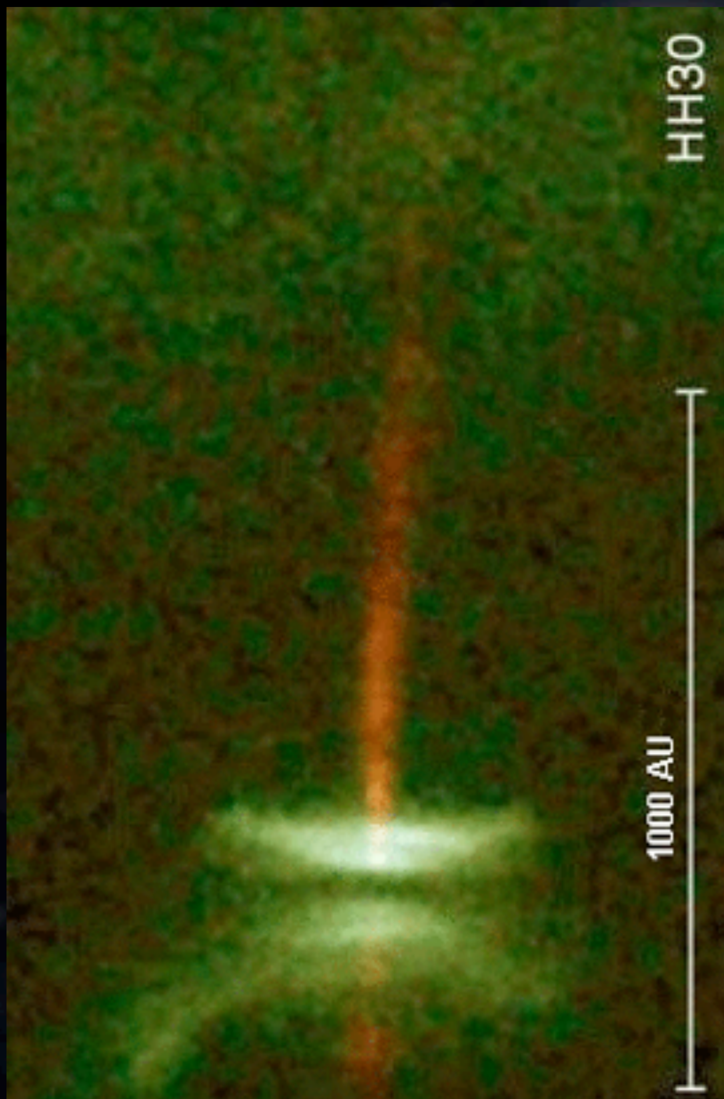
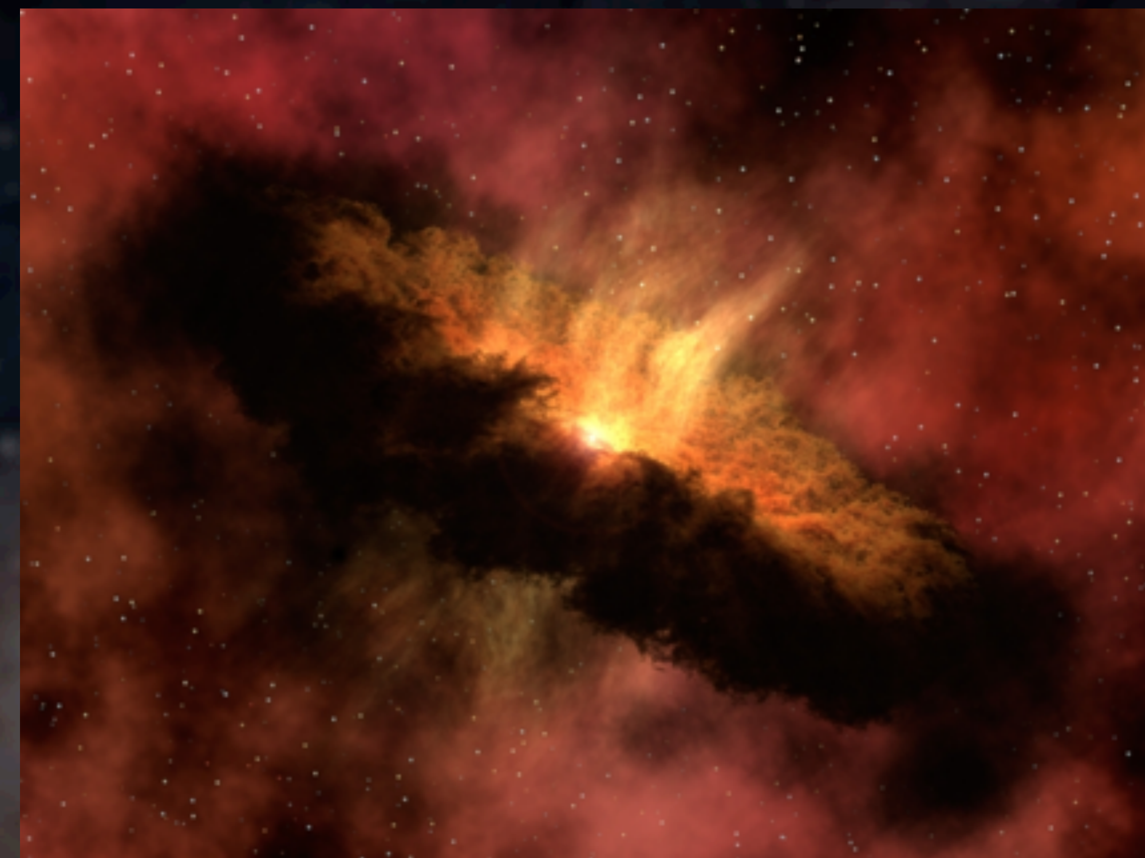
The stability of accretion disks and winds

Geoffroy Lesur (IPAG)

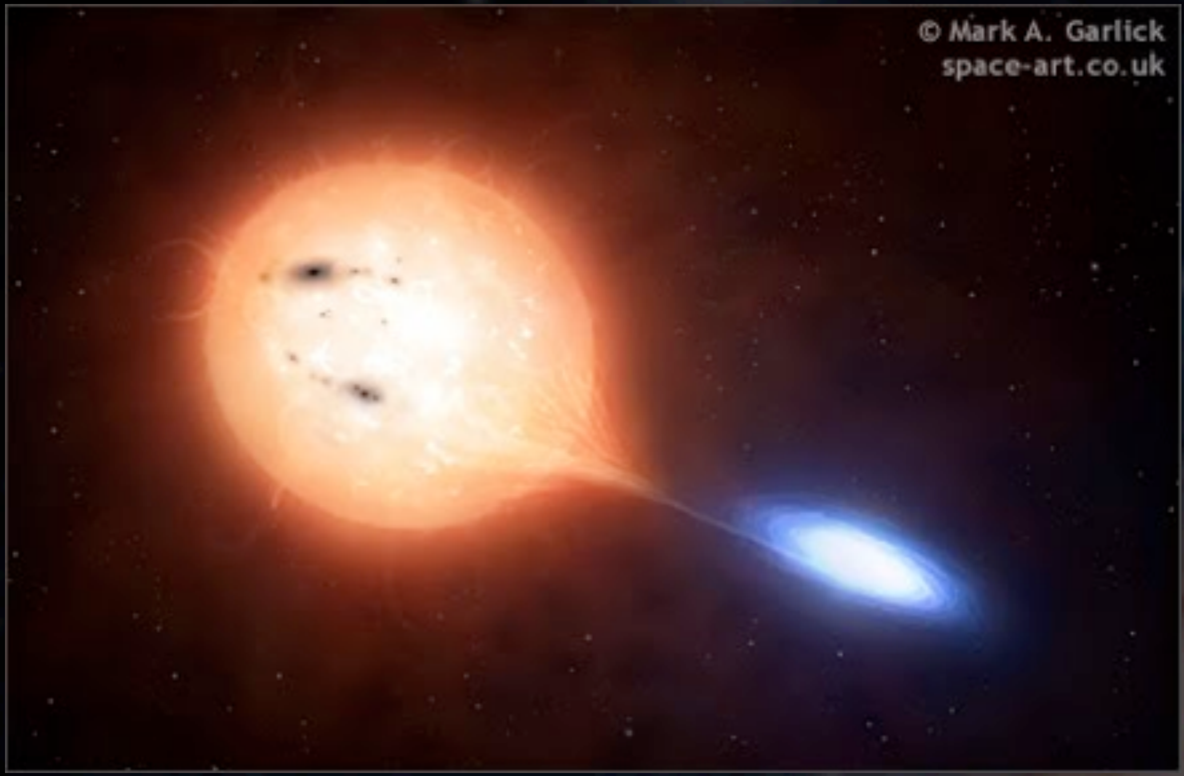
Protoplanetary disks

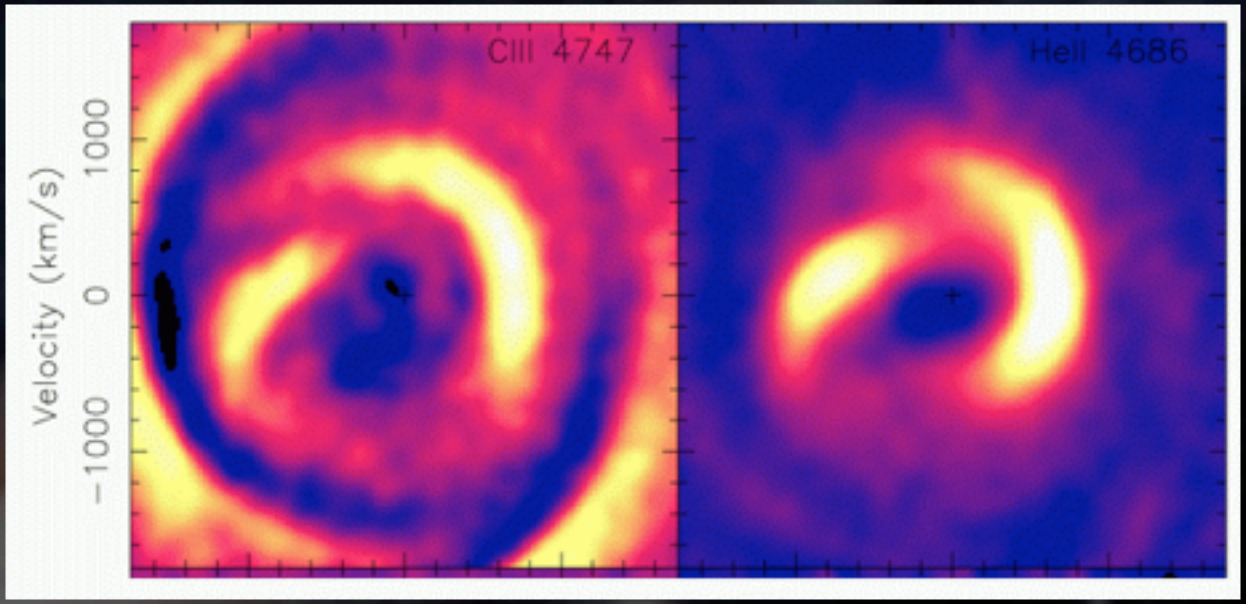




- Size: 10^{11}–10^{15} cm
- Central object: young stellar object $M \sim M_\odot$
- Temperature 10^3–10^1 K

Cataclysmic variables

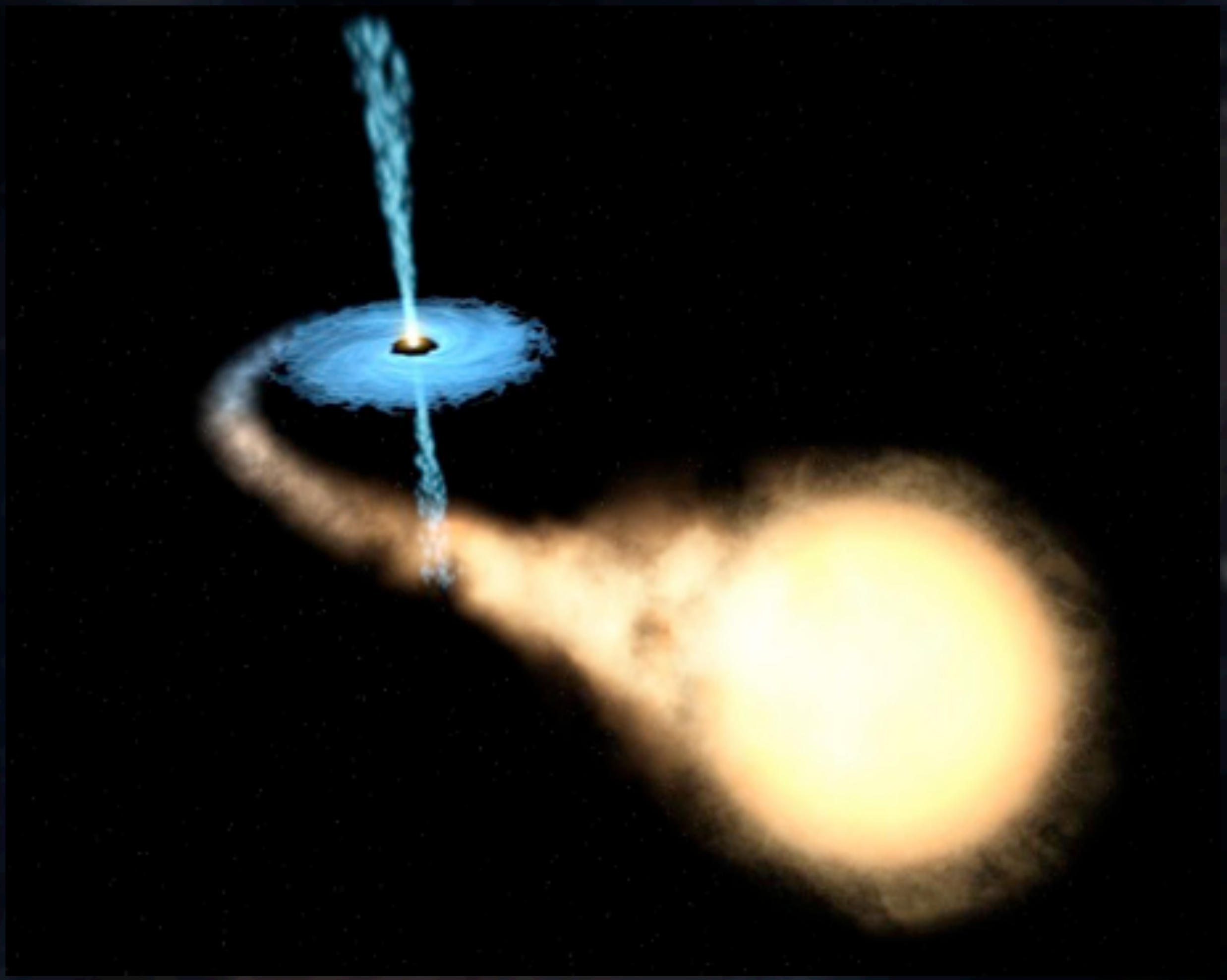




WZ SGE
Steeghs & Marsh (IAUC 7669, 7670)

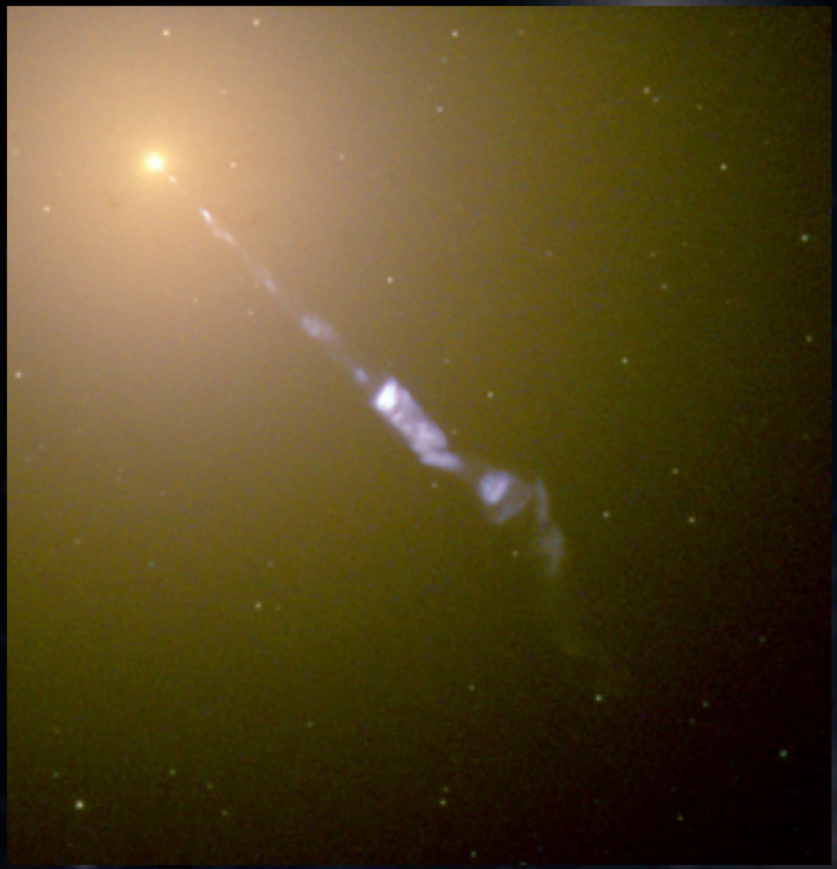
- Size: 10^9–10^{10} cm
- Central object: white dwarf $M \sim M_\odot$
- Temperature 10^5–10^3 K

X-Ray binaries

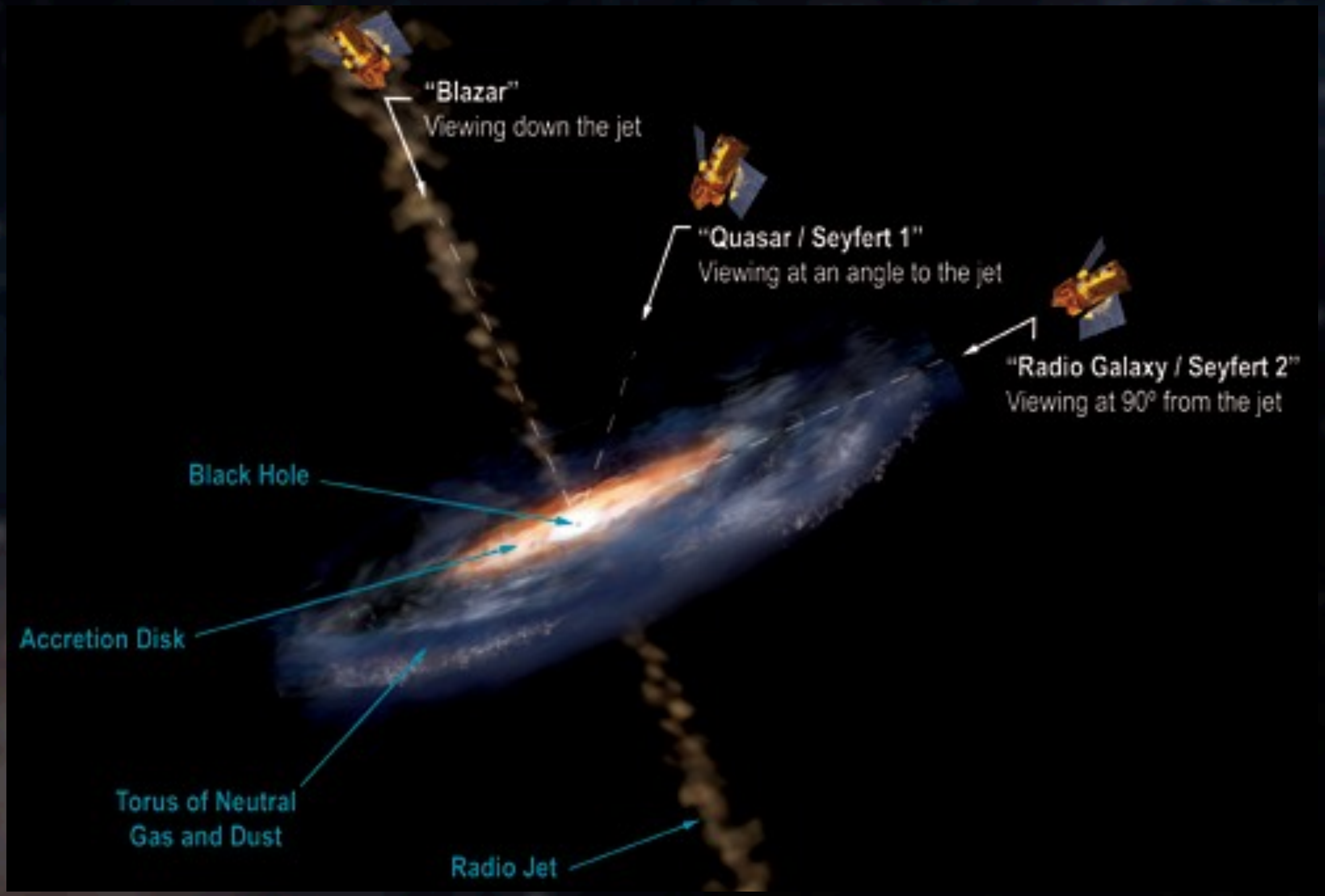


- Size: 10^6–10^{11} cm
- Central object: neutron star, black hole (1–$10\ M_\odot$)
- Temperature 10^7–10^3 K

AGNs (blazar, quasars...)

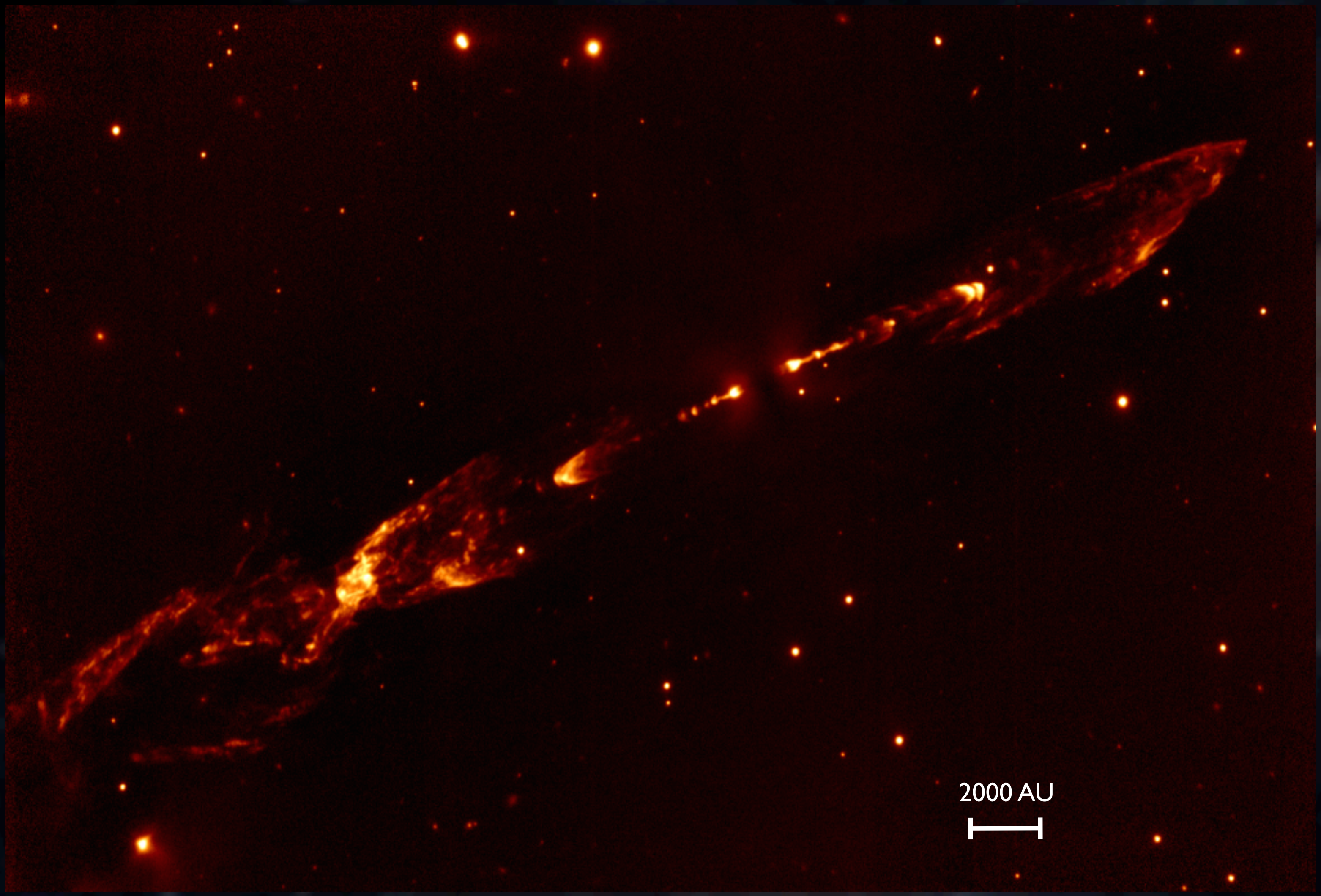


M87

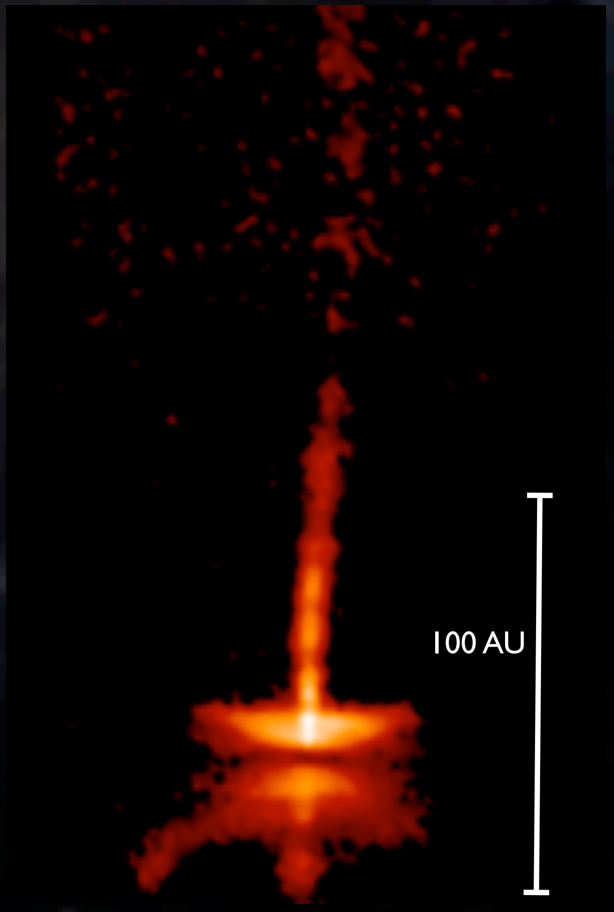


- Size: 10^6–$10^{11}\,\mathrm{cm}\left(\dfrac{M_{\mathrm{BH}}}{M_\odot}\right)$
- Central object: black hole $M_{\mathrm{BH}} = 10^6$–$10^9\,M_\odot$
- Temperature 10^5–$10^2\mathrm{K}$

