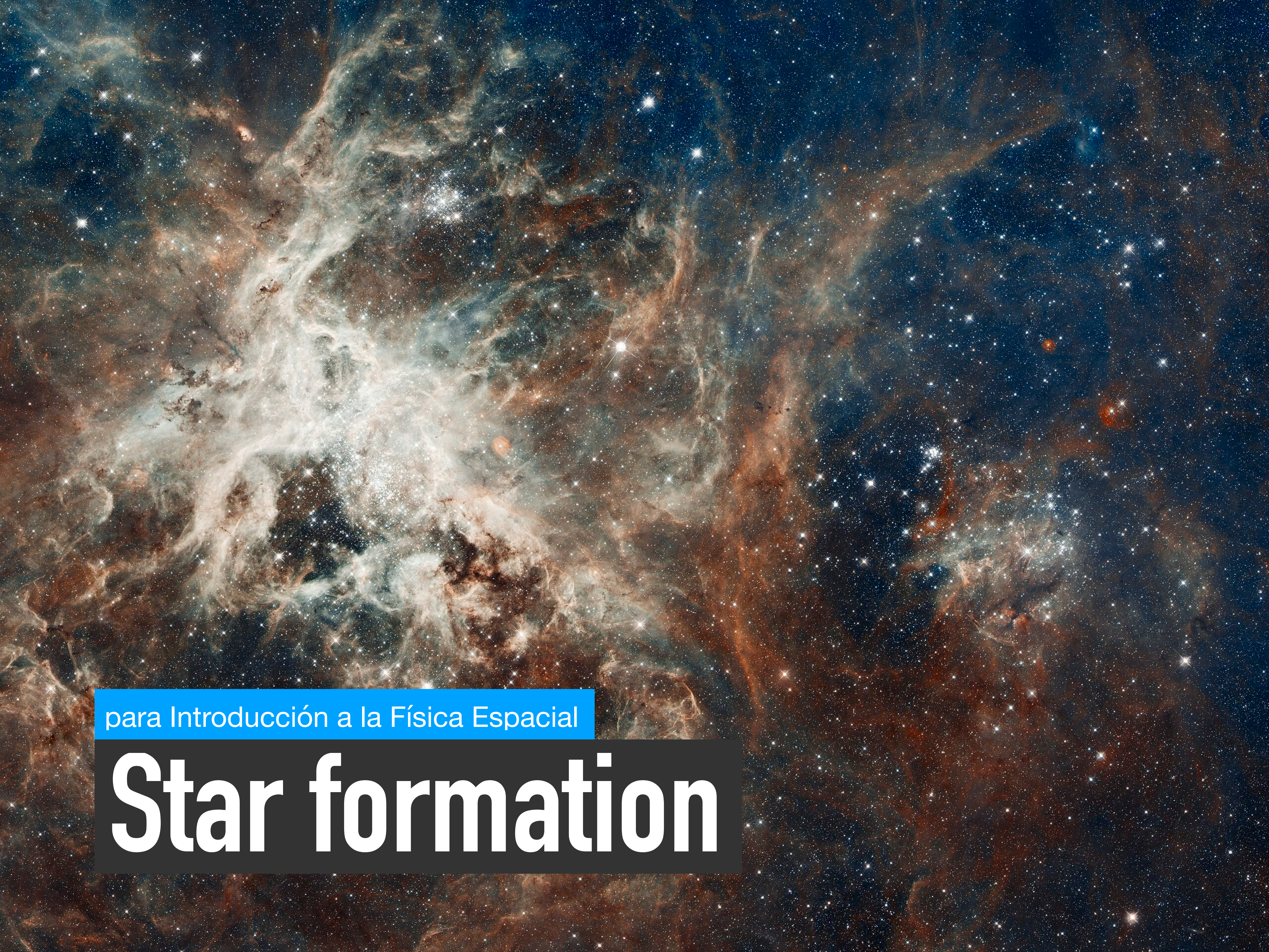
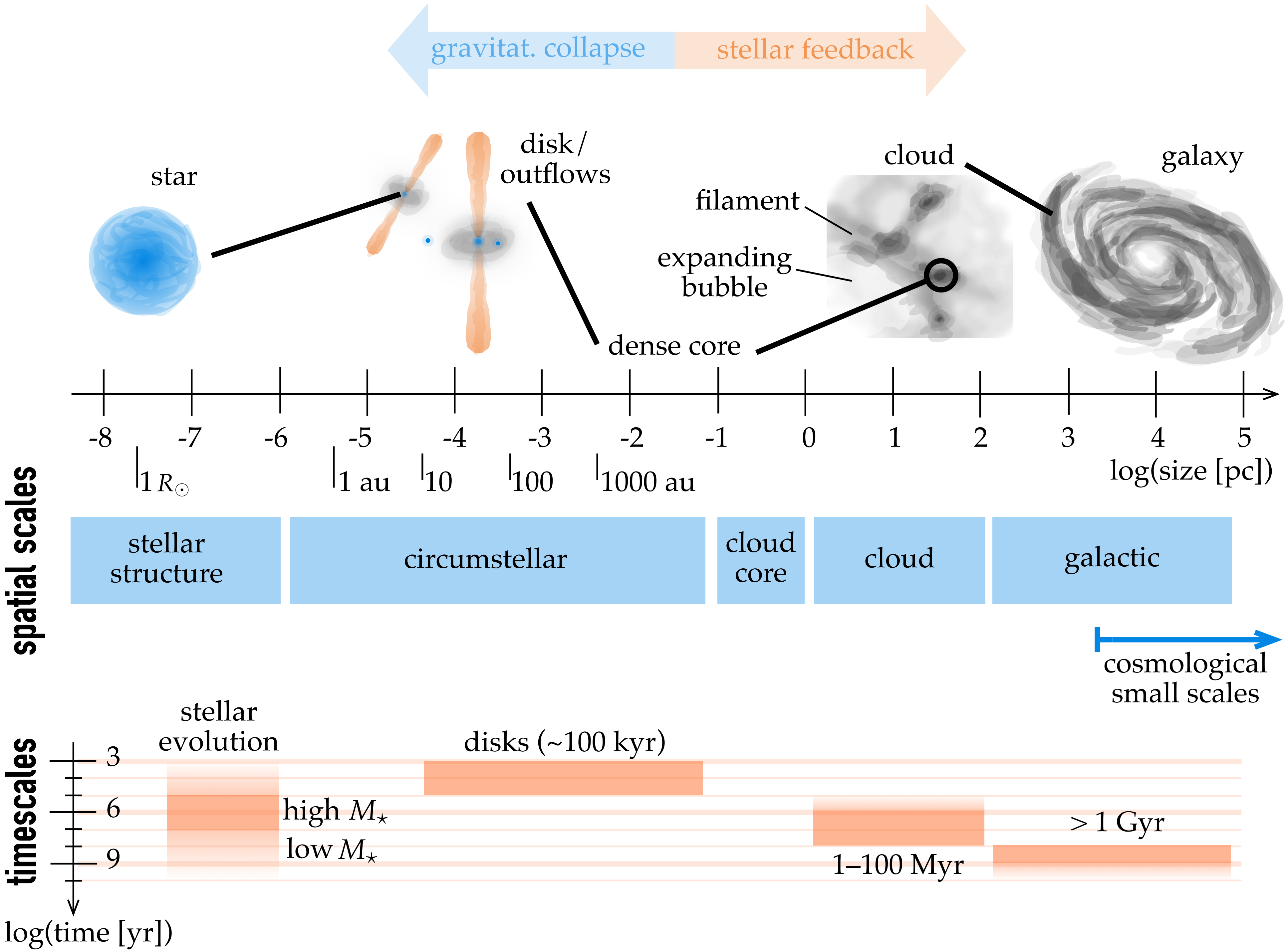


para Introducción a la Física Espacial

Star formation

Contents

Basic concepts



Scales of star formation

This diagram shows the different spatial scales where star-formation processes have an impact, as well as their duration.

	giant molecular cloud	molecular cloud	clump	cloud core	inner core	circumstellar material (disk)
Radius (pc)	20	5	2	0.1 ≈ 20 000 au	1000 au	~100 au
H_2 number density [cm^{-3}]	100	300	10^3	10^5	10^7	$\gtrsim 10^{10}$
Enclosed mass [$M_\odot$]	10^5	10^4	10^3	10^1	ongoing accretion (time-dependent)	ongoing accretion (time-dependent)
Velocity (dispersion) [km s^{-1}]	7	4	2	0.3	free-fall-like kinematics	Keplerian-like kinematics
Temperature	15	10	10	10	50	100
Speed of sound [km s^{-1}]	0.23	0.19	0.19	0.19	0.42	0.59

Typical properties of star-forming material

Bodenheimer 2011 *Principles of Star Formation*, Springer

Oliva & Kuiper 2023a,b A&A

Self-gravity of a sphere of mass M, radius R with uniform density $\rho = 3M/(4\pi R^3)$: the potential energy of a particle of mass dm outside a sphere of radius r, mass M_r is $dU = -GM_r dm/r$ (← spherical symmetry/shell theorem/Gauss's law for Newtonian gravity). If instead of a particle, we have a spherical shell of same density and thickness dr, located just at the surface of the inner sphere, $dm = 4\pi r^2 \rho dr$ (← integrating $dm = \rho dV$ along θ, ϕ). Now, $M_r = 4\pi r^3 \rho/3 = Mr^3/R^3$. In order to get the full self-gravity potential, we integrate: $U = -4\pi G \int_0^R M_r \rho r dr = \cdots = -\frac{3}{5}\frac{GM^2}{R}$.

Virial temperature: the virial theorem states that $2\langle K\rangle + \langle U\rangle = 0$. We omit the brackets. The kinetic energy is $K = \frac{3}{2}Nk_BT = \frac{3}{2}\frac{M}{\bar{m}}k_BT$. The potential energy of a self gravitating uniform sphere is $U = -\frac{3}{5}\frac{GM^2}{R}$. Substituting in the virial theorem, we get the virial temperature $T_{\text{vir}} = \frac{GM\bar{m}}{5k_BR}$. If $T < T_{\text{vir}}$, we should have collapse (U dominates). Typical numbers $\implies T_{\text{vir}} \sim [60{,}270]\,\text{K}$, so $T < T_{\text{vir}}$.

The virial parameter: The virial theorem implies $1 = -\frac{2\langle K\rangle}{\langle U\rangle} = \frac{-2\cdot 5\cdot \frac{1}{2}M\langle v^2\rangle R}{3GM^2\bar{m}}$, but $3\langle v_x^2\rangle = \langle v^2\rangle$ and then we can define $\alpha_{\text{vir}} = \frac{5\langle v_x^2\rangle R}{GM}$. The velocity dispersion (i.e., **turbulence**) works against gravity to prevent the gravitational collapse of a cloud just like the thermal pressure.

