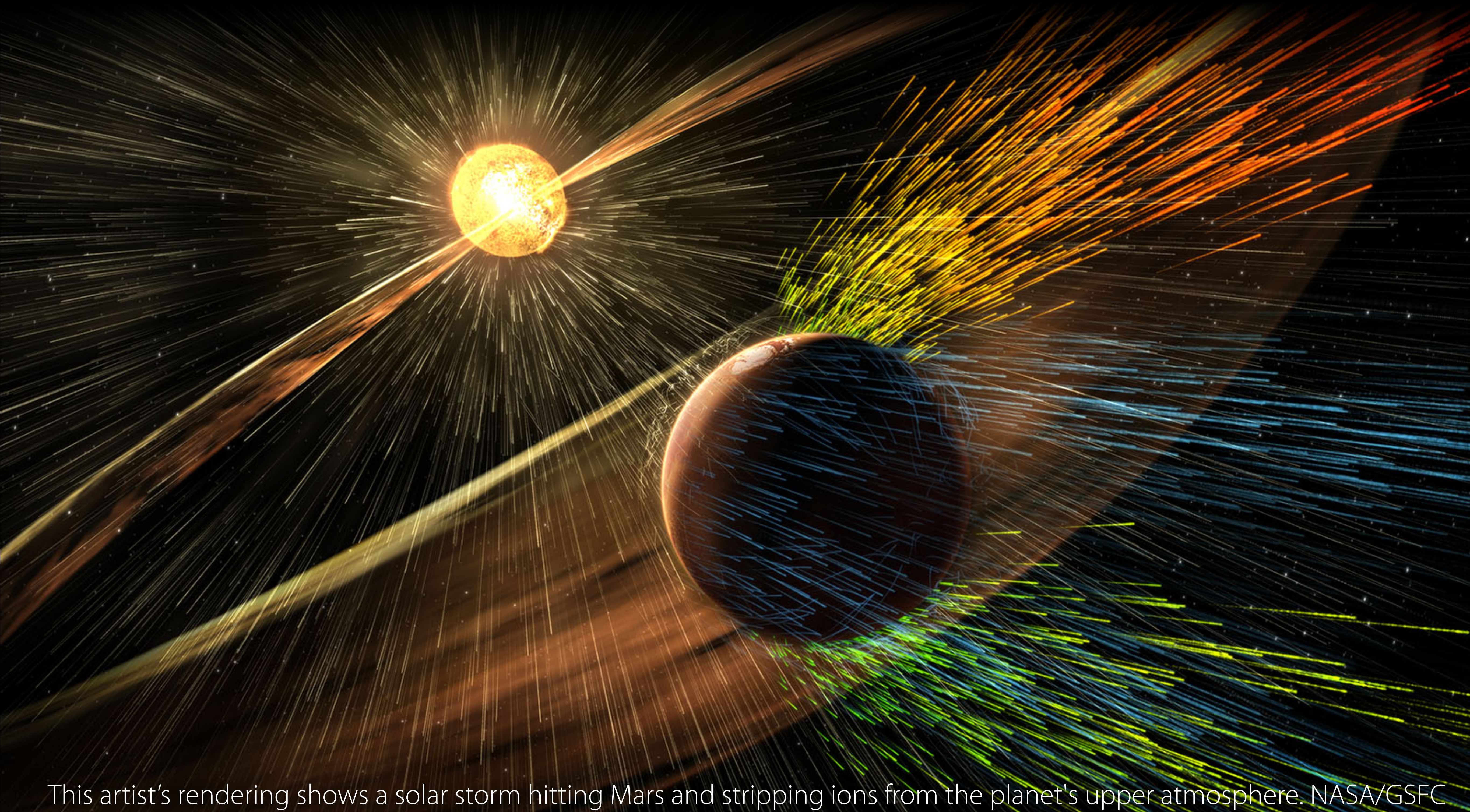


Solar Wind and Heliosphere

Stellar rotation

Solar/stellar winds

- Stream of charged particles released from the upper atmospheric layers of a star



This artist’s rendering shows a solar storm hitting Mars and stripping ions from the planet's upper atmosphere. NASA/GSFC

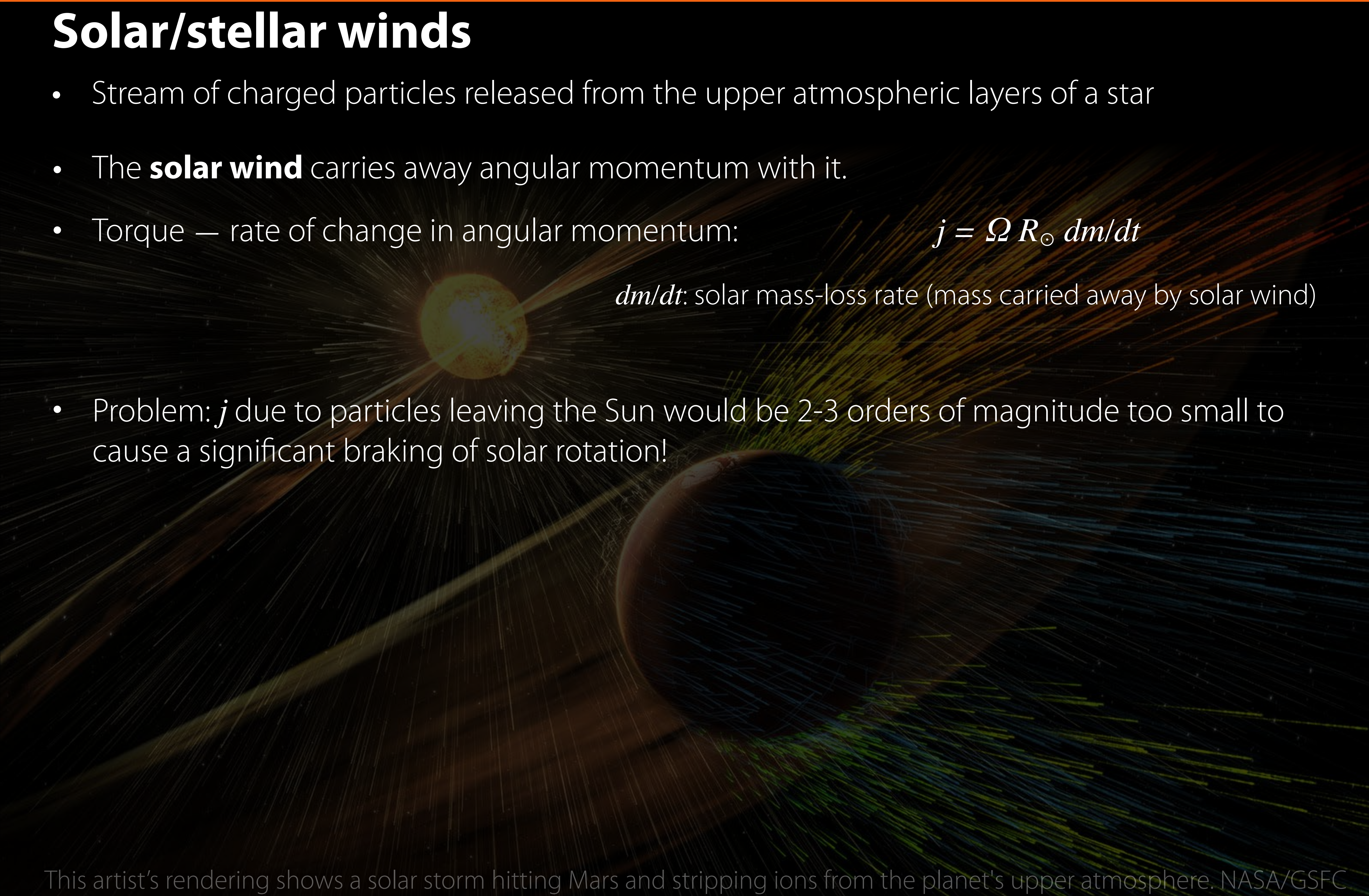
Stellar rotation

Solar/stellar winds

- Stream of charged particles released from the upper atmospheric layers of a star
- The **solar wind** carries away angular momentum with it.
- Torque — rate of change in angular momentum: $j = \Omega R_\odot \, dm/dt$

dm/dt: solar mass-loss rate (mass carried away by solar wind)

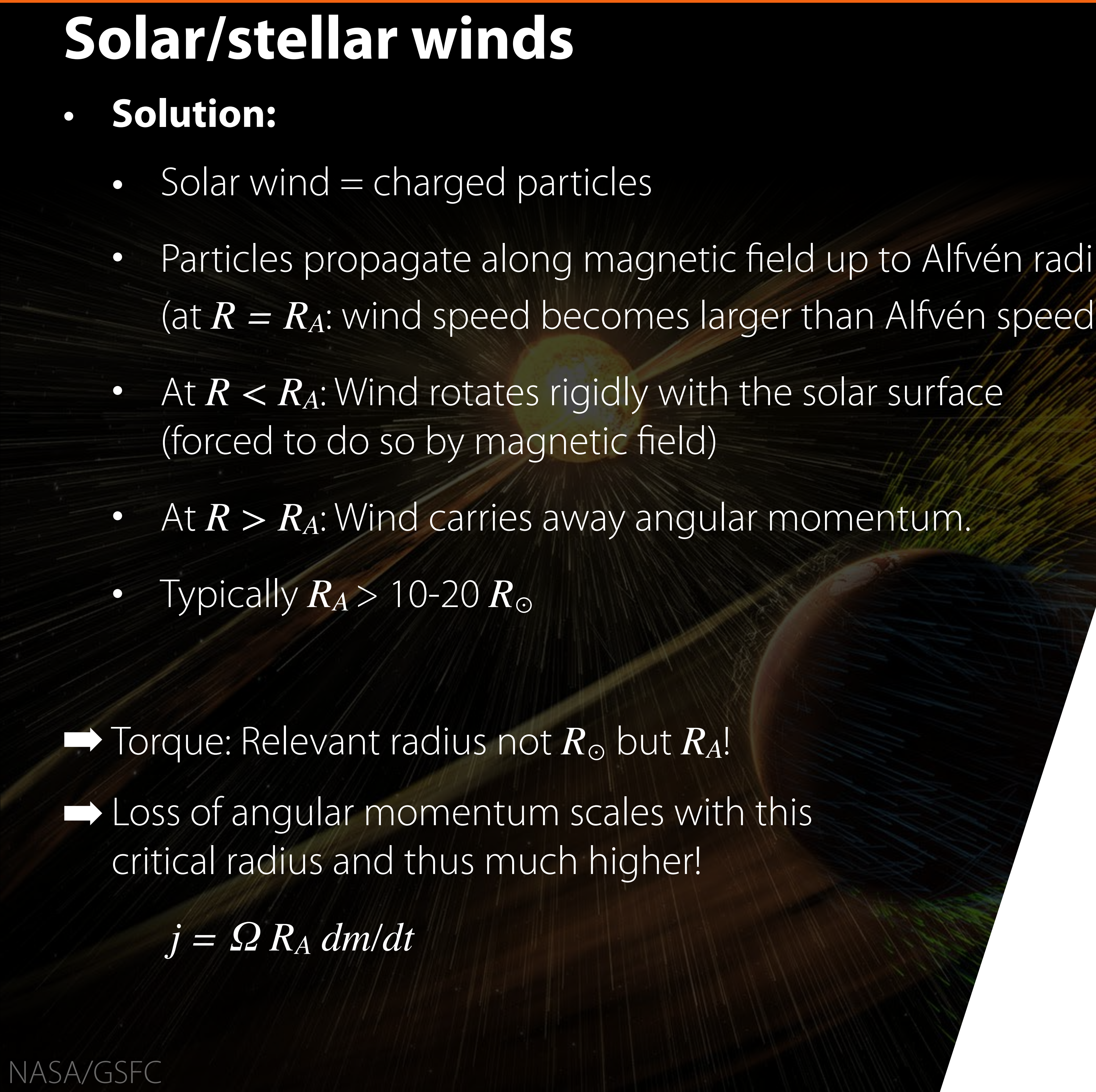
- Problem: j due to particles leaving the Sun would be 2-3 orders of magnitude too small to cause a significant braking of solar rotation!



This artist's rendering shows a solar storm hitting Mars and stripping ions from the planet's upper atmosphere. NASA/GSFC

Stellar rotation

Solar/stellar winds



- **Solution:**
  - Solar wind = charged particles
  - Particles propagate along magnetic field up to Alfvén radius R_A (at $R = R_A$: wind speed becomes larger than Alfvén speed)
  - At $R < R_A$: Wind rotates rigidly with the solar surface (forced to do so by magnetic field)
  - At $R > R_A$: Wind carries away angular momentum.
  - Typically $R_A >$ 10-20 $R_\odot$

➡ Torque: Relevant radius not $R_\odot$ but R_A!