Jets in protoplanetary disks



HH212



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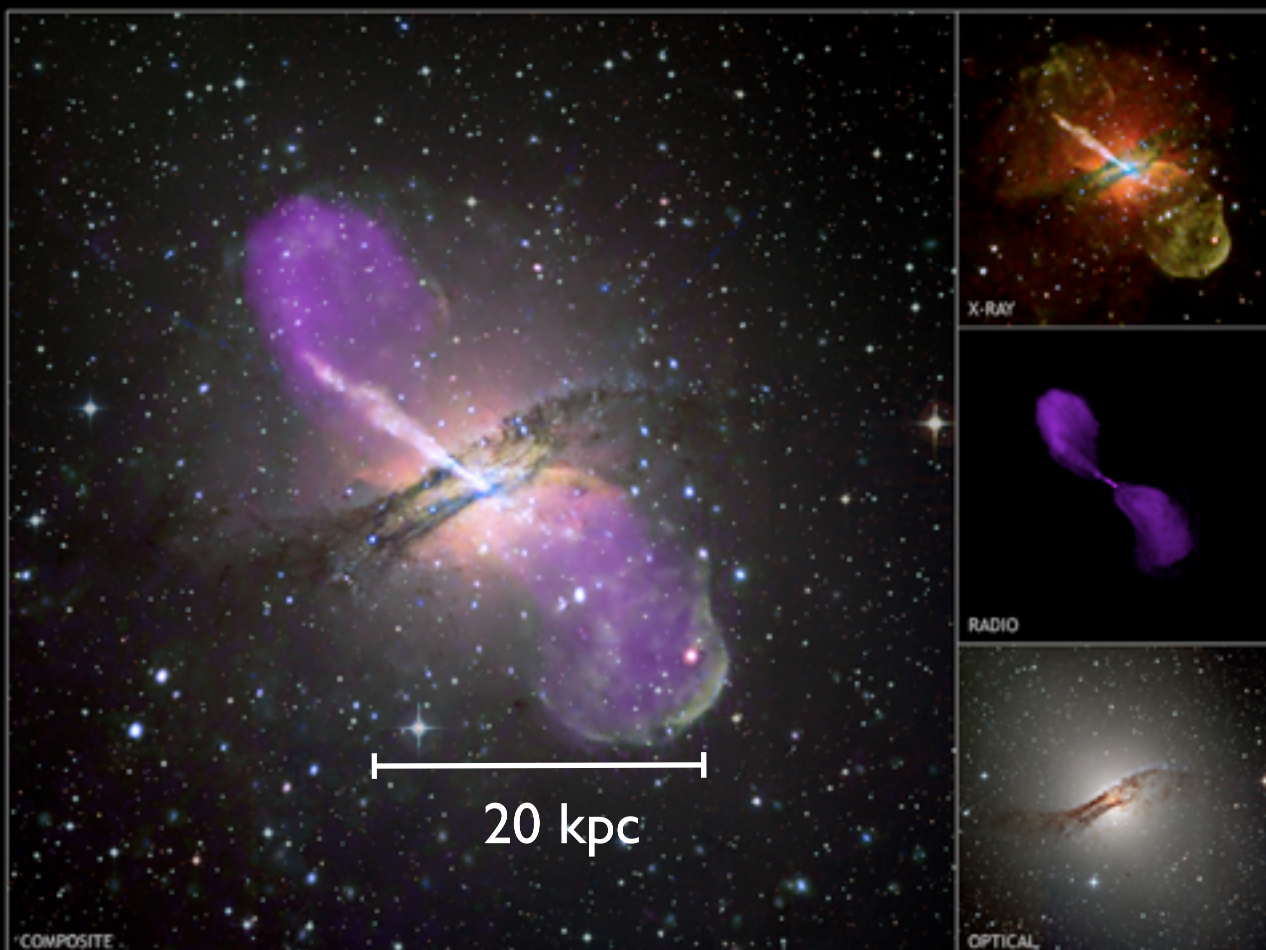
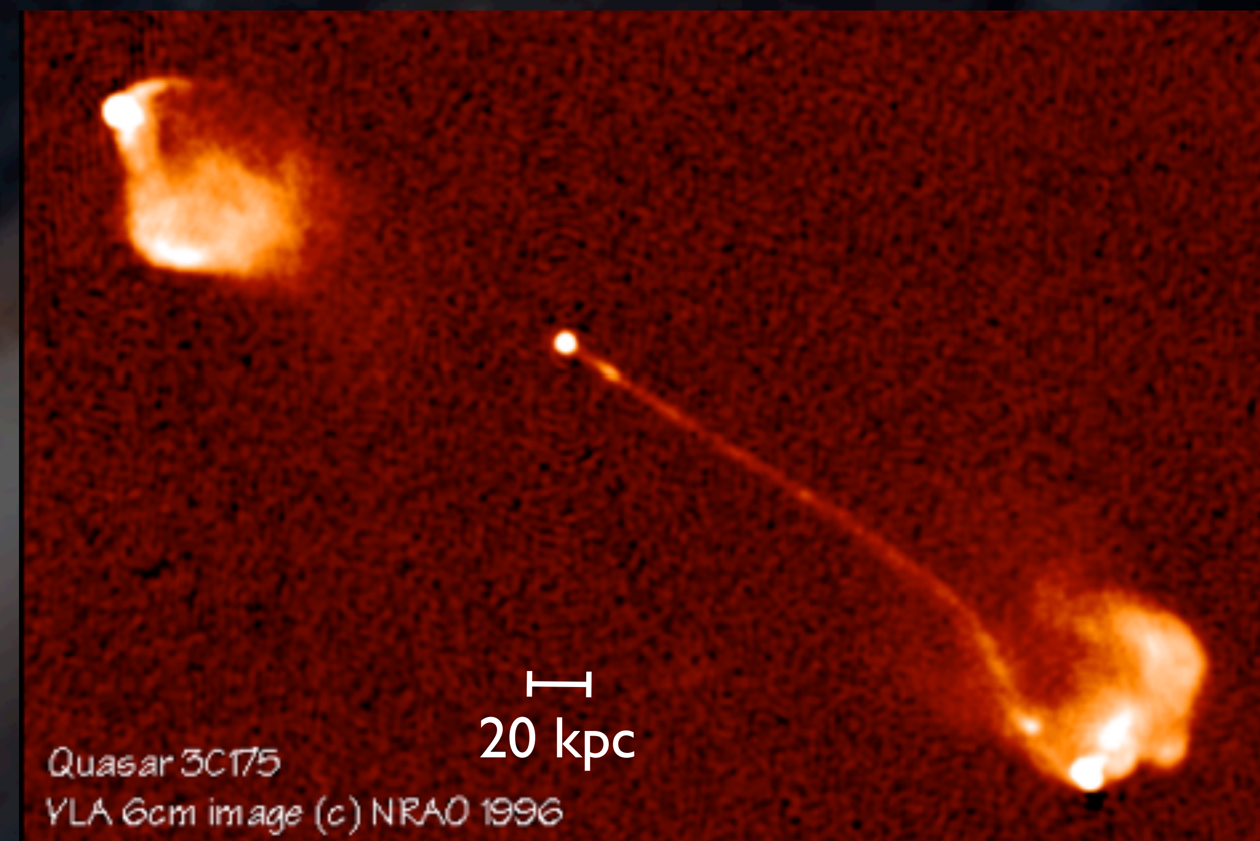
Jets in AGNs



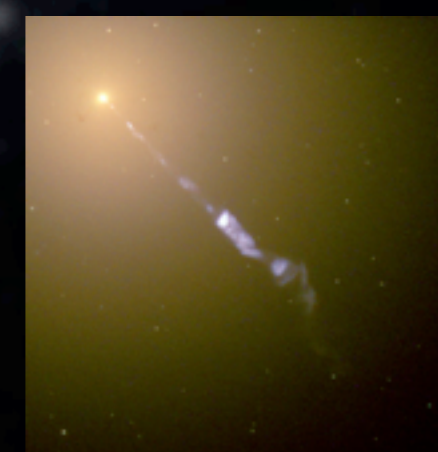
Centaurus A

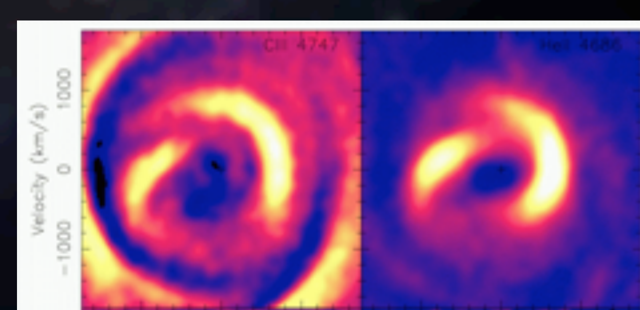


Quasar 3C175

Disks and jets in nature

- Disks are very general in nature
- Almost all of them are associated to powerful jets (with the notable exception of CVs)





Outline

- Accretion disks and jets: what are they
    - Accretion disks in nature
    - Jets in nature
    - The physics of accretion
- Turbulent disks
    - Disk Instabilities
    - The shearing box model
    - The case of the MRI
- Outflows in disks
    - Jet launching mechanisms
    - MRI & disk winds
    - disk wind stability

Disks dynamics

Radial equilibrium

$$\cancel{\frac{\partial u_r}{\partial t}} + \cancel{\boldsymbol{u}\cdot\boldsymbol{\nabla} u_r} - \frac{u_\phi^2}{R} = \cancel{\frac{\boldsymbol{B}\cdot\boldsymbol{\nabla} B_r}{\rho}} - \cancel{\frac{B_\phi^2}{\rho R}} - \frac{1}{\rho}\frac{\partial\left(\cancel{P} + \cancel{B^2/2}\right)}{\partial R} - \frac{GM_\odot}{R^2}$$

- Assume

$$v_A \ll u_\phi \qquad c_s \ll u_\phi \qquad (u_r, u_z) \ll u_\phi$$

$\longrightarrow$ $u_\phi = R\Omega(R)$ with $\Omega(R) = (GM_\odot)^{1/2} R^{-3/2}$

- Disk temporal evolution dictated by small deviations from this Keplerian profile.

$$\boldsymbol{u} = \boldsymbol{v} + R\Omega(R)\boldsymbol{e}_\phi$$

Disks dynamics

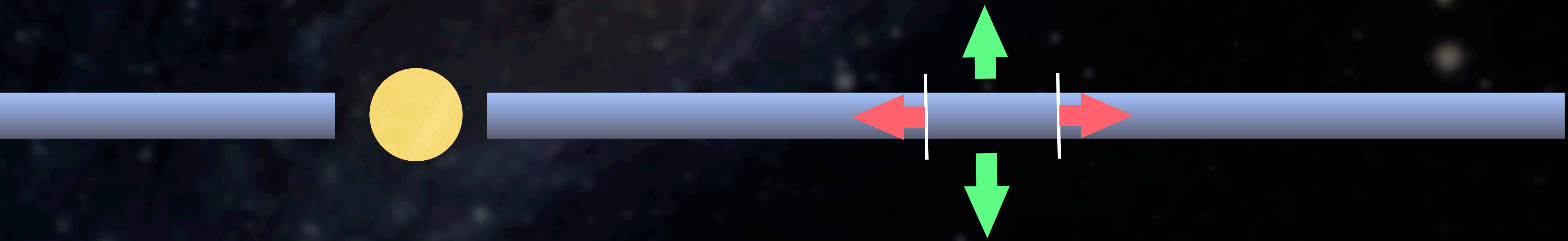
Mass conservation

$$\frac{\partial \rho}{\partial t} + \boldsymbol{\nabla} \cdot \rho \boldsymbol{u} = 0$$

Introduce the average

$$\overline{Q} = \int d\phi \int_{z=-h}^{z=+h} dz\, Q \quad \text{and} \quad \Sigma = \overline{\rho}$$

$$\frac{\partial \Sigma}{\partial t} + \frac{1}{R}\frac{\partial}{\partial R} R\overline{\rho v_r} + \left[\rho v_z\right]_{z=-h}^{+h} = 0$$



Disks dynamics

Angular momentum conservation

$$\frac{\partial(\rho R u_\phi)}{\partial t}+\boldsymbol{\nabla}\cdot\left[\rho R u_\phi \boldsymbol{u}-RB_\phi\boldsymbol{B}+R\Big(P+\frac{B^2}{2}\Big)\boldsymbol{e_\phi}\right]=0$$

- Introduce $\boldsymbol{u}=\boldsymbol{v}+R\Omega(R)\boldsymbol{e_\phi}$

$$\Omega R^2\frac{\partial\rho}{\partial t}+\cancel{\frac{\partial(\rho R v_\phi)}{\partial t}}+\boldsymbol{\nabla}\cdot\left[\rho R^2\Omega\boldsymbol{v}+\rho R v_\phi\boldsymbol{v}-RB_\phi\boldsymbol{B}+R\Big(P+\frac{B^2}{2}\Big)\boldsymbol{e_\phi}\right]=0$$

- Average & integrate vertically:

$$\Omega R^2\frac{\partial\Sigma}{\partial t}+\frac{1}{R}\frac{\partial}{\partial R}R\left(R^2\Omega\overline{\rho v_r}+R\overline{\rho v_\phi v_r}-R\overline{B_\phi B_r}\right)$$
$$+\Big[R^2\Omega\rho v_z+R\rho v_\phi v_z-RB_\phi B_z\Big]_{z=-h}^{+h}\quad=\quad 0$$

Introduce mass conservation

Disks dynamics

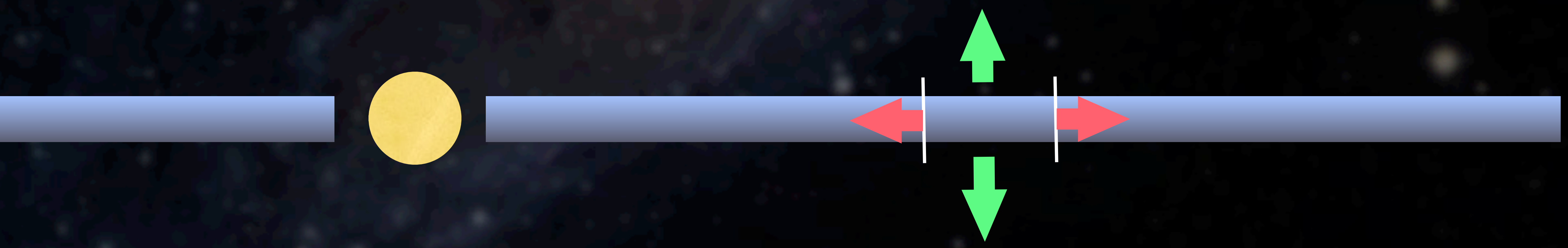
Angular momentum conservation (once mass conservation is taken into account)

$$\overline{\rho v_r}\frac{\partial}{\partial R}\Omega R^2 + \frac{1}{R}\frac{\partial}{\partial R}R^2[\overline{\rho v_\phi v_r} - \overline{B_\phi B_r}] + R\left[\rho v_\phi v_z - B_\phi B_z\right]_{z=-h}^{+h} = 0$$

Accretion

«Turbulent» torque

Wind torque



Disks dynamics

- Neglect surface terms and assume turbulent stress acts as a «viscosity»

$$\overline{\rho v_\phi v_r} - \overline{B_\phi B_r} = -\nu \Sigma R \frac{d\Omega}{dR}$$

- Angular momentum conservation (for a viscous disk)

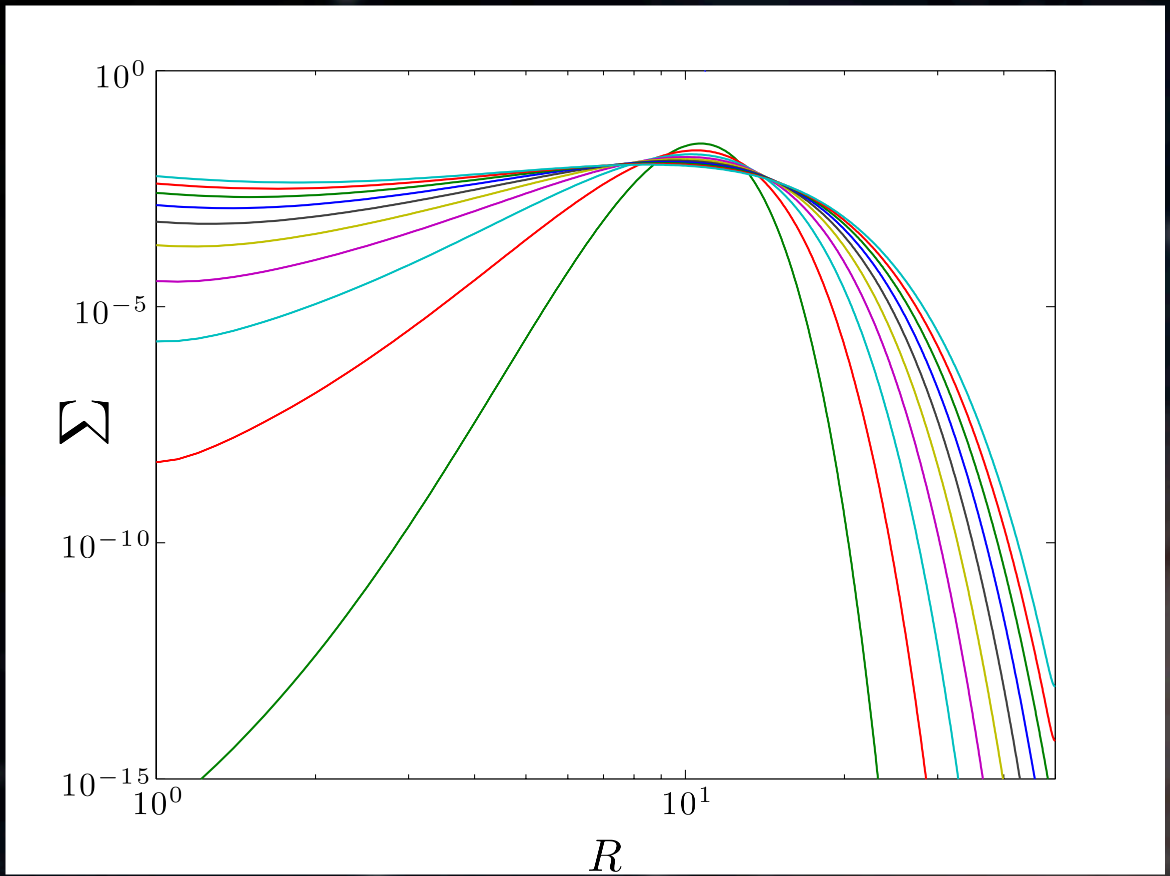
$$\overline{\rho v_r} = -\frac{3}{R^{1/2}} \frac{\partial}{\partial R}\left(R^{1/2}\nu\Sigma\right)$$

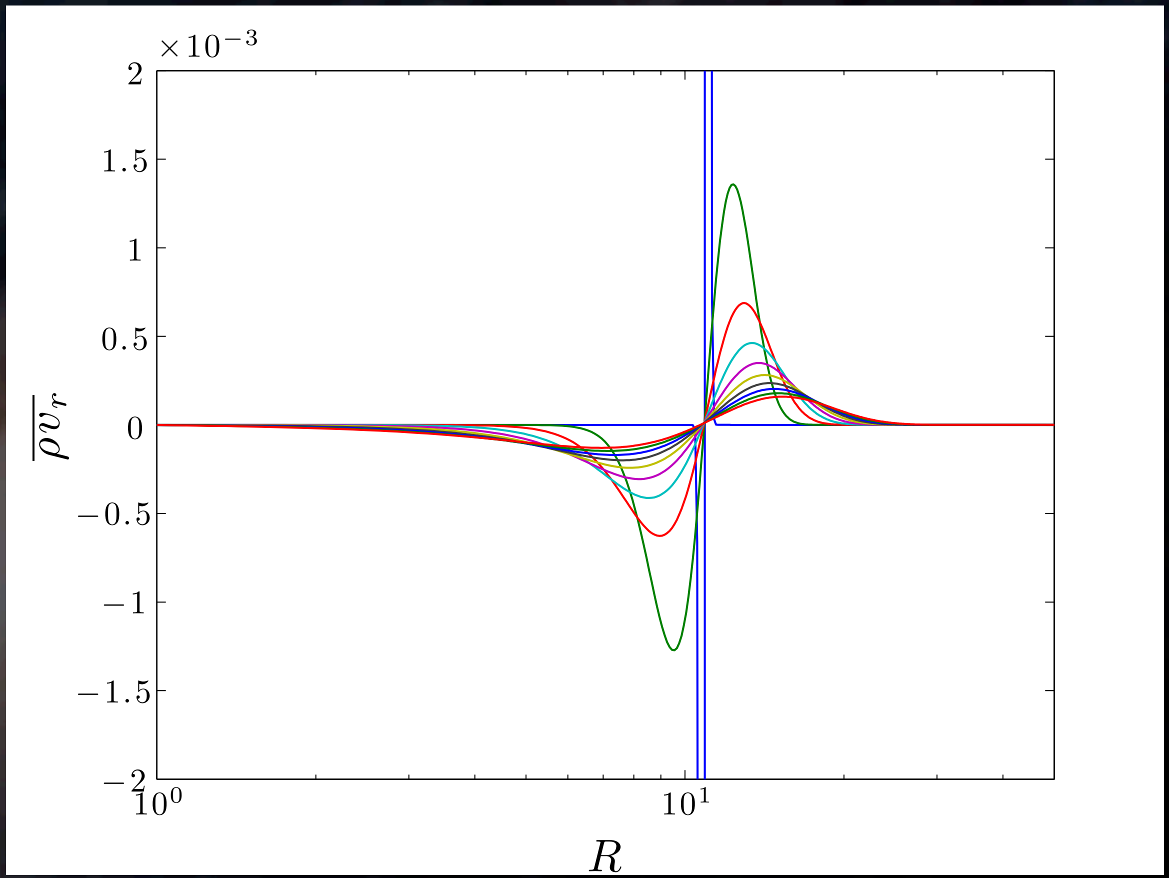
- Put it in mass conservation:

$$\frac{\partial \Sigma}{\partial t} = \frac{3}{R}\frac{\partial}{\partial R}\left(R^{1/2}\frac{\partial}{\partial R}\left(R^{1/2}\nu\Sigma\right)\right)$$

Lynden-Bell & Pringle (1974)

Viscous disk evolution





- Most of the mass accreted
- Small amount of mass excreted (evacuate the disk angular momentum)
- Viscous timescale $\tau_\nu = \dfrac{R^2}{\nu}$. How big is ν ?

What if it was pure viscosity?

$$\nu \sim u_{\rm th}\Lambda_{\rm mfp}$$

Thermal velocity — Mean free path

$$u_{\rm th} \sim 10^4\sqrt{T}\,\mathrm{cm.s^{-1}} \qquad \lambda \sim \frac{10^{16}}{n}\,\mathrm{cm}$$

- In protoplanetary disks: $T \sim 10^3\,\mathrm{K}$
  $n \sim 10^{13}\,\mathrm{cm^{-3}}$

$\nu \sim 3\times 10^8\,\mathrm{cm^2.s^{-1}}$ ➡ $\tau_\nu \sim 10^{14}\mathrm{yrs}$ >> age of the universe

- Need for an «anomalous» viscosity...

The alpha disk model

- Introduce the scaling *(Shakura-Sunyaev 1973)*

$$\nu = \alpha c_s H \qquad 10^{-4} < \alpha < 10^{-1}$$

- Estimated accretion rate: $\overline{\rho v_r} \sim -\alpha c_s \Sigma \frac{H}{R}$

New questions!

- What is responsible for this anomalous viscosity?
  - Turbulence
  - Waves
- How big is α ?
- What about the surface terms we have neglected?

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Some disk instabilities

Local instabilities:

- Magnetorotational instability (MRI): shear driven instability but requires an ionised plasma (Velikhov 1959, Chandrasekhar 1960, Balbus & Hawley 1991)
- Subcritical shear instability: probably not efficient enough, if exists (see later)
- Baroclinic instabilities: Transport due to waves. Driven by the disk radial entropy profile (see later)
- Gravitational instabilities: only for massive & cold enough disk (see later)
- Rossby wave instability: requires a local maximum of vortensity (Lovelace et. al 1999)
- Vertical convective instability: Requires a heat source in the midplane (Cabot 1996, Lesur & Ogilvie 2010)

Global instabilities:

- Papaloizou & Pringle instability: density wave reflection on the inner edge (Papaloizou & Pringle 1985)
- Accretion-ejection instability: spiral Alfvén wave reflection on the inner edge (Tagger & Pellat 1999)

Subcritical shear instabilities

The Facts:

- kelperian shear flows are *linearly stable*
- Huge reynolds numbers → nonlinear instability? (same thing as pipe flows or Couette flows)

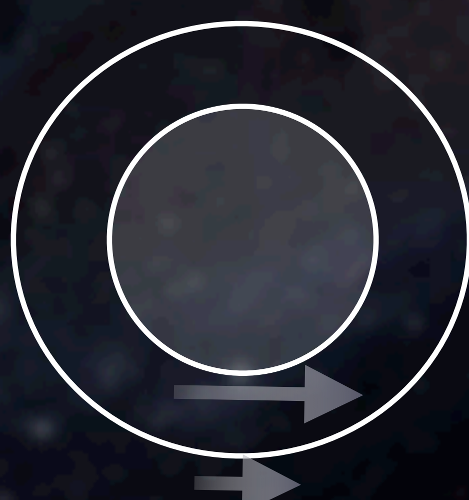
Pipe flow

Couette flow

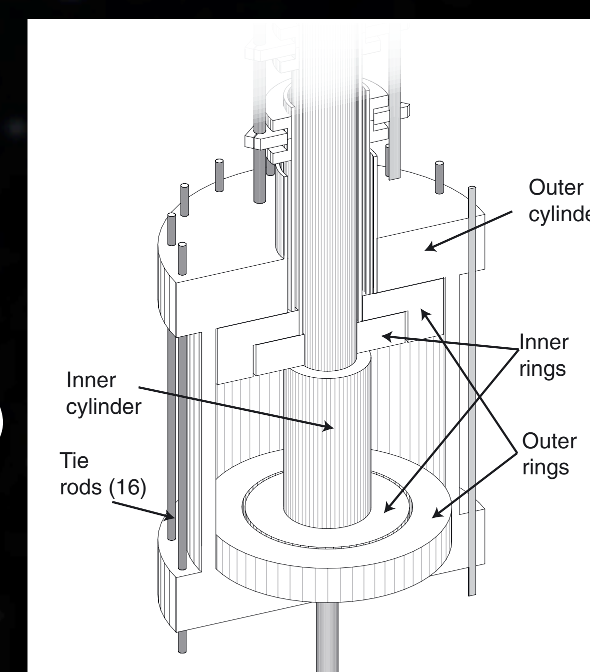
A nonlinear instability in accretion disks?

- Experimental approach: hard to «do» a disk in a lab. Boundary conditions?
- Numerical approach: high reynolds numbers unreachable.