Jeans density and mass: if the virial theorem doesn't apply (collapse), we can calculate the change in energy. $dU = -\frac{dU}{dR}dR = -\frac{3}{5}\frac{GM^2}{R^2}dR$ (minus: collapse);
$dK = dW = pdV = nk_BT4\pi R^2 dR = \frac{3Mk_BT}{\bar{m}R}dR$ and then, $dE = dU + dK = \left(-\frac{3GM^2}{5R^2} + \frac{3Mk_BT}{\bar{m}R}\right)dR$. The term inside the parenthesis should be less than zero for collapse $\implies M > \frac{5k_BTR}{G\bar{m}}$, the Jeans mass. This mass also determines a density $\rho > \frac{3}{4\pi M^2}\left(\frac{5k_BT}{G\bar{m}}\right)^3 := \rho_J$, which for the typical numbers is $n_J = \rho_J/\bar{m} \approx 4\cdot 10^6\,\text{m}^{-3}$, lower than the typical observed number density. This means that the Jeans mass is a necessary but not sufficient condition for collapse.

Fragmentation: suppose a quasi-uniform cloud, containing some areas with more density than the rest (lumps), of $M \sim M_\odot \implies n_J \sim 10^{12}\,\mathrm{m}^{-3}$, bigger than the density of the lump itself; so that the lump can't collapse. However, as the cloud as a whole collapses, the lump reaches $n > n_J$ and collapses on its own. This is a fragment.

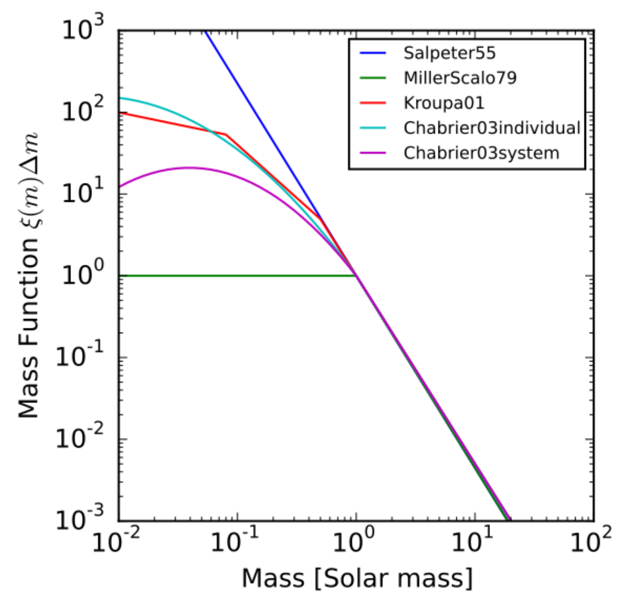
Free-fall timescale: from the hydrostatic equilibrium analysis, we get for the non-equilibrium case, $dM\frac{d^2r}{dt^2} = -dP - \frac{GM}{r^2}dM$, where the acceleration is only radial. If we turn the pressure off suddenly, we get a free fall collapse. Then, $\frac{d^2r}{dt^2} = -\frac{GM}{r^2}$. Now, we can approximate crudely, $\frac{R}{\tau_{\mathrm{ff}}^2} \sim \frac{GM}{R^2}$. This means that $t_{\mathrm{ff}} \sim \sqrt{R^3/(GM)} \sim (G\rho)^{-1/2} \sim 8 \cdot 10^{13}\,\mathrm{s} \sim 3 \cdot 10^6\,\mathrm{yr}$, lowest n; $\sim 250\,000\,\mathrm{yr}$ for highest n. The infall rate in a molecular cloud can be estimated as $\dot{M} = M/t_{ff}$.

Gas cooling: collapse $\Longrightarrow$ higher kinetic energy $\Longrightarrow$ higher temperature $\Longrightarrow$ higher pressure; this pressure tends to halt collapse. In order to continue the collapse, the cloud needs mechanisms to get rid of energy (cooling). Some mechanisms: a) energy breaks (dissociates) H_2 molecules (4.5 eV/molecule); b) atoms get ionized (H ioniz.: 13.6 eV/atom); c) electrons in atoms get excited (without ionizing them), this energy is emitted as light and escapes if the gas is optically thin; d) molecules can get to vibrational/rotational modes and then decay and emit light as well. So, if collapse is happening at this stage, the temperature can be regarded as constant and the process is isothermal.

Other factors that prevent collapse: external heating, turbulence (it's hard to get rid of turbulent kinetic energy), magnetic field (due to magnetic tension).

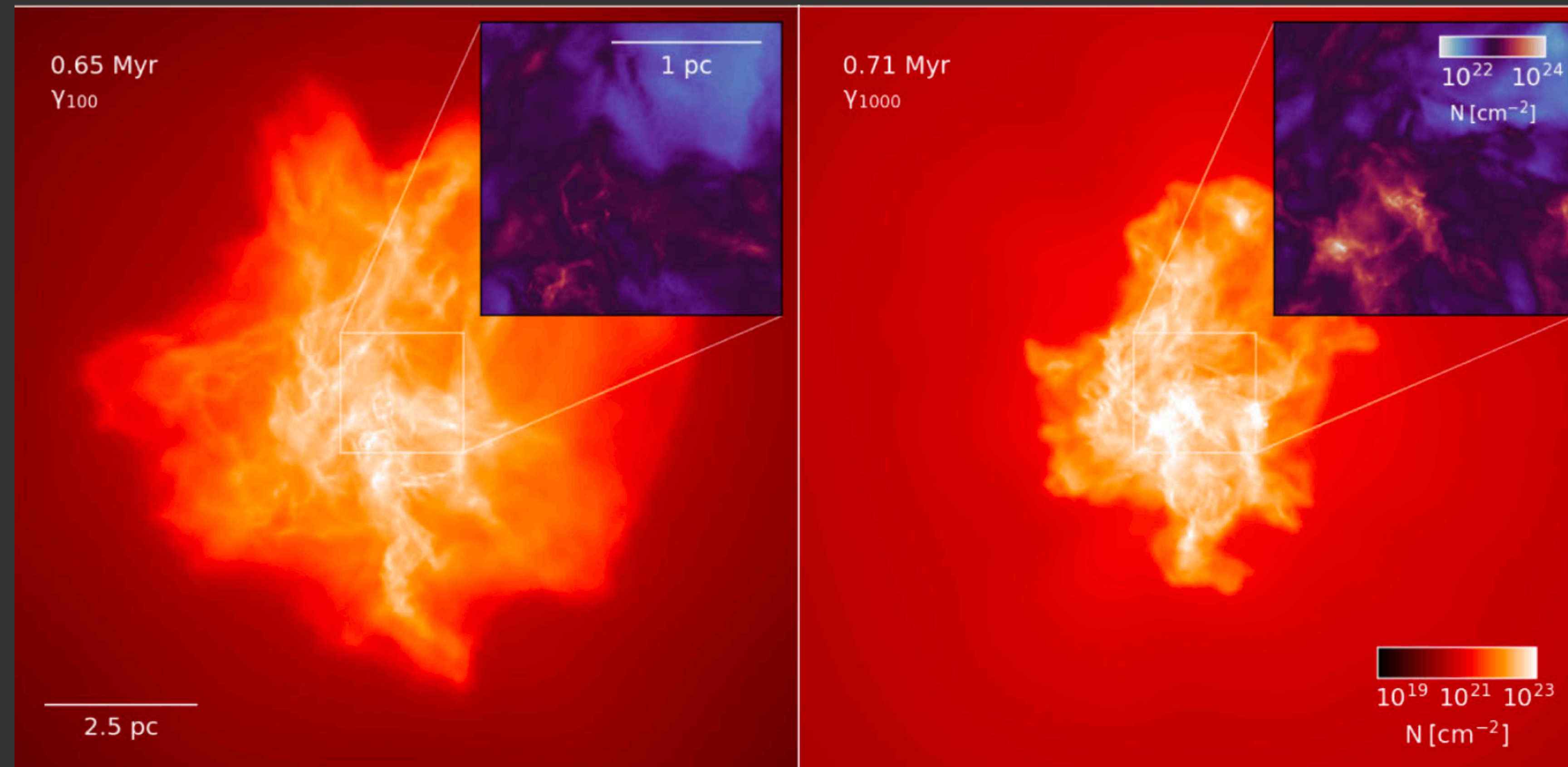
Eddington luminosity: radiation pressure can be a collapse-halting agent. Given a luminosity L, the energy flux is $L/(4\pi r^2)$; $\mathscr{D}[\text{energy flux}] = \mathscr{D}[\text{energy}]/(L^2T)$. Since a photon carries energy $h\nu$, the photon flux (number of photons/area-time) is $L/(4\pi r^2 h\nu)$. If we multiply this by the Thomson cross section σ_T ("area of one electron as seen by e.m. radiation"), we will have the number of photons/time that interact with a given electron: $L\sigma_T/(4\pi r^2 h\nu)$. Each photon carries momentum $h\nu/c$, so the force perceived by an electron is is $F_{\text{light}} = \frac{h\nu}{c} \cdot \frac{L\sigma_T}{4\pi r^2 h\nu} = \frac{2e^4L}{3m_e^2c^5r^2}$. Gravity is more significant on protons, since they are more massive, so, $F_{\text{grav}} = GMm_p/r^2$. If both forces are equal (consider a portion of gas), we get $L_E = \dfrac{3GMm_p m_e^2 c^5}{2e^4}$, the Eddington luminosity, which is the maximum luminosity for a star with a given mass. *Examples*: a) the Sun has much lower luminosity than the Eddington value, so, radiation pressure is not important; b) an O star of $60M_\odot$ and $8 \cdot 10^5 L_\odot$ has a ratio $L/L_E = 0.4$, so radiation pressure is not negligible.

Initial mass function: relative number of stars per mass interval at birth $\xi(M_\star) = \Psi(M_V)dM_V/dM_\star$ (Ψ is the initial luminosity function). This can be thought as a result of the star formation process that is useful for modeling large molecular clouds or galaxies because massive stars affect their environment (stellar feedback) and it's not the same to have 100 stars of one solar mass (no feedback, long lives) and to have 2 stars of 50 solar masses (feedback, short lives).