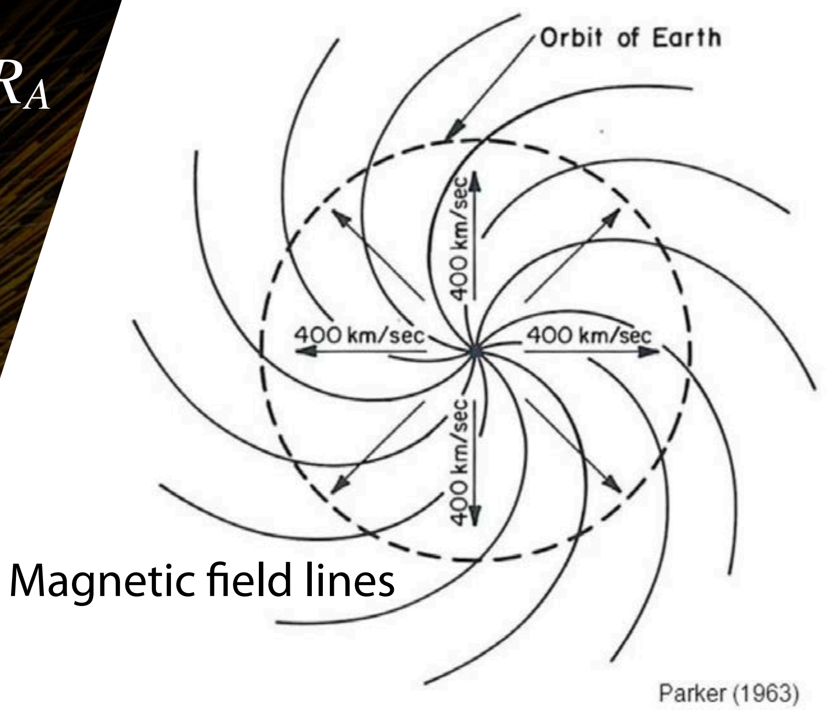
➡ Loss of angular momentum scales with this critical radius and thus much higher!

$$j = \Omega R_A \, dm/dt$$

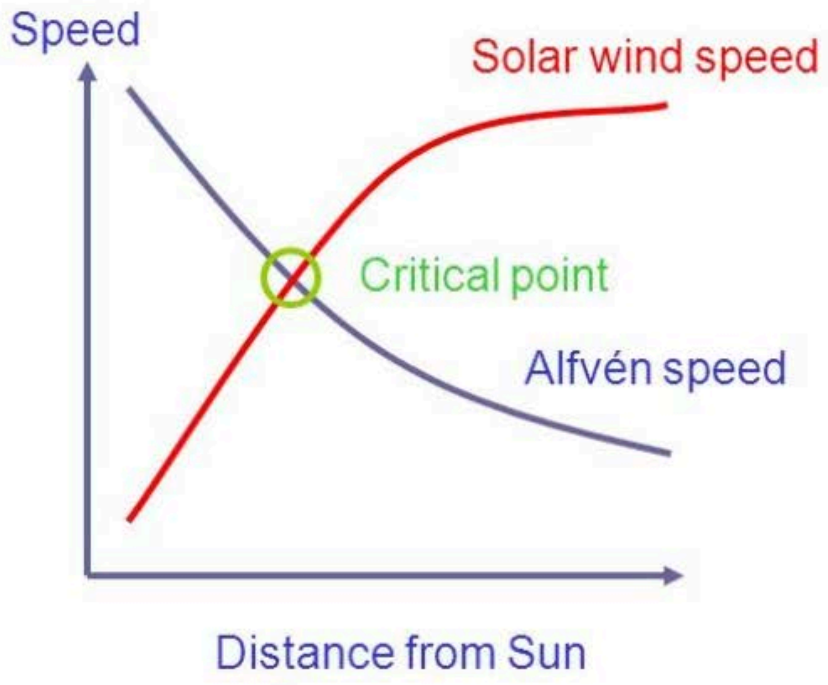
NASA/GSFC

Alfvén speed

$$v_A \equiv \frac{B}{\sqrt{\mu_0 \rho}}$$



Parker (1963)

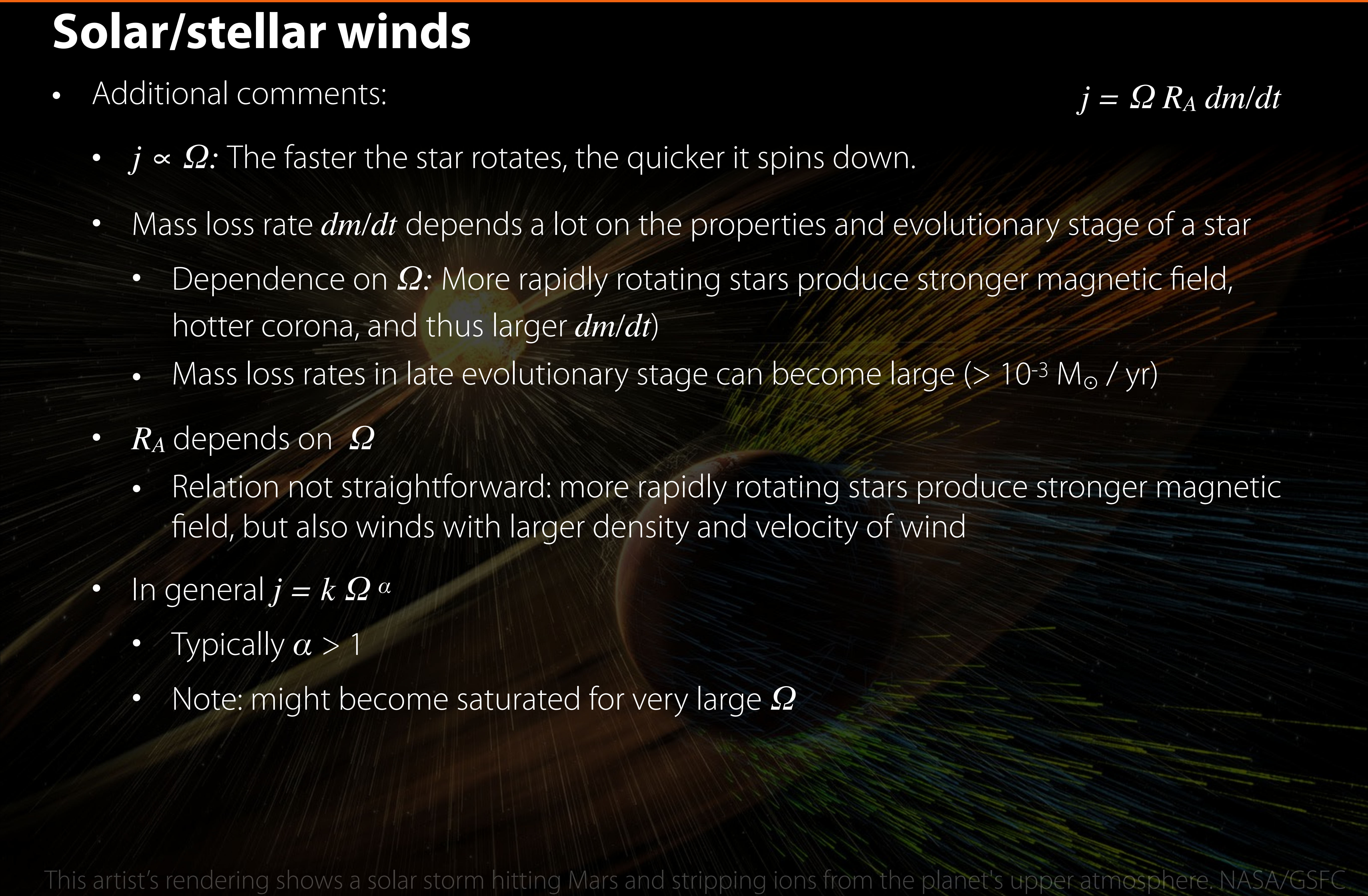


Stellar rotation

Solar/stellar winds

$$j = \Omega R_A\, dm/dt$$

- Additional comments:
  - $j \propto \Omega$: The faster the star rotates, the quicker it spins down.
  - Mass loss rate dm/dt depends a lot on the properties and evolutionary stage of a star
    - Dependence on Ω: More rapidly rotating stars produce stronger magnetic field, hotter corona, and thus larger dm/dt)
    - Mass loss rates in late evolutionary stage can become large ($> 10^{-3}\ M_\odot$ / yr)
  - R_A depends on Ω
    - Relation not straightforward: more rapidly rotating stars produce stronger magnetic field, but also winds with larger density and velocity of wind
  - In general $j = k\, \Omega^{\alpha}$
    - Typically $\alpha > 1$
    - Note: might become saturated for very large Ω

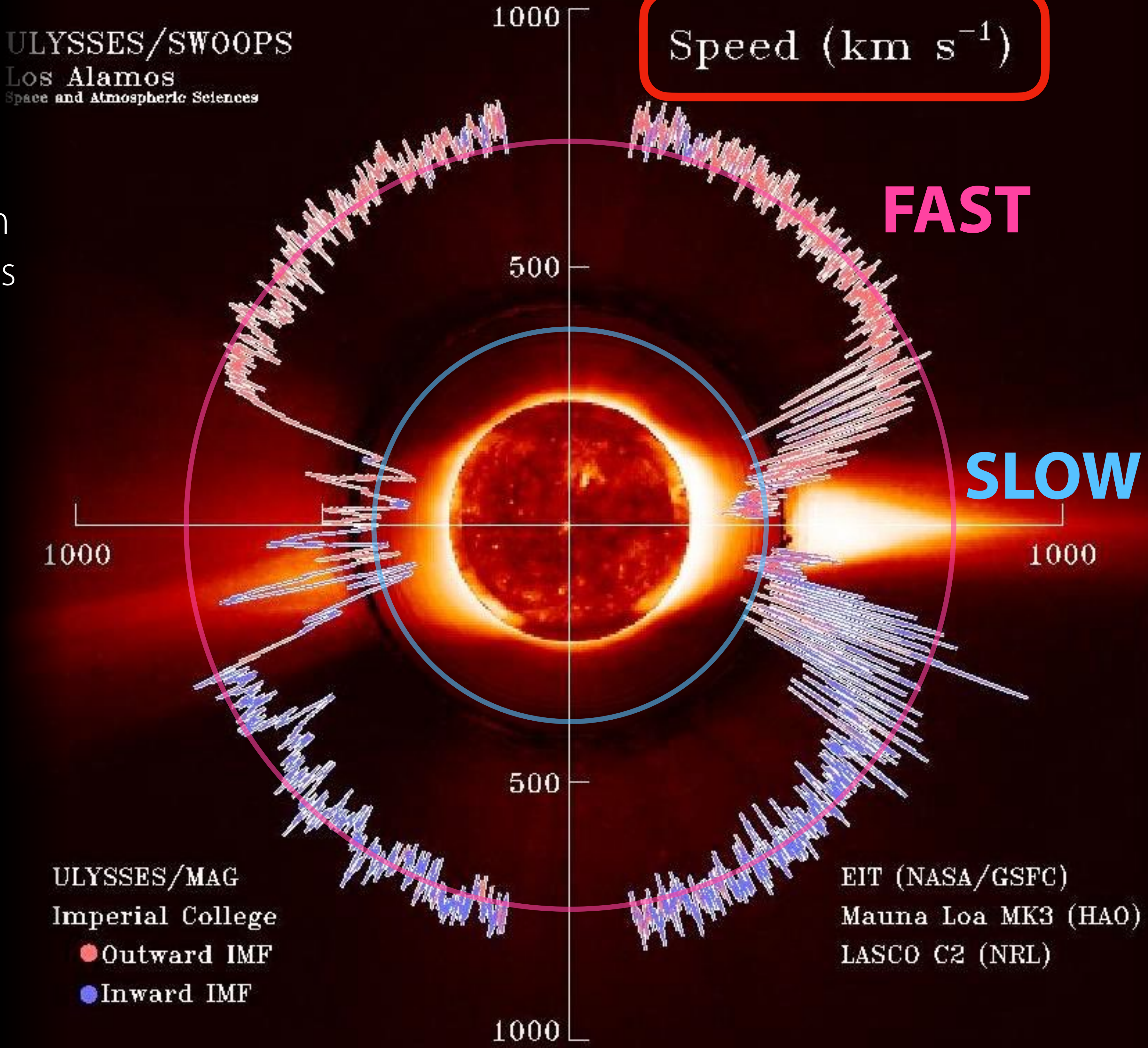


This artist's rendering shows a solar storm hitting Mars and stripping ions from the planet's upper atmosphere. NASA/GSFC

Solar Wind and Heliosphere

Components

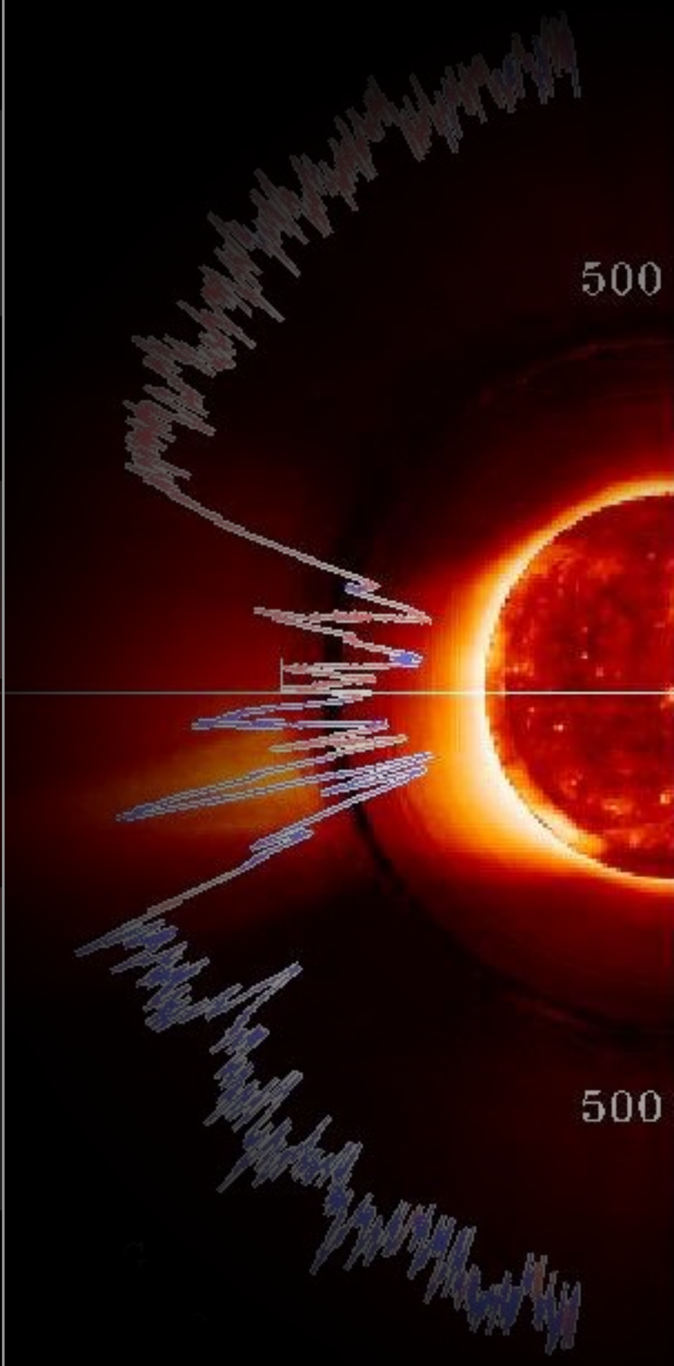
- Ulysses 1990-2009 (ESA/NASA)
- Left the ecliptic to study Sun from other inclination angles
    - Such an orbit challenging (gravity assist at Jupiter)
- Many instruments onboard, incl. SWOOPS (Solar Wind Plasma Experiment)
- 3D measurements of solar wind, ions + electrons
- Measured variation of solar wind as function of latitude
- Different components of solar wind (fast, slow)
- Solar wind varies over solar cycle



Solar Wind and Heliosphere

Properties

	Fast solar wind	Slow solar wind	Transient solar wind
Speed	> 400 km/s	<400 km/s	from < 300 km/s up to >2000 km/s
Proton density	~3 cm^{-3} homogeneous	~ 8 cm^{-3} highly variable	Often very low density
Magnetic field strength	~5 nT = 0.0005G	< 5 nT	Variable, up to 100 nT (0.01G)
Composition	95% H, 4% He	94% H, 5% He	Sometimes up to 30% He
Fluctuations	Alfvenic	Density	Often associated with interplanetary shock waves
Origin	Coronal holes coronal	Connected to coronal streamers	CMEs