Ideal Couette-Taylor



Real life Couette-Taylor (*Schartman et al. 2012*)

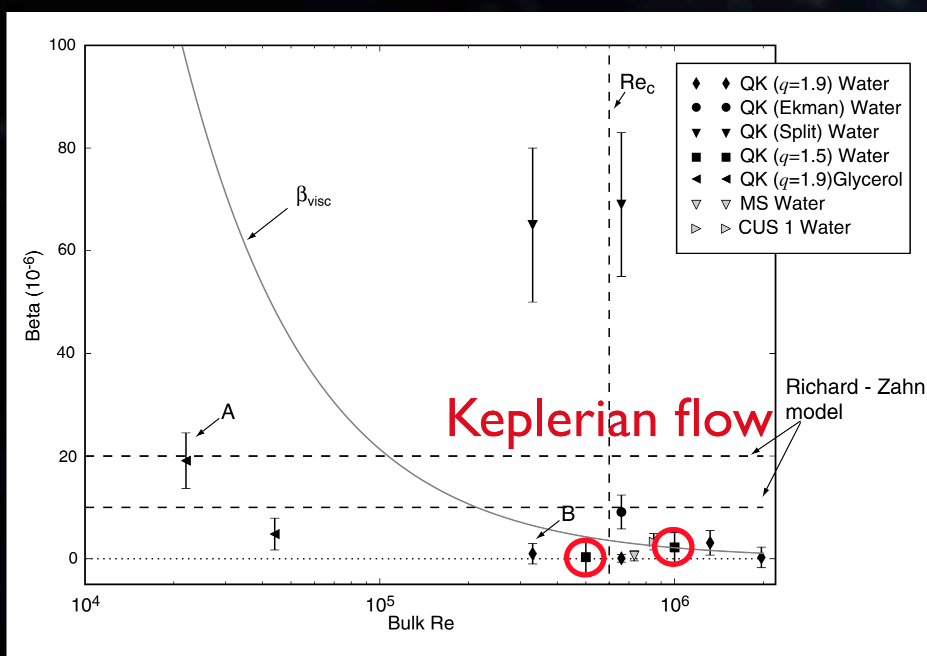


Subcritical shear instabilities

Experiments

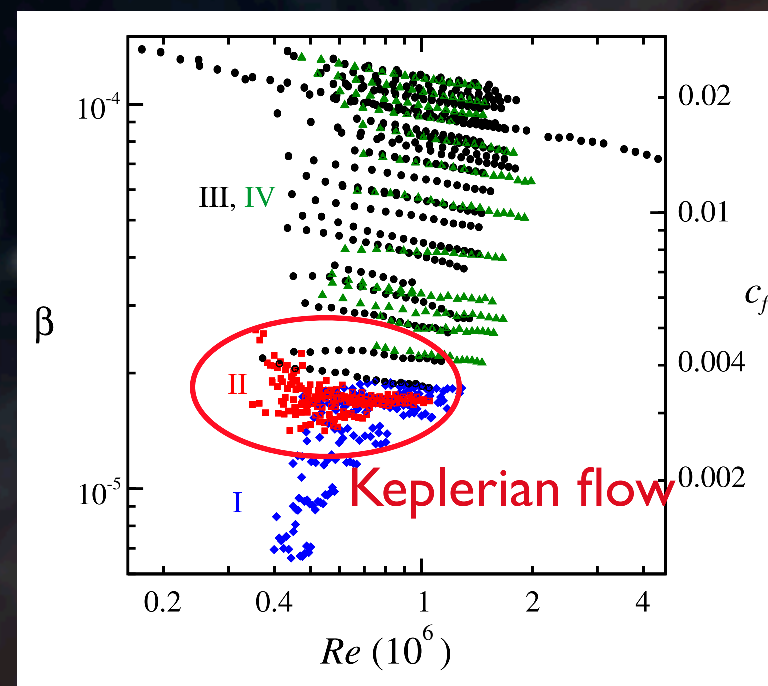
Simulations

Schartman et al. 2012



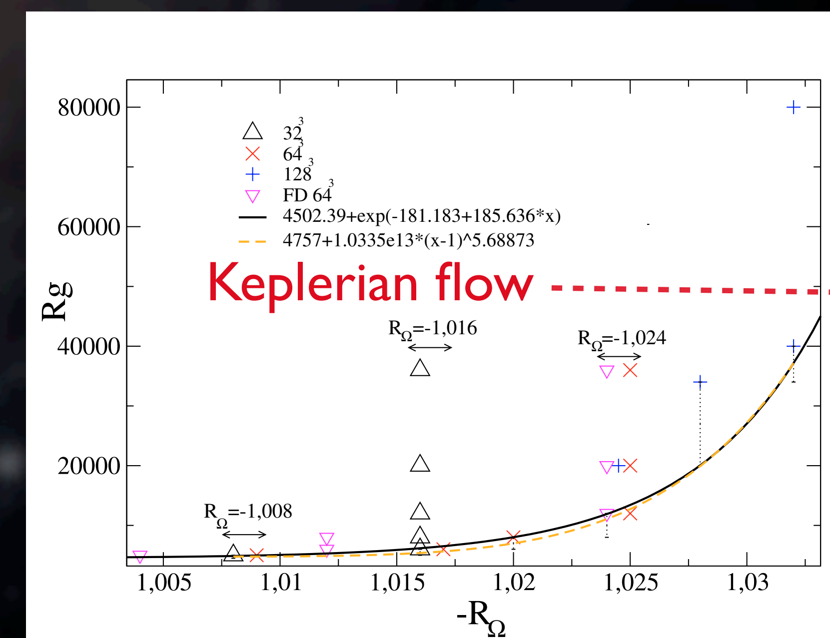
No turbulence

Paoletti & Lathrop 2012



Turbulence

Lesur & Longaretti 2005

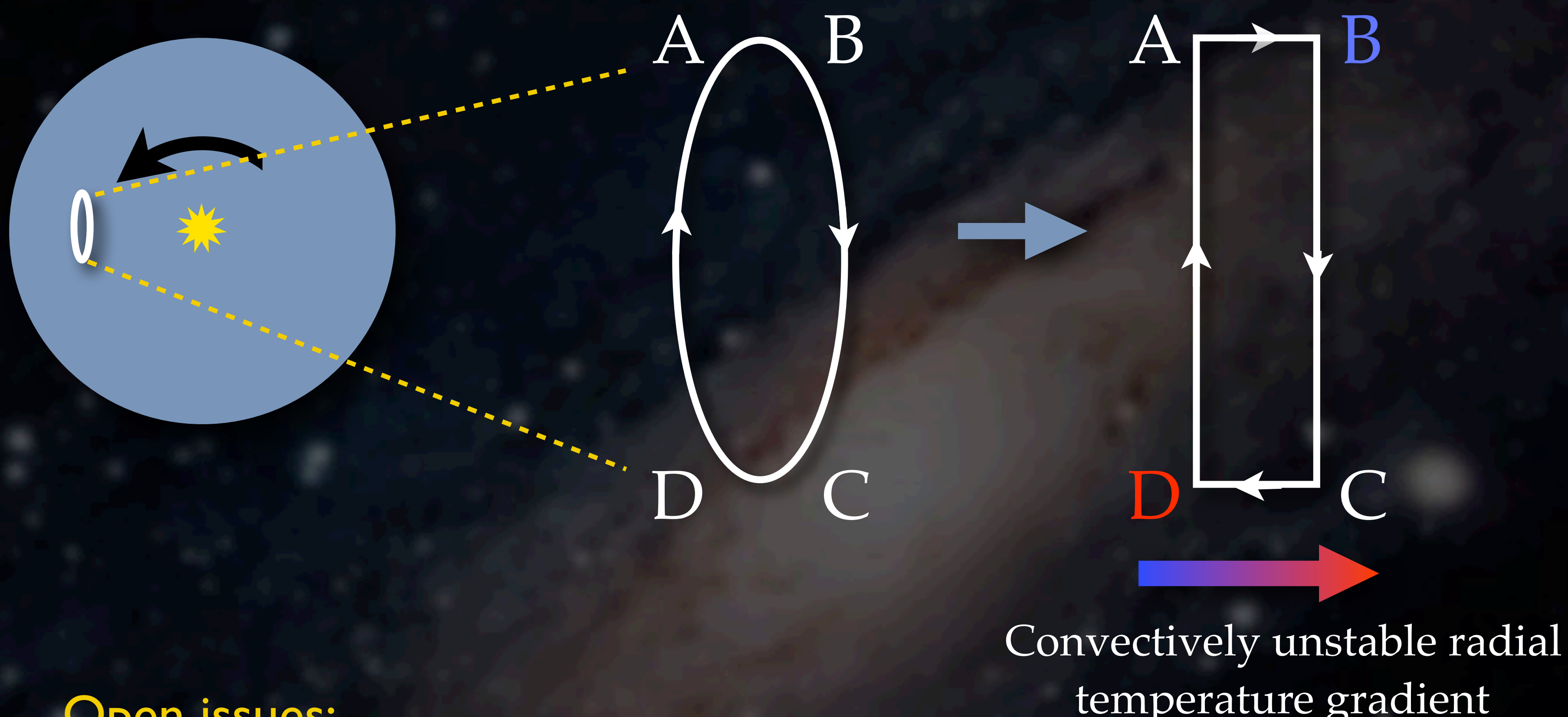


No turbulence

→ No definitive answer yet

Baroclinic instabilities

Baroclinic instabilities: nonlinear instability driven by the radial entropy gradient



Open issues:

- Vortices are unstable
- Vortices migrate
- What maintains the entropy structure?
- Instability does not survive with magnetic fields

Refs: Klahr & Bodenheimer 2003, Petersen et al. 2007a,b, Lesur & Papaloizou 2010

Gravitational instabilities

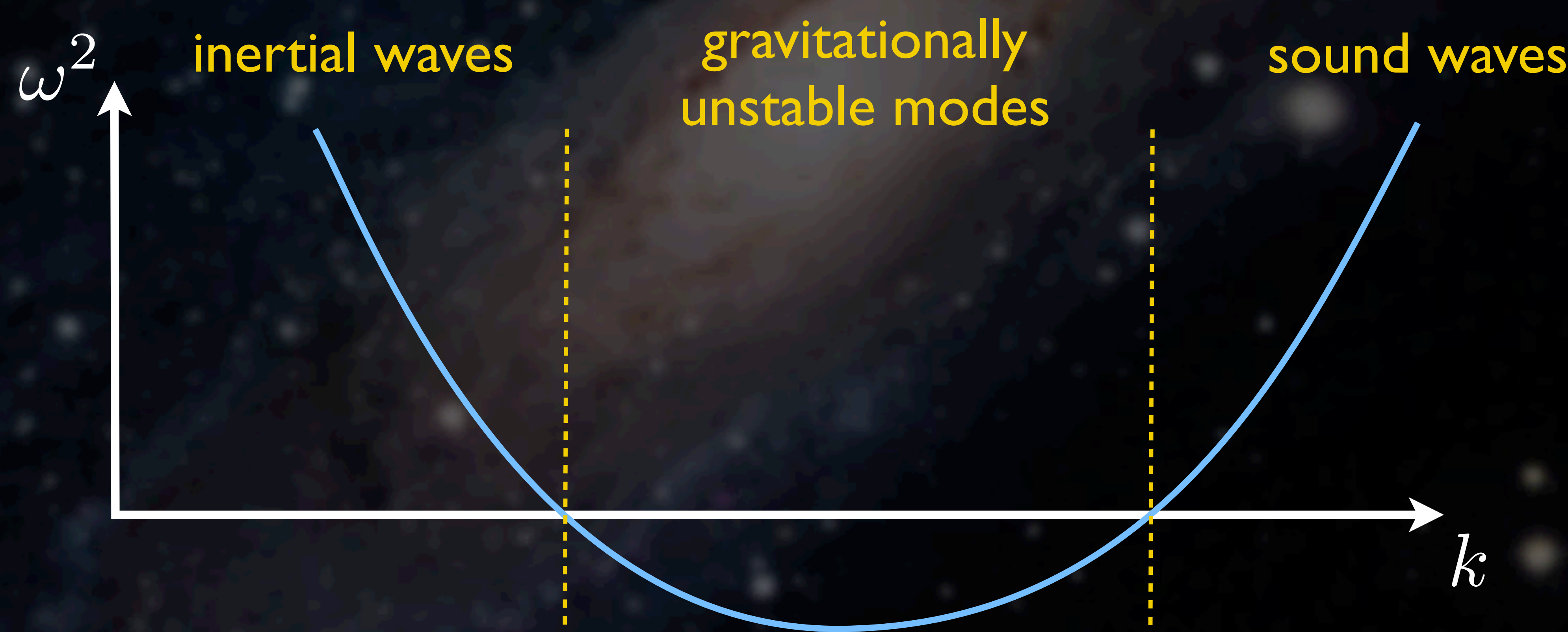
Dispersion relation for axisymmetric modes:

$$\omega^2 = k^2 c_s^2 - 2\pi G\Sigma|k| + \kappa^2 \qquad \kappa^2 = 2\Omega\left(2\Omega + \frac{d\Omega}{d\ln R}\right)$$

Thermal pressure (sound waves)

Self gravity

Local centrifugal force (epicyclic oscillations)



Criterion for instability: $\Delta' = (\pi G\Sigma)^2 - c_s^2\kappa^2 > 0 \Longrightarrow Q = \dfrac{c_s\kappa}{\pi G\Sigma} < 1$

Toomre 1964

Gravitational instabilities

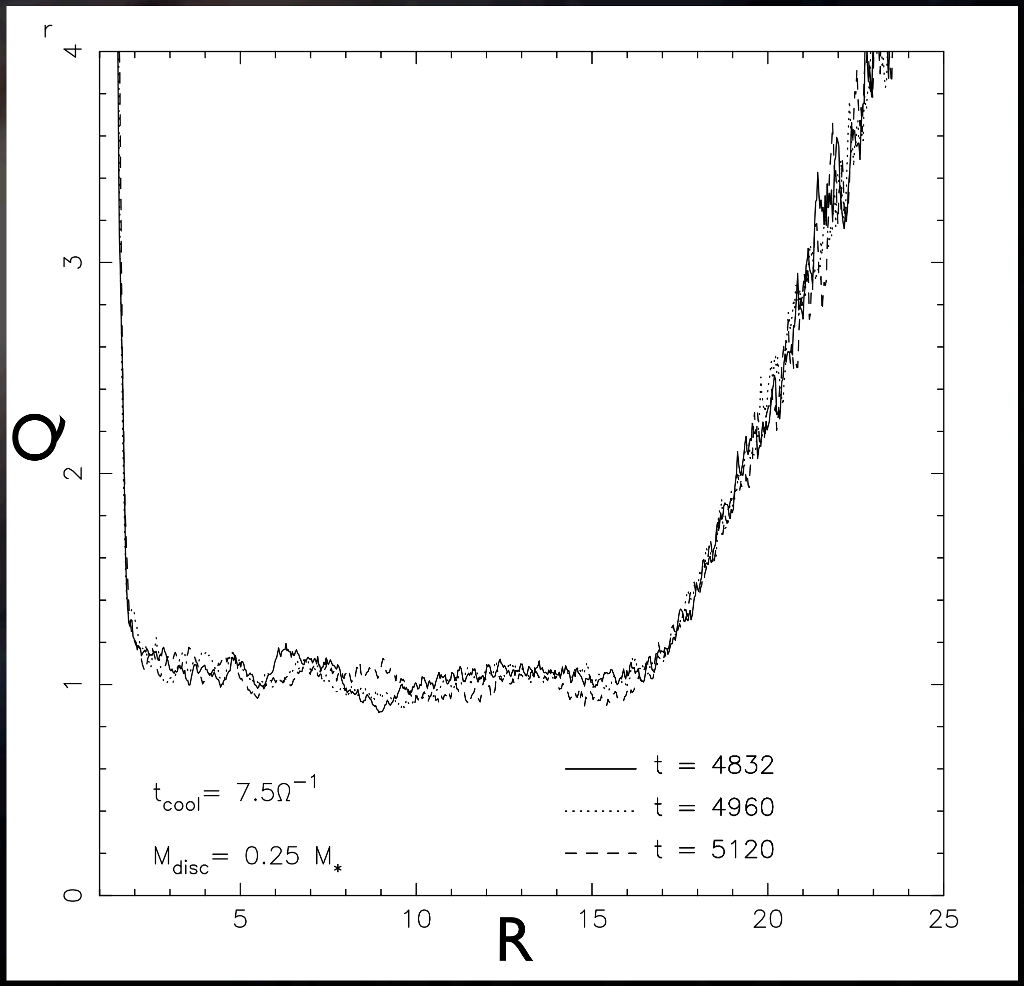
Nonlinear evolution

Consider a disk with $Q \gg 1$

- Disk cools, c_s decreases
- When $Q=1$ GI starts
- GI unstable modes produce strong shocks → heating
- c_s increases
- Q>1 GI stops

→ «gravito-turbulence»

$$Q = \frac{c_s \kappa}{\pi G \Sigma}$$

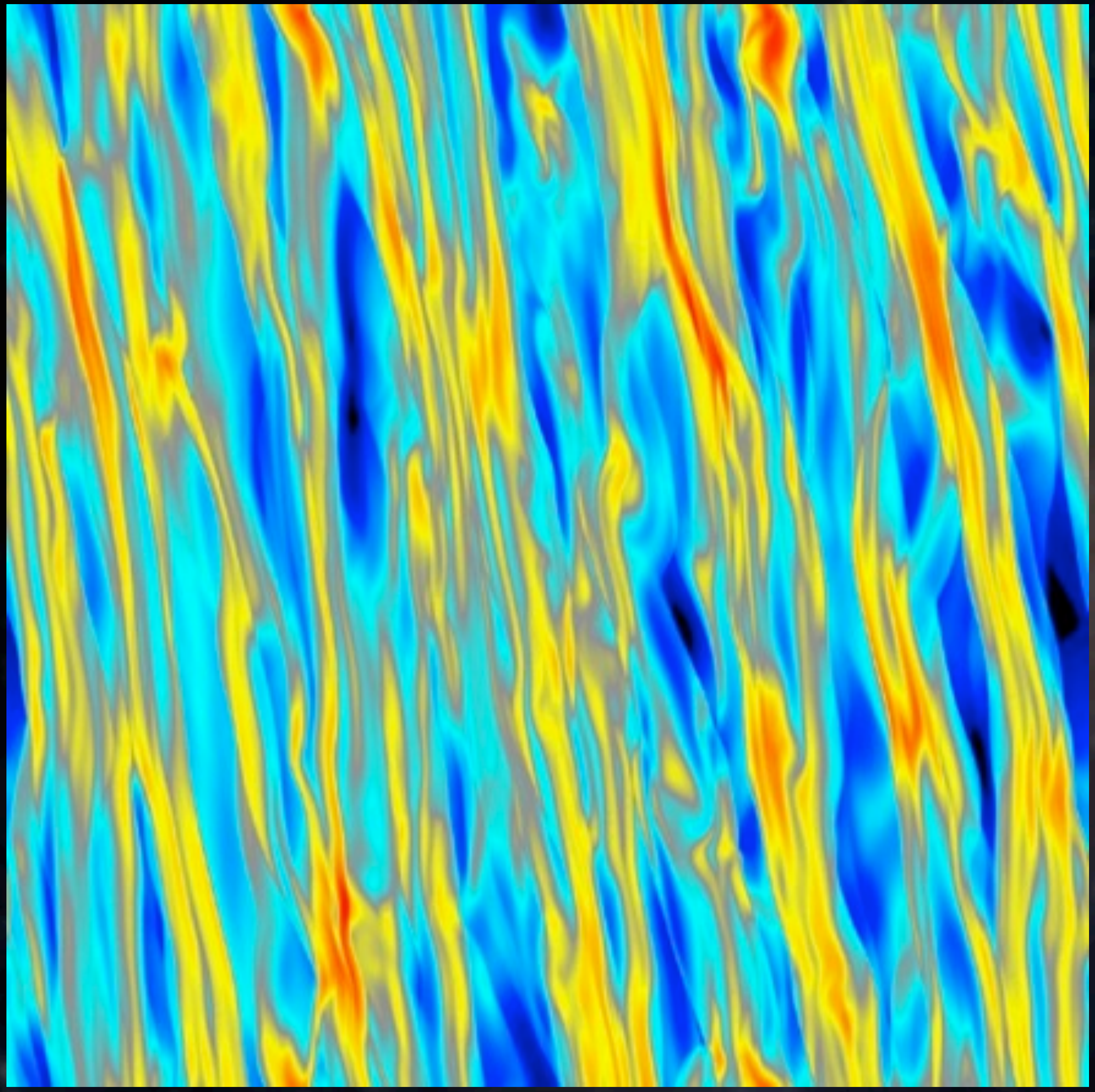


Lodato & Rice (2004)

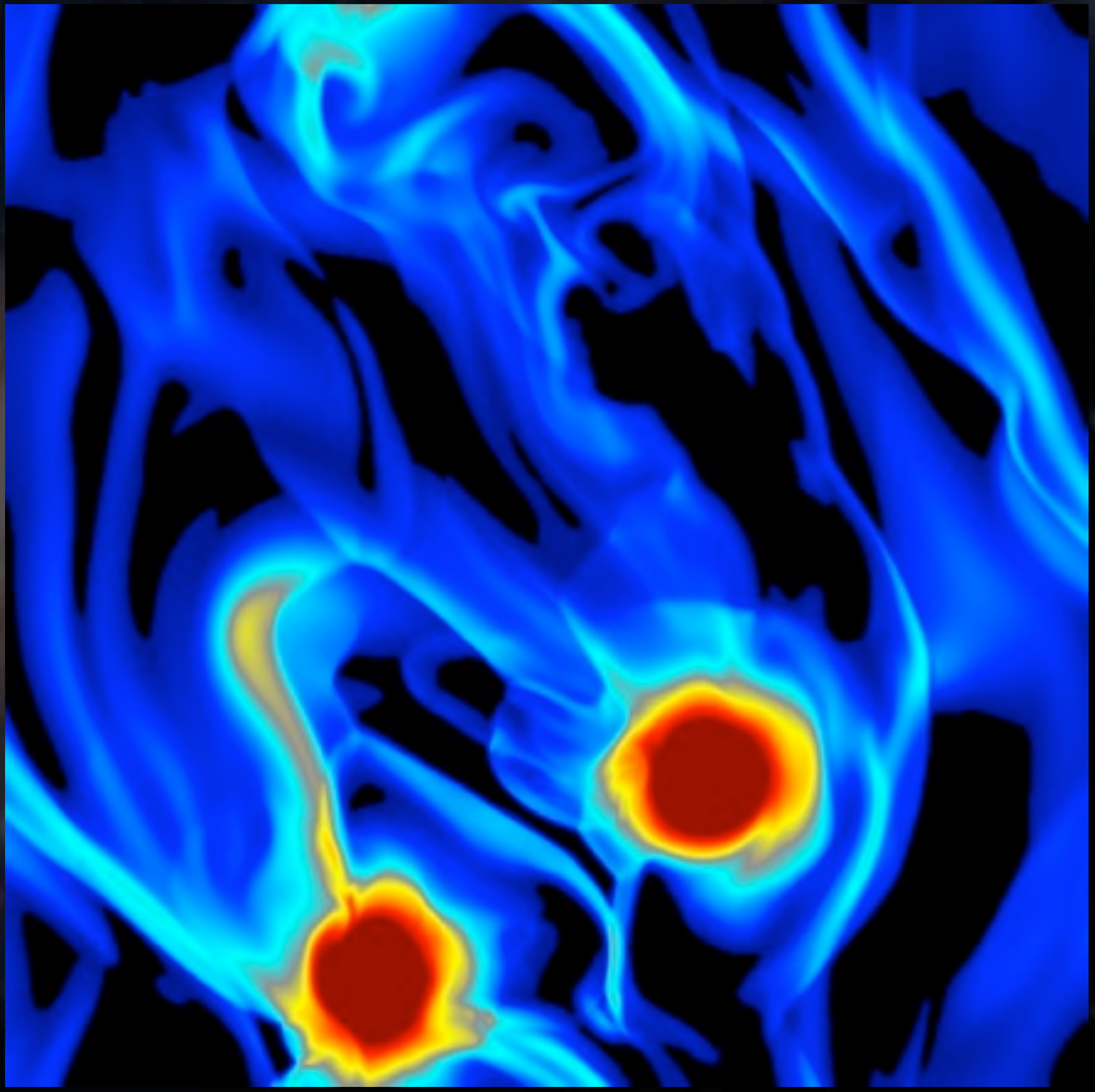
Gravitational instabilities

Well... It's not so simple!

$\tau_{\text{cool}} = 10\Omega^{-1}$

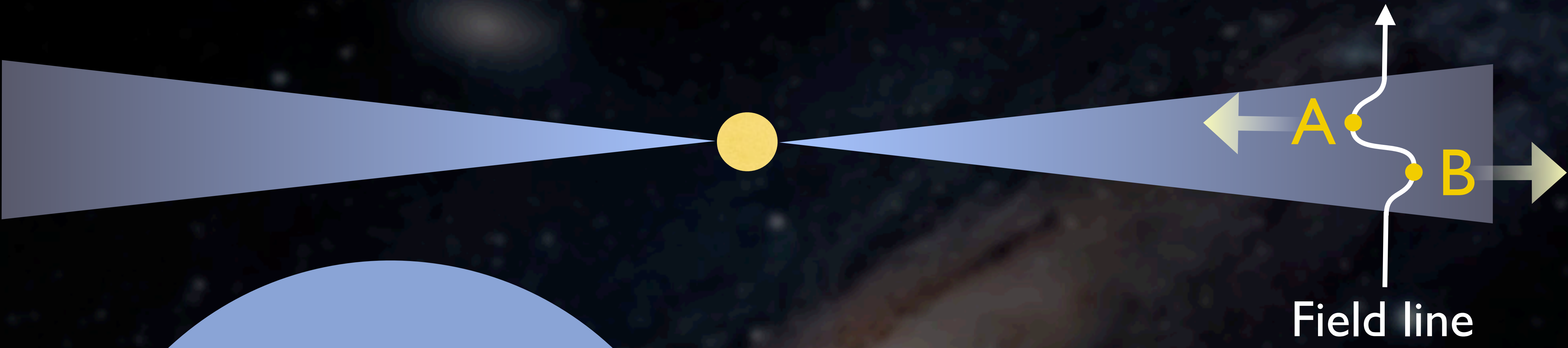


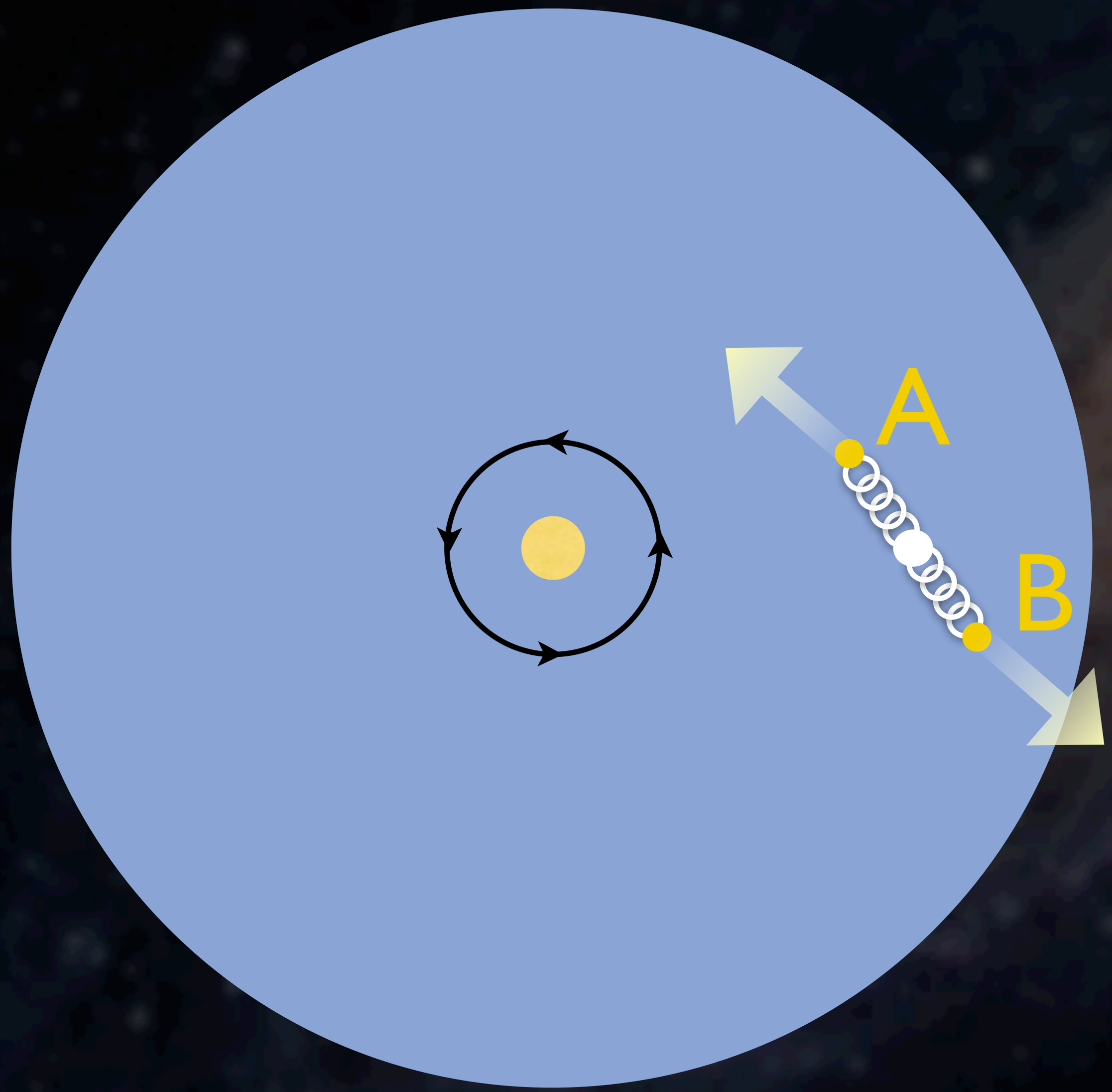
$\tau_{\text{cool}} = 2\Omega^{-1}$



The outcome depends on the cooling time, but also on numerical schemes and resolution... See Paardekooper (2012), Meru & Bate (2012)

Magnetorotational instability



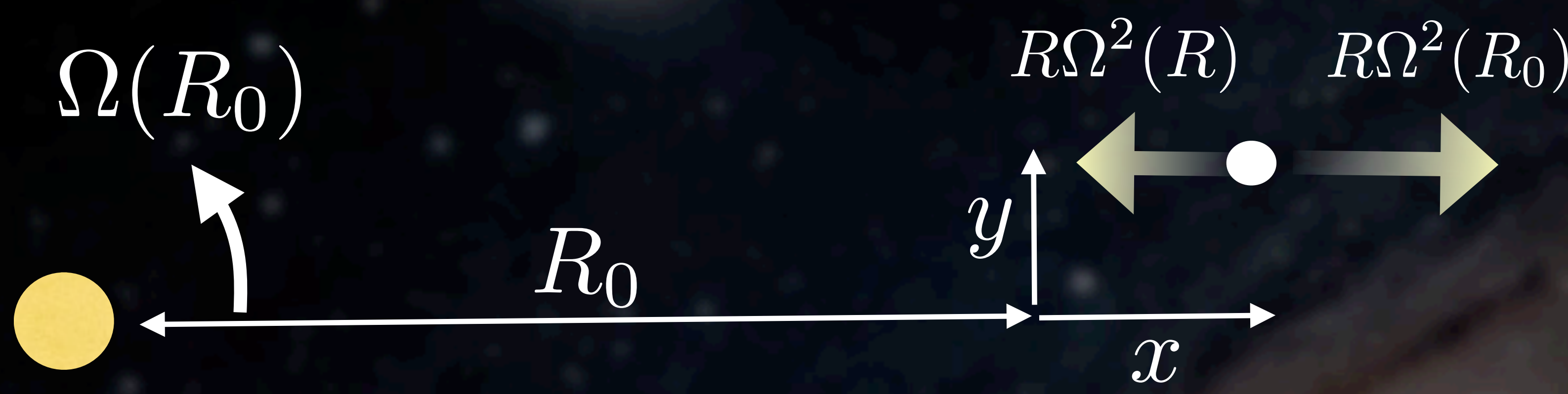


Main properties

- Due to an interaction between magnetic tension and epicyclic motions
- Not too strong magnetic fields required («weak field instability»)
- Need a sufficiently high ionization fraction

Balbus & Hawley 1991, Balbus 2003

Magnetorotational instability



Effective (tidal) radial acceleration: $-x\dfrac{d\Omega^2}{d\ln R}$

Resulting equation of motion for a fluid particle:

$$\begin{aligned}\ddot{x} - 2\Omega\dot{y} &= -\frac{d\Omega^2}{d\ln R}x \\ \ddot{y} + 2\Omega\dot{x} &= 0\end{aligned}$$

→ Epicyclic oscillations at frequency $\kappa = \left(4\Omega^2 + \dfrac{d\Omega^2}{d\ln R}\right)^{1/2}$