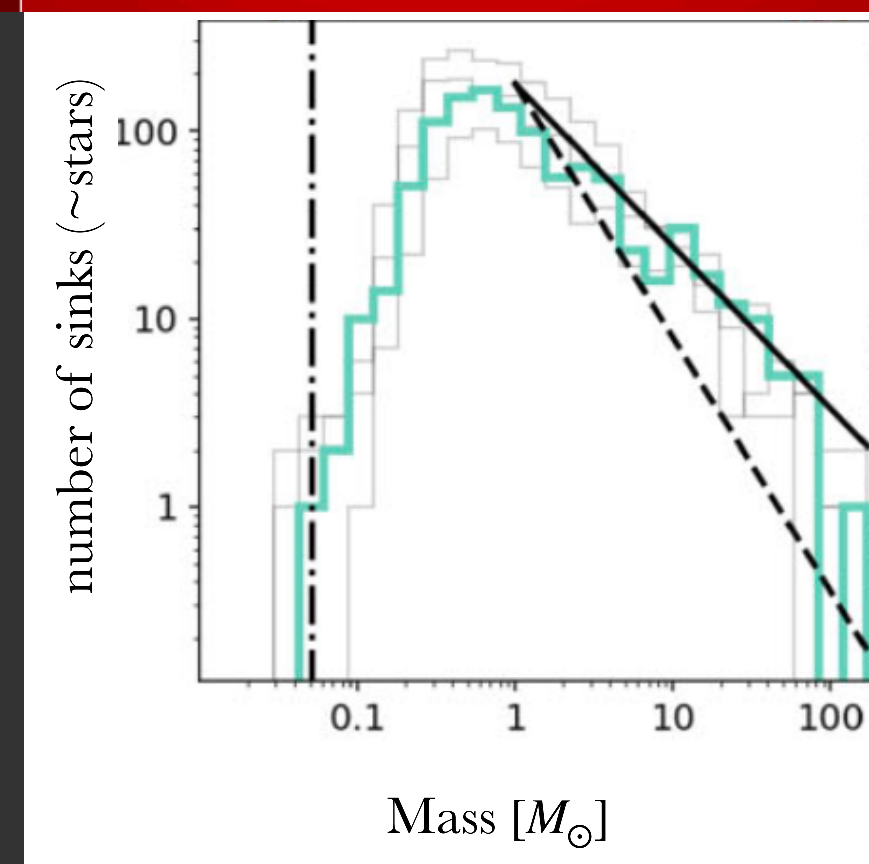
Star formation rate: the amount of mass that is converted into stars ($M_\odot$/yr). This is useful for modeling galaxies.

Metallicity: the proportion of "metals" present in an environment, that is, the amount of elements heavier than helium. Nucleosynthesis of metals occurs inside of stars (light elements are created in low-mass stars, while heavy elements can only be created inside of massive stars, supernovae explosions and neutron star mergers). Metallicity increases as successive generations of stars die and enrich the environment. Primordial star formation is star formation in the early universe, where the primary difference is the almost zero metallicity.



Cloud core formation simulations

These simulations study the formation of cores in a molecular cloud via cloud fragmentation. The cloud initially has a turbulent velocity field. The plot on the right shows the number of sinks formed (roughly equivalent to the number of stars formed) with which masses (=initial mass function). The aim of the paper is to study star formation in *starburst galaxies* (galaxies with an unusually high star formation rate).



Cusack, Clark, Glover 2025 A&A

Inside of a forming star

First Larson core: in a collapsing core, the point in which the dust becomes opaque to its own radiation (radiation cannot escape).

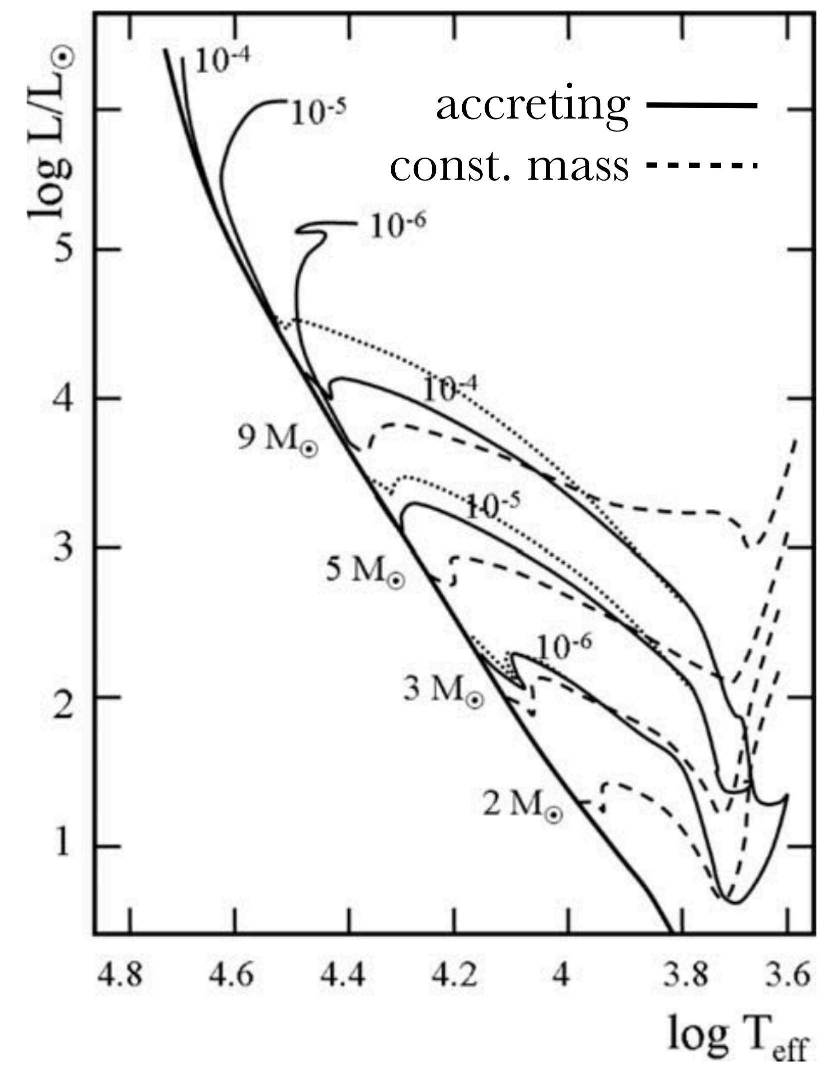
Second Larson core: molecular gas is made from molecular hydrogen. At a temperature of ≈ 2000 K, H_2 dissociates into H. Gravitational energy from the collapse is used to dissociate hydrogen instead of increasing thermal pressure, and then the gravitational collapse continues until it is stopped by an atomic hydrogen core. This is the second Larson core

Introduction to the HR diagram: A Hertzprung-Russel diagram is simply a plot of stellar luminosity against its surface temperature. For historical reasons, the temperature increases from right to left. We will encounter this when studying stellar evolution.

Protostellar evolutionary tracks:

Initial parameters: mass, chemical composition, l/H (=mean free path of largest convective elements/pressure scale height), initial structure (e.g. a polytrope with index 2.5).

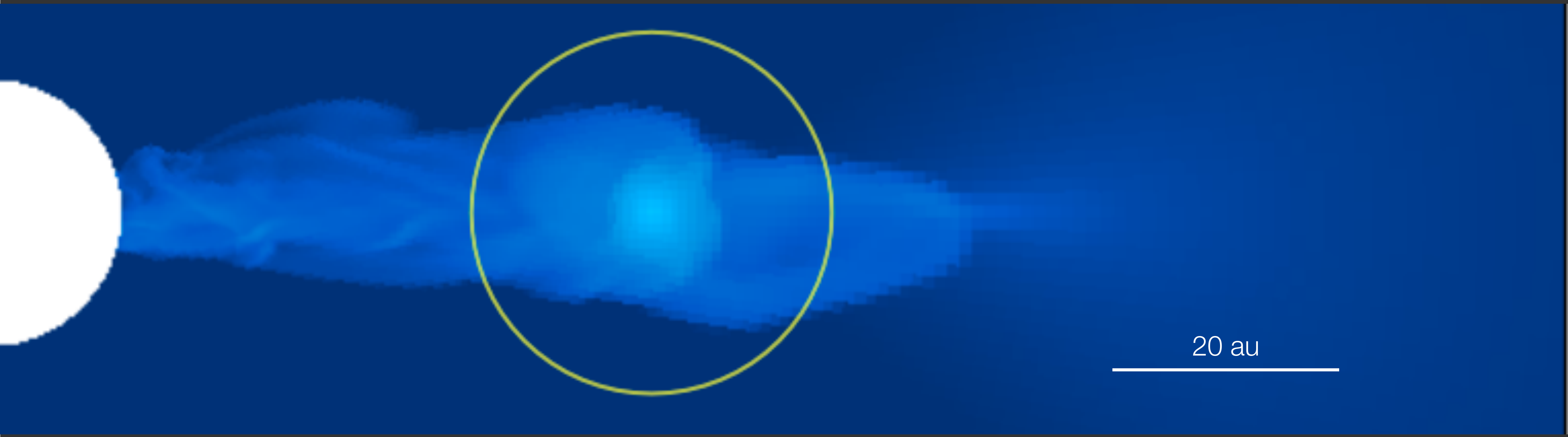
Hayashi track ($<0.5\ M_{\odot}$): Fully convective phase. The evolution in the HR diagram is vertical and downwards. Energy transport is quite efficient in the interior; the rate of energy loss is determined by a thin radiative layer at the surface. The evolution in the track assumes no accretion but instead a contraction of the initial mass.



Accreting vs non-accreting tracks: A non-accreting protostar (equilibrium assumption) has a forbidden zone at the right of the Hayashi track in the HR diagram, but an accreting protostar has no such limitation. In the diagram, the solid line is the birth line of accreting protostars, and the dashed lines are tracks assuming direct contraction from an initial, constant mass. (From Maeder 2009, Springer)

Henyey track (>0.5 $M_\odot$): Radiative phase. The evolution in the HR diagram is relatively horizontal. Contraction $\Longrightarrow$ T interior increases $\Longrightarrow$ opacity decreases (see Kramer's opacity law), the interior becomes radiative (convection stops). Luminosity is no longer controlled by the surface layer, but by the opacity of the whole radiative region. Sharp bend to the left in the evoluc. track. During the track: short contraction times, approximately the Kelvin-Helmholtz timescale (relatively constant luminosity).

Transition to the main sequence: When H starts burning (main sequence), the contraction stops (and for high masses, the luminosity increases gradually, for a solar mass track, the luminosity decreases slightly).