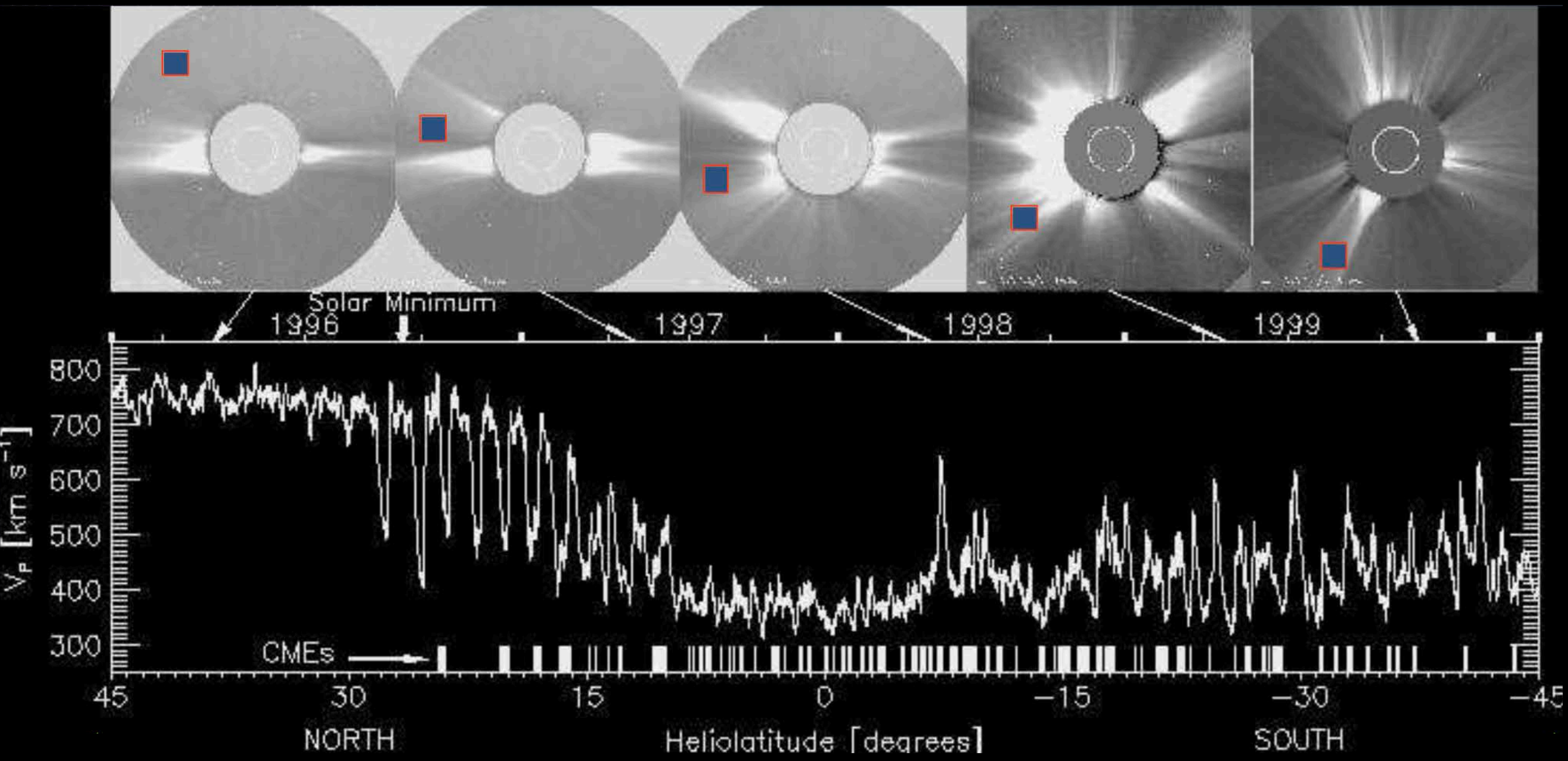


Solar Wind and Heliosphere

Variation with solar cycle

Solar activity
minimum
(8/1996)

Increasing activity,
towards
solar maximum



Solar Wind and Heliosphere

Solar Wind generation

- Still many open questions (one of the key objectives of Solar Orbiter!)
- Possible explanation for the solar wind:
  - Plasma confined lower solar atmosphere in closed magnetic field
  - Magnetic reconnection events between open and closed magnetic field

  ➡ Dynamical released into open field

- Fast Wind generation in Coronal Holes:
- Many possible physical processes (yet to be explained):
    - Reconnection events in the low corona
    - MHD turbulence
  - MHD waves
  - Low-frequency Alfvén waves
  - Heavy ion velocity filtration
  - . . .
  - Some include generation of different types of MHD waves, ion cyclotron waves, fast collisionless shocks, cascades …

Solar Wind and Heliosphere

Solar Wind generation

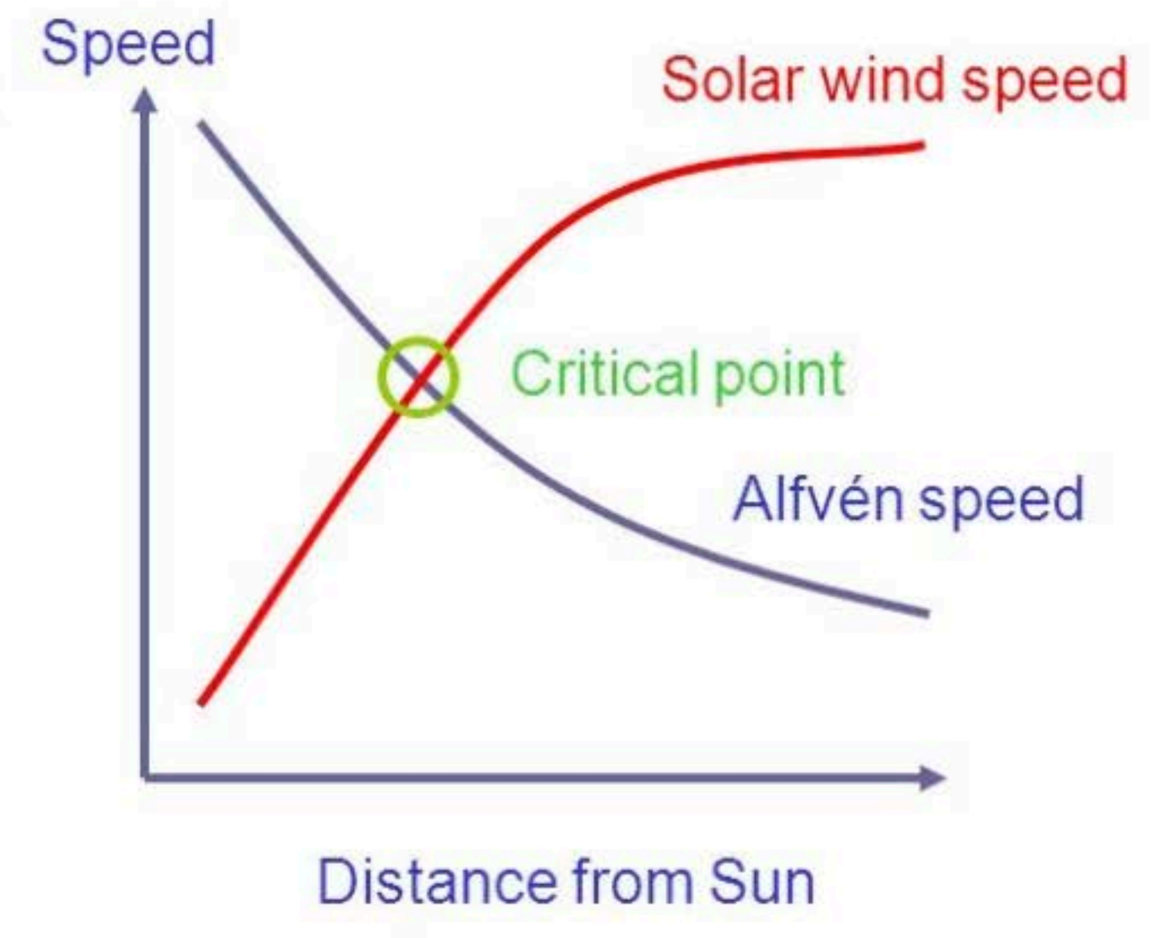
- Solar wind is channeled by magnetic field up to the Alfven radius $\mathbf{R}_A$, i.e. point where wind speed > Alfven speed

  ➡ Critical point

Alfvén speed

$$v_A \equiv \frac{B}{\sqrt{\mu_0 \rho}}$$

- Important:
  - At which height is the energy deposited that is needed for accelerating the solar wind?
  - Where are waves reflected and damped? (radial gradient of Alfvén speed!)
  - Energy deposited below or above the Parker critical point?



Solar Wind and Heliosphere

Parker's theory of the solar wind

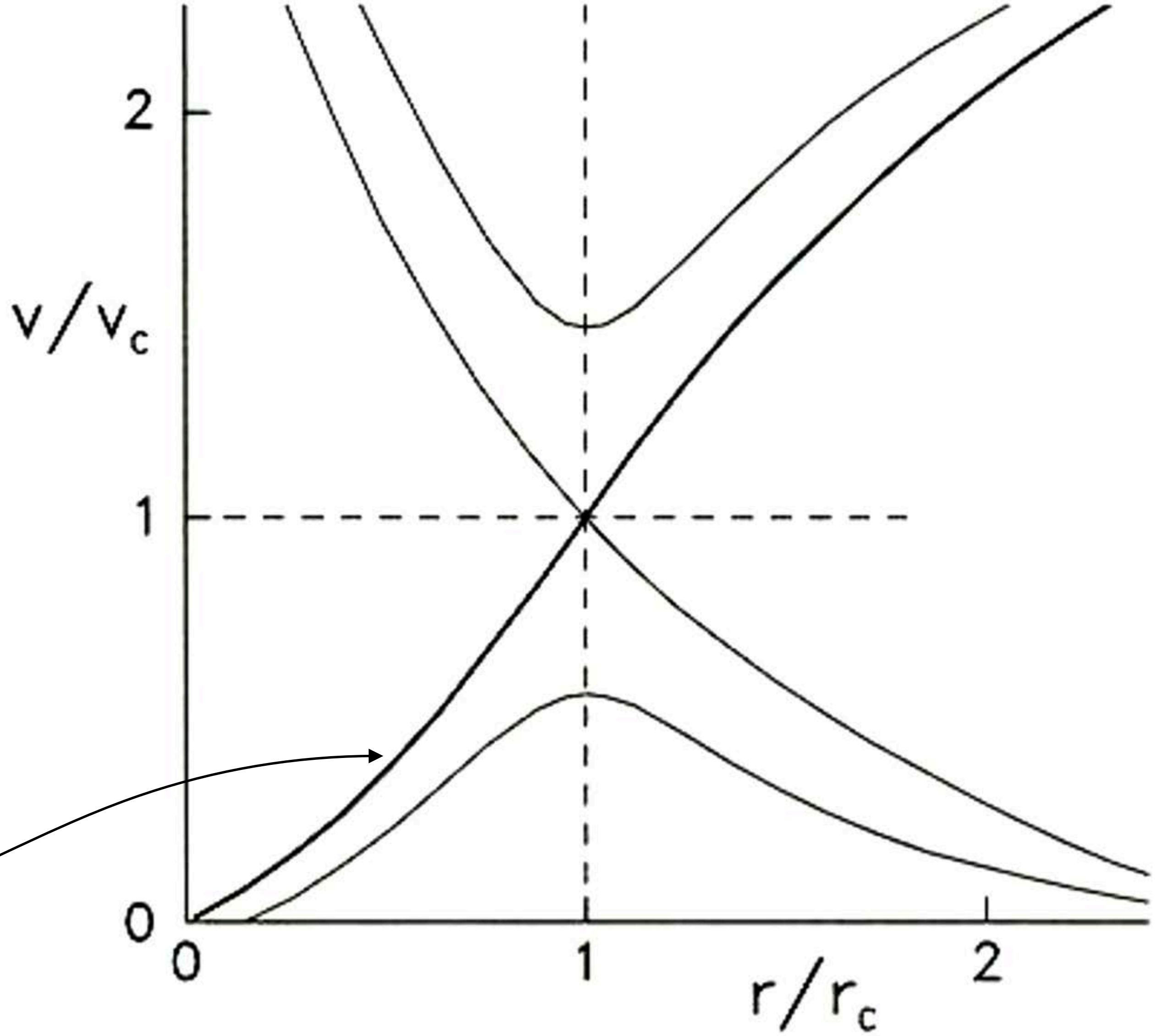
- **Basic idea:** Dynamic equilibrium between hot corona and interstellar medium.
- Mass and momentum balance equations:

$$\frac{d}{dr}(\rho r^2 v) = 0 \qquad v\frac{dv}{dr} = -\frac{1}{\rho}\frac{dP}{dr} - \frac{GM}{r^2}$$

➡ Parker's Equation for solar wind speed (assuming isothermal atmosphere)

$$\frac{1}{v}\frac{dv}{dr}(v^2 - c_s{}^2) = \frac{2c_s{}^2}{r} - \frac{GM}{r^2}$$