Balbus 2003

Magnetorotational instability

Induction equation for a displacement $\boldsymbol{\xi}$
and a spatial dependence $\propto \exp(ikz)$

$$\delta \boldsymbol{B} = i(\boldsymbol{k}\cdot\boldsymbol{B})\boldsymbol{\xi}$$

The magnetic tension force is then

$$\frac{i(\boldsymbol{k}\cdot\boldsymbol{B})}{\rho}\delta\boldsymbol{B} = -(\boldsymbol{k}\cdot\boldsymbol{v_A})^2\boldsymbol{\xi}$$

Resulting equation of motion for a fluid particle:

$$\begin{aligned} \ddot{x} - 2\Omega\dot{y} &= -\Big(\frac{d\Omega^2}{d\ln R} + (\boldsymbol{k}\cdot\boldsymbol{v_A})^2\Big)x \\ \ddot{y} + 2\Omega\dot{x} &= -(\boldsymbol{k}\cdot\boldsymbol{v_A})^2 y \end{aligned}$$

Introduce:

$$x = x_0\exp(i\omega t)$$
$$y = y_0\exp(i\omega t)$$

$\Rightarrow$ Dispersion relation

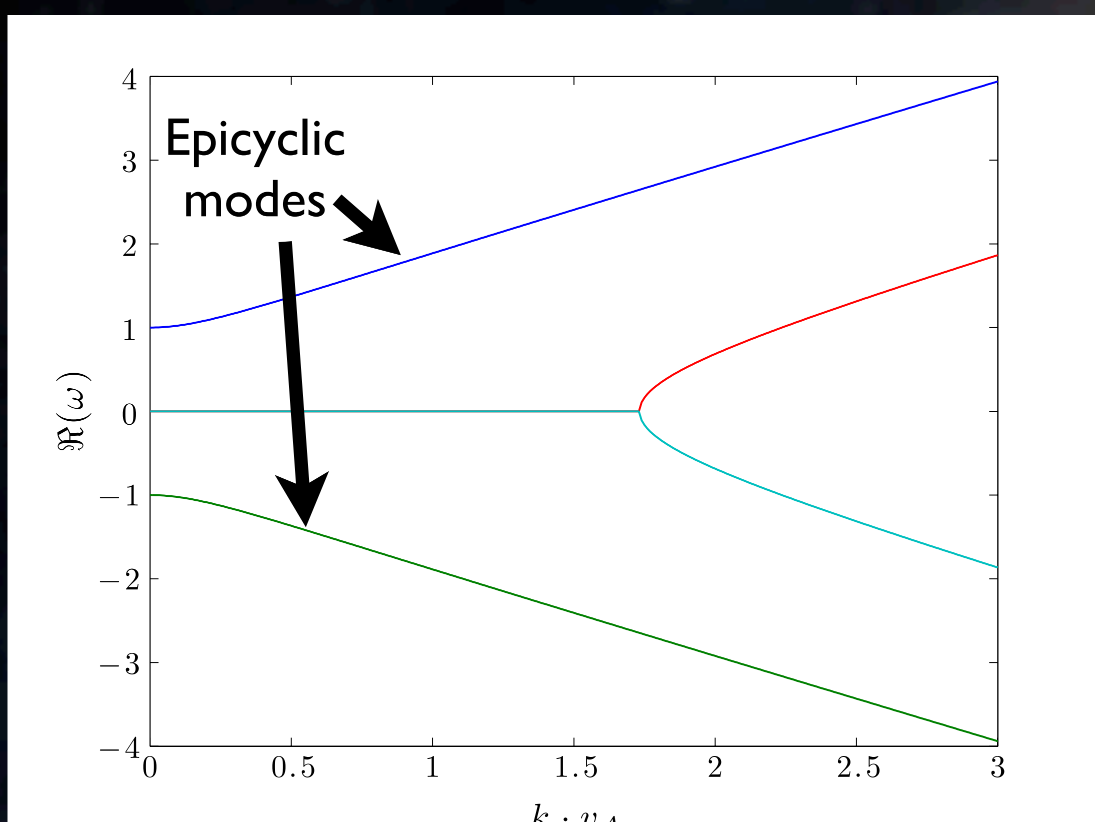
Magnetorotational instability

$$\omega^4 - \omega^2[\kappa^2 + 2(\boldsymbol{k}\cdot\boldsymbol{v_A})^2] + (\boldsymbol{k}\cdot\boldsymbol{v_A})^2\left[(\boldsymbol{k}\cdot\boldsymbol{v_A})^2 + \frac{d\Omega^2}{d\ln R}\right] = 0$$

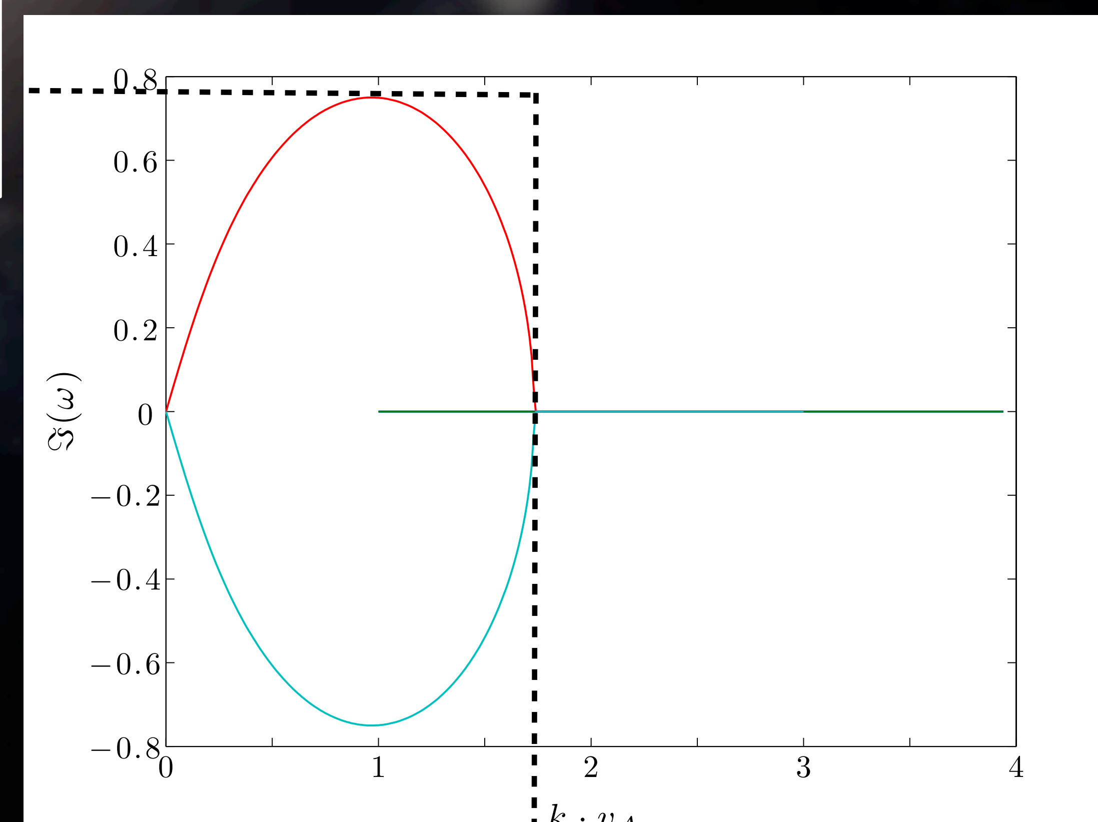
Stabilizing Destabilizing

$$\Omega \propto R^{-q} \implies \frac{d\Omega^2}{d\ln R} = -2q\Omega^2$$

For an accretion disk (q=3/2)



$$\frac{1}{2}\left|\frac{d\Omega}{d\ln R}\right|$$



$$\left(\left|\frac{d\Omega^2}{d\ln R}\right|\right)^{1/2}$$

Magnetorotational instability

Main properties:

- Fast instability ($\gamma \sim \Omega$)
- Condition for instability $\frac{d\Omega}{dR} < 0$ satisfied in disks
- Also works for different field topologies (eg toroidal field: Balbus & Hawley 1992)
- Need weak enough fields ($\boldsymbol{k} \cdot \boldsymbol{v}_A \lesssim \Omega$)

→ Can it explain the anomalous transport needed in disks?

The shearing box model

Problem:

- Computing a full disk is computationally expensive

→ Local resolution is poor

→ Boundary conditions

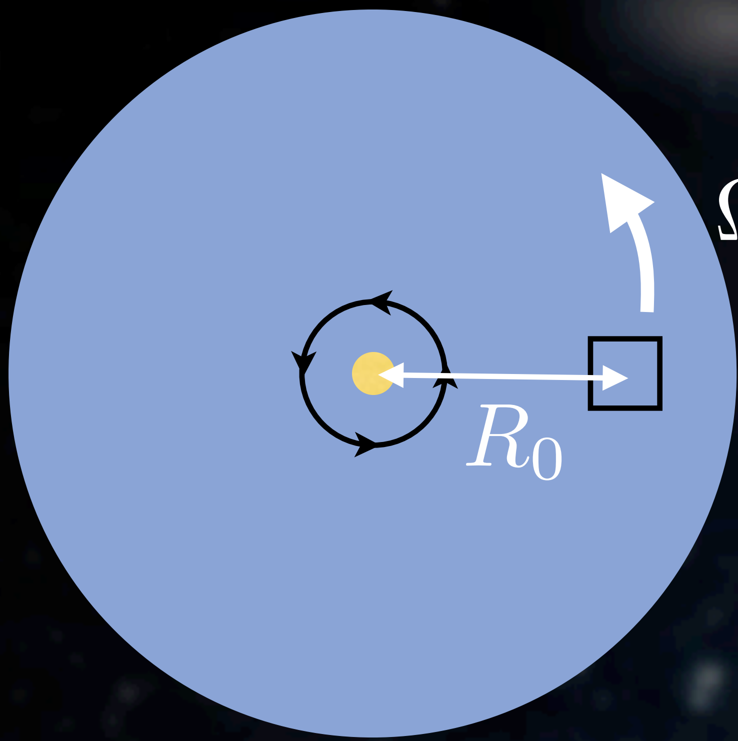
Goal:

- Define a simplified setup which mimics the local properties of an accretion disks

→ Simplifies numerical simulations & boundary conditions

→ Better convergence properties

The shearing box model



$\Omega(R_0)$

$$
\begin{aligned}
\partial_t \rho + \boldsymbol{\nabla} \cdot \rho \boldsymbol{u} &= 0 \\
\partial_t \rho \boldsymbol{u} + \boldsymbol{\nabla} \cdot (\rho \boldsymbol{u} \otimes \boldsymbol{u}) &= -\boldsymbol{\nabla} P + \boldsymbol{J} \times \boldsymbol{B} \\
&\quad -2\rho \boldsymbol{\Omega} \times \boldsymbol{u} - \rho \boldsymbol{\nabla} \psi + \rho \nu \Delta \boldsymbol{u} \\
\partial_t \boldsymbol{B} &= \boldsymbol{\nabla} \times (\boldsymbol{u} \times \boldsymbol{B}) + \eta \Delta \boldsymbol{B}
\end{aligned}
$$

With the effective potential: $\psi = -q\Omega^2 x^2 + \dfrac{1}{2}\Omega^2 z^2$

This set of equations admits a simple solution (isothermal EOS):

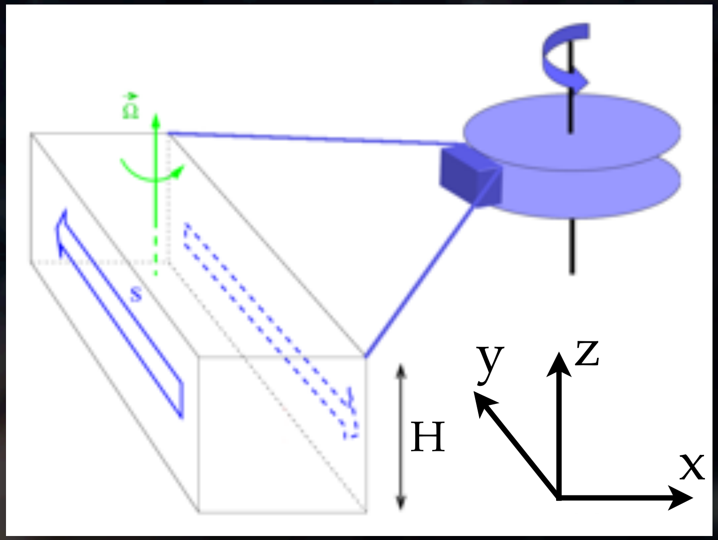
$\boldsymbol{u} = -q\Omega x \boldsymbol{e_y}$ → Mean keplerian shear

$\rho = \rho_0 \exp\left(-\dfrac{z^2}{2H^2}\right)$ → Disk scale height $H = \dfrac{c_s}{\Omega}$

The shearing box model

Separate the mean shear from the fluctuations:

$$\boldsymbol{u} = -q\Omega x \boldsymbol{e_y} + \boldsymbol{v}$$



Shearing box equations:

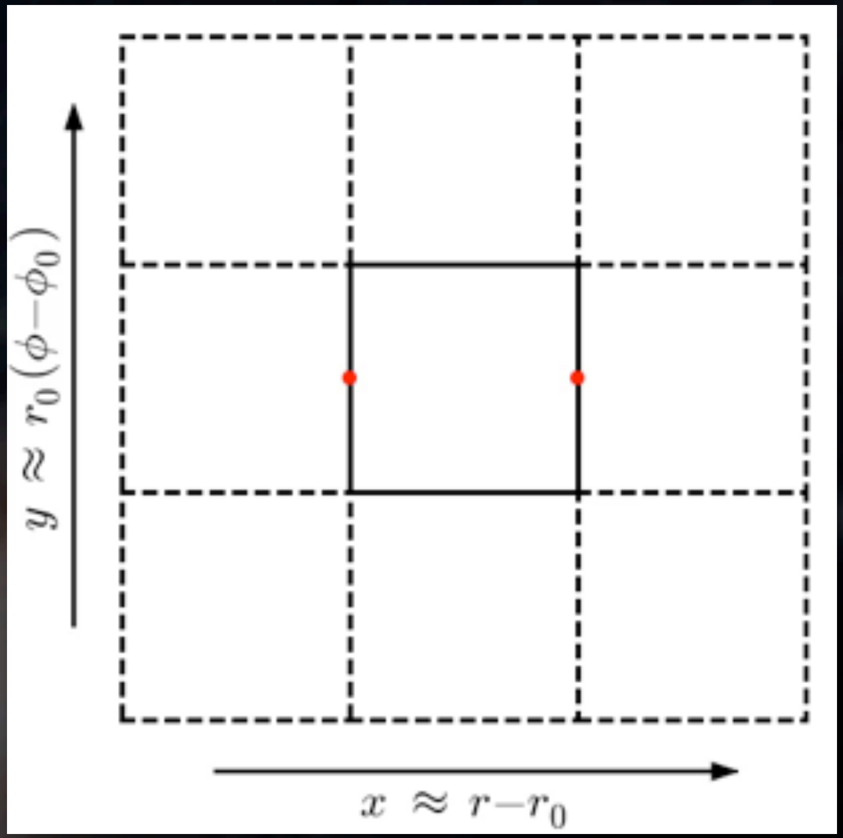
$$\begin{aligned}
\partial_t \rho - q\Omega x \partial_y \rho + \boldsymbol{\nabla} \cdot \rho \boldsymbol{v} &= 0 \\
\partial_t \rho \boldsymbol{v} - q\Omega x \partial_y \rho \boldsymbol{v} + \boldsymbol{\nabla} \cdot (\rho \boldsymbol{v} \otimes \boldsymbol{v}) &= -\boldsymbol{\nabla} P + \boldsymbol{J} \times \boldsymbol{B} - 2\rho \boldsymbol{\Omega} \times \boldsymbol{v} \\
&\quad + \rho q \Omega v_x \boldsymbol{e_y} - \rho \Omega^2 z \boldsymbol{e_z} + \rho \nu \Delta \boldsymbol{v} \\
\partial_t \boldsymbol{B} - q\Omega x \partial_y \boldsymbol{B} &= \boldsymbol{\nabla} \times (\boldsymbol{v} \times \boldsymbol{B}) - q\Omega B_x \boldsymbol{e_y} + \eta \Delta \boldsymbol{B}
\end{aligned}$$

→ Now, solve that...

Boundary conditions

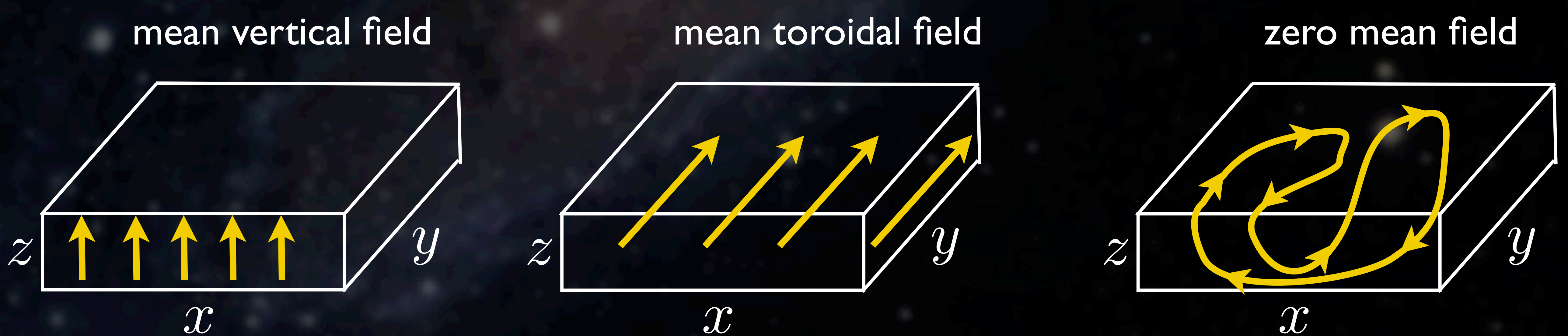
Boundary conditions

- Use shear-periodic boundary conditions= «shearing-sheet»
- Allows one to use a sheared Fourier Basis
- periodic in *y* and z (non stratified box)



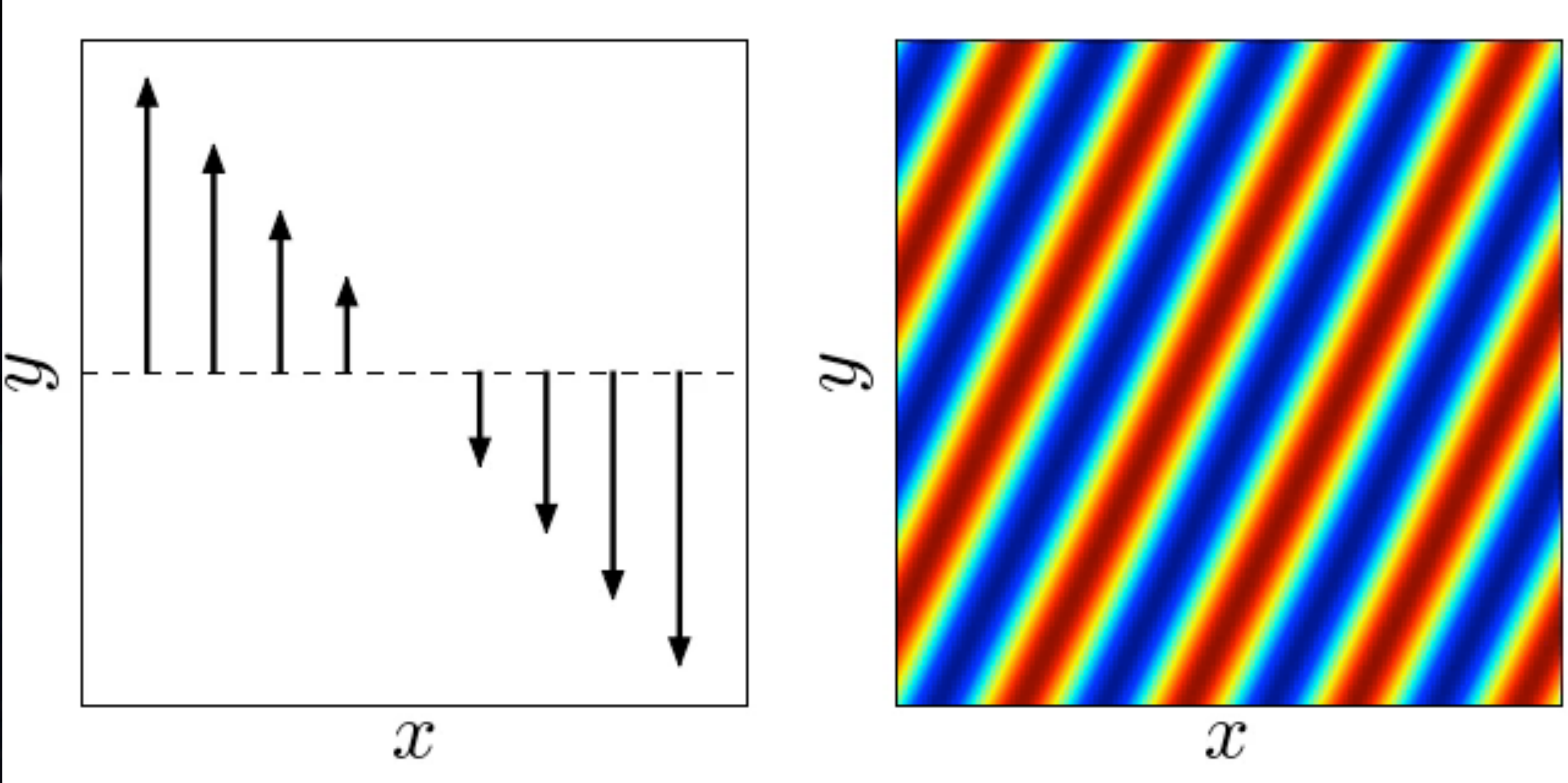
Courtesy T. Heinemann

Vertical and toroidal total magnetic flux conserved



Spectral methods for shearing boxes

Shearing wave



Courtesy T. Heinemann

Spectral methods for shearing boxes

The shearing box involves equations of the type:

$$\frac{\partial Q}{\partial t} - q\Omega x \frac{\partial Q}{\partial y} = H(Q)$$

Assume Q can be decomposed into:

$$Q(t, \boldsymbol{x}) = \tilde{Q}(t) \exp\left[i\boldsymbol{k}(t) \cdot \boldsymbol{x}\right]$$

One has:

$$\frac{\partial Q}{\partial t} = \left[\frac{d\tilde{Q}}{dt} + i\tilde{Q}\frac{d\boldsymbol{k}}{dt} \cdot \boldsymbol{x}\right] \exp\left[i\boldsymbol{k}(t) \cdot \boldsymbol{x}\right]$$

$$\rightarrow \quad \frac{d\tilde{Q}}{dt} + i\tilde{Q}\frac{d\boldsymbol{k}}{dt} \cdot \boldsymbol{x} - iq\Omega x k_y = \widetilde{H(Q)}$$

Cancel explicit *x* dependency:

$$\frac{dk_x}{dt} = q\Omega k_y \quad \rightarrow \quad \boldsymbol{k} = \boldsymbol{k}_0 + q\Omega k_y t \boldsymbol{e_x}, \qquad \frac{d\tilde{Q}}{dt} = \widetilde{H(Q)}$$

The Snoopy code
a spectral method for sheared flows

- MHD equations solved in the sheared frame
- Compute non linear terms using a pseudo spectral representation
- 3rd order low storage Runge-Kutta integrator
- OpenMP and/or MPI parallelization
- Written in C
- Available online http://ipag.osug.fr/~glesur/snoopy.html

Advantages:

- Shearing waves are computed exactly (natural basis)
- Exponential convergence when resolution is increased
- Magnetic flux conserved to machine precision
- Sheared frame & incompressible approximation: no CFL constrain due to the background sheared flow/sound speed.
- Very weak numerical dissipation: tight control on physical dissipation processes

Disadvantages:

- Slower than finite differences for the same resolution (number of real grid points)
- Shocks/diskontinuities can't be treated spectrally (Gibbs oscillations)
- Strongly parallel spectral codes are not very efficient

MRI simulations
Dimensionless numbers

- Mean field amplitude:

$$\beta = \frac{\Omega^2 H^2}{\langle v_A \rangle^2}$$

- Reynolds number

$$Re = \frac{\Omega H^2}{\nu}$$

- Magnetic Reynolds number

$$Rm = \frac{\Omega H^2}{\eta}$$

- Magnetic Prandtl number

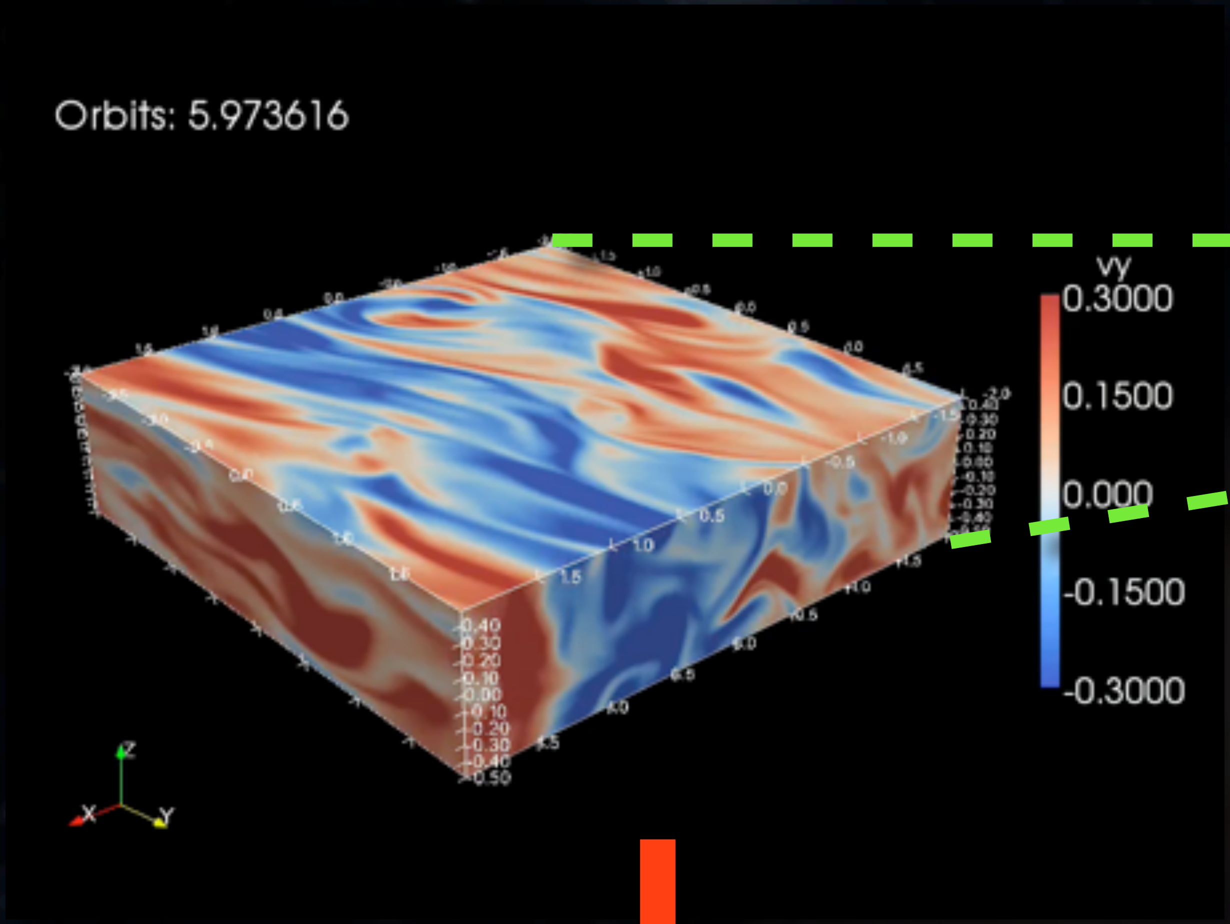
$$Pm = \frac{\nu}{\eta}$$

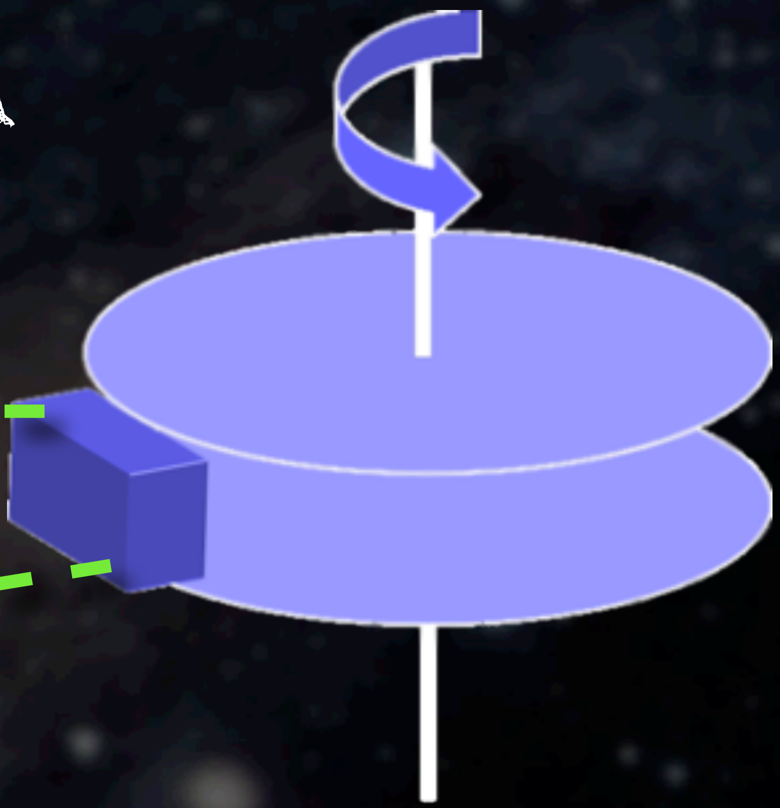
- Turbulent transport of angular momentum