Formation of a First Larson core

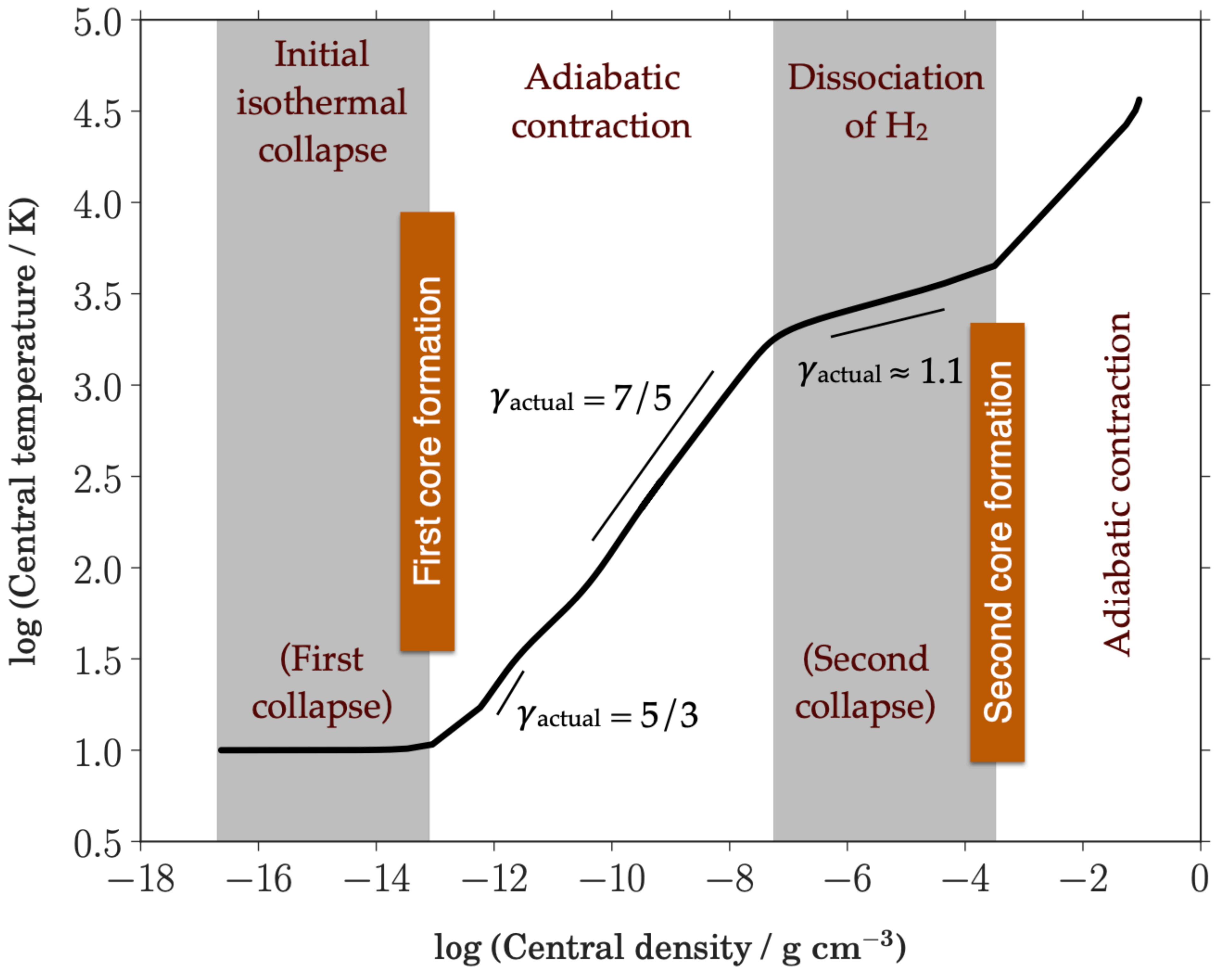
This simulation shows the formation of a spherical-like overdensity in an accretion disk that becomes gravitationally bound and collapses into a first Larson core.

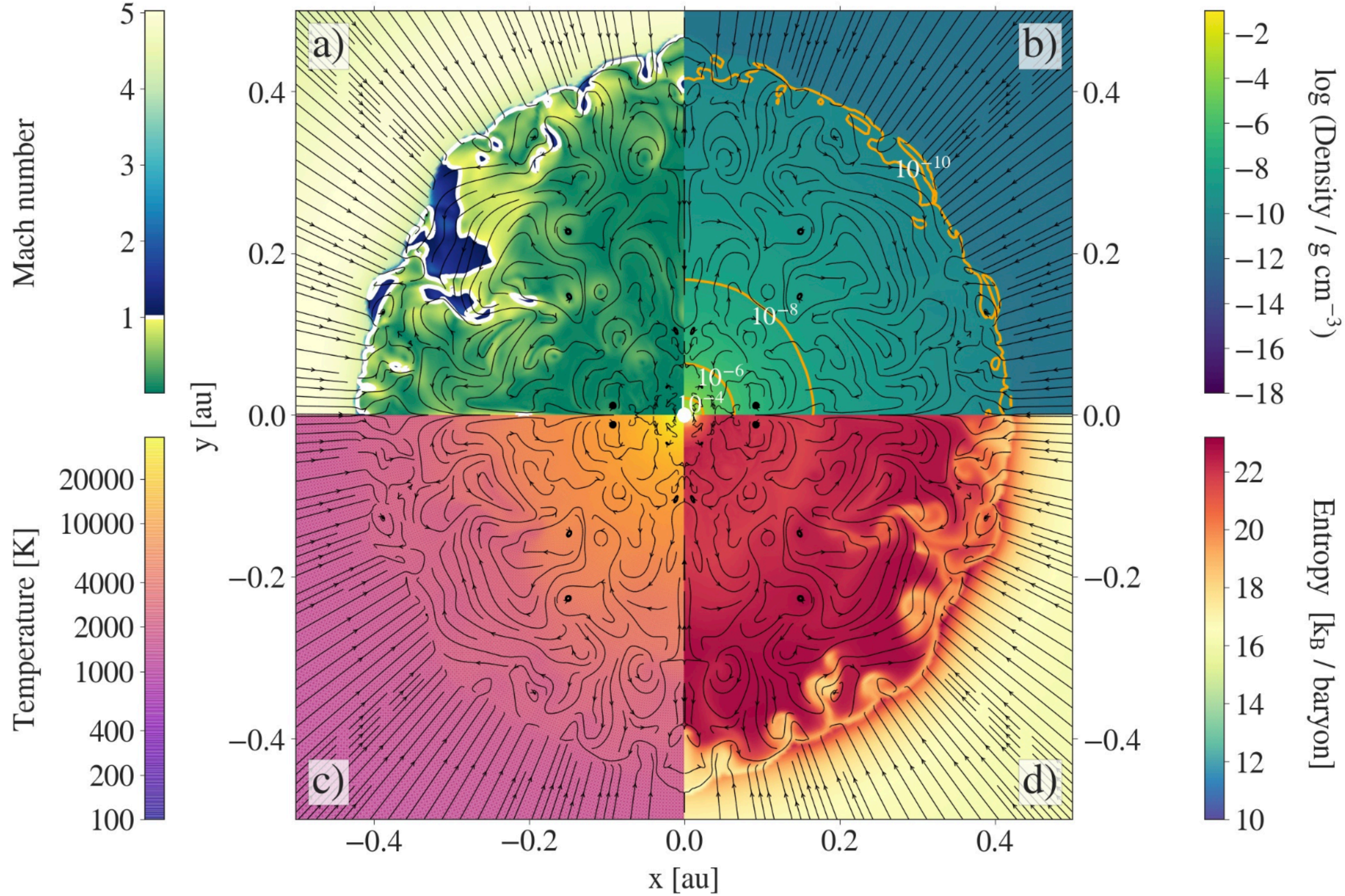
Oliva & Kuiper 202 A&A

Thermodynamics during the collapse

The simulations in this paper allowed for the calculation of the polytropic index of the thermodynamical processes during the formation of the first and second Larson cores. At first, the cloud has a constant temperature of ~10 K (efficient cooling). When the first core forms, it becomes opaque (radiation is trapped), so the process is adiabatic. During the dissociation of hydrogen, the core is able to collapse because temperature is not rising quick enough for the thermal pressure to stop gravity. Finally, the collapse continues adiabatically during the pre-main-sequence.