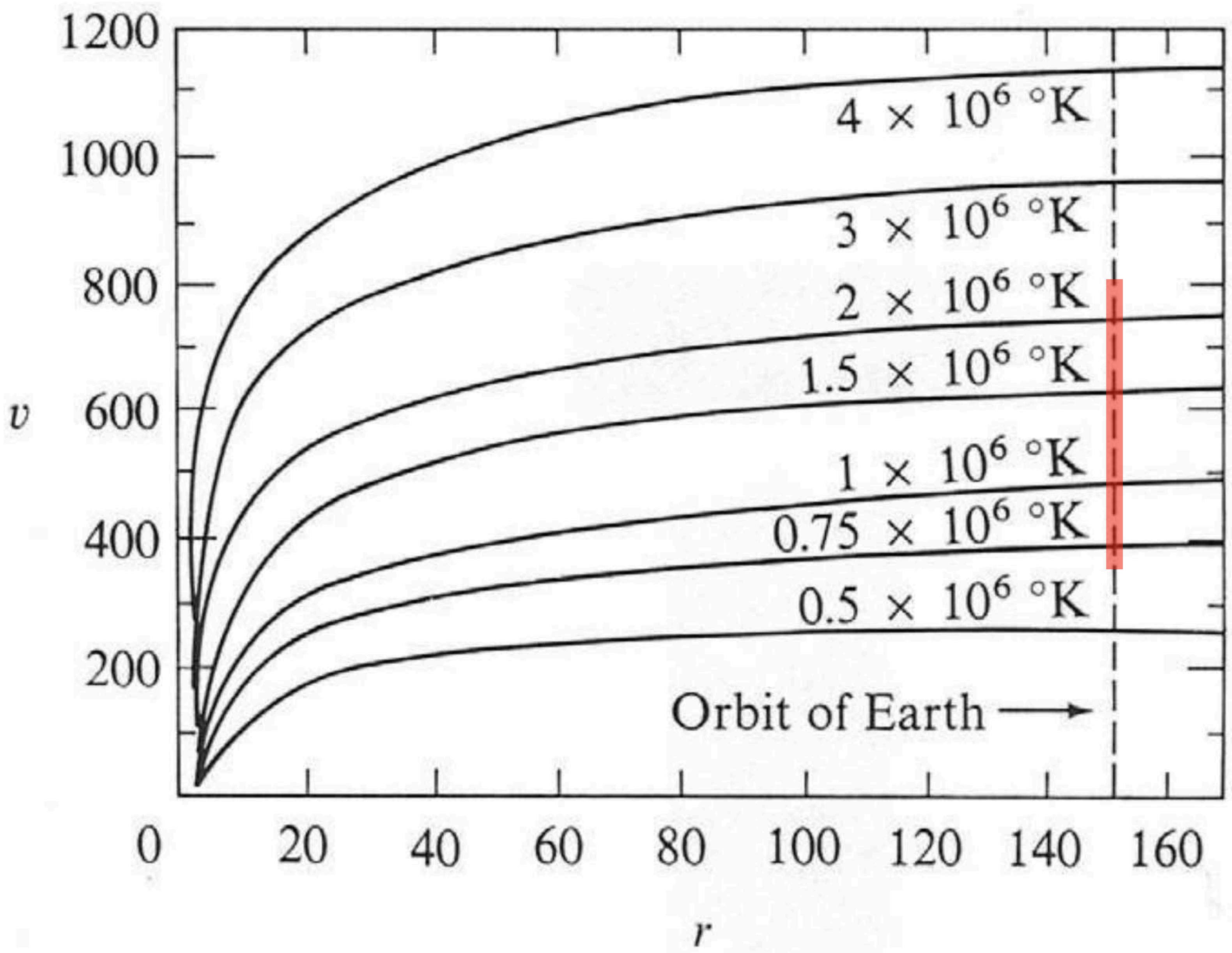
- Different (theoretically) possible solutions but some not valid for various reasons as in contrast to observations (some imply supersonic at solar surface or produce insufficient pressure against interstellar medium)
- One left: Correct solution (?!)

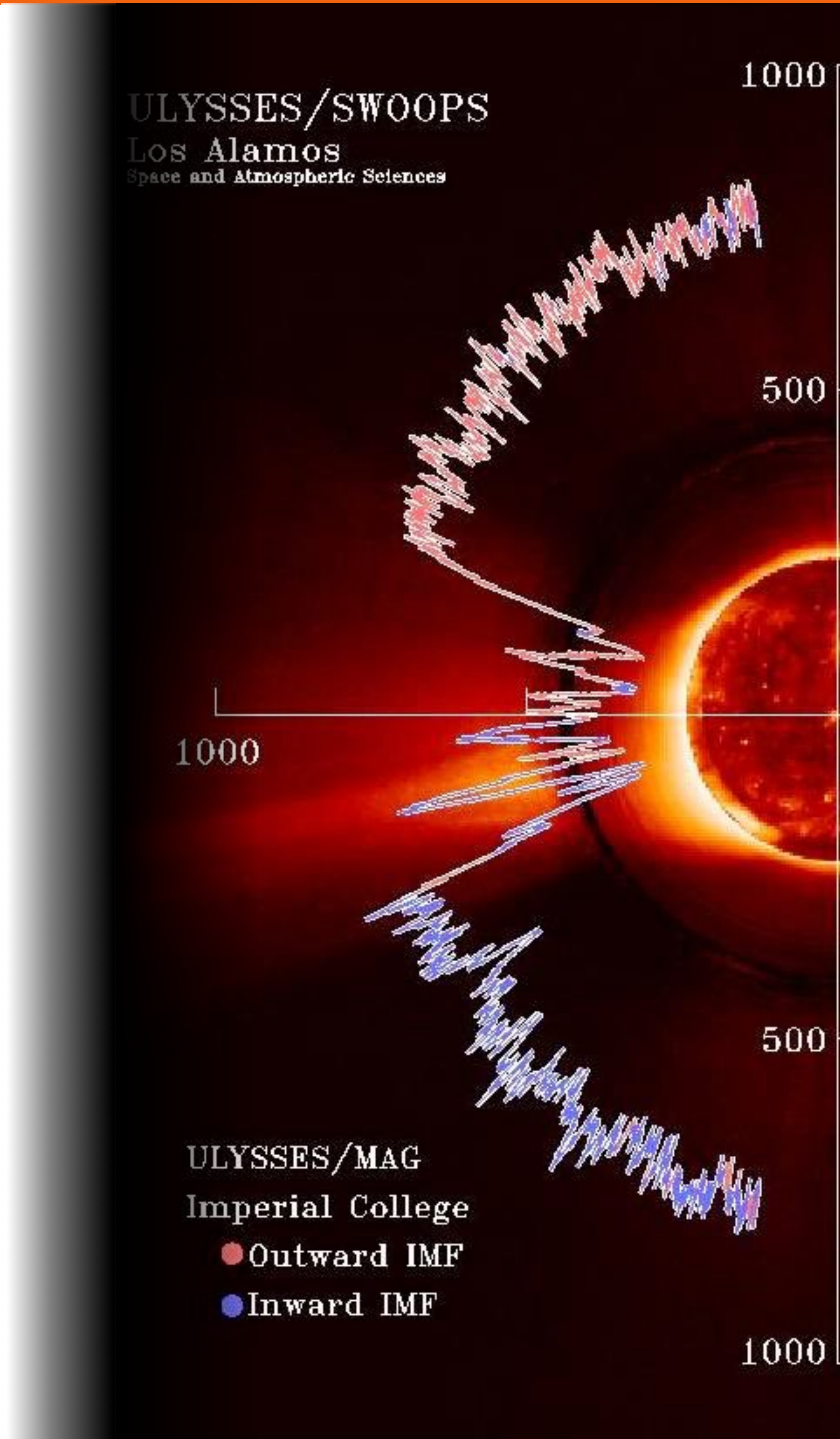


Solar Wind and Heliosphere

Parker's theory of the solar wind

- Speed of solar wind predicted by Parker's model for different coronal temperatures
  (simplified, isothermal case; no magnetic field)
- Consistent with measured velocities and coronal temperatures





Solar Wind and Heliosphere

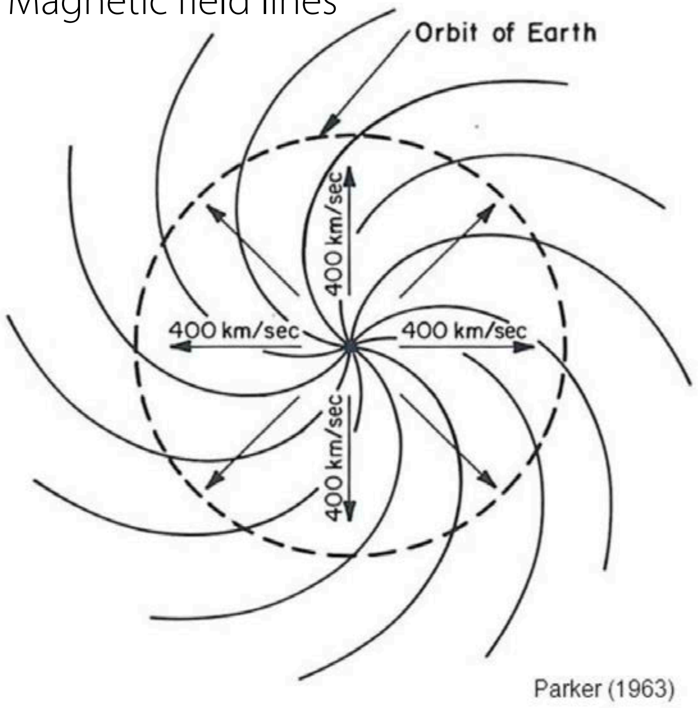
Stop here?

Parker spirals

- Magnetic field of the Sun extends into interplanetary space (throughout solar system)
- Field tethered to Sun while Sun rotates, magnetic tension

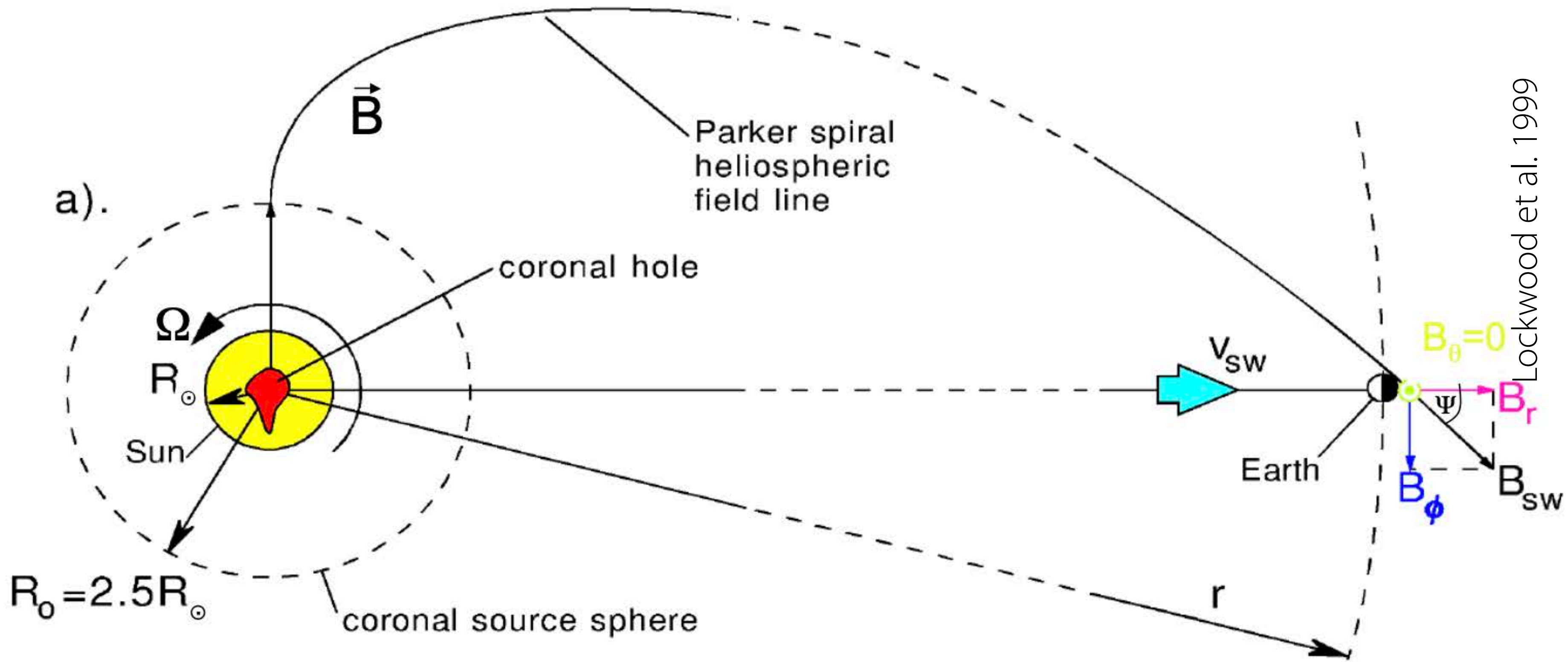
  ➡ **Parker spirals**

Magnetic field lines



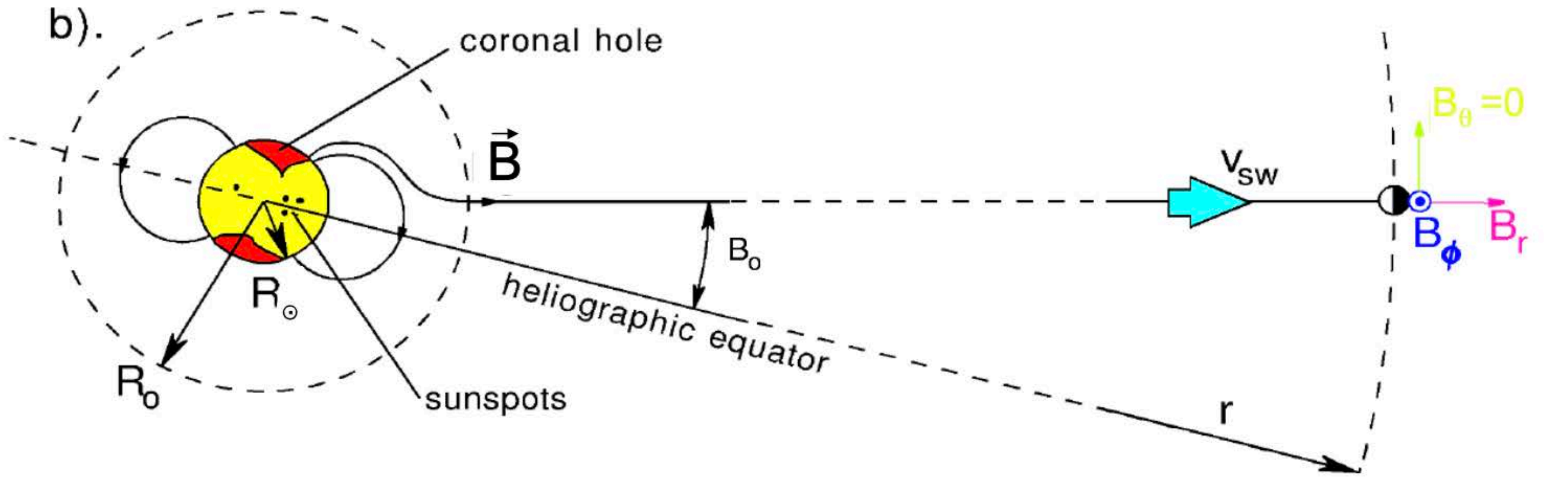
Parker (1963)