$$\alpha = \frac{\langle \rho v_x v_y - B_x B_y \rangle}{\rho \Omega^2 H^2}$$

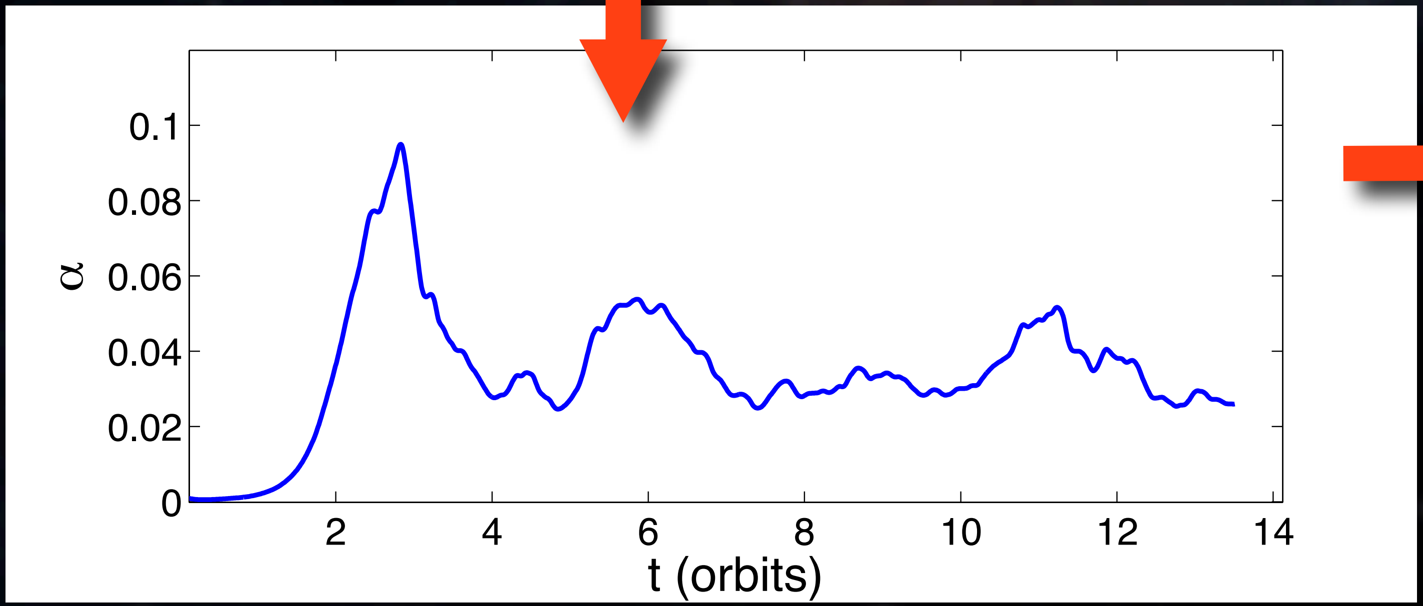
MRI simulations

Typical simulation





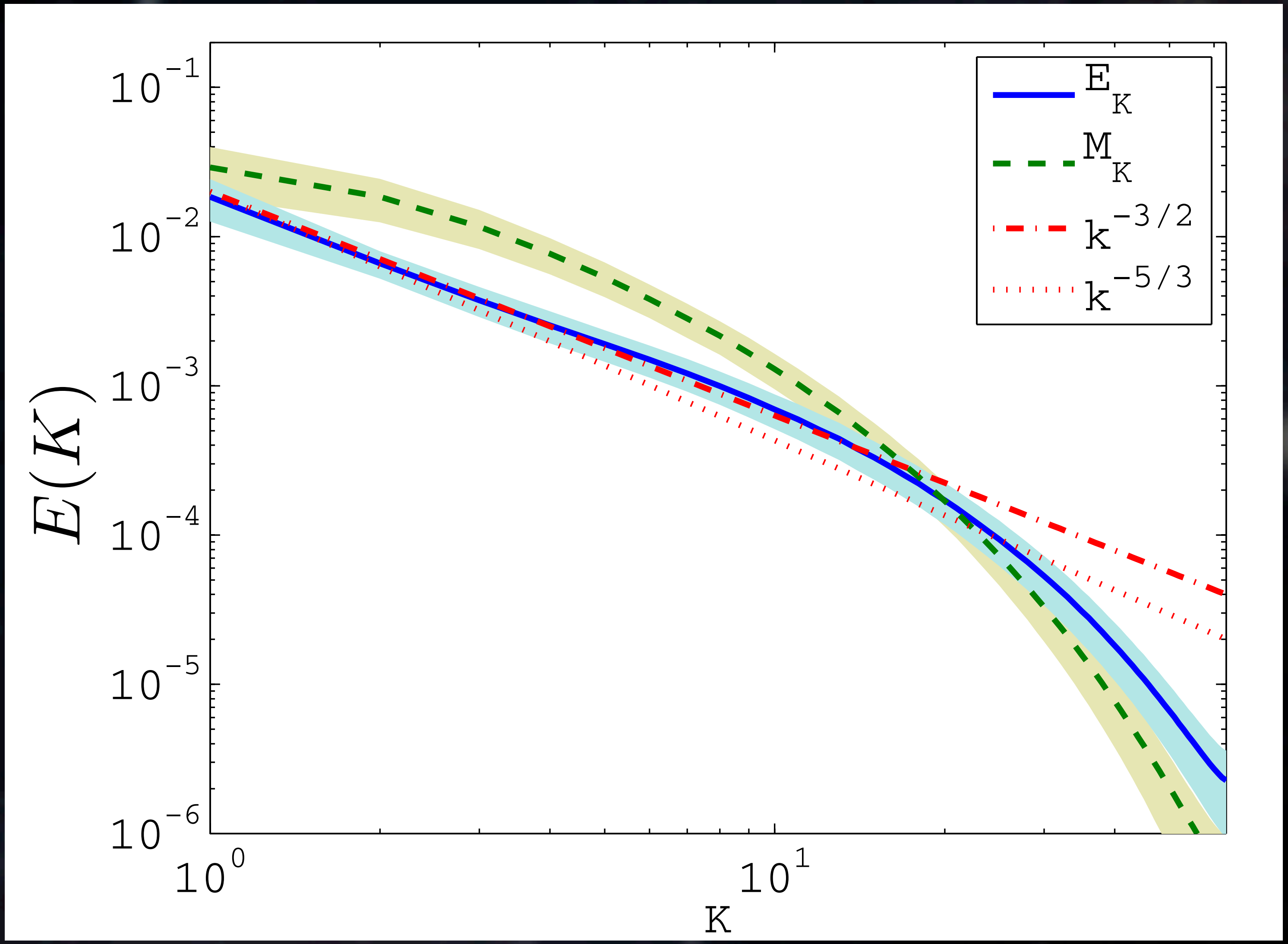
Simulation parameters: Re=1000, Pm=1, β=1000
3D map of vy (azimuthal velocity)



It works!

Is it the end of the story?

MRI simulations
Typical spectrum

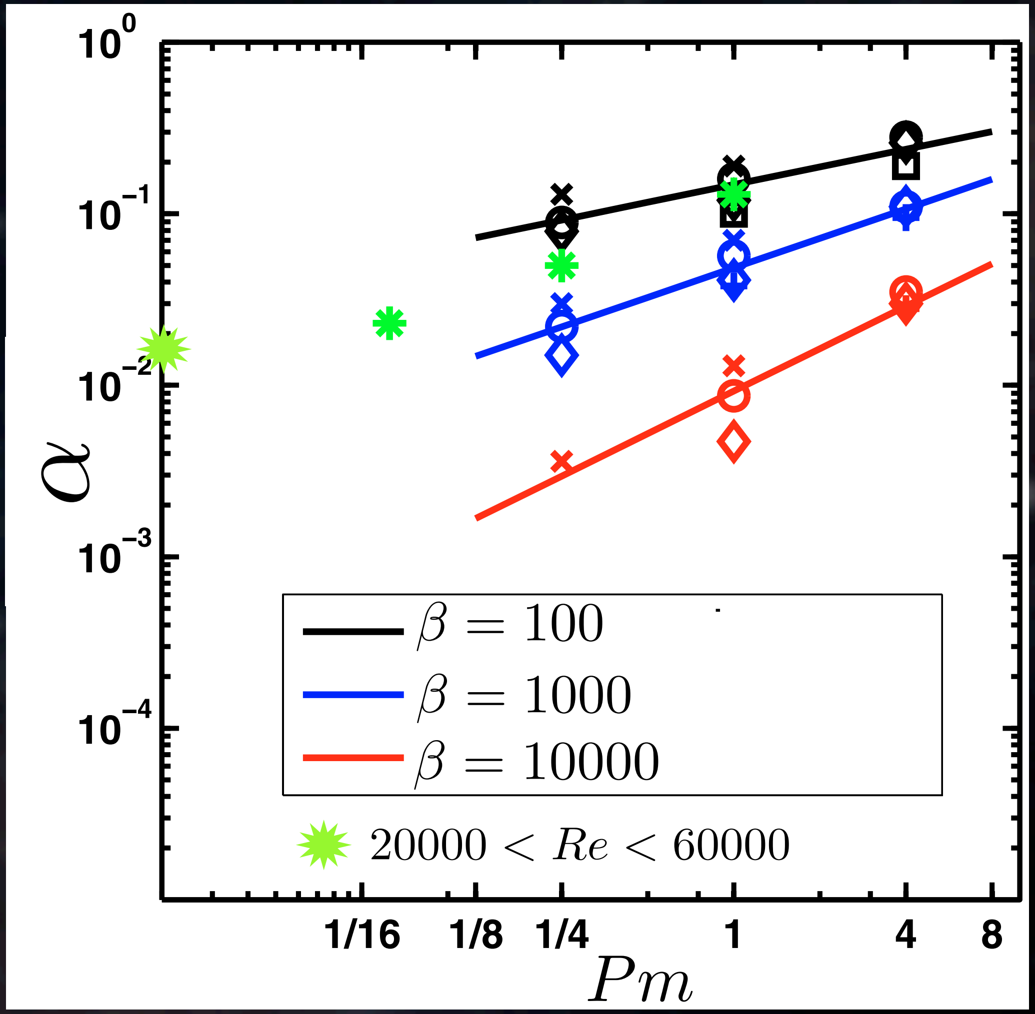


768x384x192
~500 turnover times

$K^{-3/2}$ for kinetic energy?

MRI simulations

Simulations with a mean vertical field



Longaretti & Lesur (2010)

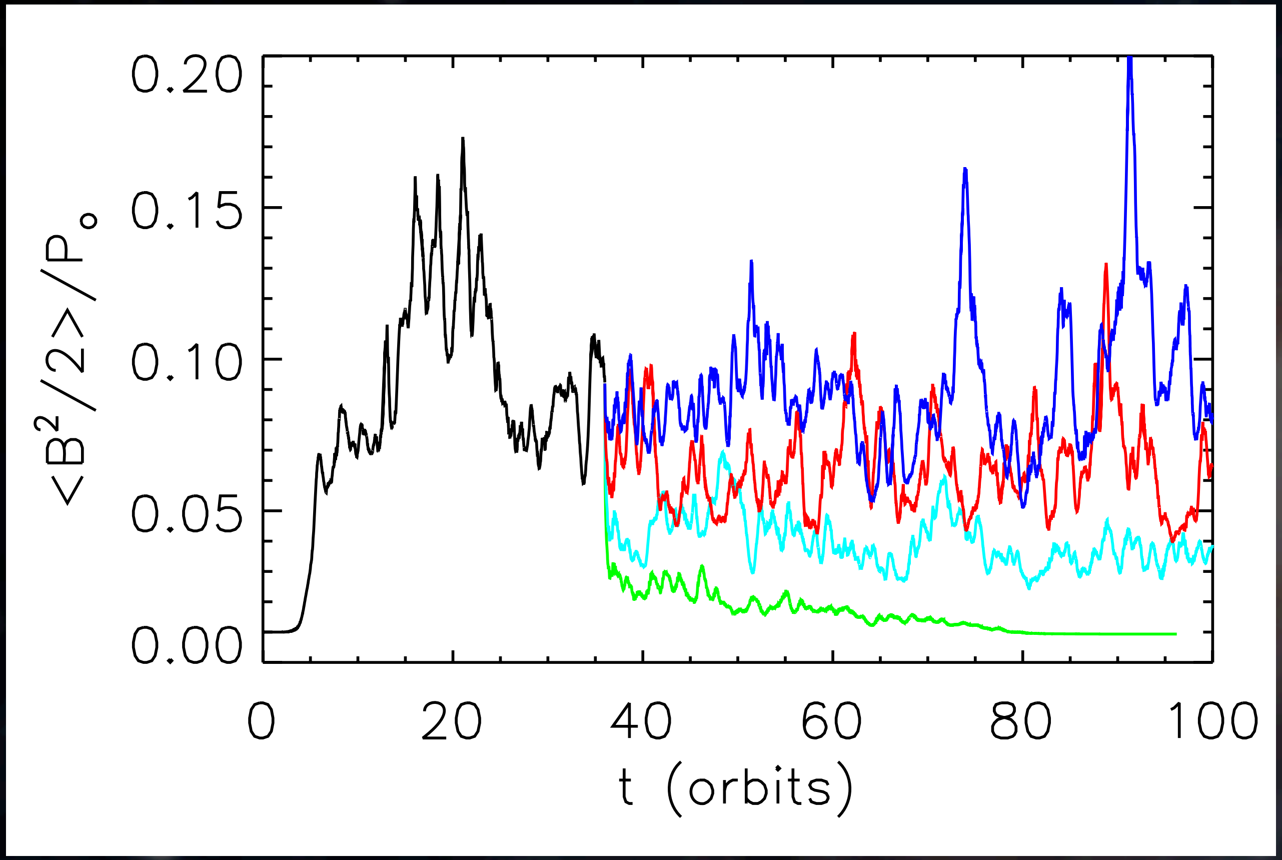
Turbulent transport varies by 2 order of magnitude!

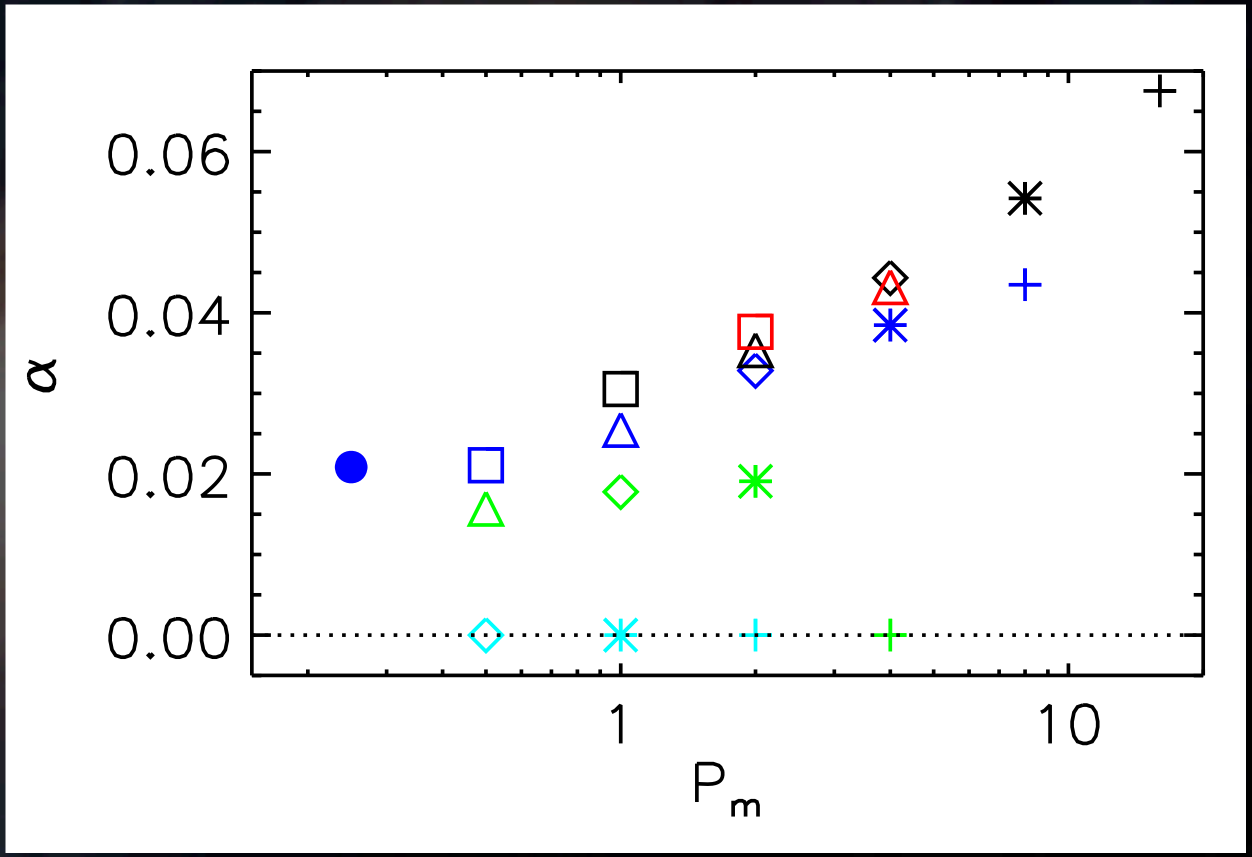
$$\alpha = \alpha(Pm, \beta)$$

MRI simulations

Simulations with a mean toroidal field

Simon & Hawley (2009)





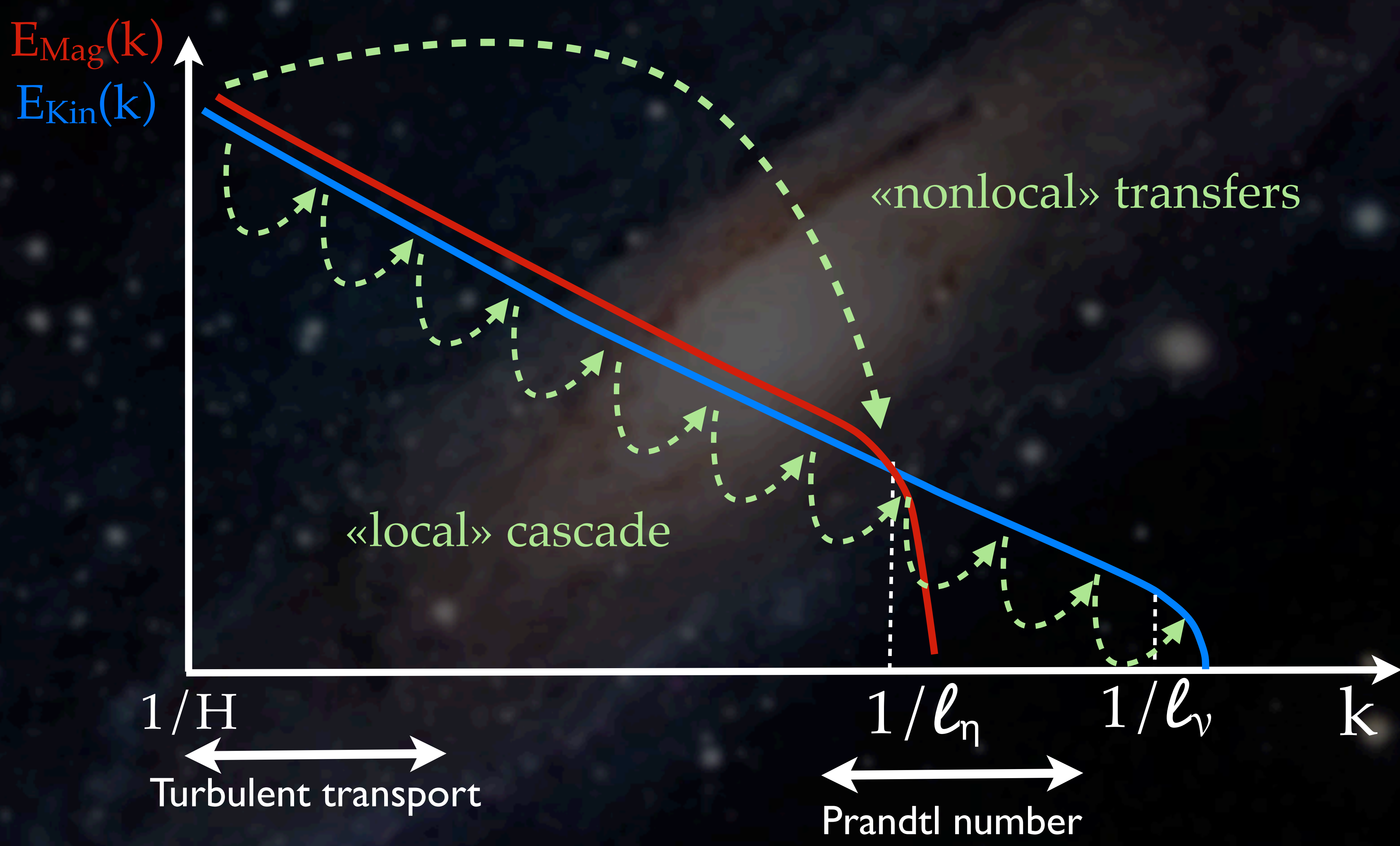
$\beta = 100$

- Weaker transport with a mean toroidal field
- Same trend with *Pm*

$$\alpha = \alpha(Pm, \beta, \text{topology})$$

MRI simulations

Simulations with a mean field

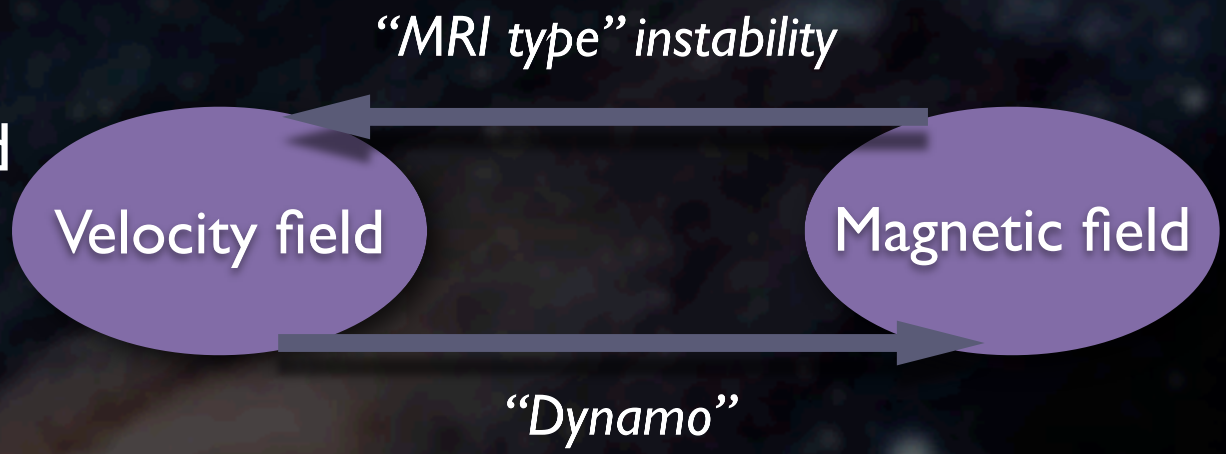


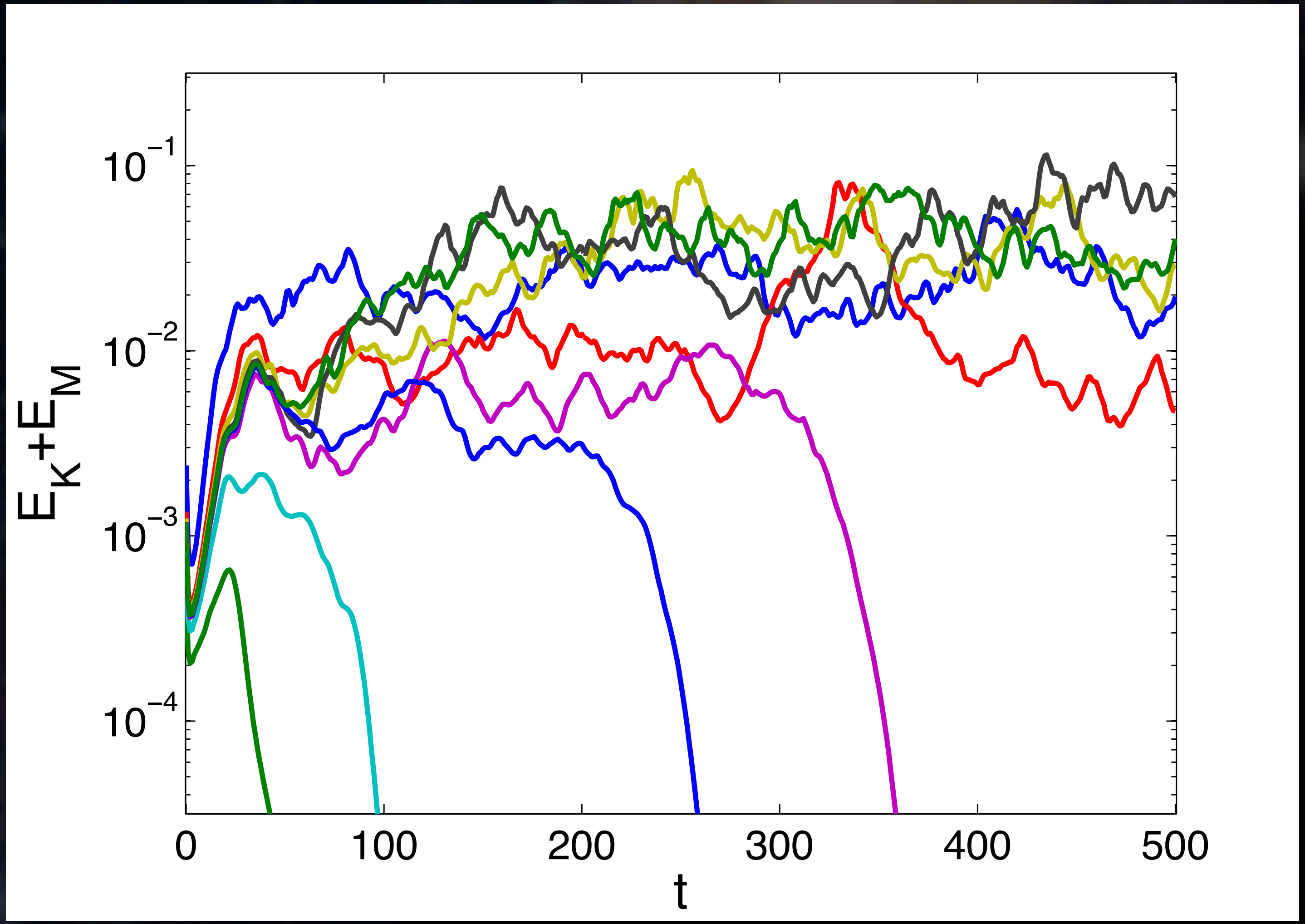
→ Pm effect shows nonlocal transfers are important in numerical simulations

MRI simulations

Simulations with no mean field

No mean field is a peculiar situation: the field has to be regenerated by a dynamo action. Often called «MRI dynamo»

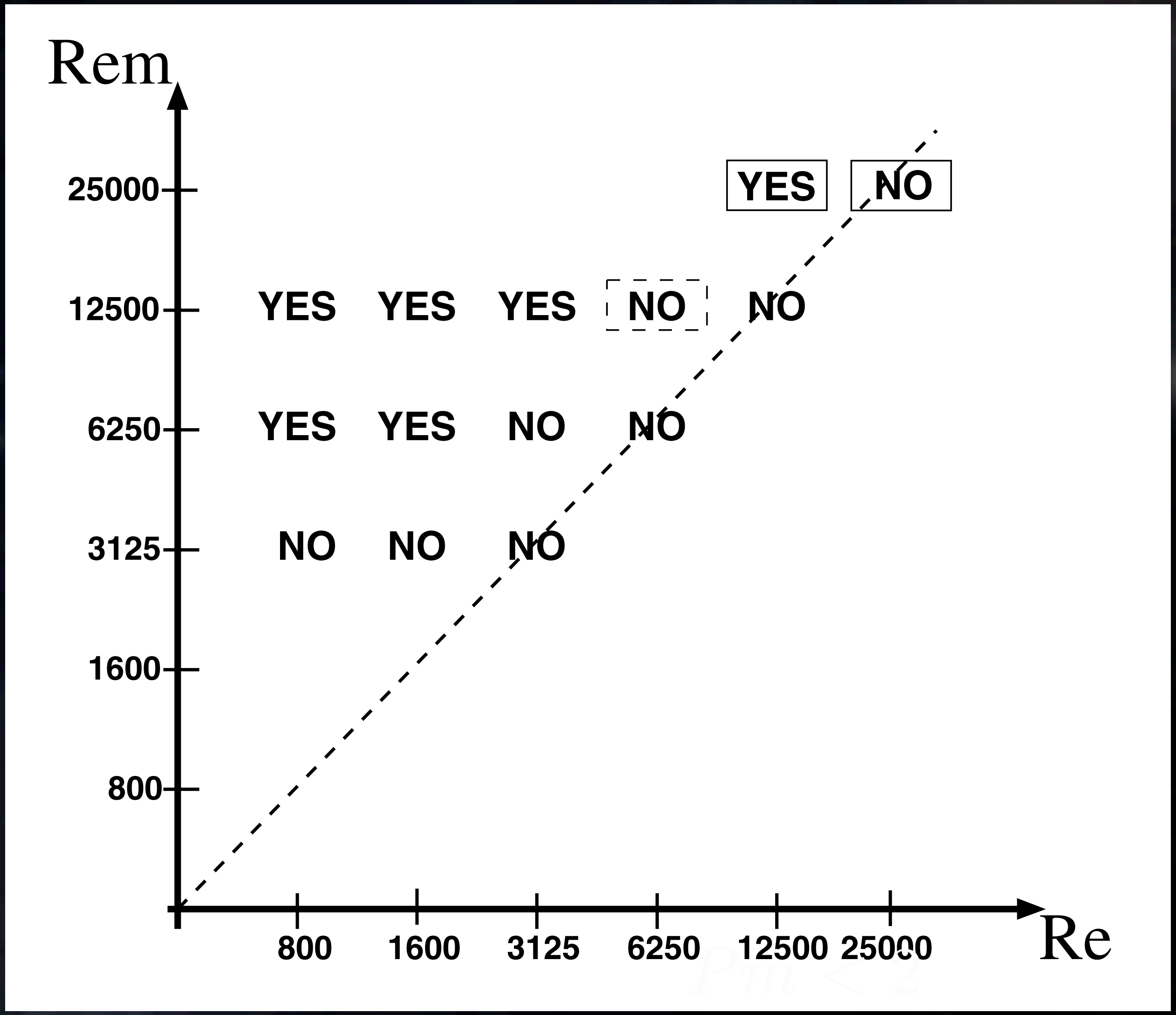




Subcritical instability!