Bhandare et al 2018 A&A

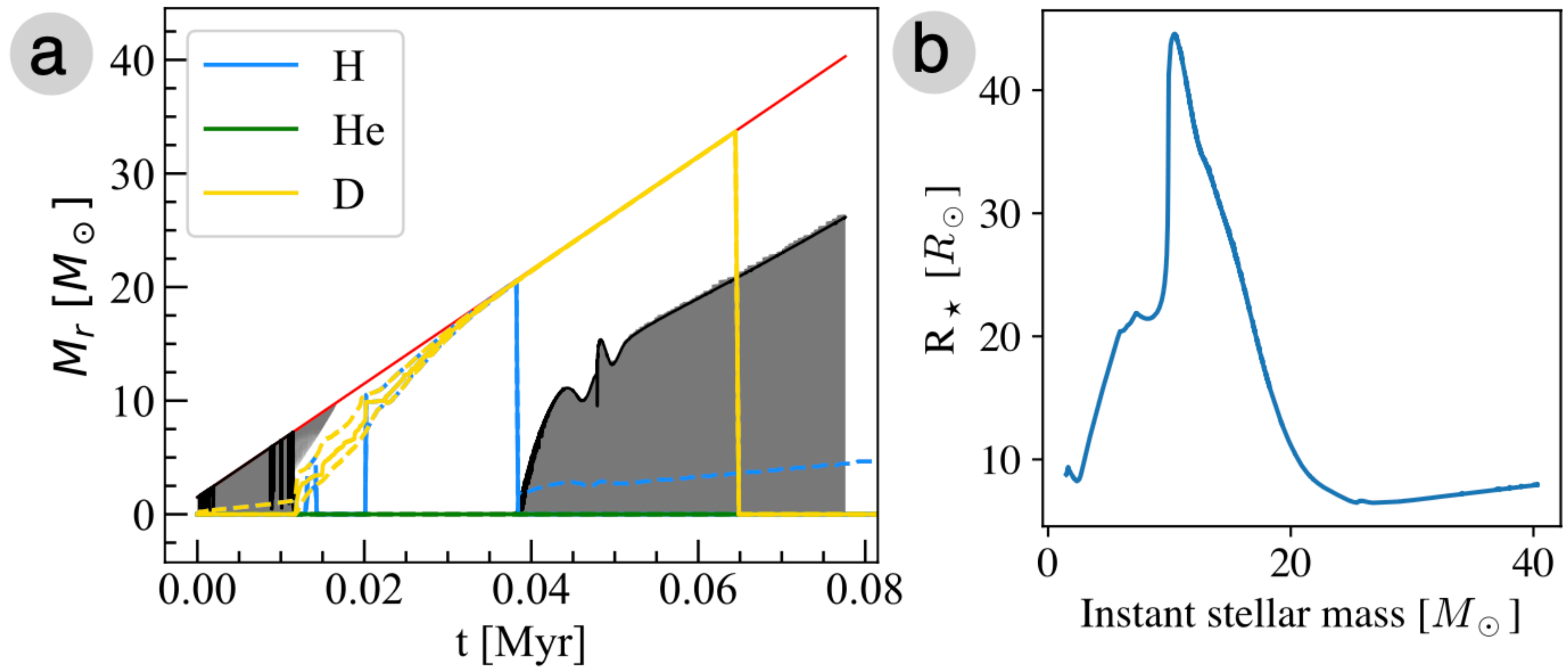




Structure of the second Larson core

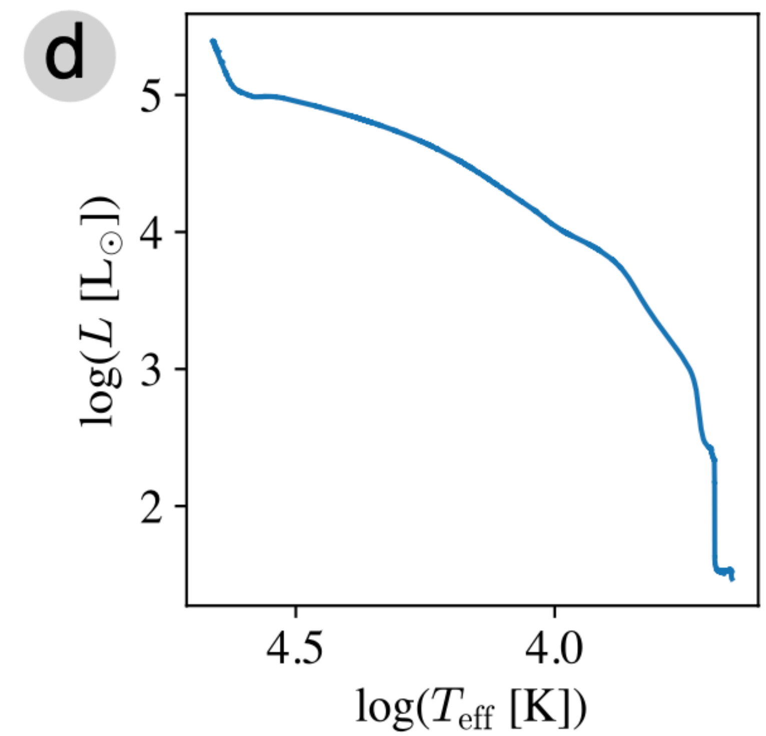
This simulation describes the interior of a second Larson core. It starts from the collapse of a cloud.

Bhandare et al 2020 A&A



The pre-main sequence evolution of a massive star

(a) Shows the interior of a forming massive star accreting from a disk. The vertical axis is the enclosed mass (the total mass is the red line, increasing in time as accretion goes on). The shadowed parts indicate a convective structure and the unshadowed parts a radiative structure. First, deuterium burning starts (D, yellow), and then hydrogen (H, green), point at which the star reaches the main sequence. Low mass stars stop their accretion at the arrival to the main sequence, but accretion rates are so high in massive star formation that accretion continues. (b) When deuterium starts burning, the stellar envelope is heated and the radius of the star expands. After the excess heat is radiated away, the radius decreases (arrival to the main seq.). The *central* values of density and temperature (not shown) increase all the time during this process. (d) HR diagram of the pre-main seq. (birth line).



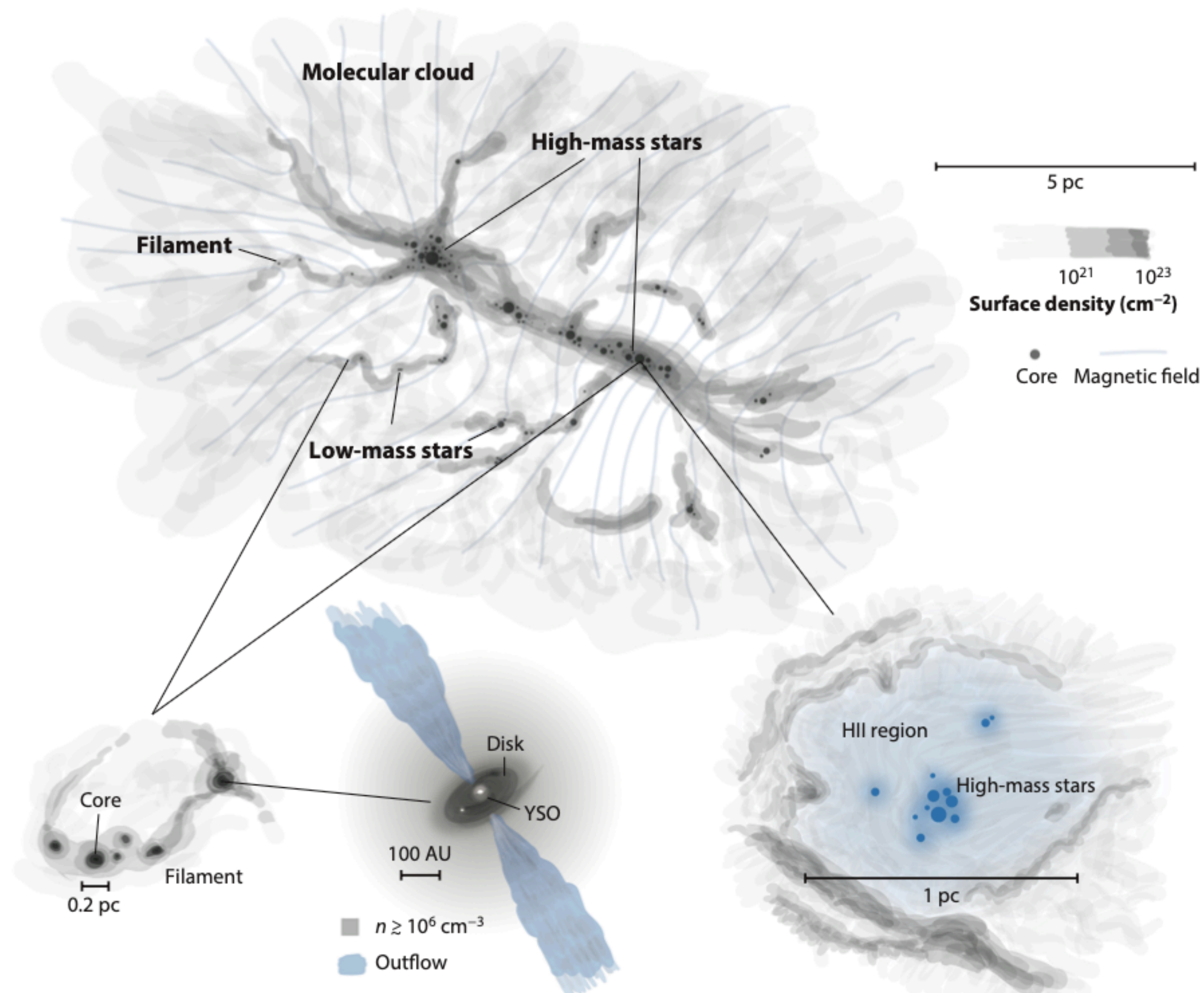
Oliva et al (in prep.)

Low vs high-mass star formation

Differences between massive and low-mass star formation: From Beuther, Kuiper, Tafalla 2025:

1. Turbulence is not different between low and high mass star formation (massive star forming regions are not more turbulent; the assumption was commonly made for explaining how large amounts of mass were able to accumulate to form massive stars and because of observations of regions with ongoing SF, but not applicable to initial conditions for SF)
2. Mean densities for massive SF regions are higher $\Longrightarrow$ smaller Jeans length $\Longrightarrow$ smaller core fragmentation separation. This means that multiplicity strongly increases from low- to high- mass stars
3. The biggest difference is that massive protostars produce strong ionizing radiation that impacts the environment through constructive and destructive feedback.

Term	Mass	Observations
Low-mass star	$< 8M_\odot$	Solar-like star
Massive star	$> 8M_\odot$	OB star, massive enough to produce a type II supernova. (O,B0, B1)
Very massive star	$[10^2,10^3]M_\odot$	R136a1 (291 $M_\odot$), BAT99-98 (226 $M_\odot$), R136a2 (195 $M_\odot$), etc.
Ultramassive star	$[10^3,10^4]M_\odot$	Unlikely to be formed in the present-day universe (probably in early universe)
Supermasive star	$[10^4,10^8]M_\odot$	equilibrium dominated by radiation pressure. Collapse due to GR instability. Also unlikely to be formed in the present epoch but maybe in the early universe.



Stellar feedback: radiation pressure

The radiation pressure classical problem:

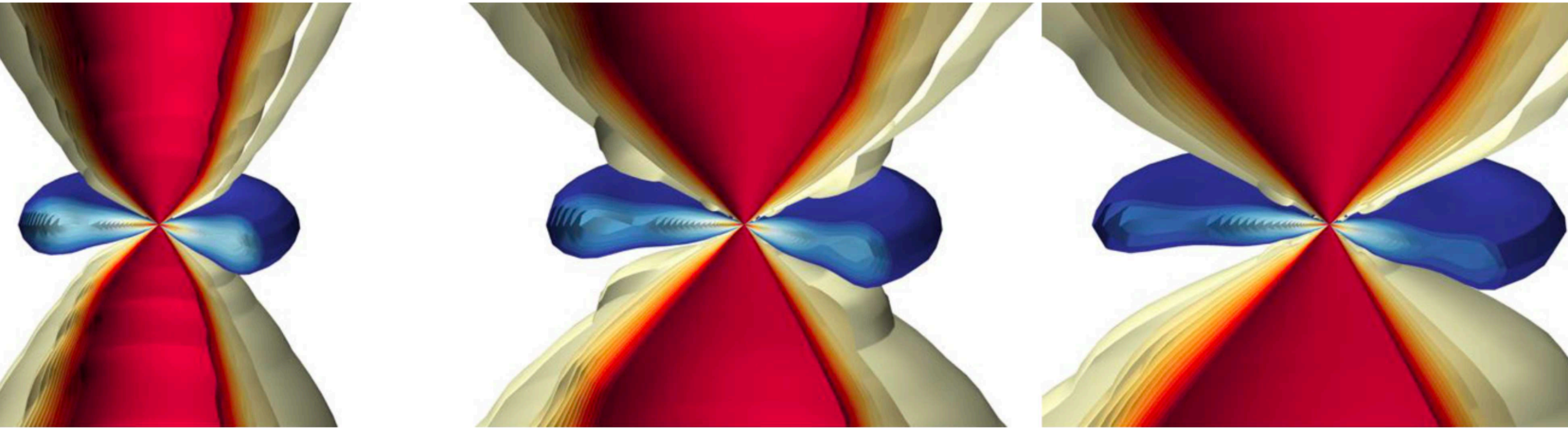
When the main sequence is reached, radiation from the core surface shifts to the UV region, producing ionizing photons. These photons may ionize the dusty infalling gas, and this effect must be taken into account since it slows down infall. Opacity of the dust in the infalling envelope: {UV, optical: high; IR: lower}, however, dust species evaporate above 1500 K (dust destruction front; eg. ~ 8 AU for $L \sim 1000 L_\odot$, scales as $L^{1/2}$).