Poleward view



Lockwood et al. 1999

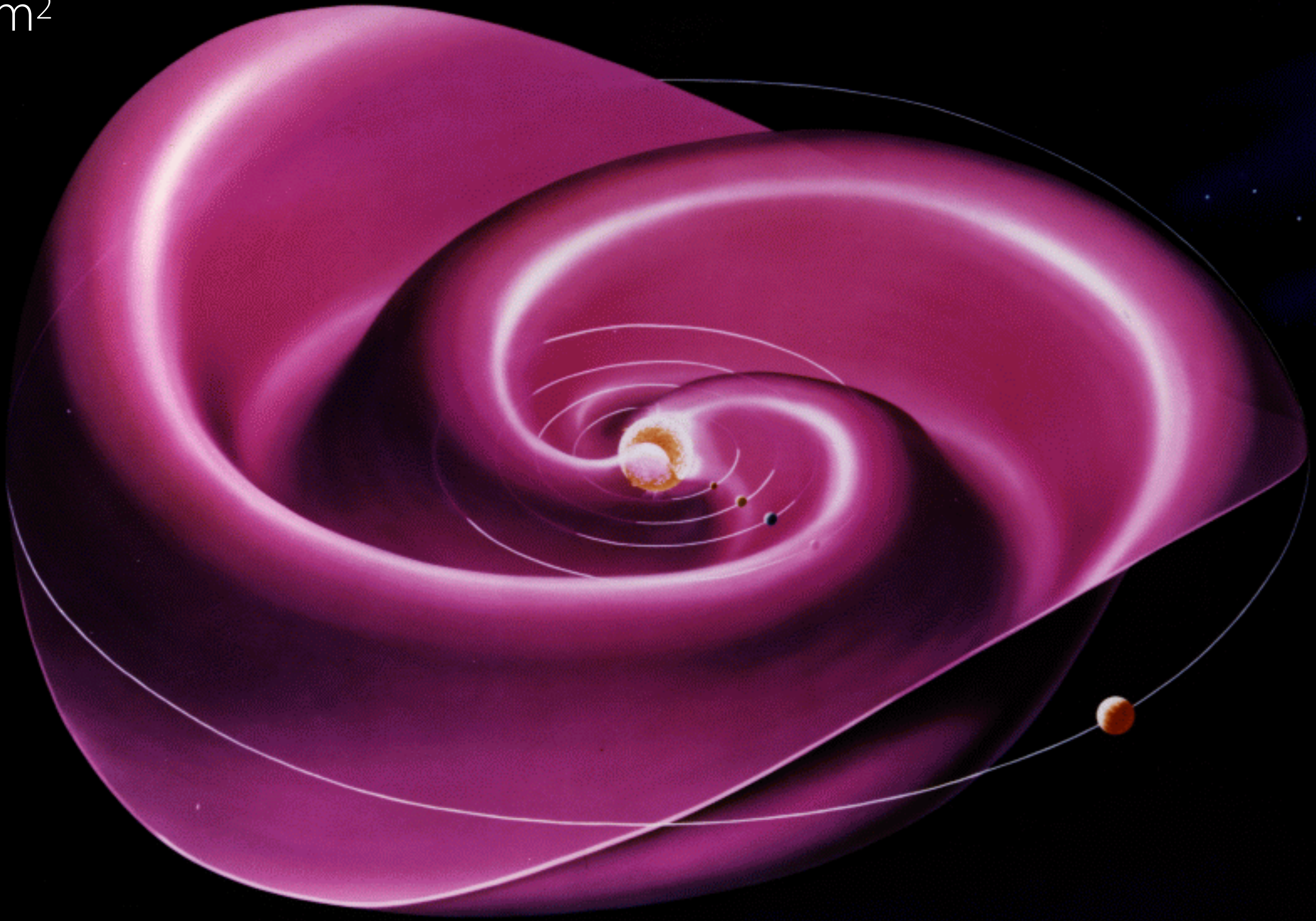
View from ecliptic



Solar Wind and Heliosphere

Heliospheric current sheet

- As a result:
  Heliospheric/interplanetary **current sheet** = surface where polarity of solar magnetic field changes
  - Electric current only ~10^{-10} A/m^2
  - Thickness of current sheet ~10 Mm at 1AU



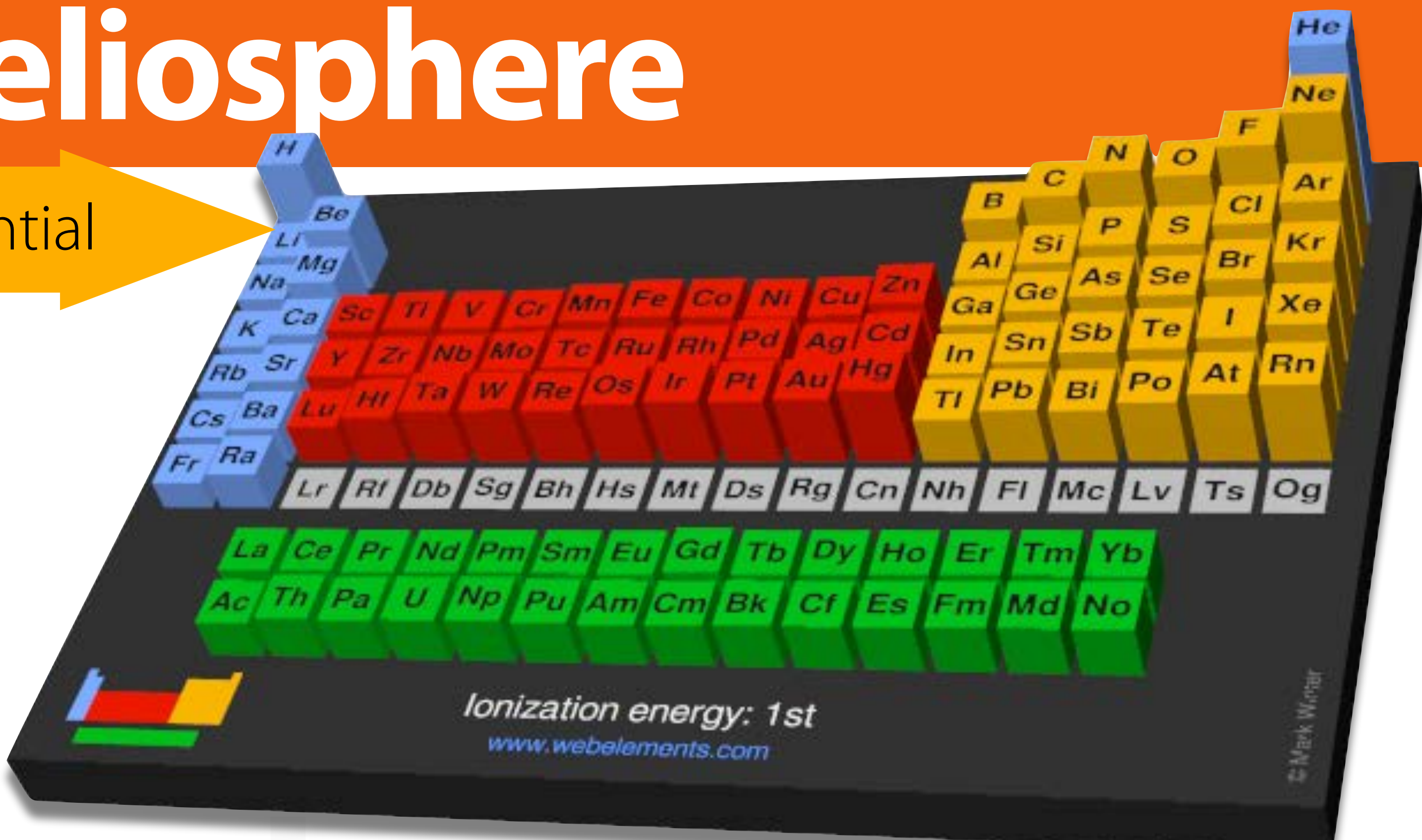
Heil / L. Svalgaard and J.M. Wilcox, Nature 262, p. 766, 1978.
K. Schatten, J.M. Wilcox, and N. Ness, Solar Physics, 6, p. 442, 1969.

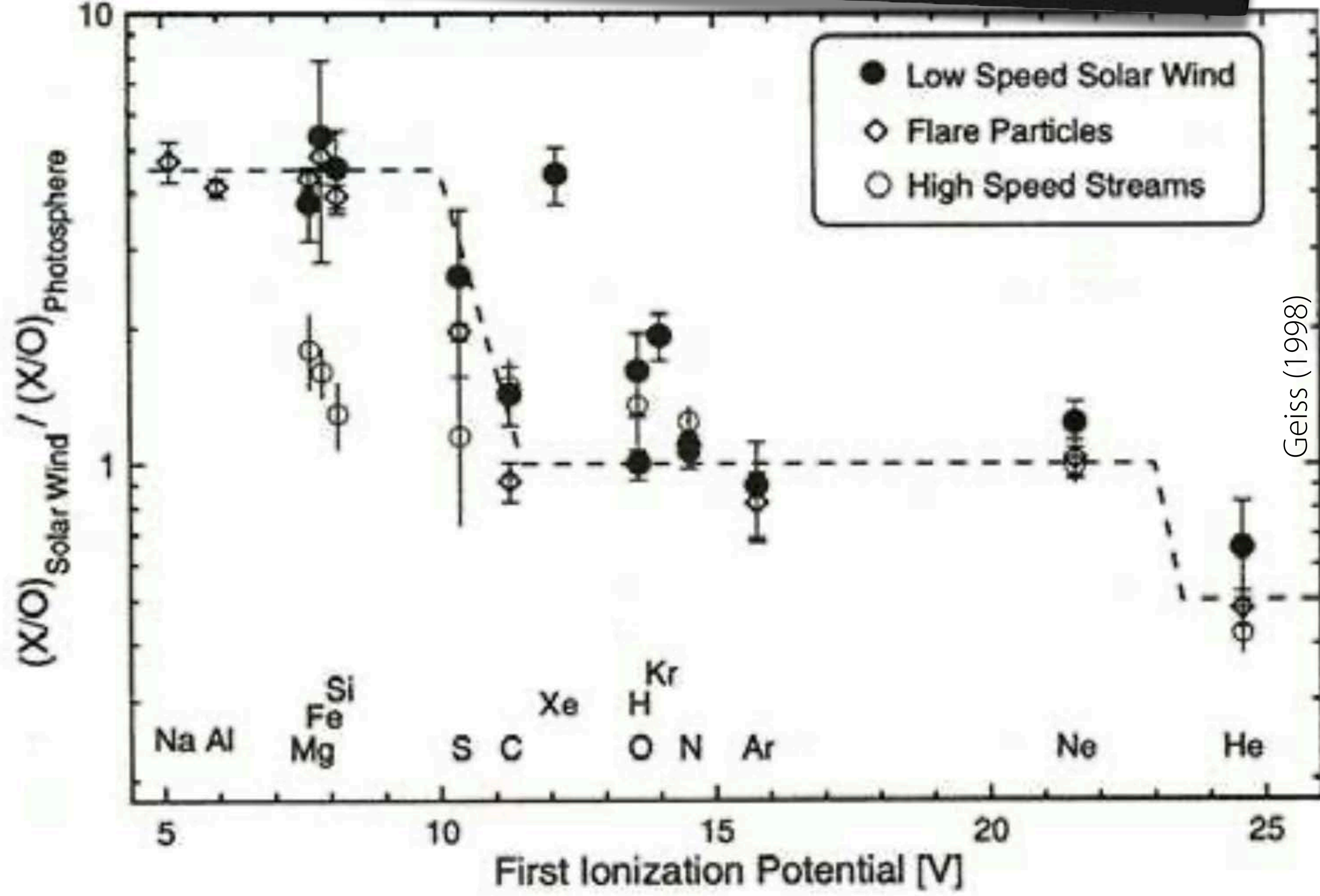
Solar Wind and Heliosphere

FIP effect

FIP = First Ionisation Potential

- Transition region, corona, solar wind: **Element abundances** differ from photospheric values!
- Elements with low FIP (< 10 eV): up to six times more abundant than elements with FIP> 10 eV.
- Fast solar wind: only weak (or absent) FIP effect
- Slow solar wind: FIP effect very pronounced
- Fast CME ejecta: Higher fractionation
- Differences imply dependence on the conditions in the region from where the wind plasma is originating in the first place





Geiss (1998)

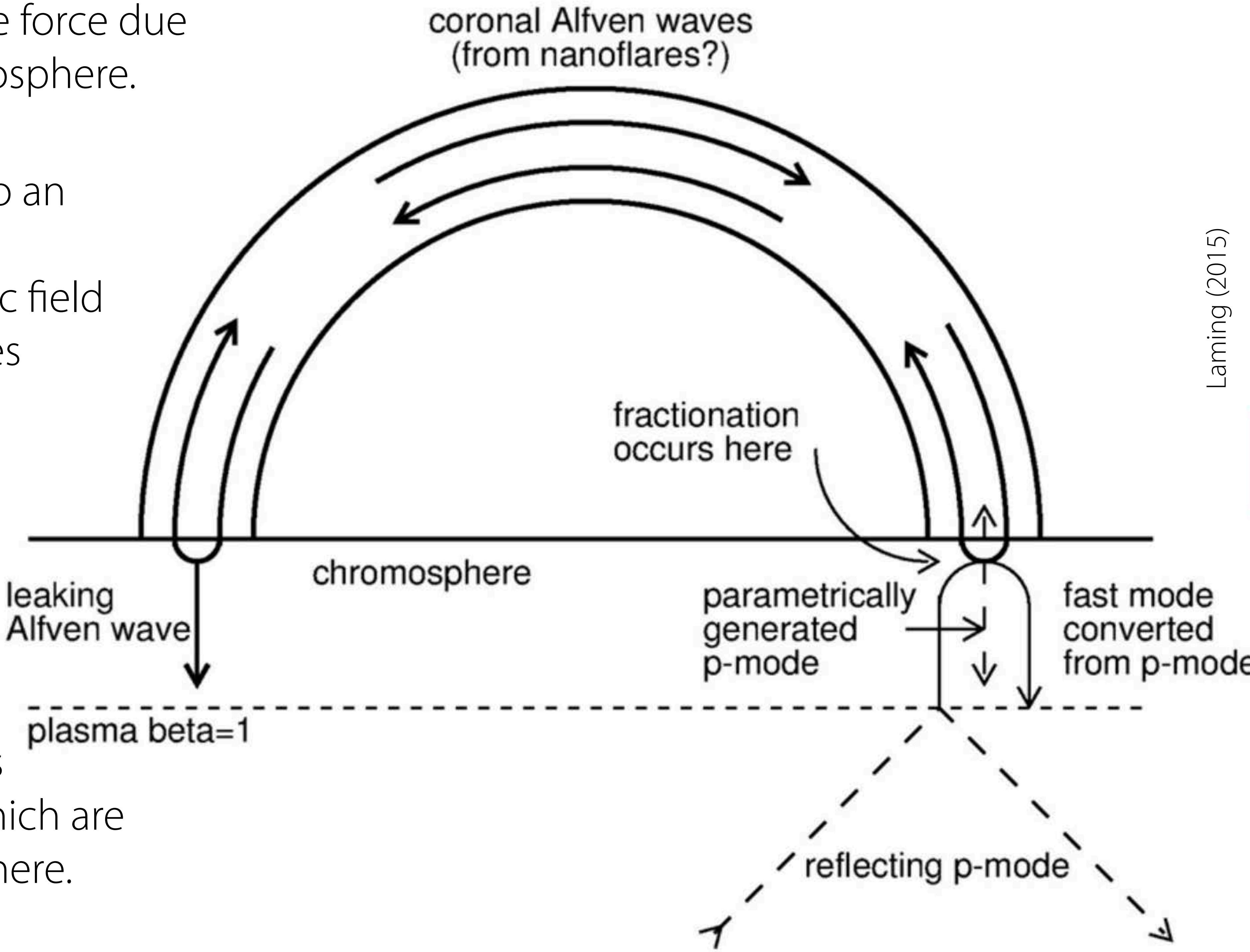
Solar Wind and Heliosphere

FIP effect

- Explanation: Acceleration of particles in chromosphere more efficient for low-FIP elements (ionised in chromosphere, high-FIP elements not ionised)
- Caused by ponderomotive force due to Alfvén waves in chromosphere.
- **Ponderomotive force** = net non-linear force due to an **inhomogeneous** (!) oscillating electromagnetic field acting on charged particles

➡ Charged particle moves towards area of weaker field strength (instead of oscillating around an initial point as in a homogeneous field)

➡ Acts on charged particles and thus on elements which are ionised in the chromosphere.