MRI simulations

Simulations with no mean field

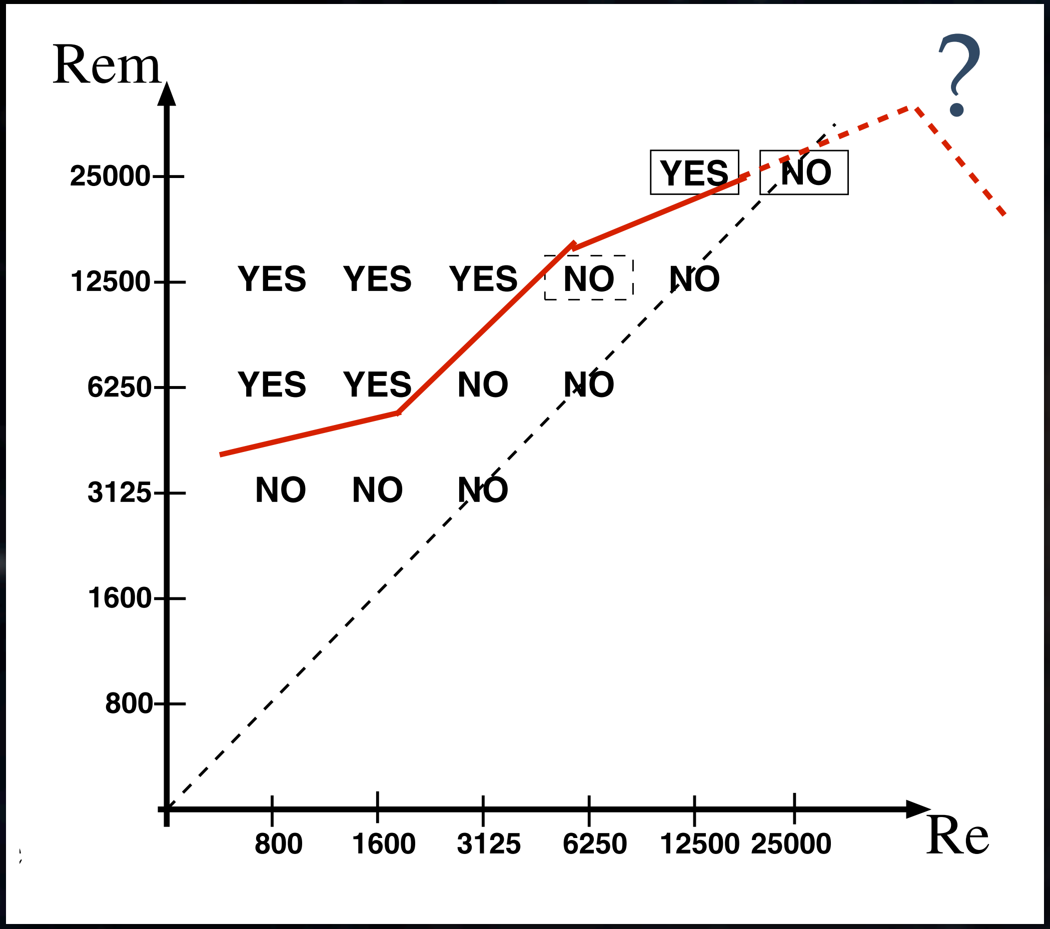


Fromang *et al.* 2007

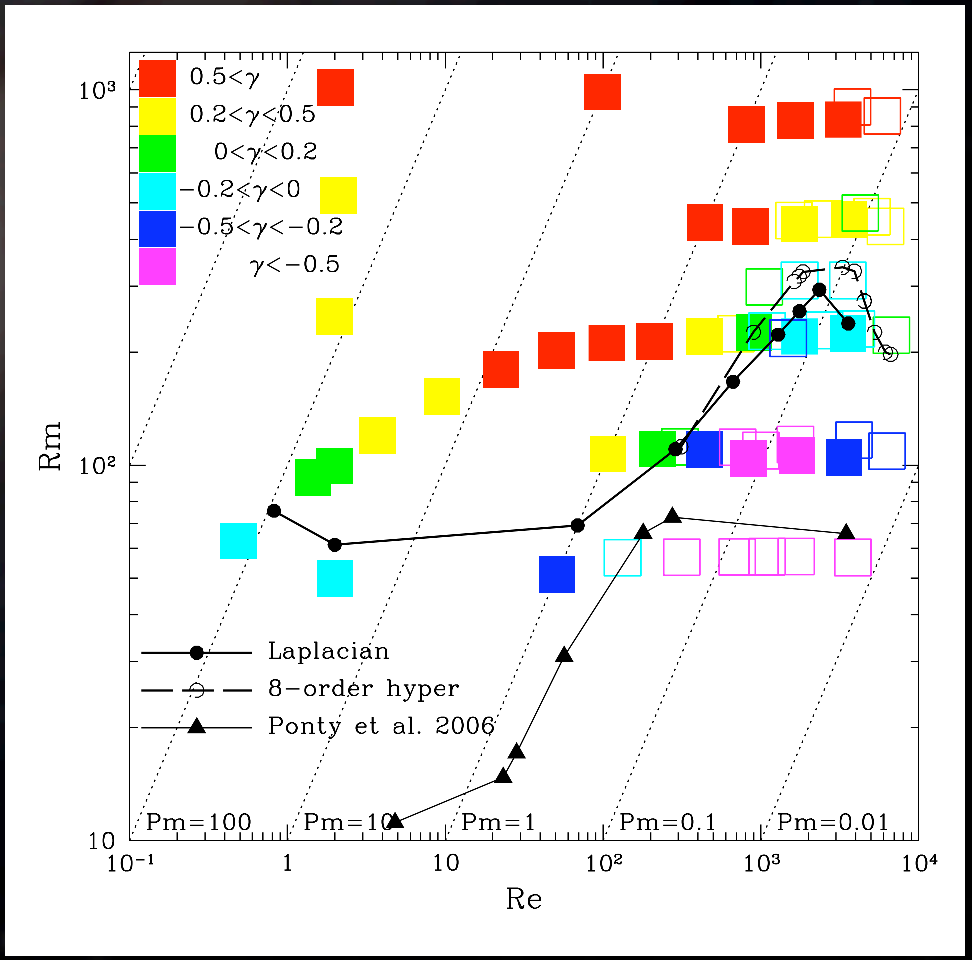
MRI simulations

Simulations with no mean field

Zero net flux MRI
(Fromang et al. 2007)



Small scale dynamo
(Schekochihin et al. 2006)



- Similar behaviour in the limit of small Pm ?

MRI simulations

Simulations with no mean field

Understanding the mechanism underlying the MRI dynamo

Brute force approach

Use the biggest computer with the largest resolution possible

Cross your fingers

«try to be smart» approach

Simplify the system so that only relevant physical processes are in the simulation

Is it a good representation of reality?

«Do we care?» approach

Set

$$Re = Rm = \infty$$

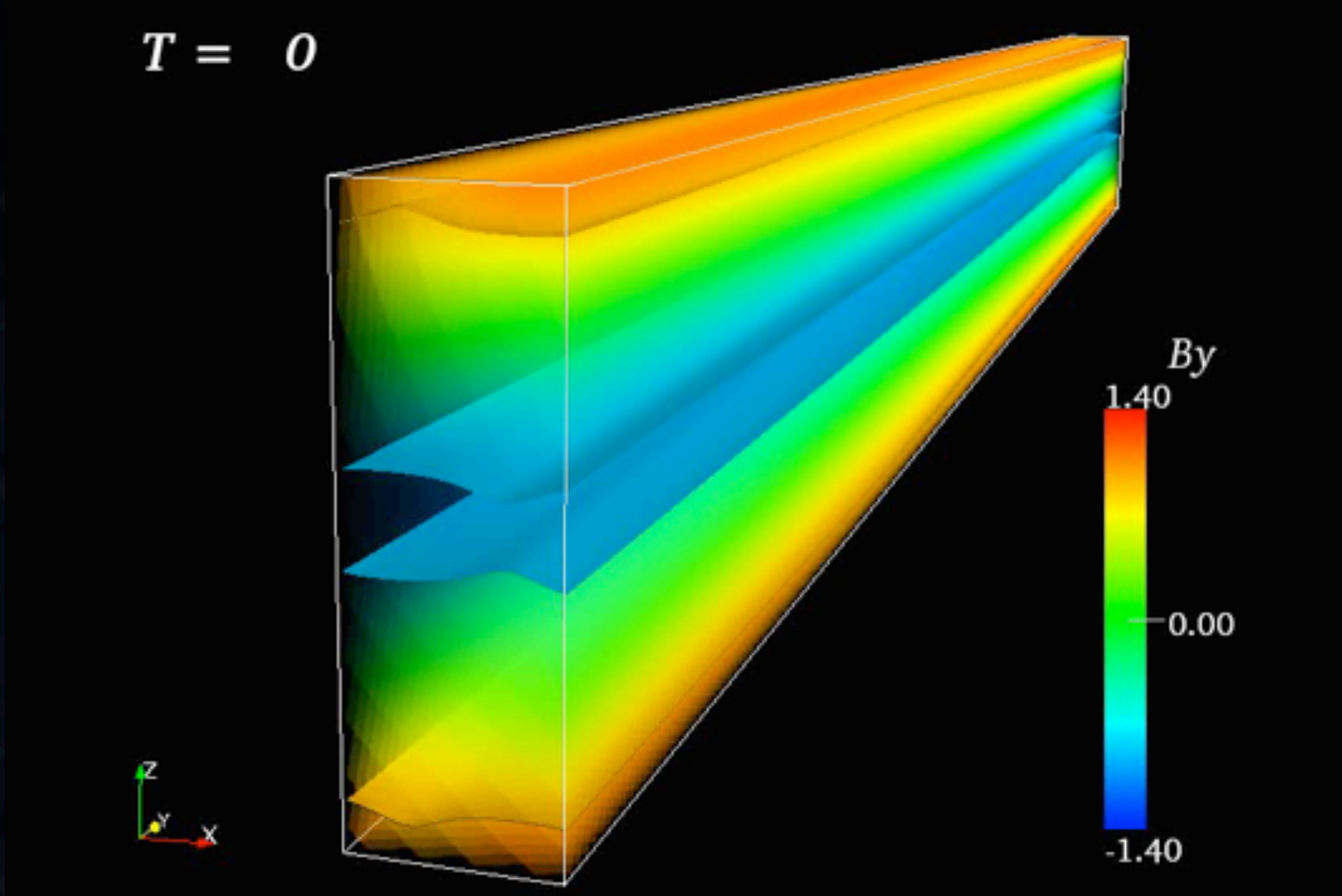
in your simulation

Problem solved (but your simulation is dominated by numerical artefacts...)

MRI simulations

Simulations with no mean field

Try to be smart approach



Herault et al. 2011

- Identify dynamo cycles and characterise their mechanism
- Turbulent transport can be described as a sum on these cycles

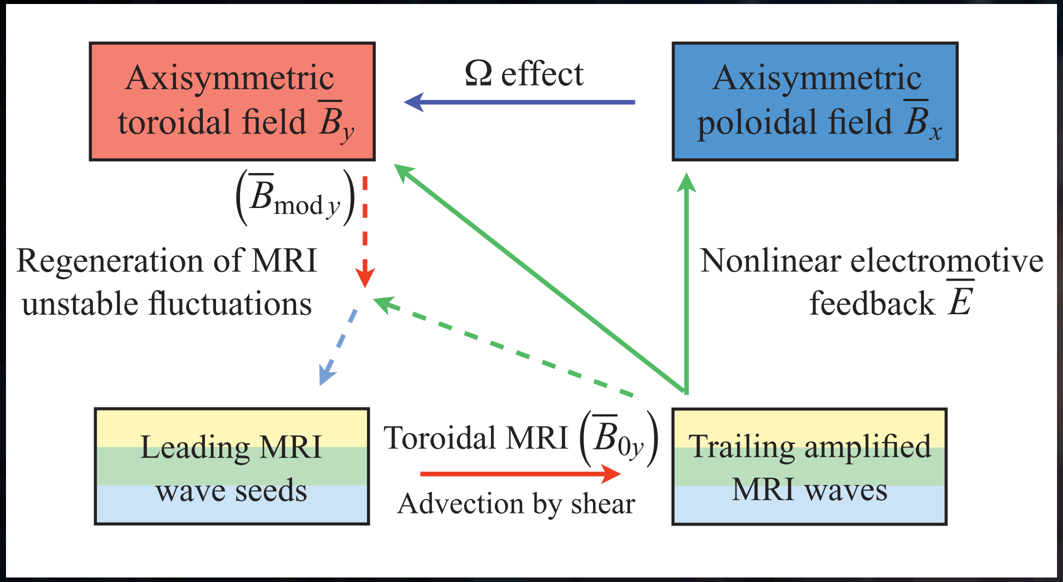
→ A systematic way to characterise the transport due to the MRI?

MRI simulations

Simulations with no mean field

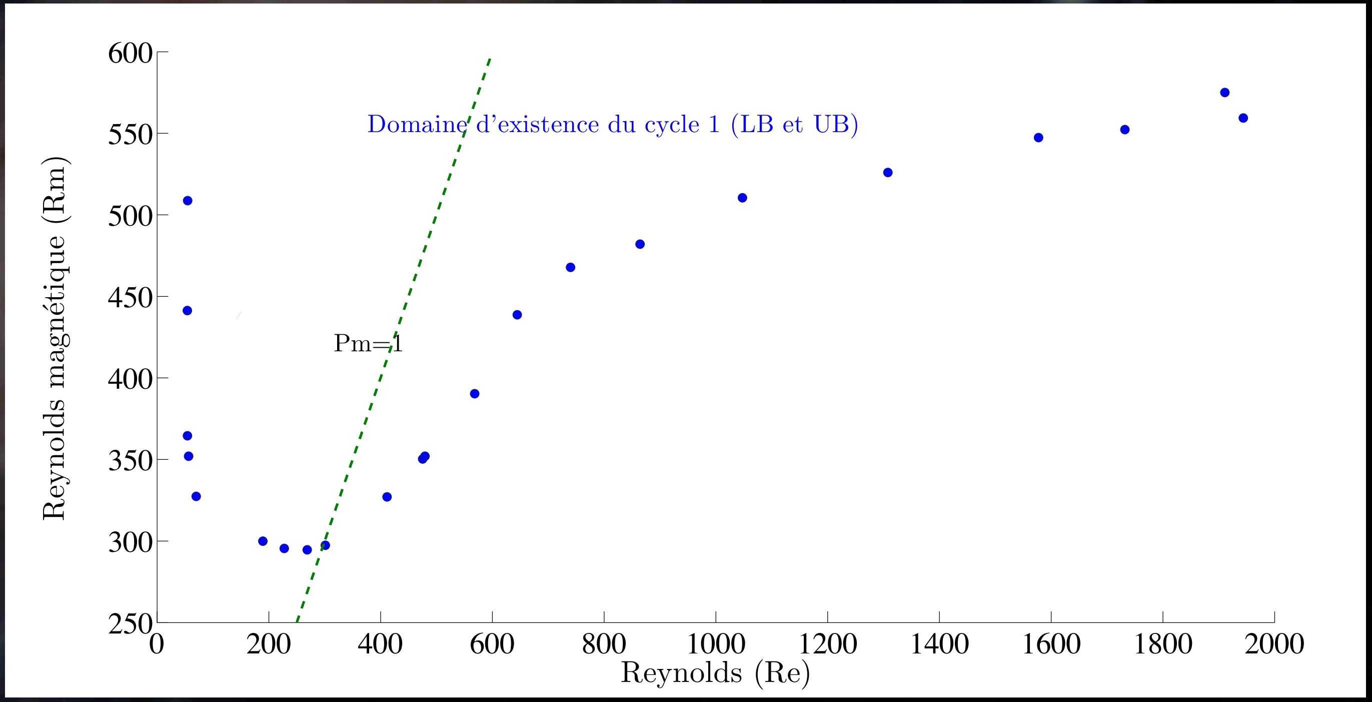
Dynamo cycle

Describe a dynamo cycle



Herault et al. 2011

Follow a dynamo cycle in (Re,Rm) plane



See A. Riols poster

MRI simulations
Conclusion

How big is α-MRI ?

- It depends on the field strength
- It depends on the field topology
- It depends on the magnetic Prandtl number

$$\rightarrow \quad 0 \lesssim \alpha \lesssim 10^{-1}$$

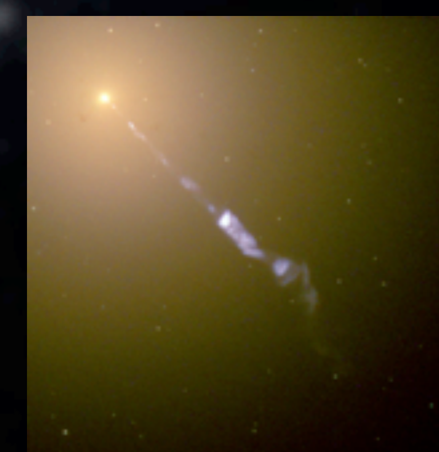
Overall, we don't really know...

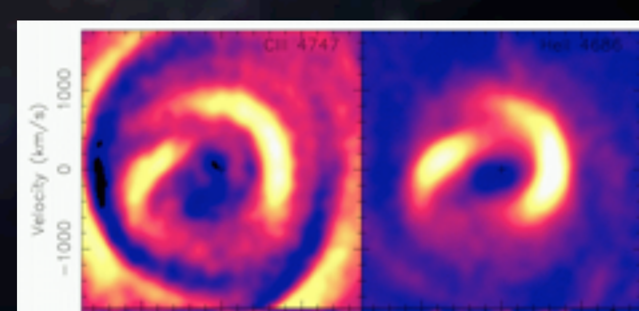
Outline

- Accretion disks and jets: what are they
  - Accretion disks in nature
  - Jets in nature
  - The physics of accretion
- Turbulent disks
  - Disk Instabilities
  - The shearing box model
  - The case of the MRI
- Outflows in disks
  - Jet launching mechanisms
  - MRI & disk winds
  - disk wind stability

Remember: disks and jets in nature

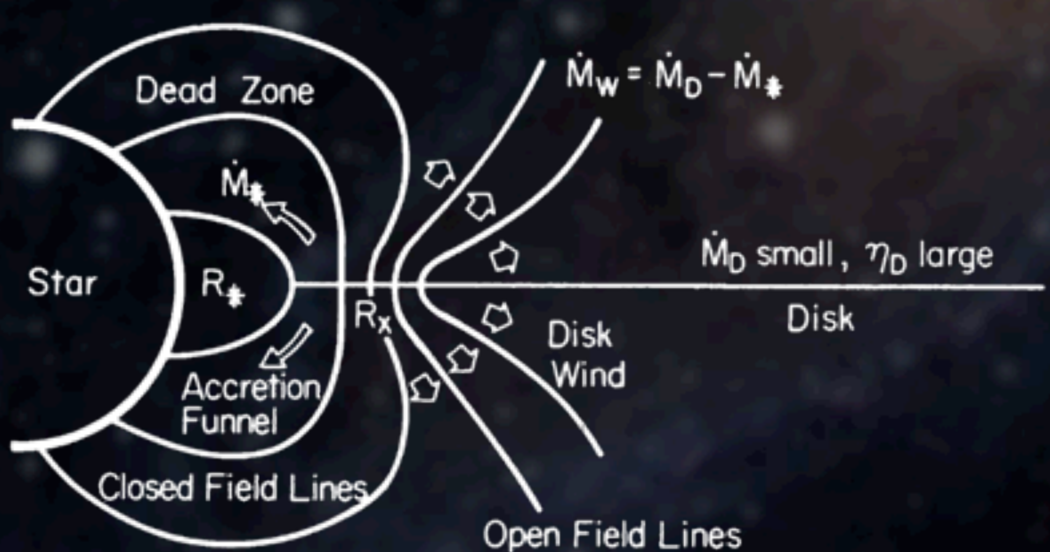
- Disks are very general in nature
- Almost all of them are associated to powerful jets (with the notable exception of CVs)



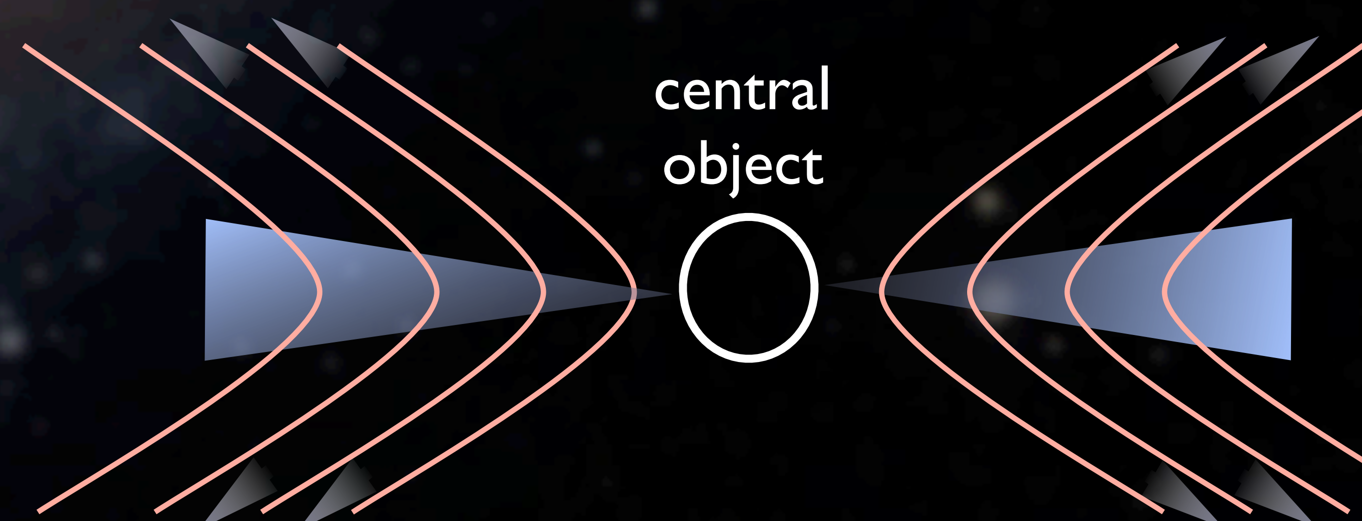


Launching a jet

- Blandford-Znajek: energy extracted by the magnetic field surrounding a rotating black hole (Blandford & Znajek 1977)
- X-wind: ejection from a reconnection point in the disk (Shu et al. 1994)
- Disk wind: ejection due to a magnetocentrifugal acceleration process in the disk (Blandford & Payne 1982)
- Stellar wind

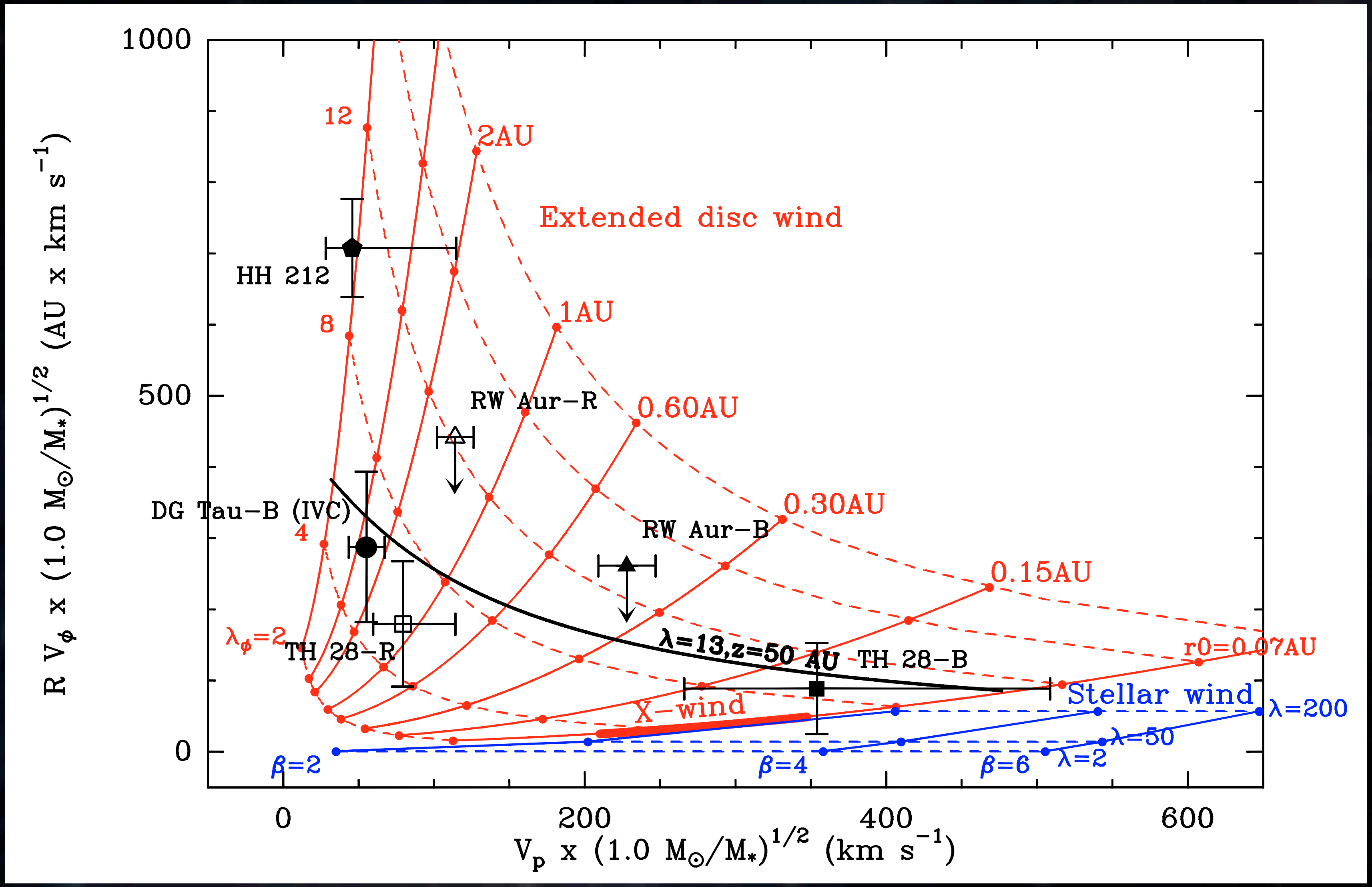


X-wind



Disk wind

Which Jet launching mechanism in young stars?