Opacity: We define the mean free path of light in a given frequency as $\ell_\nu = 1/(\kappa_\nu \rho)$, where κ_ν is called opacity $\mathscr{D}[\kappa] = L^2 M^{-1}$. This is analogy to particle collisions, where $\ell = 1/(\sigma n)$, where σ is the scattering cross section and n the number density of particles. Also, from [Notes on Astrophysics § Radiative transfer ¶ Radiation pressure], we know that $dE_\nu = F_\nu dA dt$, which implies that the momentum is $dE_\nu / c = F_\nu dA dt / c$. Pressure $=$ momementum$/(dAdt)$ $= F_\nu / c$.

Gravity against radiation pressure:
$\dfrac{GMdm}{r^2} = dPdA \implies \dfrac{GM}{r^2} = \dfrac{dP\cancel{dA}}{\rho dr \cancel{dA}}$. Taking dr as the free path of a light ray, we have $\dfrac{GM}{r^2} = \dfrac{\bar{\kappa} F}{c}$, where $\bar{\kappa} = \dfrac{1}{F} \displaystyle\int_0^\infty \kappa_\nu F_\nu d\nu$

Mass limit for a star in a dust envelope: since $L_\star = 4\pi r^2 F$, we have $\dfrac{L_\star}{M_\star} = \dfrac{4\pi G c}{\bar{\kappa}}$. For dust at $T \sim [300{,}1000]\,\mathrm{K}$, $\bar{\kappa} \approx 0.8\ \mathrm{m/kg}$, which gives us $L_\star / M_\star = 0.31429$, or expressed in solar units, $(L_\star / L_\odot)/(M_\star / M_\odot) \approx 1\,600 L_\odot / M_\odot$ which corresponds to a main-sequence star of about $15 M_\odot$. According to this calculation, no massive stars should be formed! Disregarding dust, the mass limit, which becomes the Eddington limit, becomes $\sim 200 M_\odot$. If the star is accreting, the Eddington limit also limits the accretion rate. Comparing the spherical accretion luminosity $L_{\mathrm{acc}} = GM\dot{M}/R$ to the Eddington limit yields $\dot{M} \approx 10^{-2}\,\mathrm{M_\odot/yr}$ for $100\,M_\odot$.

The solution to the radiation pressure problem are accretion disks, which enable continued accretion while allowing radiation pressure to escape through the polar regions.



Feedback escapes through the poles while the star accretes from the disk

These simulations show how the outflows caused by the radiation from the star (stellar feedback) escape through the rotation axis while accretion continues through the midplane. This enables protostars to grow to high masses.

Kuiper & Hosokawa 2018 A&A

Kuiper, Klahr, Beuther & Henning 2014 Springer

Stellar feedback: photoionization

HII regions: a UV photon with $\lambda < 912\,\text{Å}$ can ionize a hydrogen atom from the ground level ($n = 1$). O-B stars (high enough temperature), emit UV photons, and ionize the ISM around them, forming HII regions. Typical temperatures: 6000 K.

Strömgren spheres: consider a sphere of radius R_S (Strömgren sphere) where there is ionization and recombination. Recombination rate $\propto n_p$ and $\propto n_e$, so, recombination rate = $\alpha n_p n_e$, where α is the recombination coefficient ($n = N/V$). In a steady state, number of ionizations = number of recombinations, and so, the total number of ionizations within the Strömgren sphere is $(4\pi/3)R_S^3 \alpha n_p n_e$. The star emits N_γ ultraviolet photons that produce the ionizations, so, $N_\gamma = (4\pi/3)R_S^3 \alpha n_p n_e$.

Emission from recombination: if in the recombination, the free electron is captured and jumps to $n = 1$, then a UV photon is emitted. However, this can also happen in stages: first, the free electron is captured to $n > 1$, and then it jumps again to $n = 1$. This produces radiation in visible light and even in radio. Hot gas from HII regions also emits bremsstrahlung.

Hypercompact HII regions: when a massive protostar gains enough mass to go into the main sequence it's still accreting, but starts emitting ionizing radiation. The ionized region is small, because of the infalling gas. Typical scales: size < 0.01 pc (=2000 au); number density $\sim 10^6\ \text{e}^-/\text{cm}^3$. It can be roughly obtained by thinking of the gravitational radius (radius for which sound waves can't escape) $r_g = GM_\star/c_i^2$, where $c_i \approx 10\,\text{km/s}$ is the sound speed of the ionized gas.

Ultracompact HII regions: as the collapse from the cloud progresses, the density of the infalling envelope decreases (ram pressure decreases) and the star, having gained more mass, increases the ionizing flux. Typical scales: size $\sim 0.1\,\text{pc}$ (size of the cloud core), number density $\sim 10^4\ \text{e}^-/\text{cm}^3$. These scale can be derived from the Strömgren radius.

Later stages: when $R_S \gg r_g$, the HII region is not gravitationally bound and it begins to ionize the molecular cloud material in the neighborhood. The disk becomes photoionized and starts to photoevaporate, which may have an impact on the final mass of the star. If the HII region is too big (escape from the cloud core), the high temperatures (~8000 K) expand the gas further and rapid mass loss from the cloud can occur, even destroying the star-forming region.

Primordial star formation

Formation of the first stars, hydrogen: in the very early universe, the environment conditions are very different because dust grains were not present. The formation and excitation of H_2 was a crucial cooling mechanism to lower the Jeans mass enough to form stellar-mass objects. For the first stars, the hydrogen gas was recombined (=neutral). Hydrogen molecules were not formed by direct attachment of two H atoms, but by leftover reactions from the recombination epoch such as $H^- + H \rightarrow H_2 + e^-$ (although free electrons were rare and only a small quantity of molecular hydrogen was produced this way below densities of 10^8 cm^{-3}). Above the density of 10^8 cm^{-3}, three-body interactions are possible and H_2 can be formed by the interaction of 3 H atoms. Above 2000 K (high enough density), molecular hydrogen dissociates and cooling occurs.

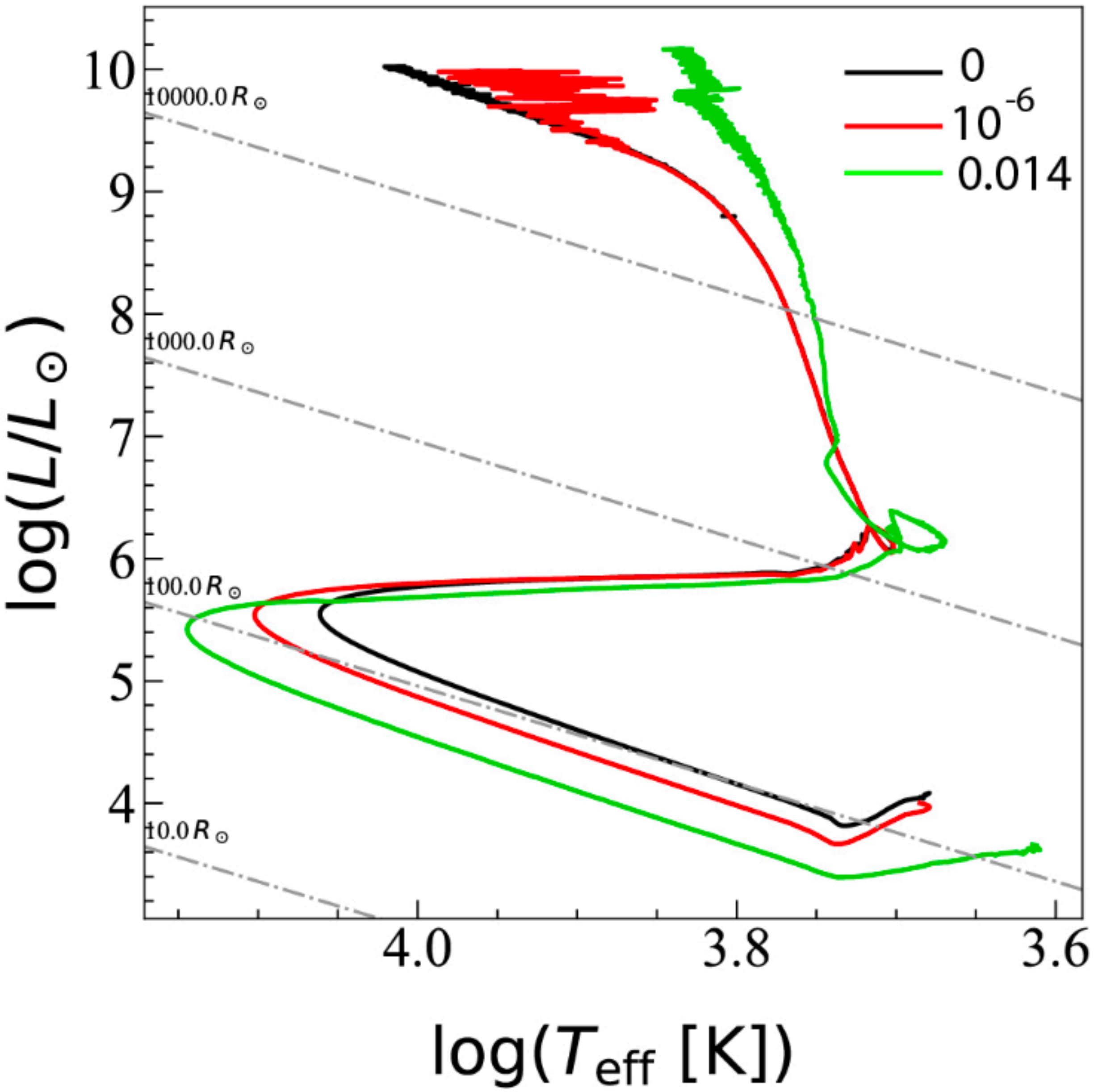
Formation of the first stars, fragmentation: the Jeans mass for the conditions of the formation of the first stars ($T \approx 200\,\mathrm{K}$, $n \approx 10^4\,\mathrm{g\,cm^{-3}}$) is about 1000 $M_\odot$. A key question in primordial star formation is whether the cloud core formed this way will fragment or form stars of such masses.