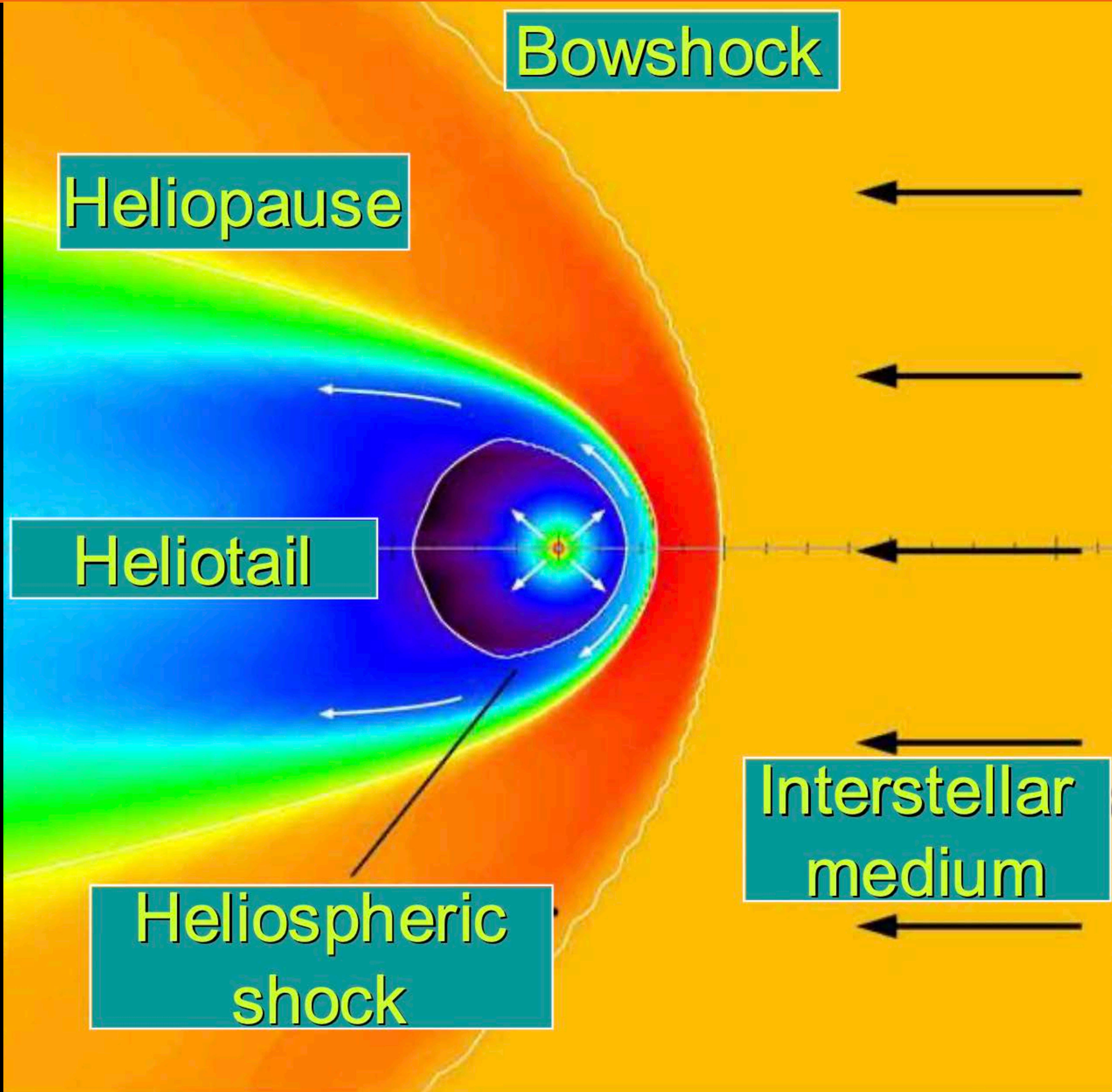


Laming (2015)

Solar Wind and Heliosphere

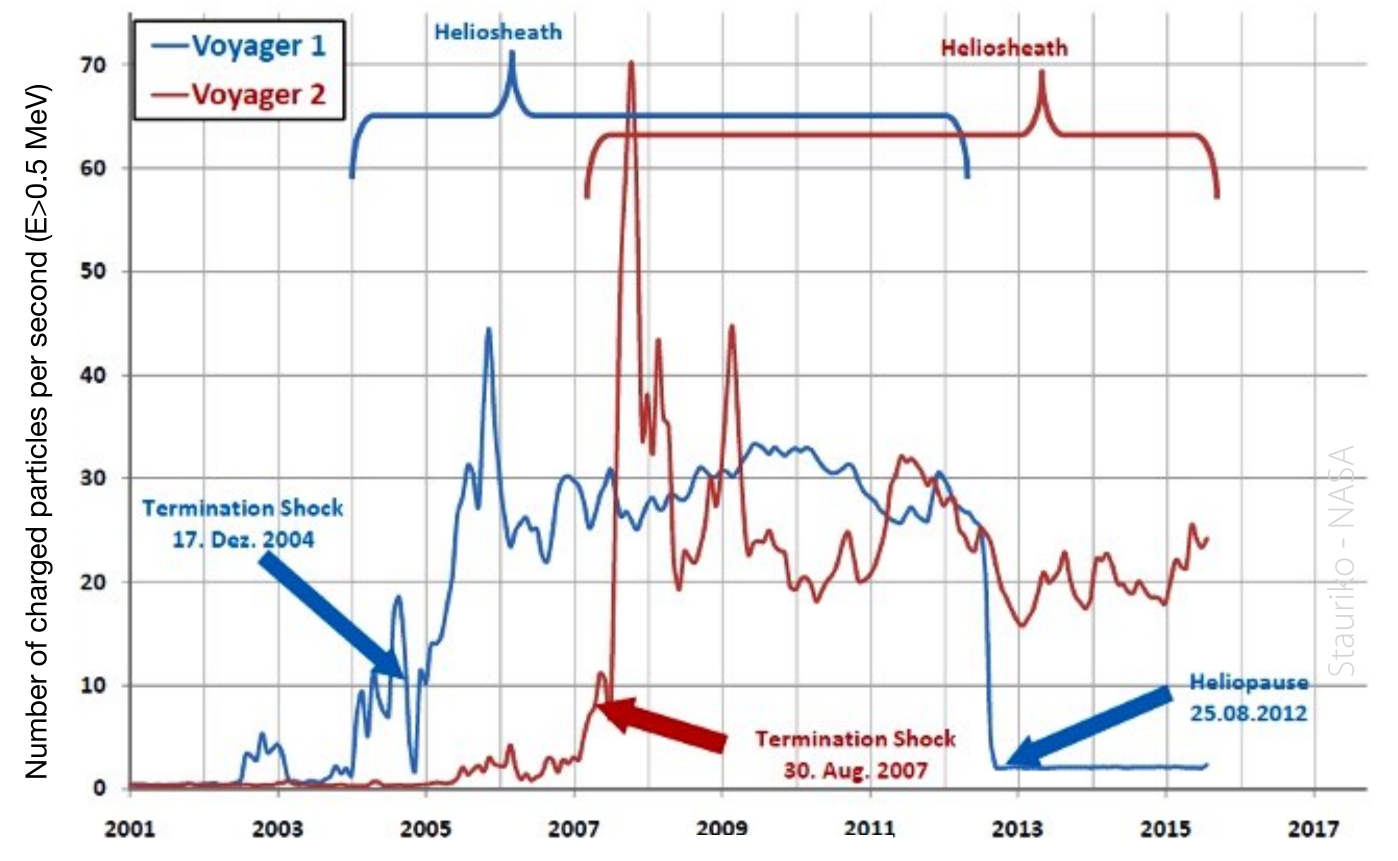
Heliosphere

- Definition: Region where solar wind and solar magnetic field dominate over interstellar medium and galactic magnetic field
- "Bubble" embedded in interstellar medium, produced by outflowing solar wind
- **Heliopause:** boundary of the heliosphere
- **Bow shock:** interstellar medium is slowed relative to the Sun.
- **Heliospheric shock:** solar wind is decelerated relative to Sun

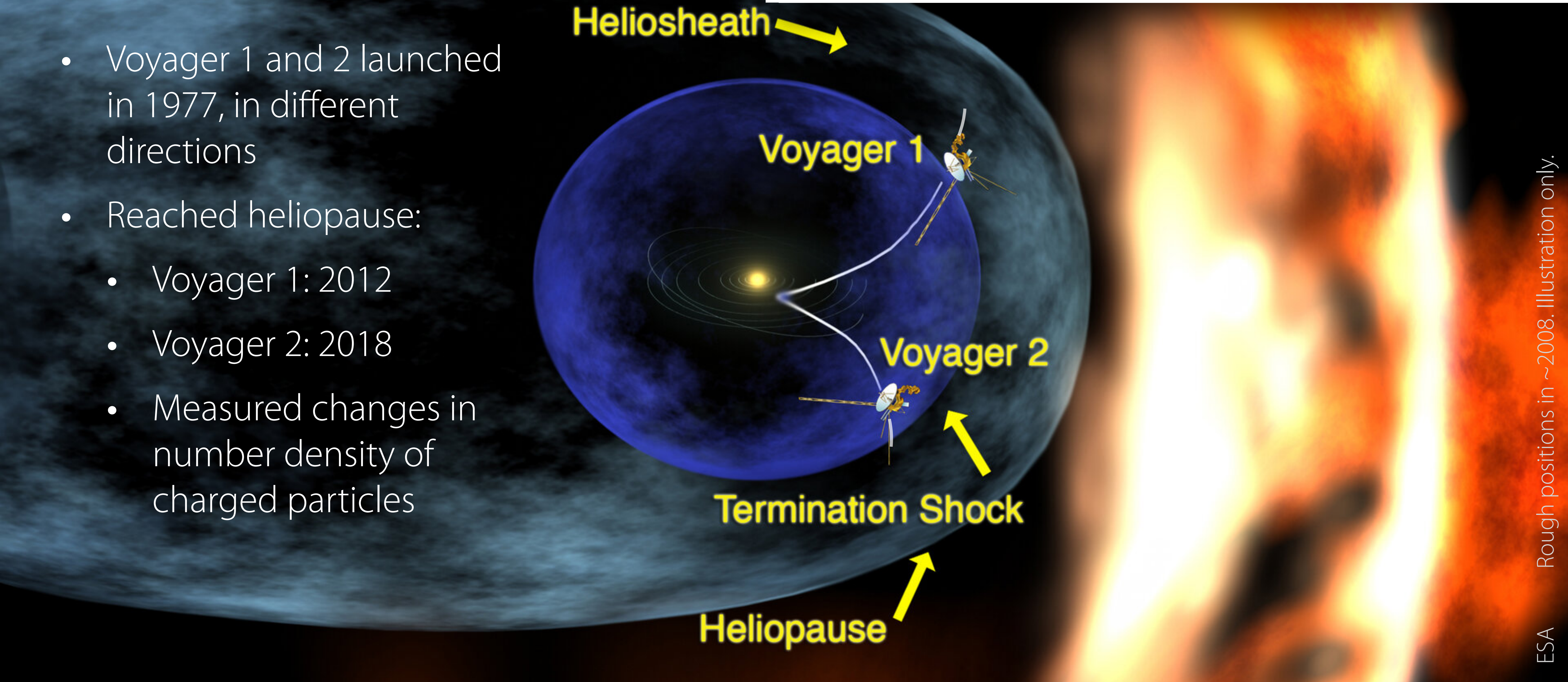


Solar Wind and Heliosphere

Heliosphere



- Voyager 1 and 2 launched in 1977, in different directions
- Reached heliopause:
  - Voyager 1: 2012
  - Voyager 2: 2018
  - Measured changes in number density of charged particles