Ferreira et al. 2006

Jets are rotating → disk winds are favoured

Launching a jet

The Blandford & Payne launching mechanism

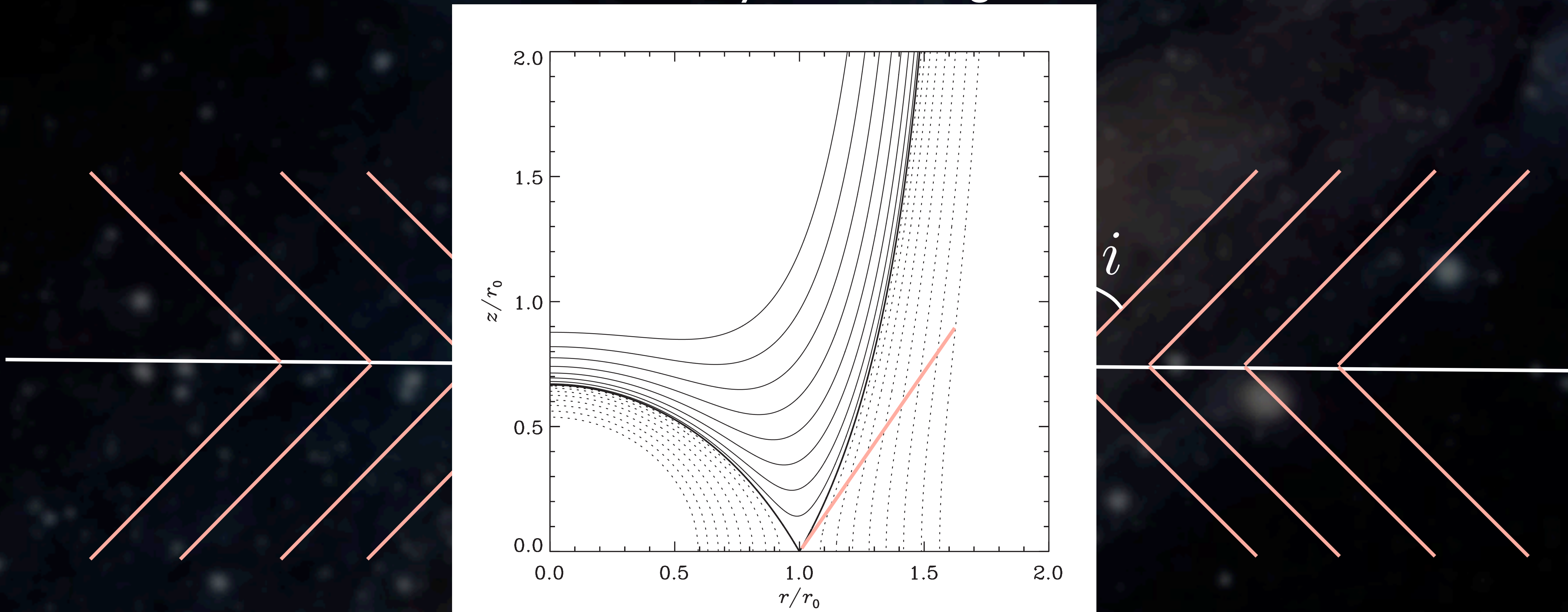


Figure 1. Contours of the effective potential experienced by matter that is forced to rotate with the Keplerian angular velocity at radius r_0. The contour values are unequally spaced. Dotted contours correspond to values lower than that of the saddle point.

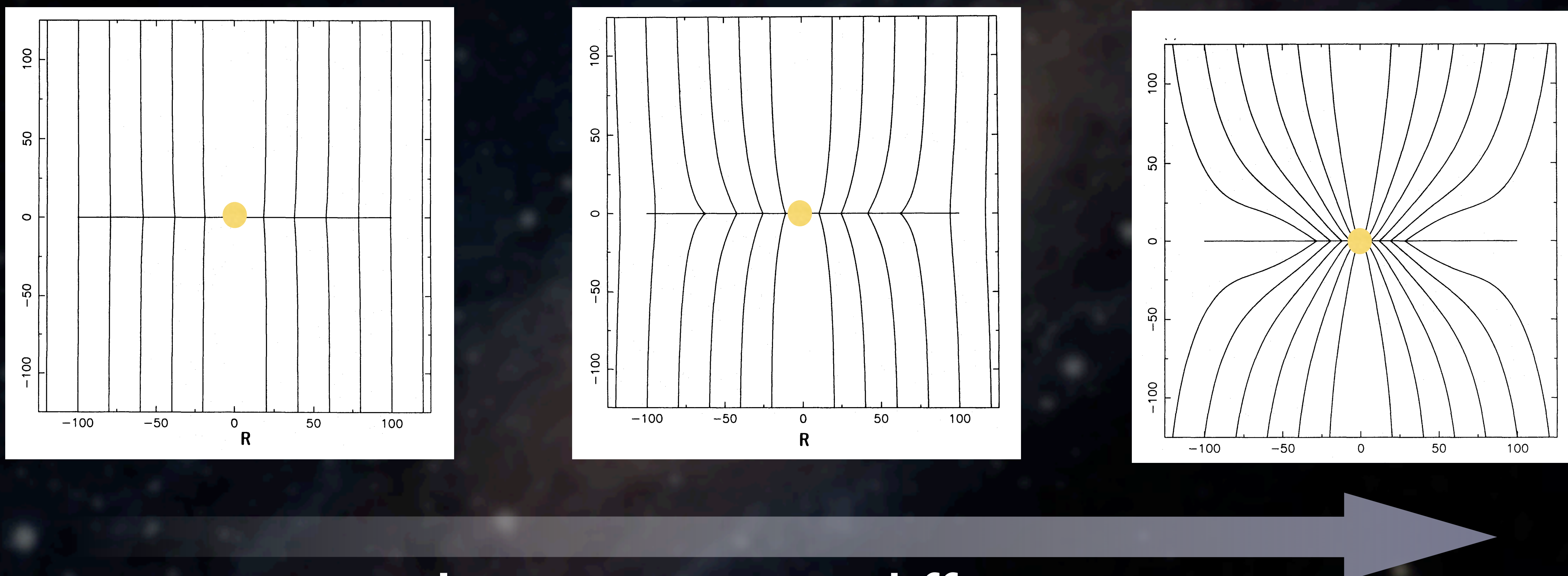
Launching a particle at rest from z=0 requires

$$i > 30^\circ$$

Large scale magnetic configuration

- Ejection requires relatively inclined field lines (i>30°)
- How is it sustained?

Lubow et al. 1994

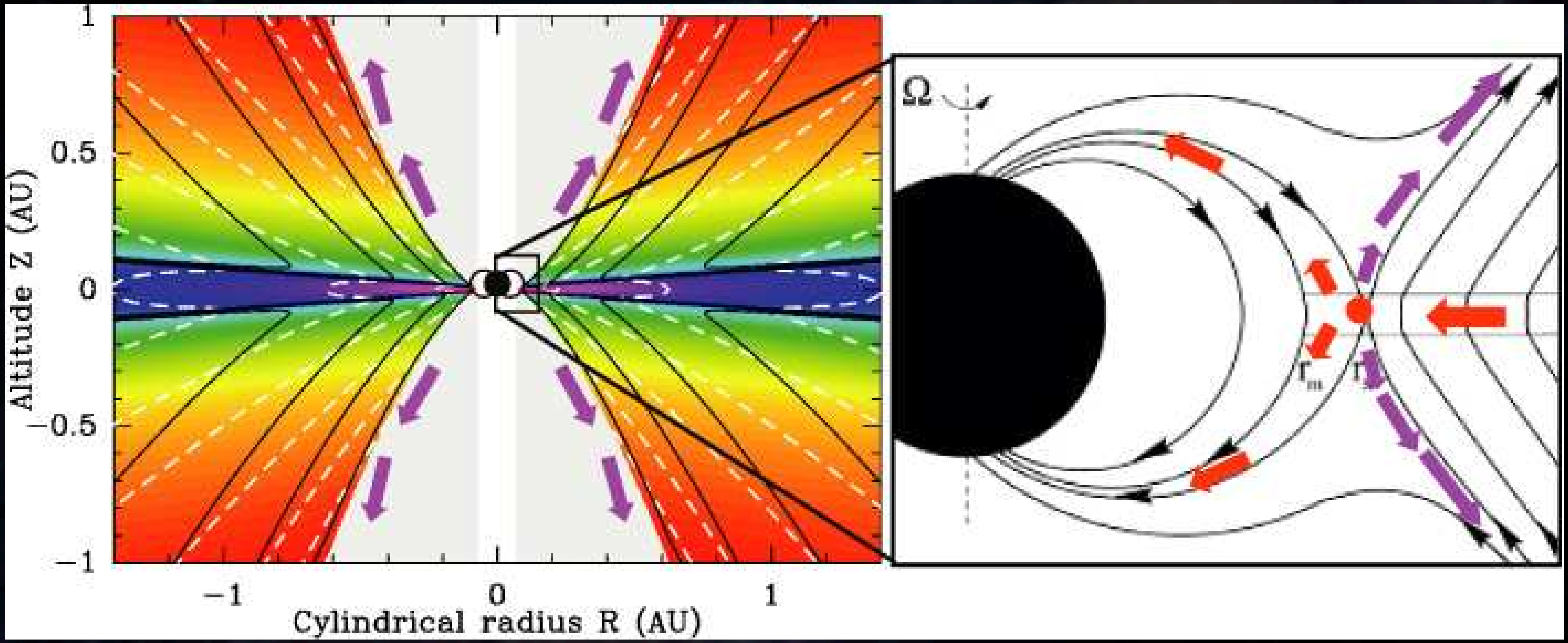


Less magnetic diffusion

- Need magnetic diffusion (so that matter can be accreted)
- Not too much ! (otherwise i=0°)

Launching a jet

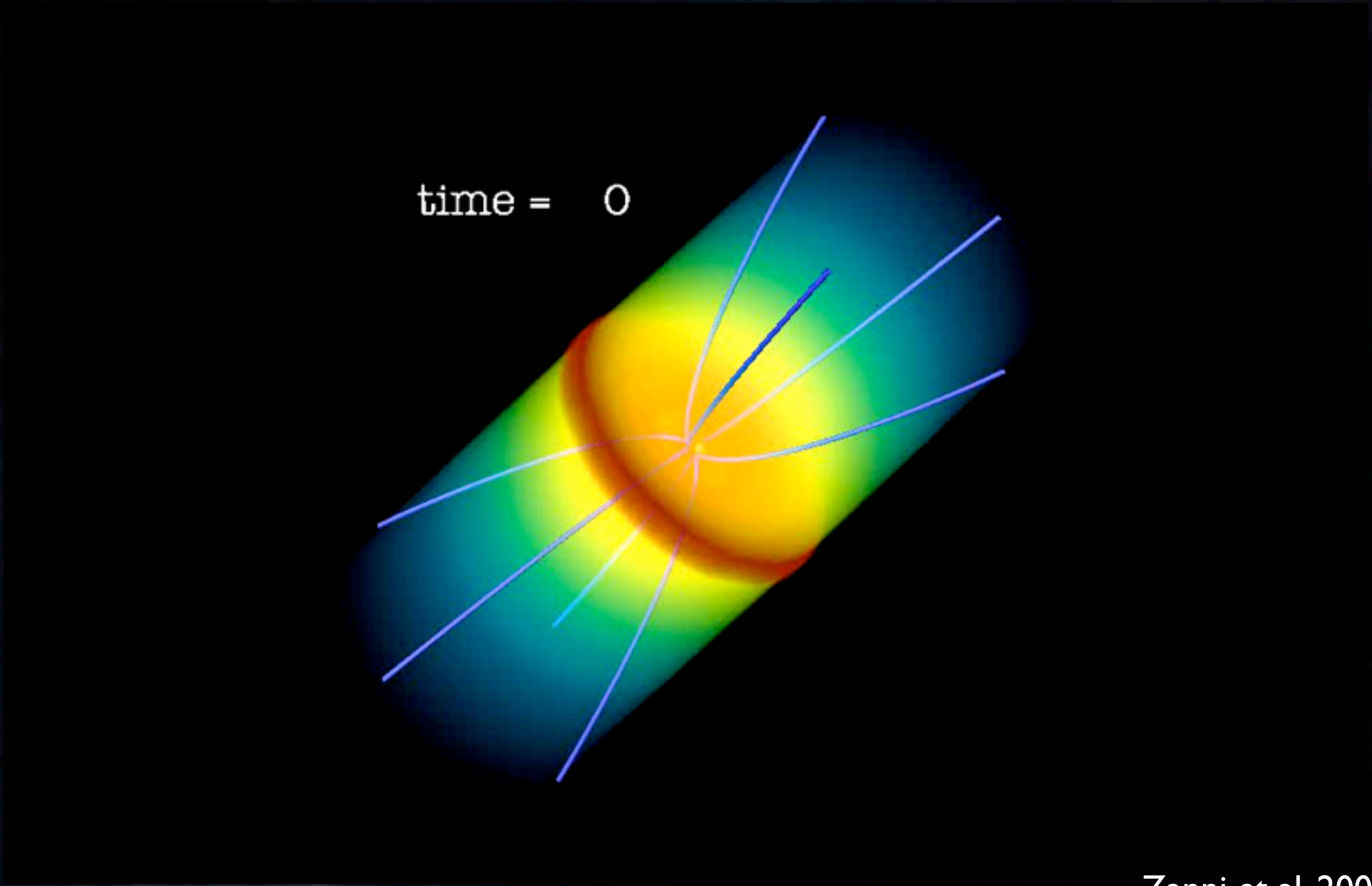
Global (2D) Self-similar solutions



Ferreira 1997

Launching a jet

Global simulation (2.5 D) with magnetic diffusion



Zanni et al. 2007

Condition of existence

Disk wind scenario

- Magnetic diffusion $\alpha_M \sim 1$ → Turbulence?
- Poloidal field at equipartition $\beta \sim 1$

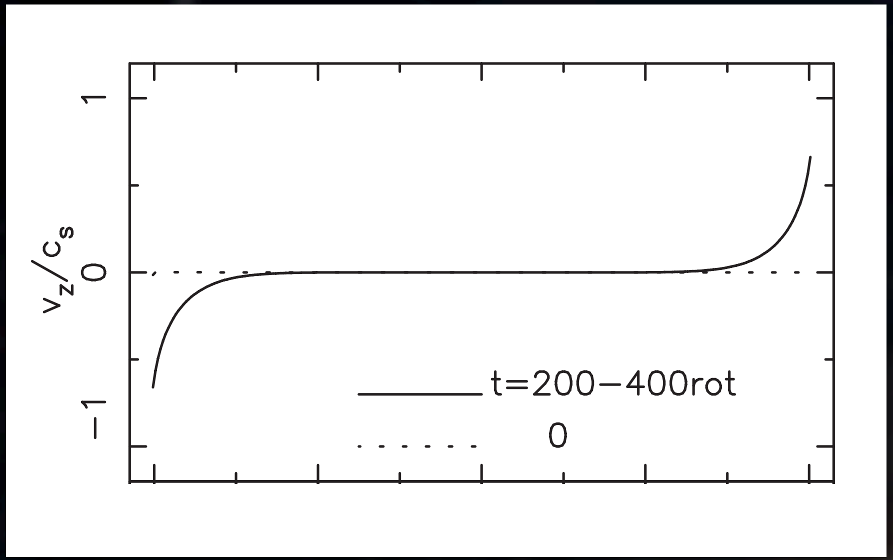
Refs: Ferreira 1997, Casse & Ferreira 2000, Zanni et al. 2007

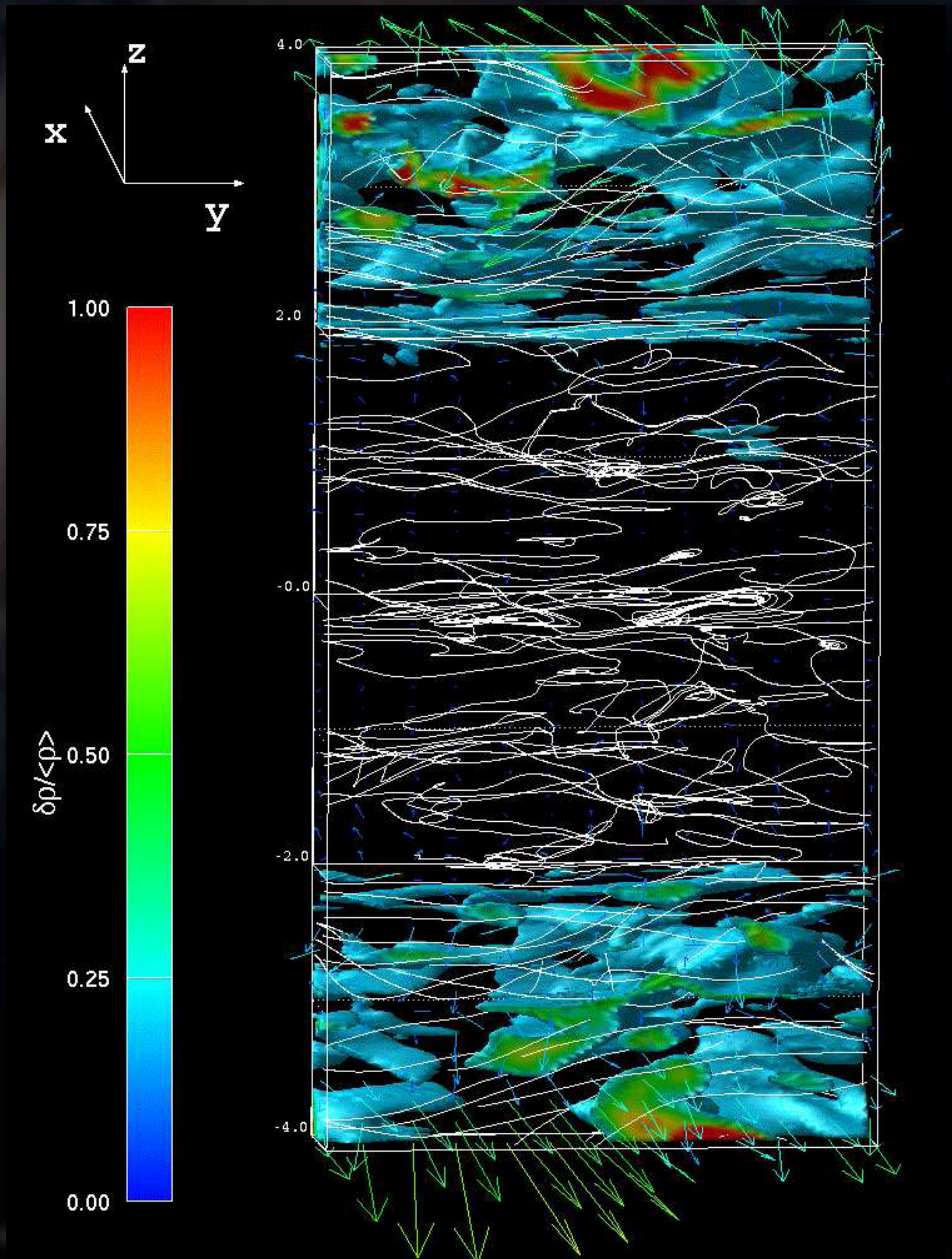
One need to understand the interplay between the MRI and jet launching mechanisms

→ Outflows in a shearing box

Turbulent driven «evaporation»

$$\beta = 10^4 — 10^5$$





- MRI with outflow boundary conditions
- «Escape» speed~sound speed
- No wind driving mechanism (matter ultimately falls back?)

Suzuki & Inutsuka (2009)

MRI in a strongly magnetised & stratified box

MRI dispersion relation:

$$\omega^4 - \omega^2[\kappa^2 + 2(\boldsymbol{k}\cdot\boldsymbol{v_A})^2] + (\boldsymbol{k}\cdot\boldsymbol{v_A})^2\left[(\boldsymbol{k}\cdot\boldsymbol{v_A})^2 + \frac{d\Omega^2}{d\ln R}\right] = 0$$

MRI is quenched when: $k\cdot v_A = \sqrt{3}\Omega$

Assuming $k \sim H^1$ and using $c_s = \Omega H$: $\beta_{\mathrm{quench}} \sim 1$

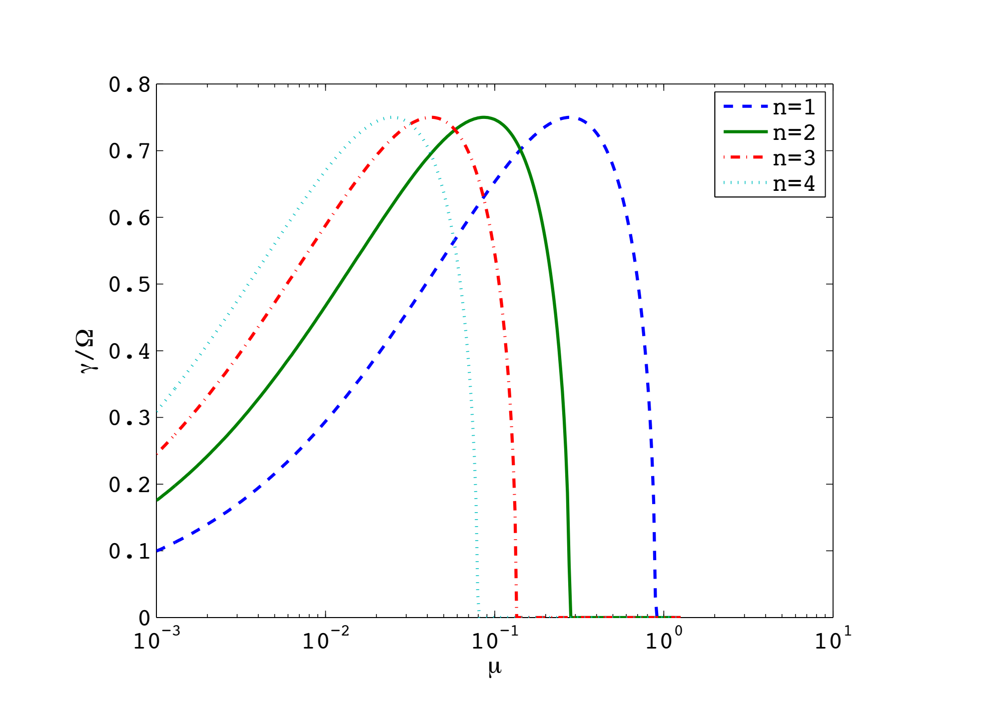
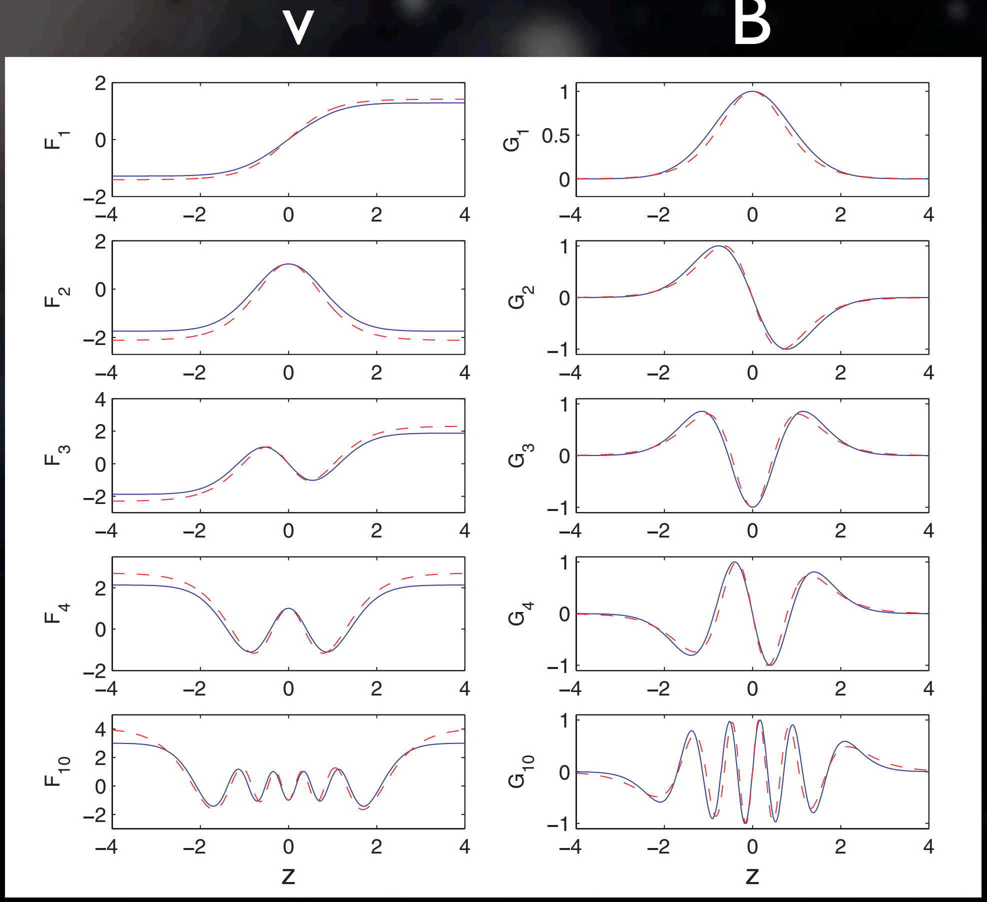


Fig. 2. Growth rates of the largest stratified MRI eigenmodes as a function of μ. These growth rates were deduced from eq. (18) in Latter et al. (2010).

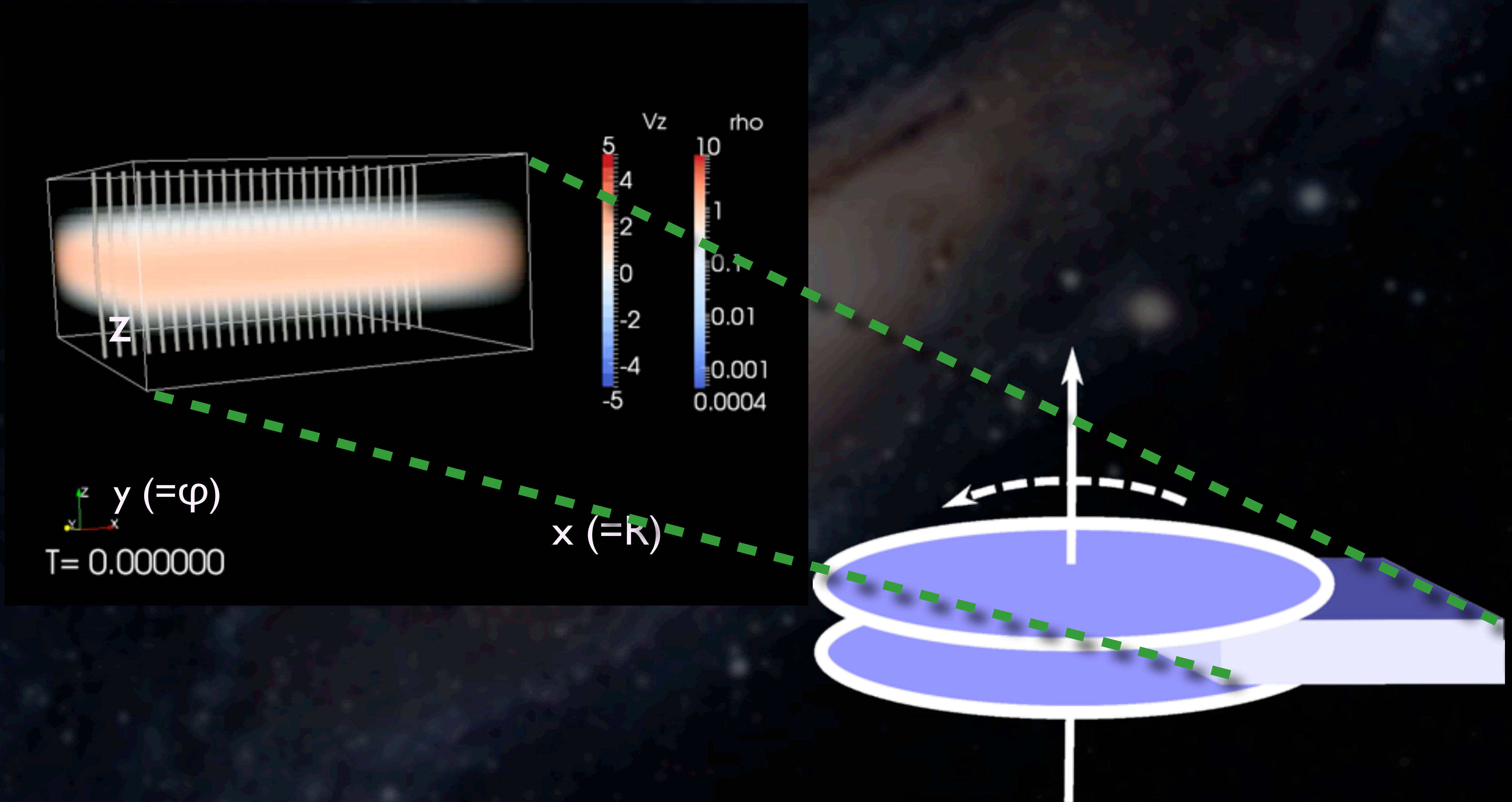
Growth rates



Eigenmodes

Strongly magnetised MRI modes (II)

$$\beta = 10$$



Outflow streamline configuration

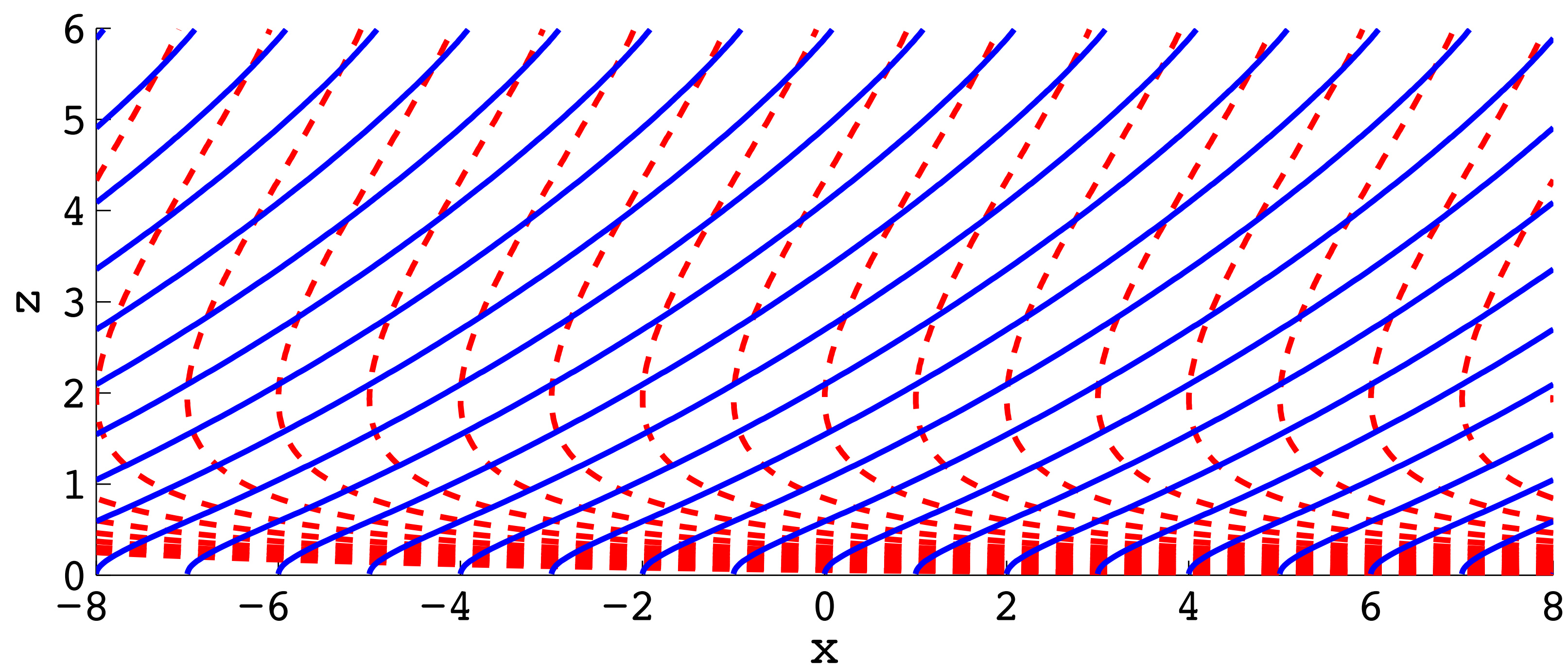


Fig. 3. Streamlines (red dashed lines) and field lines (blue plain lines) of our steady solution obtained at $t = 95$.