Formation of the first stars, stars: the 1000 $M_\odot$ baryonic matter cloud core is surrounded by a dark-matter halo. After the gravitational collapse is triggered (assuming no fragmentation), and H_2 is dissociated, matter free-falls and temperatures reach $\sim 20\,000\,\mathrm{K}$ (formation of a core). Accretion rates are of the order of a few $10^{-2}\,\mathrm{M_\odot\,yr^{-1}}$. In the core, hydrogen starts to ionize and outside it remains neutral H (although a small H_2 remains to provide enough cooling through dissociation). Accretion continues. Because of the lack of dust (and therefore lower opacity), the Eddington limit is higher than in the solar-metallicity case and higher masses and accretion rates are possible. Opacity is dominated by electron scattering. On the right, a plot of the limiting accretion rate (based on the Eddington limit) as a function of the enclosed mass obtained from simulations of the first stars. The different lines types are estimations by different authors.

HR diagram of a supermassive star

The authors compute the evolution of a supermassive star from its formation with the Geneva stellar evolution code for three different metallicities (zero for the early universe, low metallicity and solar metallicity). They assume that the star never ends accretion. In the present day universe, stellar feedback is strong in part because of the radiation pressure felt by dust particles. With no dust in the early universe, maybe stars grow continuously throughout their life until they become supermassive stars.

Nandal, Zwick et al 2024 A&A

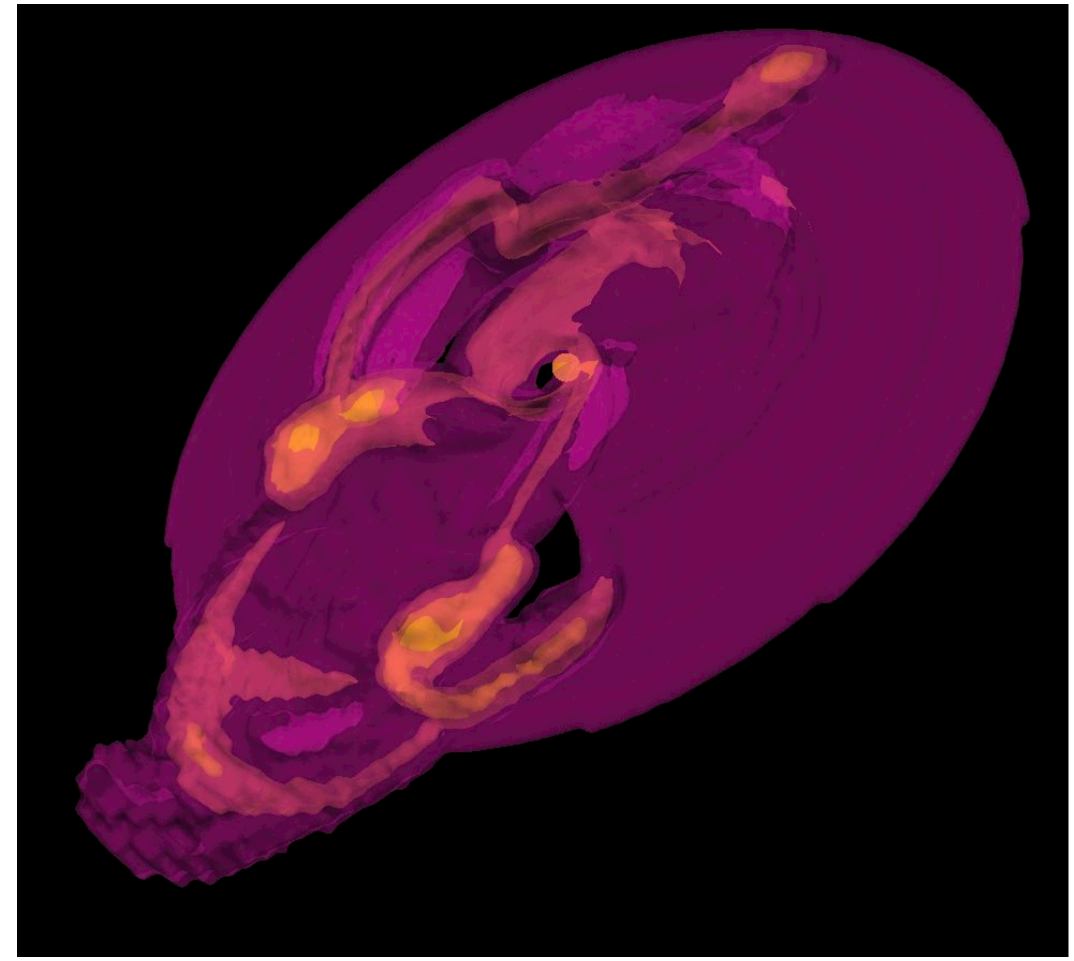


Disks and jets

Disks: when a slowly rotating molecular cloud collapses under the action of gravity, density is redistributed (the "radius" R of the cloud decreases and the center mass increases although the total mass M is constant). This means that the moment of inertia $I \propto MR^2$ of the cloud decreases. By the conservation of angular momentum ($I\Omega = \text{const}$), this means that the material starts rotating faster. When the angular velocity reaches the Keplerian value $\Omega_k = \sqrt{GM/R^3}$, the material starts moving in a circular orbit and forms a disk.

Accretion disks: a purely Keplerian disk would not transport material inwards (Keplerian orbits are circular and stable). Then, additional mechanisms for removing angular momentum from the disk and allowing the inwards transport of material are needed to feed the star. Jets and winds can provide the angular momentum transport but are not the only possible mechanisms in all phases of star and planet formation (self-gravity and several HD and MHD instabilities could also do the job).

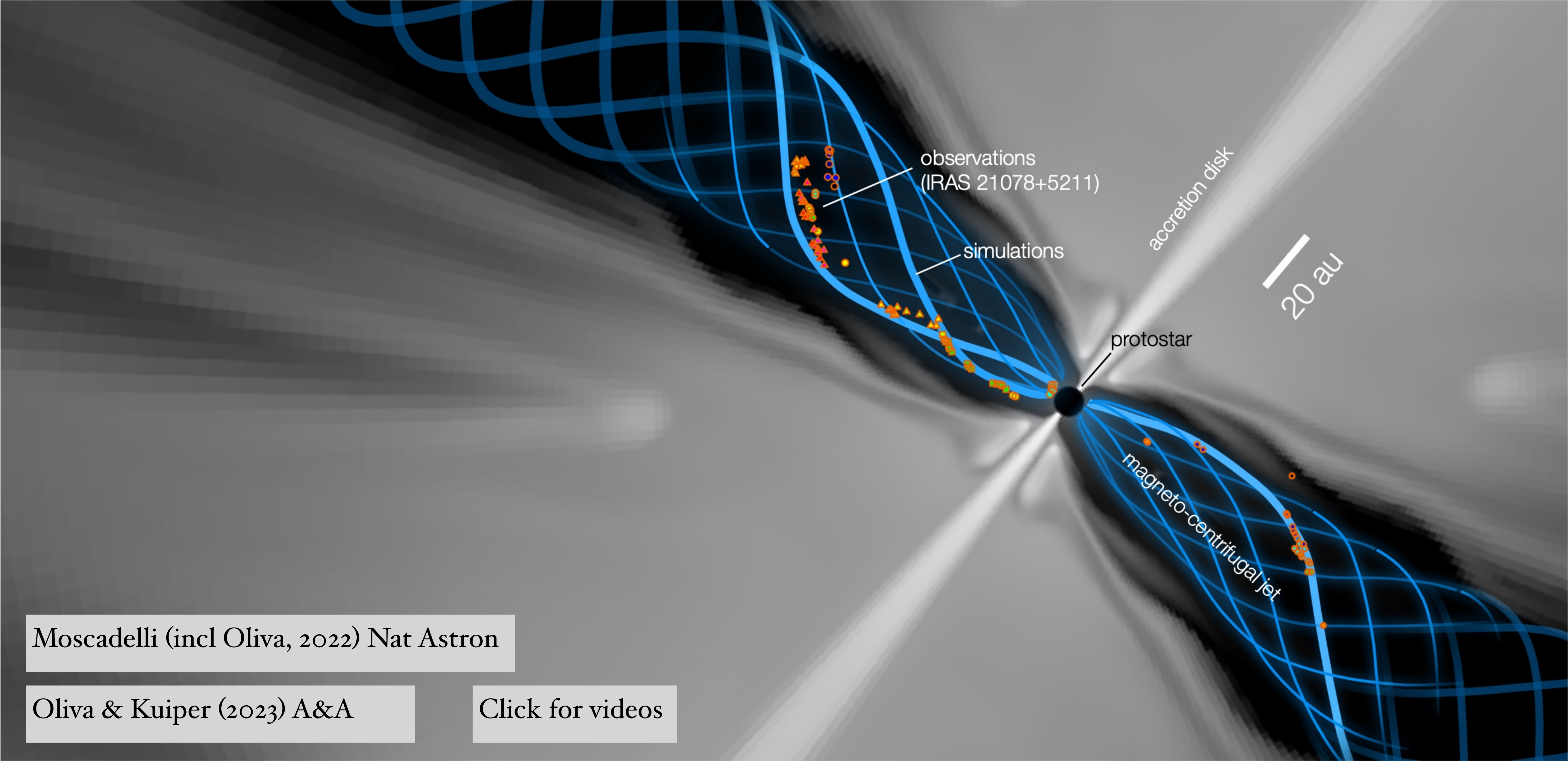
Terminology: around a forming star: *circumstellar disk*. Around a binary: *circumbinary disk.* Planet-forming disks: *protoplanetary disk*. Around a forming planet: *circumplanetary disk*.



Oliva & Kuiper (2020)

Click to see 3D view

This is a simulation of an accretion disk undergoing disk fragmentation (density is shown; lighter color is denser). The central (proto)star is being fed by the disk. The gravitational instability has broken down the disk into fragments that will become companion stars



Magnetically-driven outflows: magnetic field lines present in the cloud twist as the gravitational collapse goes on. The twisting of the field lines creates a magnetic pressure gradient that overpowers gravity and drives a large-scale *magnetic tower flow*. Very close to the protostar, another mechanism, the *magnetocentrifugal mechanism* drives a high-speed jet.