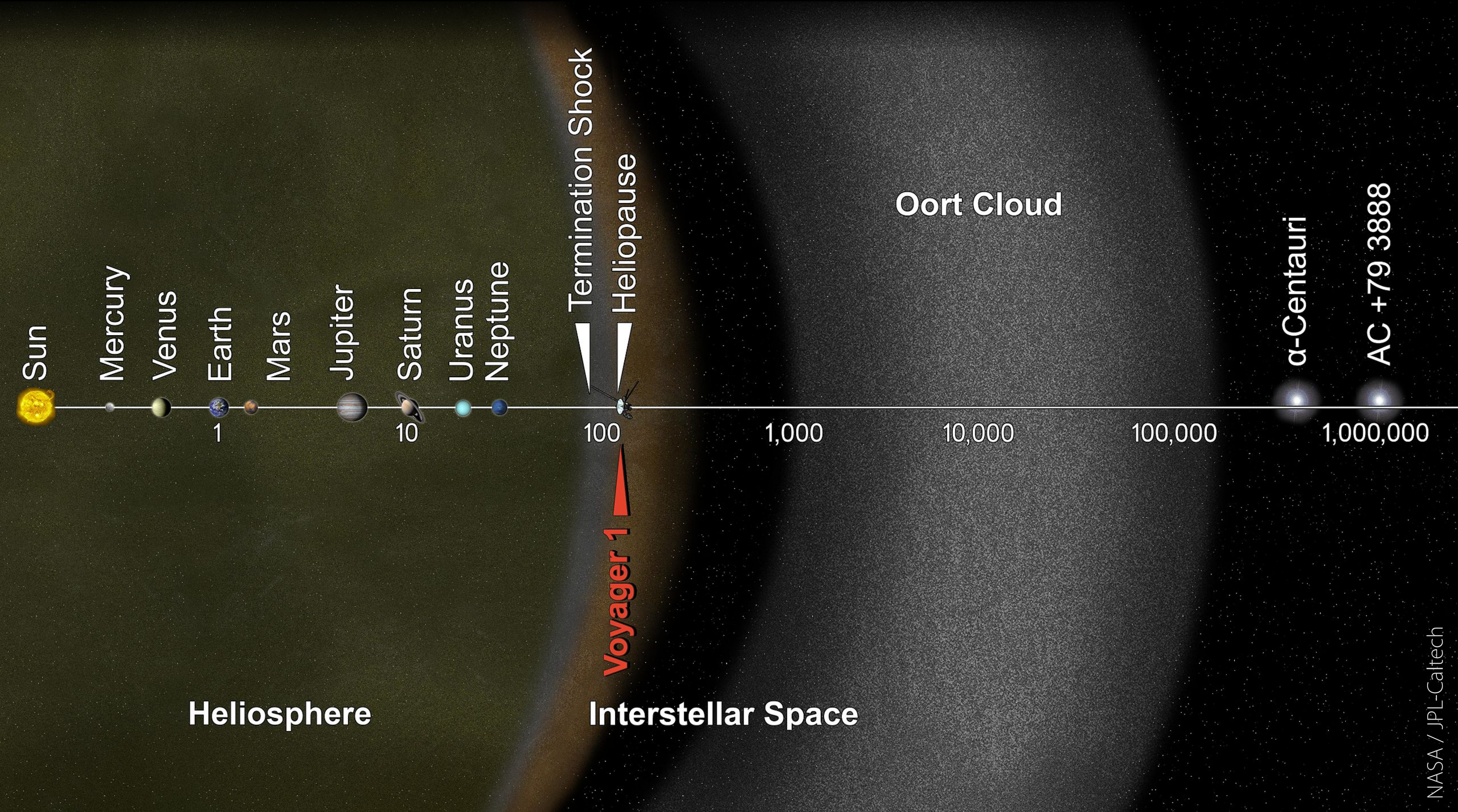
Rough positions in ~2008. Illustration only.

ESA

Solar Wind and Heliosphere

Heliosphere

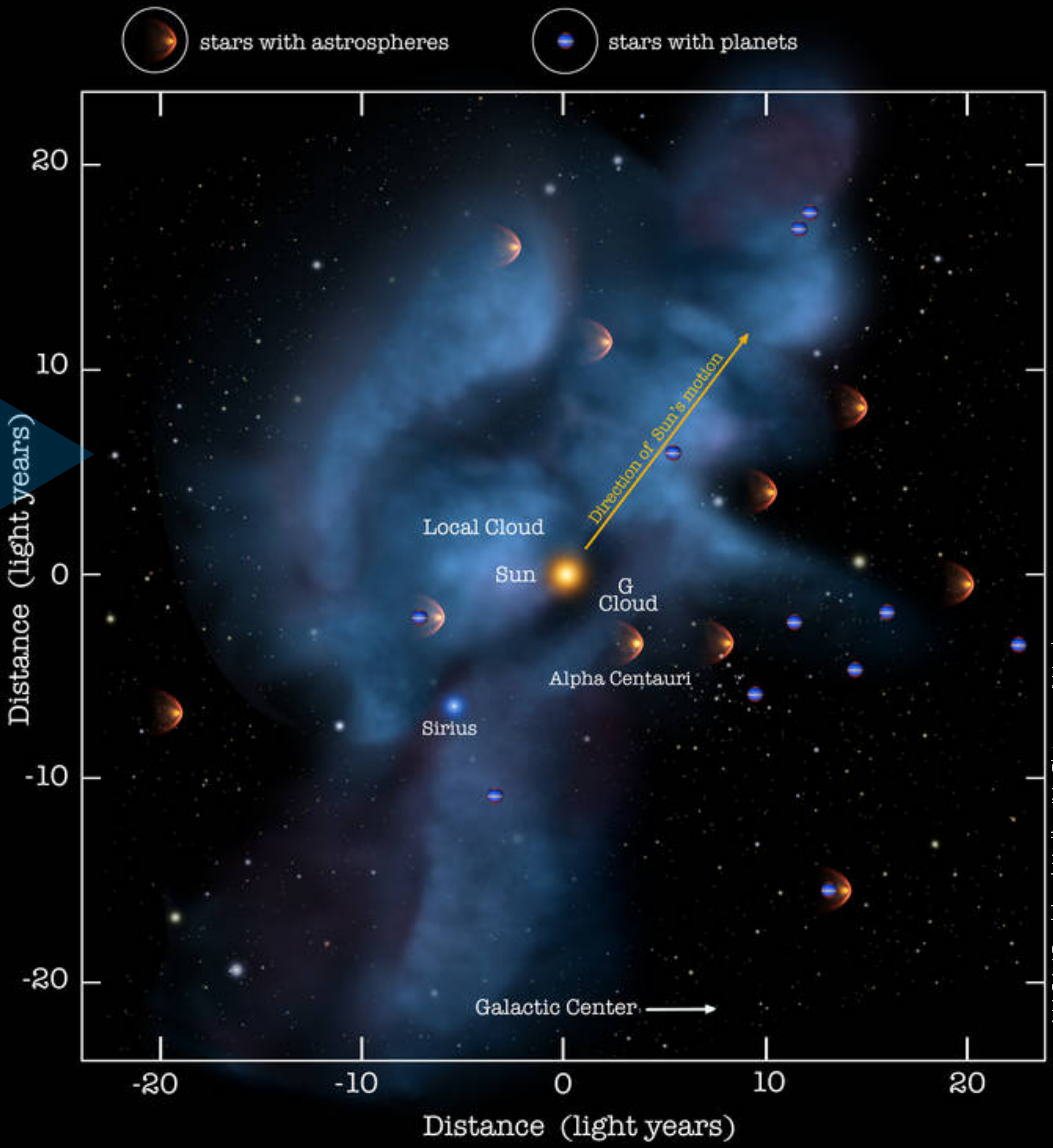
- **Heliosphere:** protective shielding against Galactic Cosmic Rays (hazardous for life)



NASA / JPL-Caltech

Beyond the heliosphere

- Astrosphere = like heliosphere but around other stars
- Detected around some other (nearby) stars
- Astrosphere around α Centauri (nearest star)!
- More candidates
- Many systems likely analogous to our solar system with astrospheres shielding a planetary system.



NASA/Goddard/Adler/U. Chicago/Wesleyan

Stellar Evolution

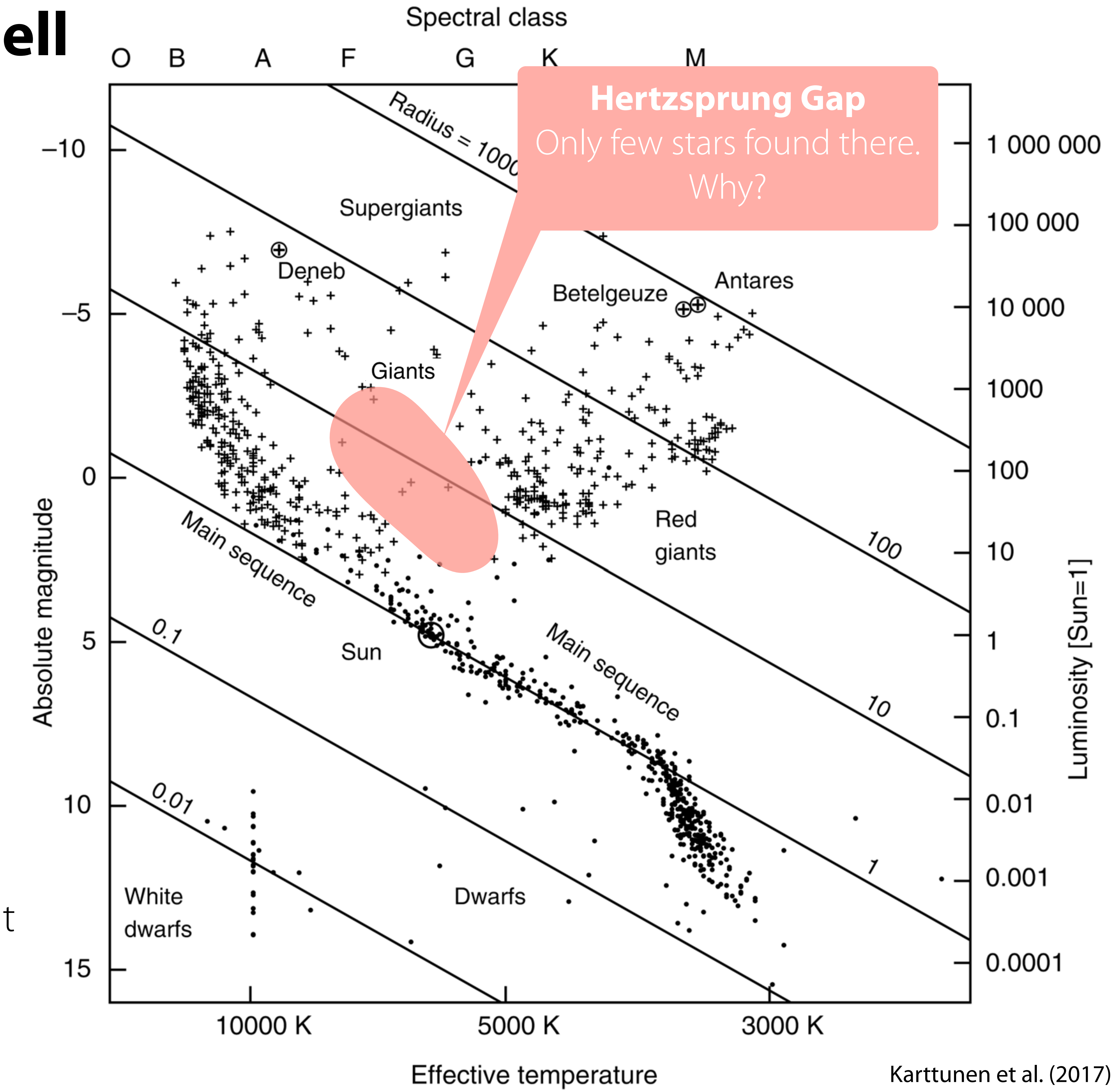
Stellar evolution

Hertzsprung-Russell diagram

- Stars clearly not distributed randomly
- Connections between
    - T_{eff}/Spectral class
    - Luminosity/absolute brightness
    - Radius

➡ Constraints for models of structure and evolution of stars

- Theory shows: Groups of stars represent different stages of their evolution but depends on mass (and chemical composition)



Hertzsprung Gap
Only few stars found there.
Why?

Karttunen et al. (2017).