The blue lines show a simulation and the orange dots show observations. In this study, it was shown that protostellar jets are driven by the magnetocentrifugal mechanism. The material around the protostar, close to the rotation axis is both sub-Alfvénic and super-Keplerian (it moves with a speed larger than needed for a circular orbit). This means that the plasma wants to move into a larger orbit (in the direction of the disk) but it is instead forced outwards by the magnetic field lines (following a helical and slightly conical path).

Planet formation

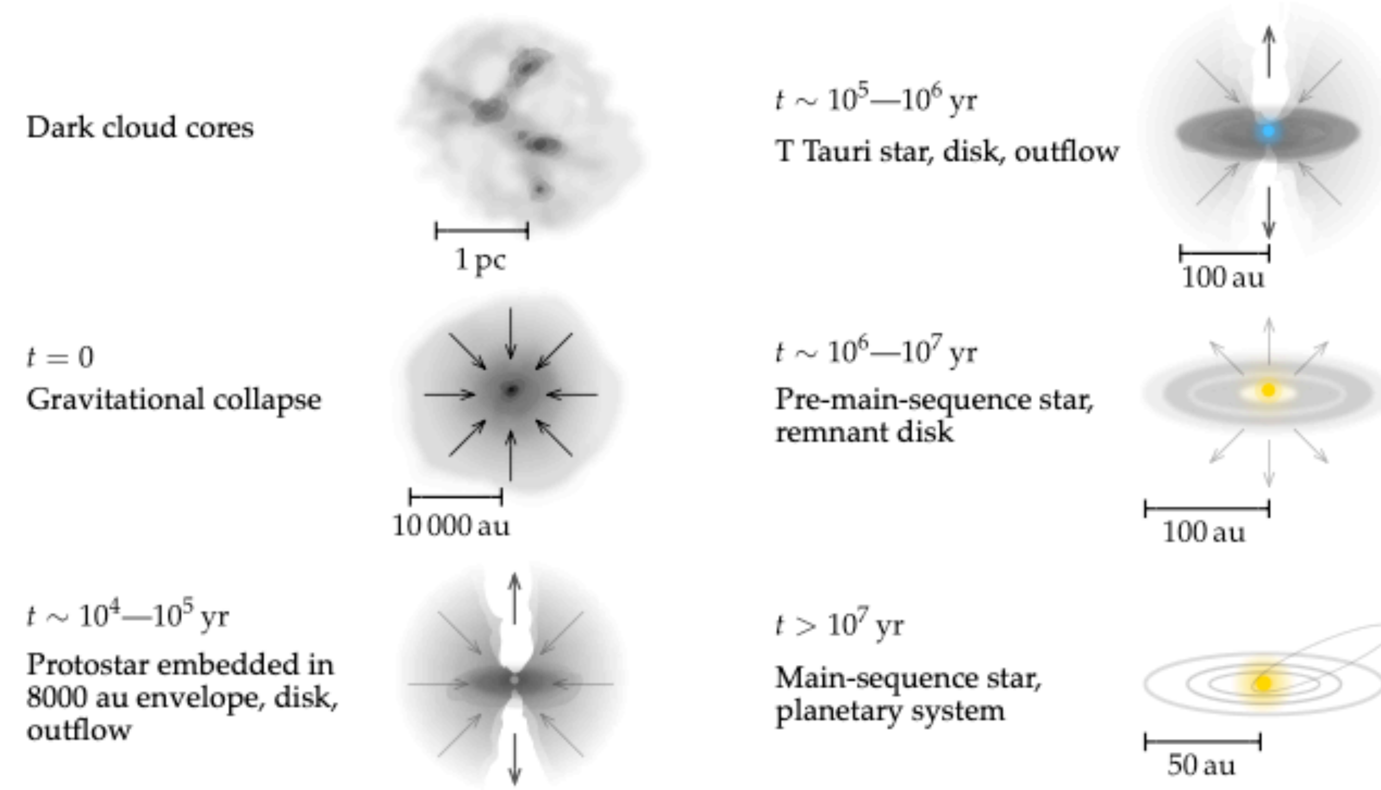
Stages of low-mass star and planet formation: traditionally, and motivated by differences in observed spectra, the different stages of star formation are:

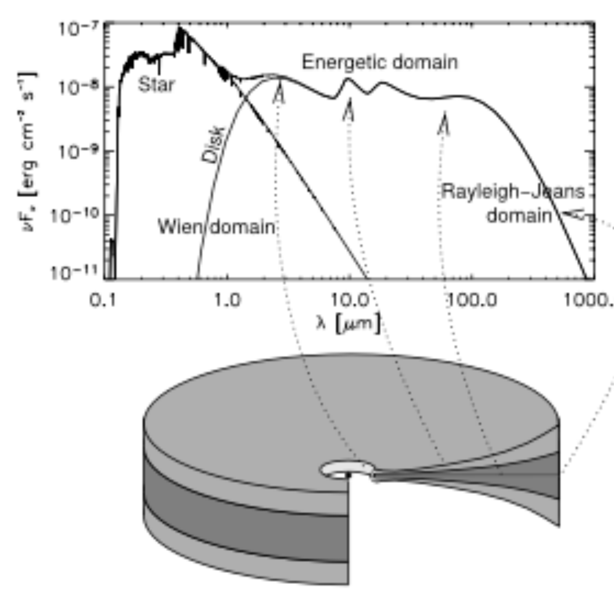
Class-0: collapse and disk and jet formation, very deeply embedded in the cloud.

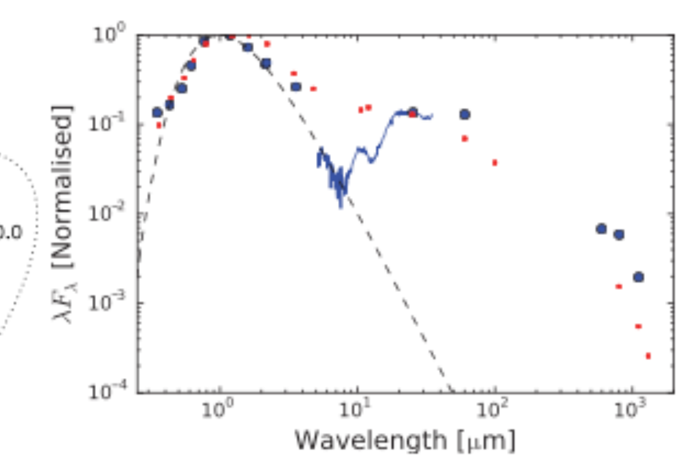
Class I: spectra contain emission from dust grains and show signs of X-ray photoevaporation. This means that in class I disks there is already a young star.

Class II: spectra contain emission from dust grains, a big inner hole and accretion. Current theories of planet formation propose those are remnant disks that contain forming planets. Those planets form gaps. Planets can migrate inwards and outwards in the disk over time.

Class I and class II disks are called *transition disks* (see on the right a diagram of the spectrum and a real spectrum of a protoplanetary disk).



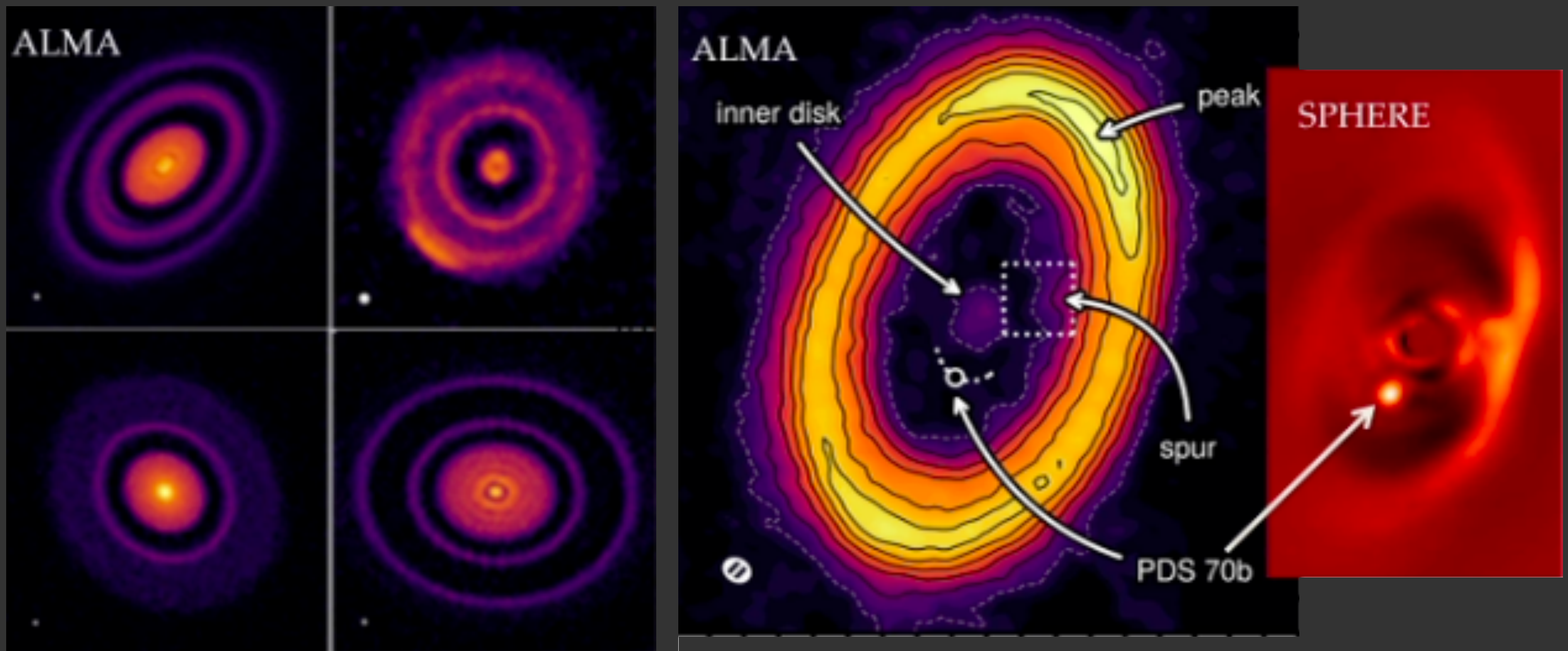




Dulemond et al 2007

Owen 2016

From the introduction of Thomas Rometsch's PhD thesis



Observations of protoplanetary disks

These are disks observed with ALMA (two figures on the left) and the VLT (on the right). Notice the gaps opened by planets. In the case of PSD 70, two giant planet have been directly imaged, while still accreting and still embedded in their protoplanetary disk.