Stellar evolution

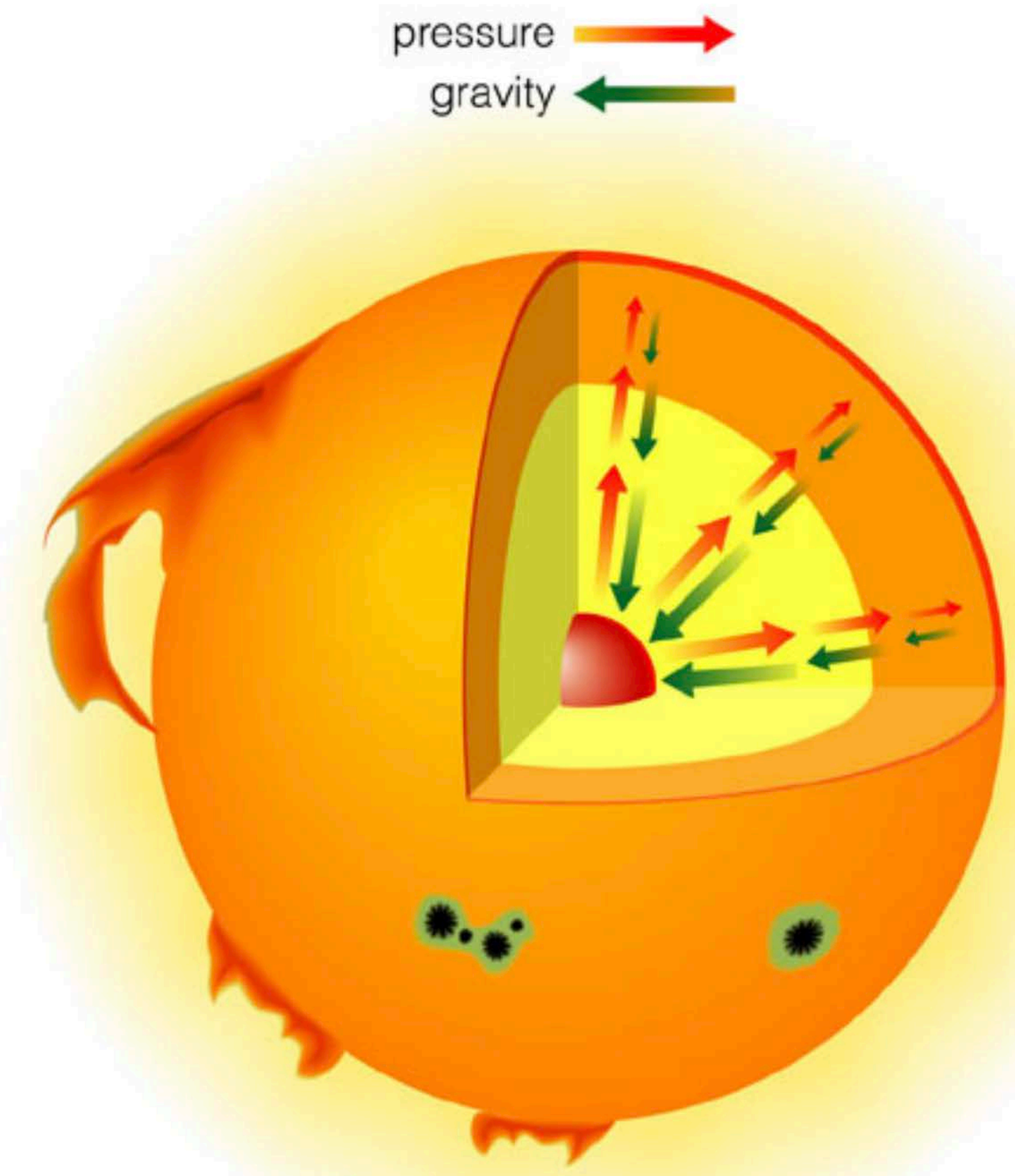
Introduction

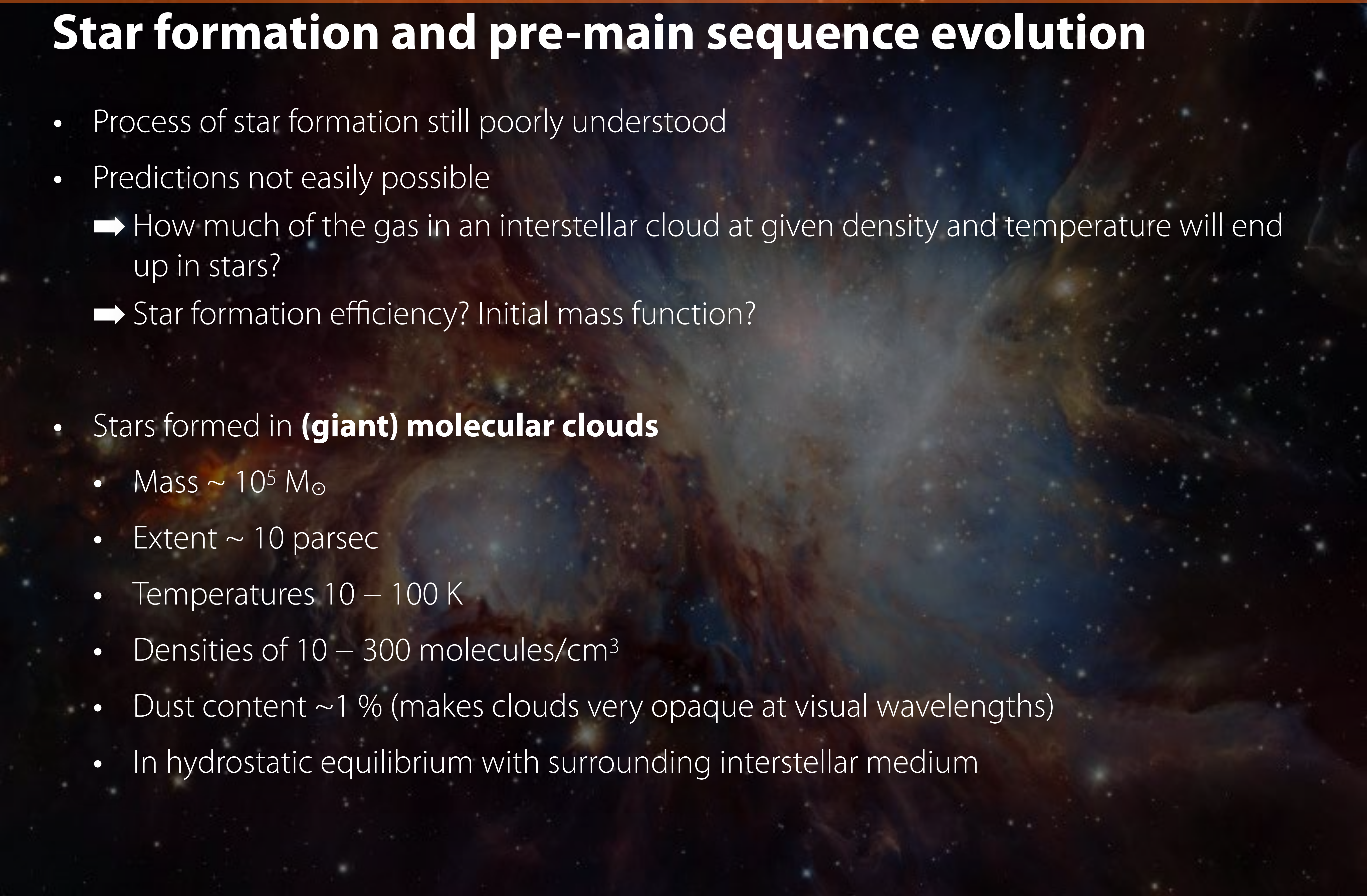
- Stellar evolution equations highly non-linear

➡ Complicated solutions that cannot always be anticipated based on fundamental principles, much depends on detailed numerical computations

- Simplifications and approximations usually needed (e.g., for convection, rotation, magnetism,…)

- **Basic principles**
  - Hydrostatic equilibrium required to keep a star stable and shining for extended time
  - The battle against gravity — energy source needed to keep the star stable
  - Slow adjustments of structure during evolution to maintain stability, some phase with more rapid changes
  - Time scales important — determine which processes matter the most at a given evolution stage





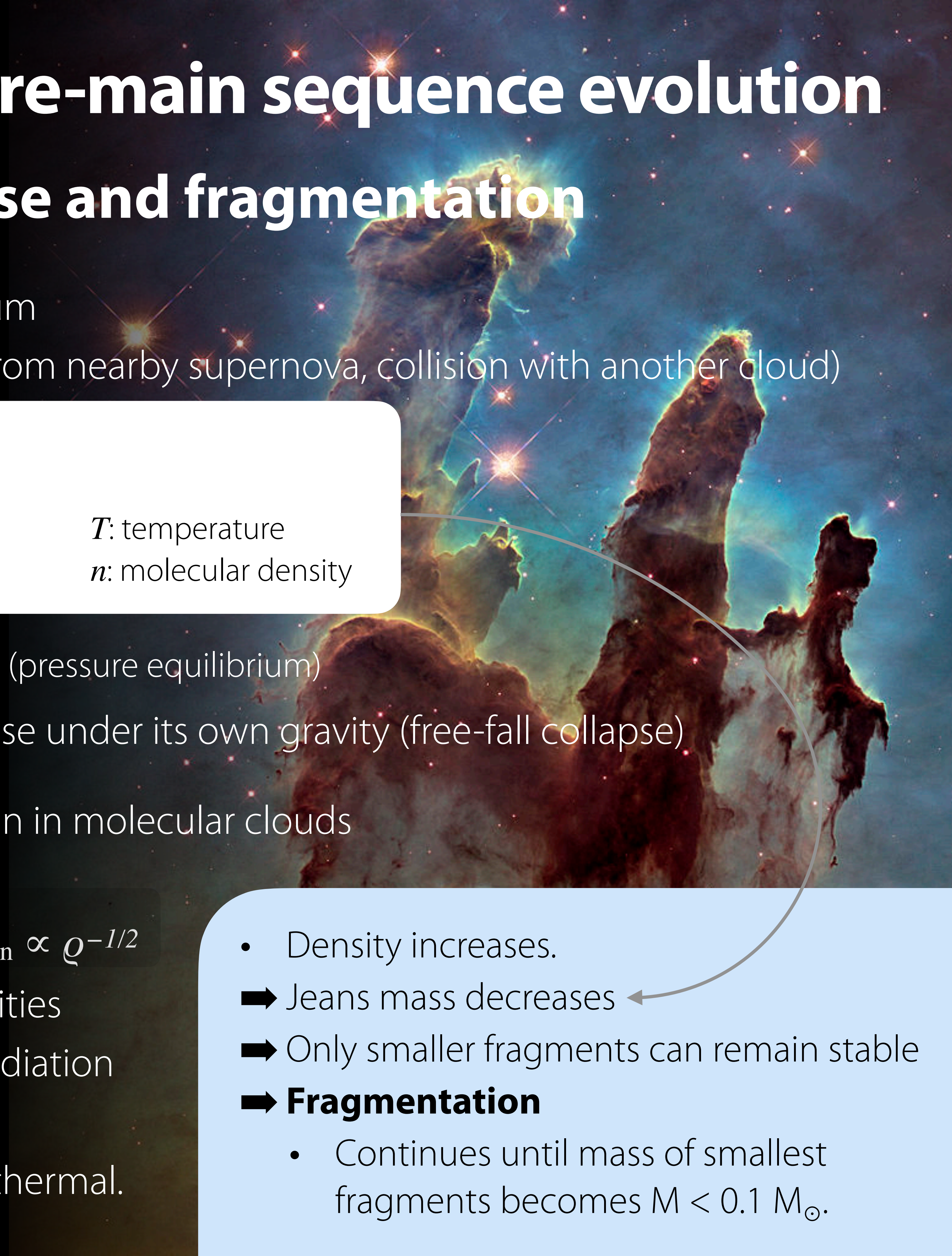
Star formation and pre-main sequence evolution

- Process of star formation still poorly understood
- Predictions not easily possible
  ➡ How much of the gas in an interstellar cloud at given density and temperature will end up in stars?
  ➡ Star formation efficiency? Initial mass function?

- Stars formed in **(giant) molecular clouds**
  - Mass ~ 10^5 $M_\odot$
  - Extent ~ 10 parsec
  - Temperatures 10 – 100 K
  - Densities of 10 – 300 molecules/cm^3
  - Dust content ~1 % (makes clouds very opaque at visual wavelengths)
  - In hydrostatic equilibrium with surrounding interstellar medium

Star formation and pre-main sequence evolution

Interstellar cloud collapse and fragmentation



- Cloud initially in hydrostatic equilibrium
- Now: Perturbation (e.g., shock wave from nearby supernova, collision with another cloud)
- (Critical) **Jeans mass**

$$M_{\mathrm{J}} \approx 4 \times 10^4\, M_\odot \left(\frac{T}{100\,\mathrm{K}}\right)^{3/2} \left(\frac{n}{\mathrm{cm}^{-3}}\right)^{-1/2}$$

T: temperature
n: molecular density

  - Cloud with $M < M_{\mathrm{J}}$ remains stable (pressure equilibrium)
  - Cloud with $M > M_{\mathrm{J}}$ starts to collapse under its own gravity (free-fall collapse)
  - $M_{\mathrm{J}} \sim 10^3 - 10^4\, M_\odot$ for typical T and n in molecular clouds

- **Collapse** — dynamical timescale $\tau_{\mathrm{dyn}} \propto \varrho^{-1/2}$
  - ~millions of years due to low densities
  - Cloud transparent in far-infrared radiation
  - ➡ Cools efficiently.
  - ➡ Early stages of the collapse are isothermal.

- Density increases.
- ➡ Jeans mass decreases
- ➡ Only smaller fragments can remain stable
- ➡ **Fragmentation**
  - Continues until mass of smallest fragments becomes $M < 0.1\, M_\odot$.

Star formation and pre-main sequence evolution

Formation of a protostellar core and disk

- Density increases.

➡ Gas becomes increasingly opaque in IR, cools less.

➡ Temperature and gas pressure increase

➡ Cloud core establishes hydrostatic equilibrium

➡ Dynamical collapse slowed down to slow contraction

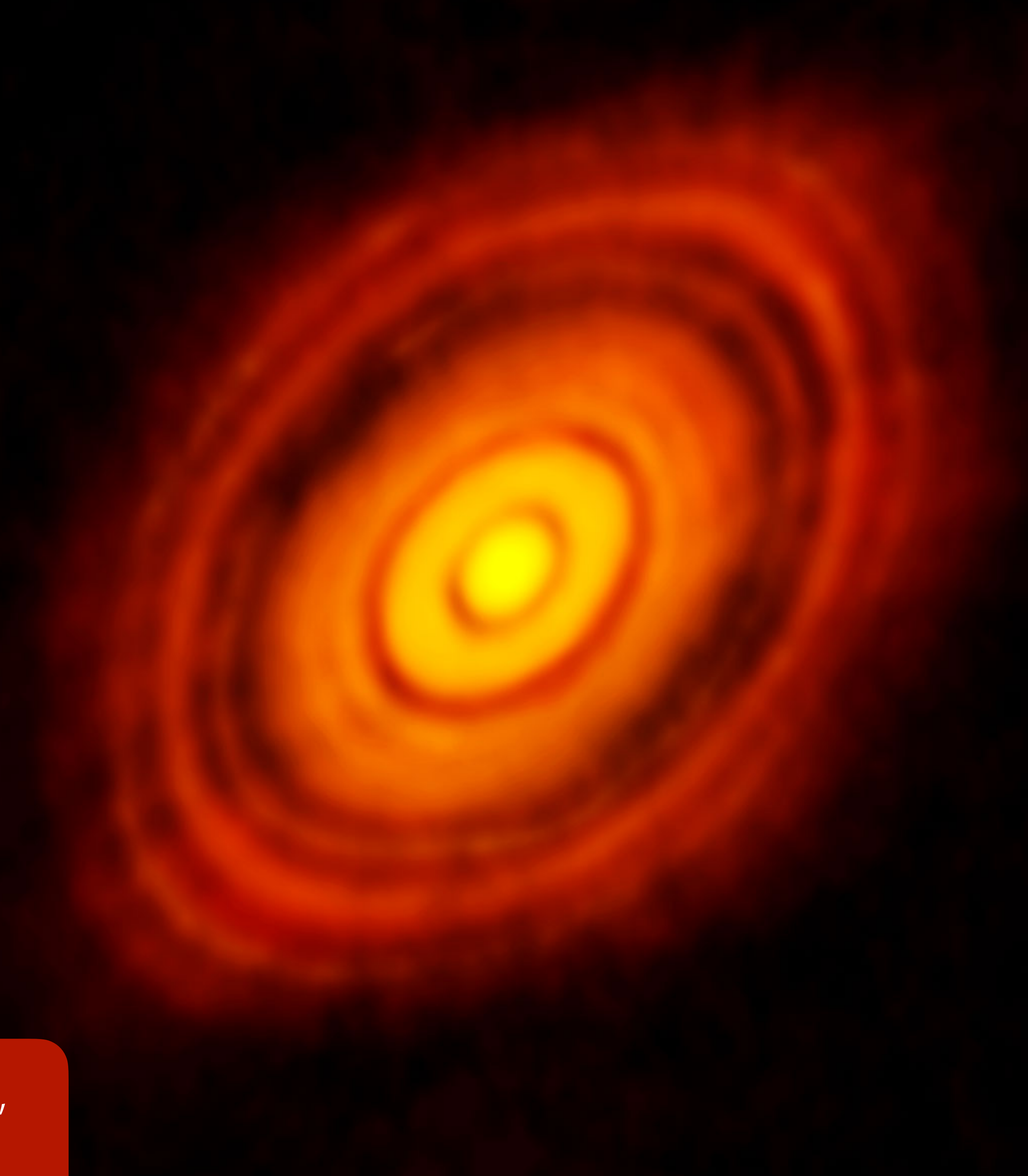
➡ **A protostar if formed!**

- Conservation of angular momentum of infalling gas

➡ **Accretion disk** around protostar is formed

- Can be observed at infrared to sub-millimeter wavelengths

T Tauri stars: pre-main sequence (age ~ 10^6 yrs), accretion disk still optically thick, large IR excess observed



HL Tauri ALMA (ESO/NAOJ/NRAO)