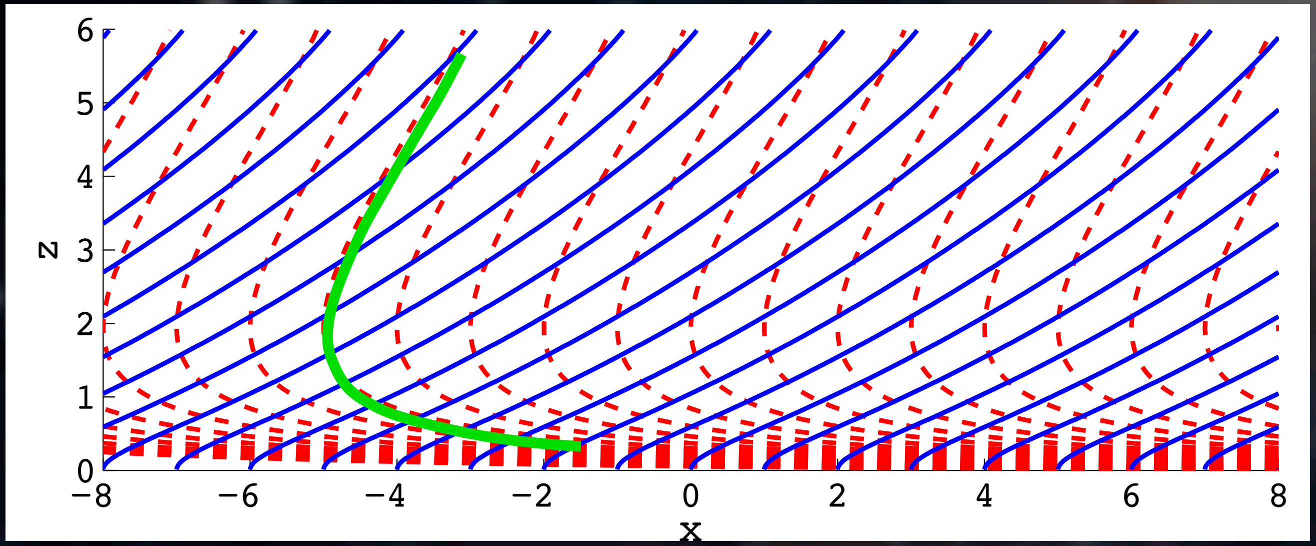
- Streamlines inclination~30° (Satisfies Blandford & Payne criterion)
- Poloidal ***v*** and ***B*** not aligned → magnetic flux is advected.

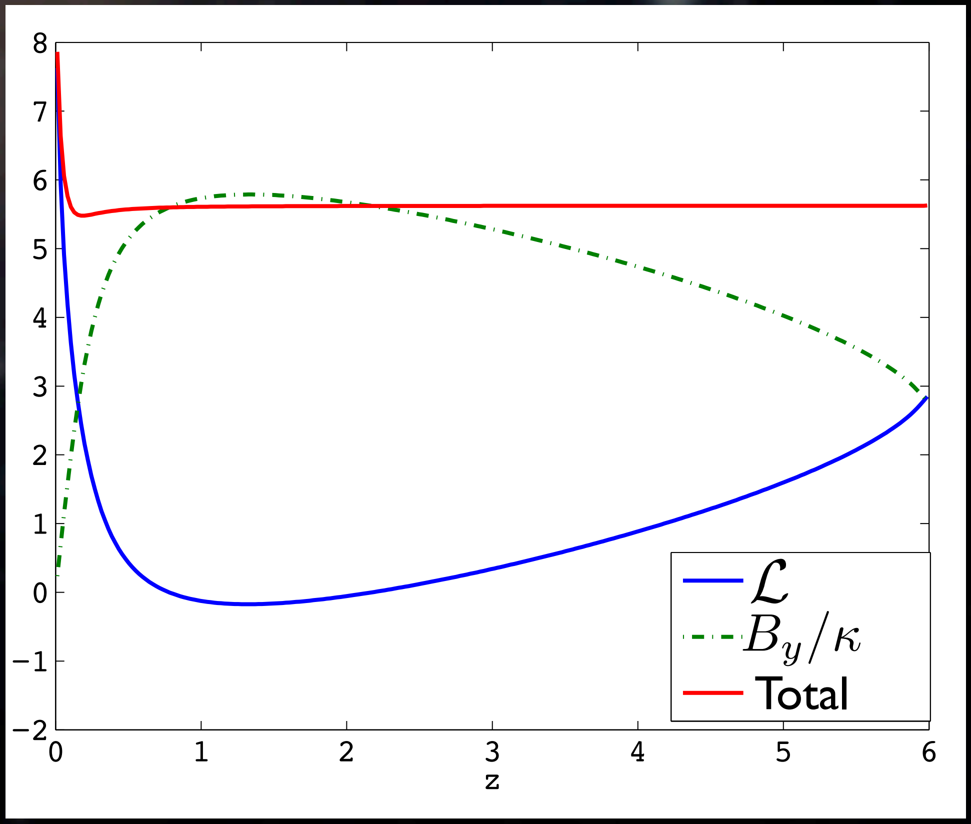
Angular momentum conservation

- Angular momentum conservation equation reads:

$$\rho \boldsymbol{v_p} \cdot \boldsymbol{\nabla} \left[\mathcal{L} - \frac{B_y}{\kappa} \right] = 0$$

- «Effective» angular momentum conserved along a streamline

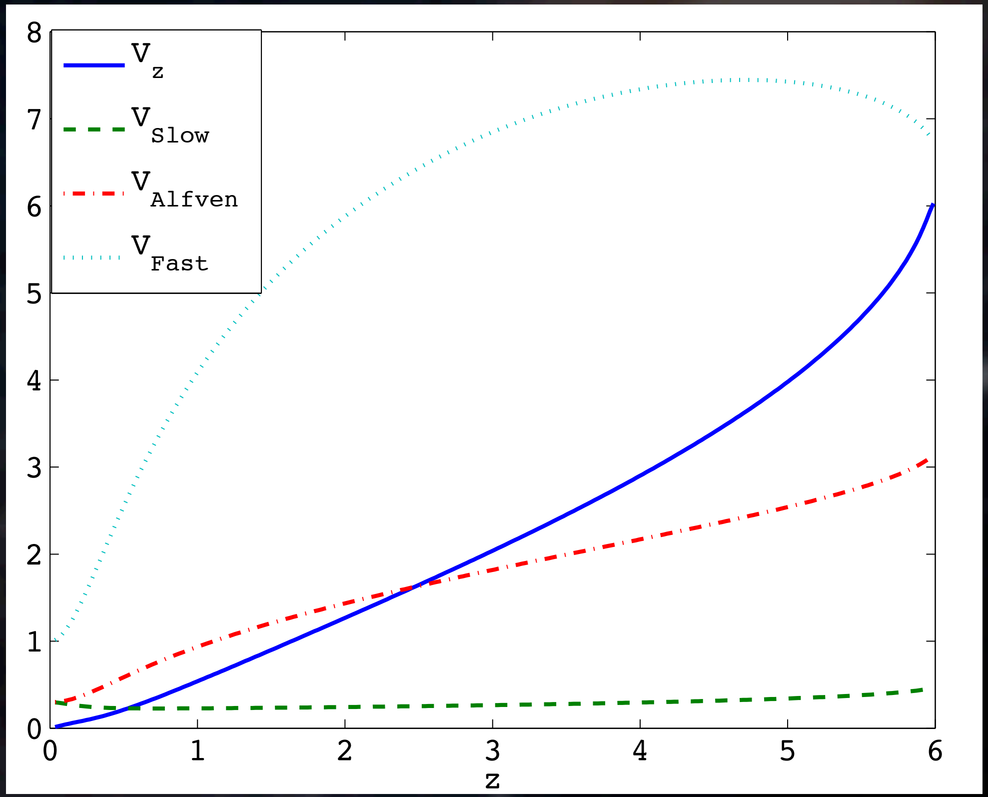




- Demonstrates the magnetocentrifugal effect is driving the outflow

Critical points

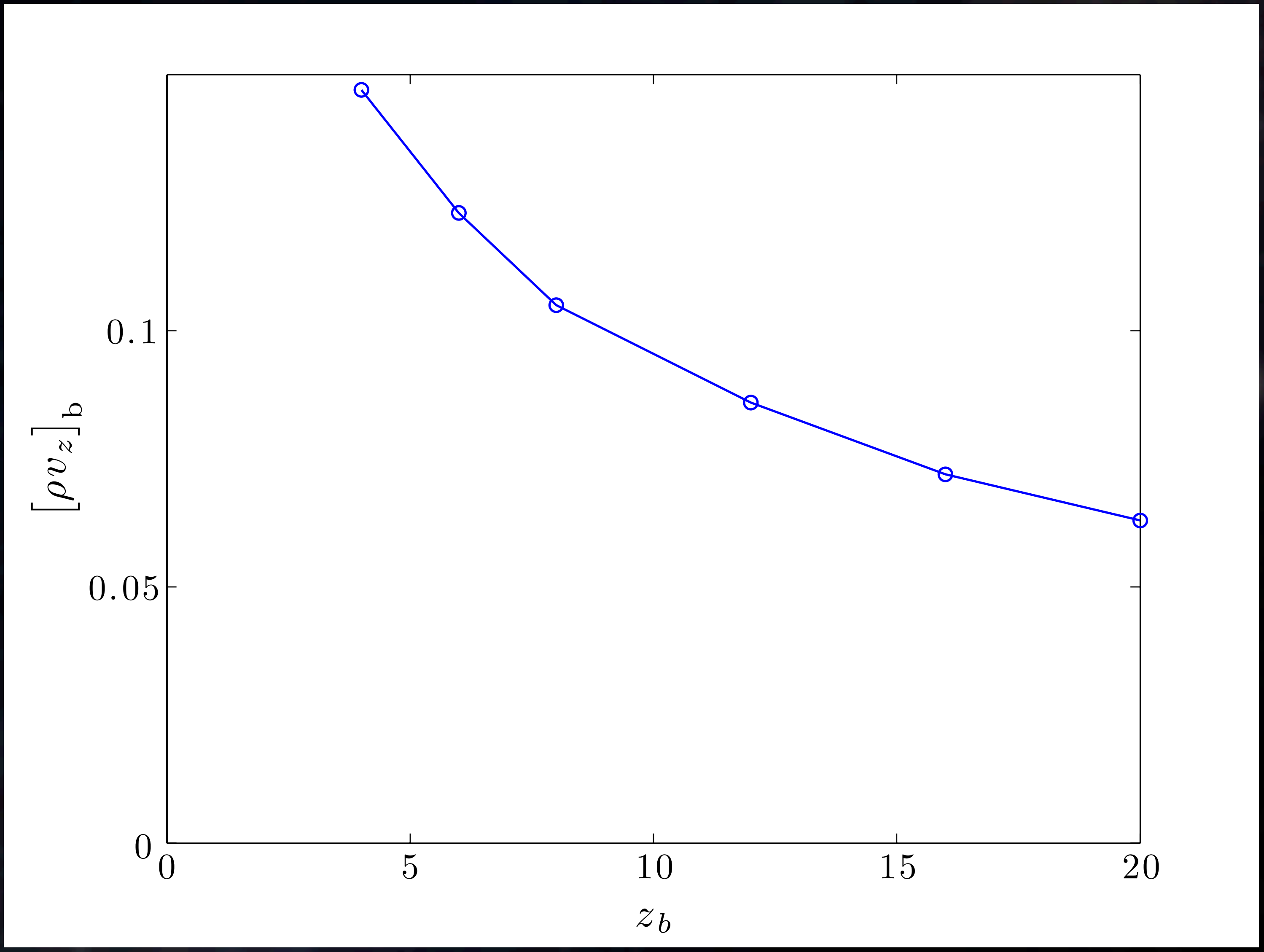
- Critical points are points (planes) where v_z is equal to a wave speed (slow magnetosonic, Alfvén and fast magnetosonic points)
- An outflow is fully determined by its launching conditions once it has crossed all the critical points



The MRI outflow is sub-fast: some quantities are boundary condition dependent.

Mass loss rate

Mass loss rate as a function of the altitude of the boundary condition



Mass loss rate is boundary condition dependent !

Refs: Fromang et al. (2013), Lesur et al. (2013)

MRI & Disk winds

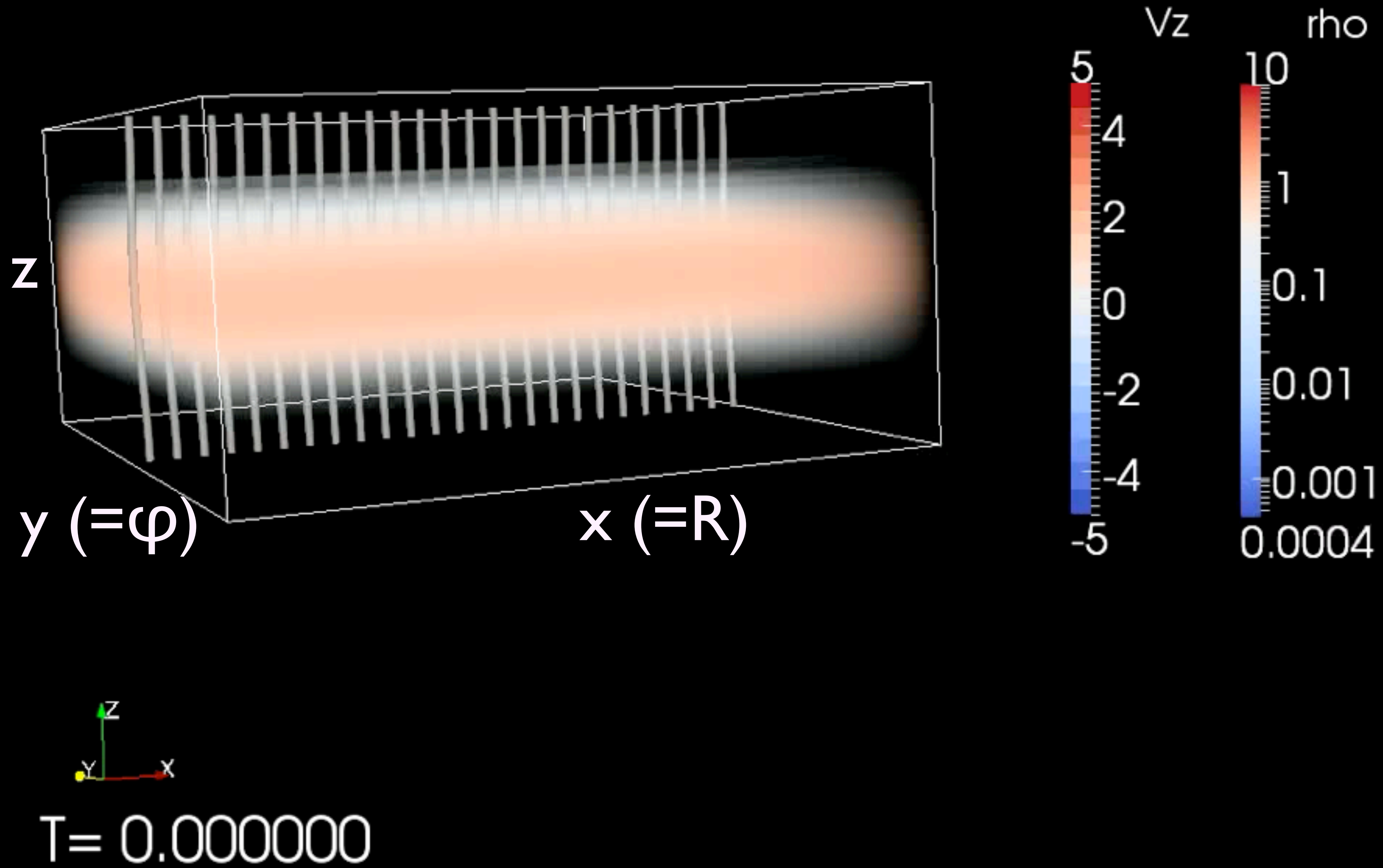
For strong enough fields ($\beta \lesssim 10$) MRI modes spontaneously evolve into disk wind

- Driven by magneto-centrifugal acceleration *à la* Blandford & Payne (1982)
- Super-Alfvenic but sub-fast

Limitations (due to shearing box)

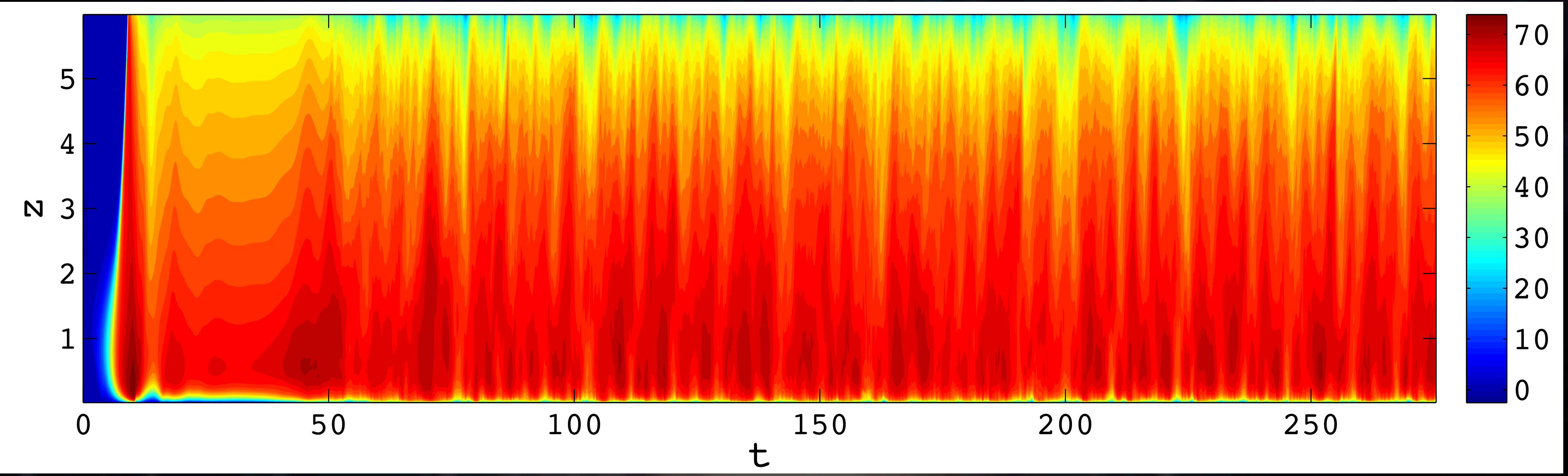
- Mass-loss rate depends on boundary conditions
- Magnetic field is dragged inward (unrealistic for a quasi steady solution)
- Is the outflow really escaping the system?

Outflow stability

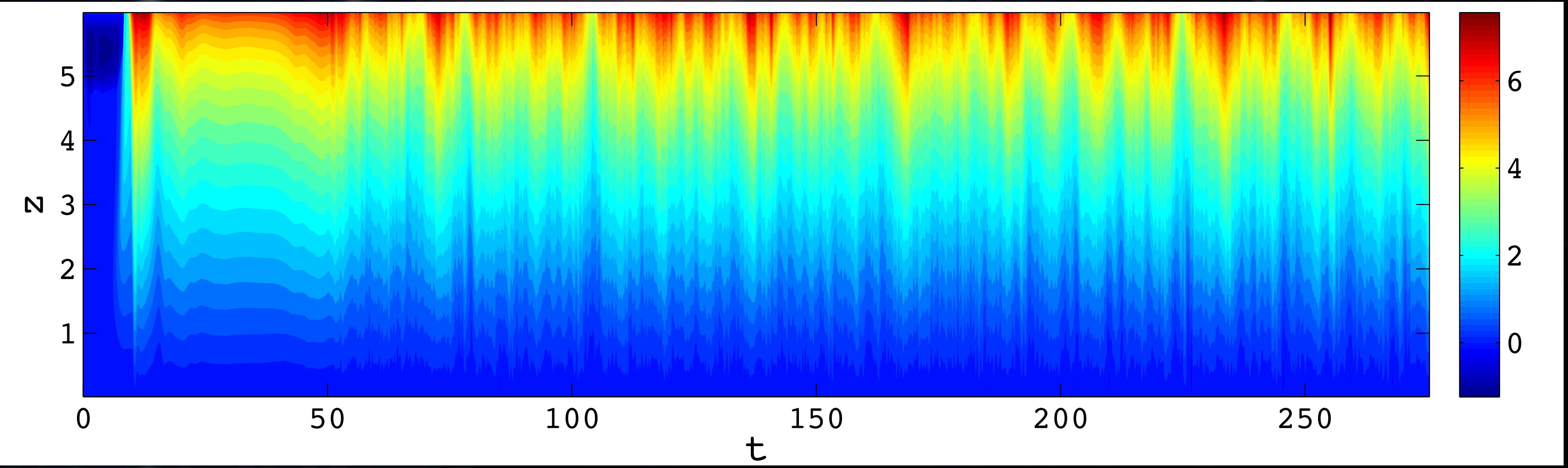


Long term evolution

Poloidal field inclination



Vertical velocity



Lesur et al. 2013

Outflow instability properties

- Outflow instability localised around z~0.6

- Vorticity profile peaked at the same location

➡ Shear driven instability? (Kelvin-Helmholtz type)

➡ Lubow et al. 1994 instability?

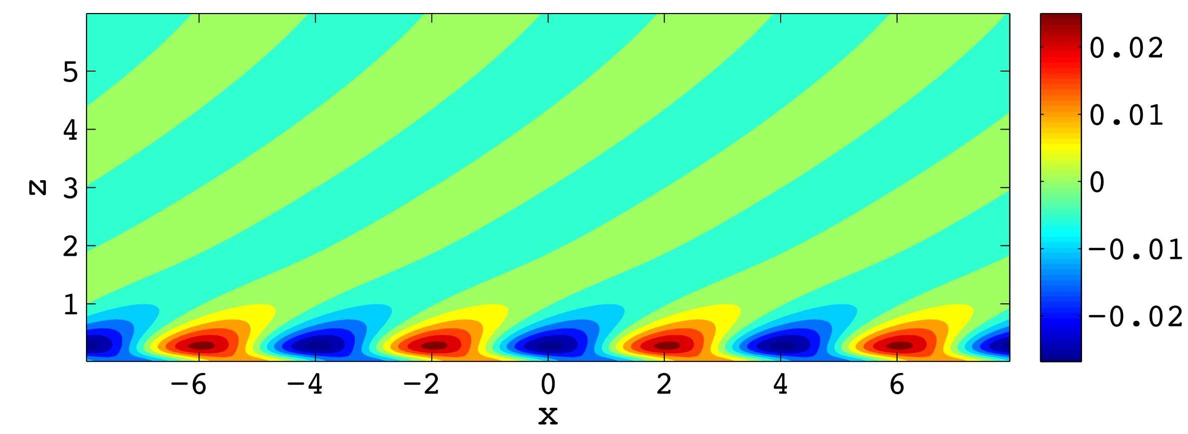


Fig. 14. Density fluctuations corresponding to the $n = 4$ eigenmode.

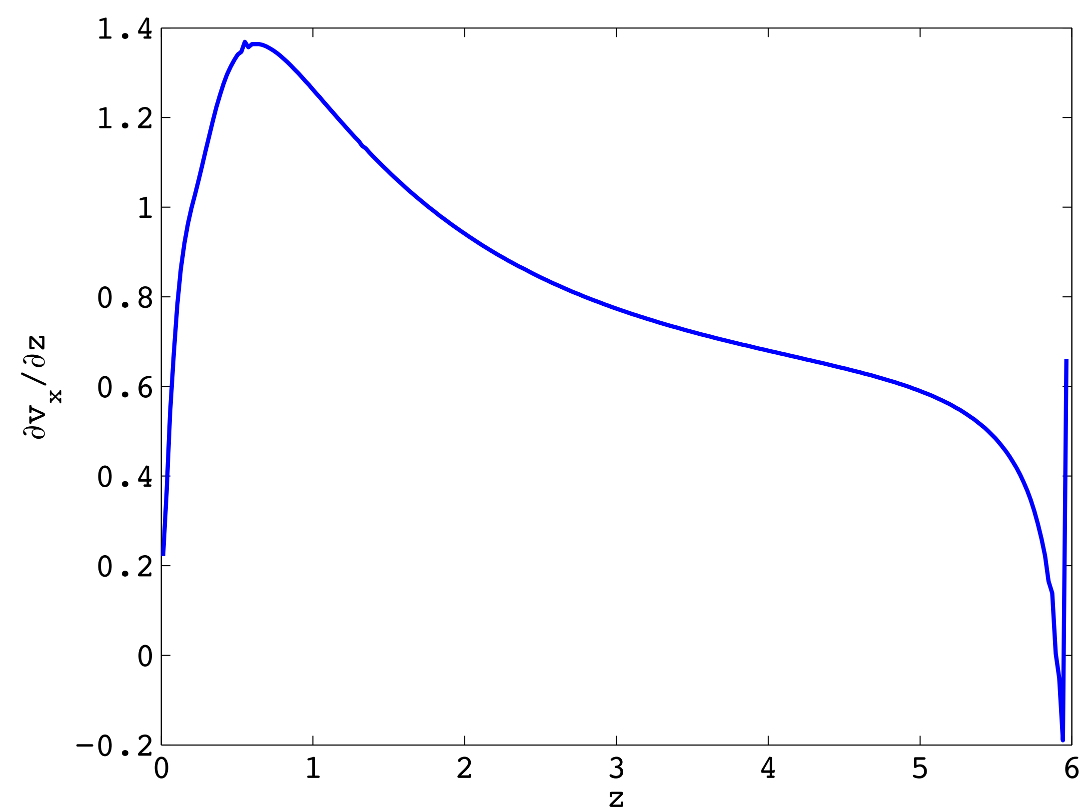
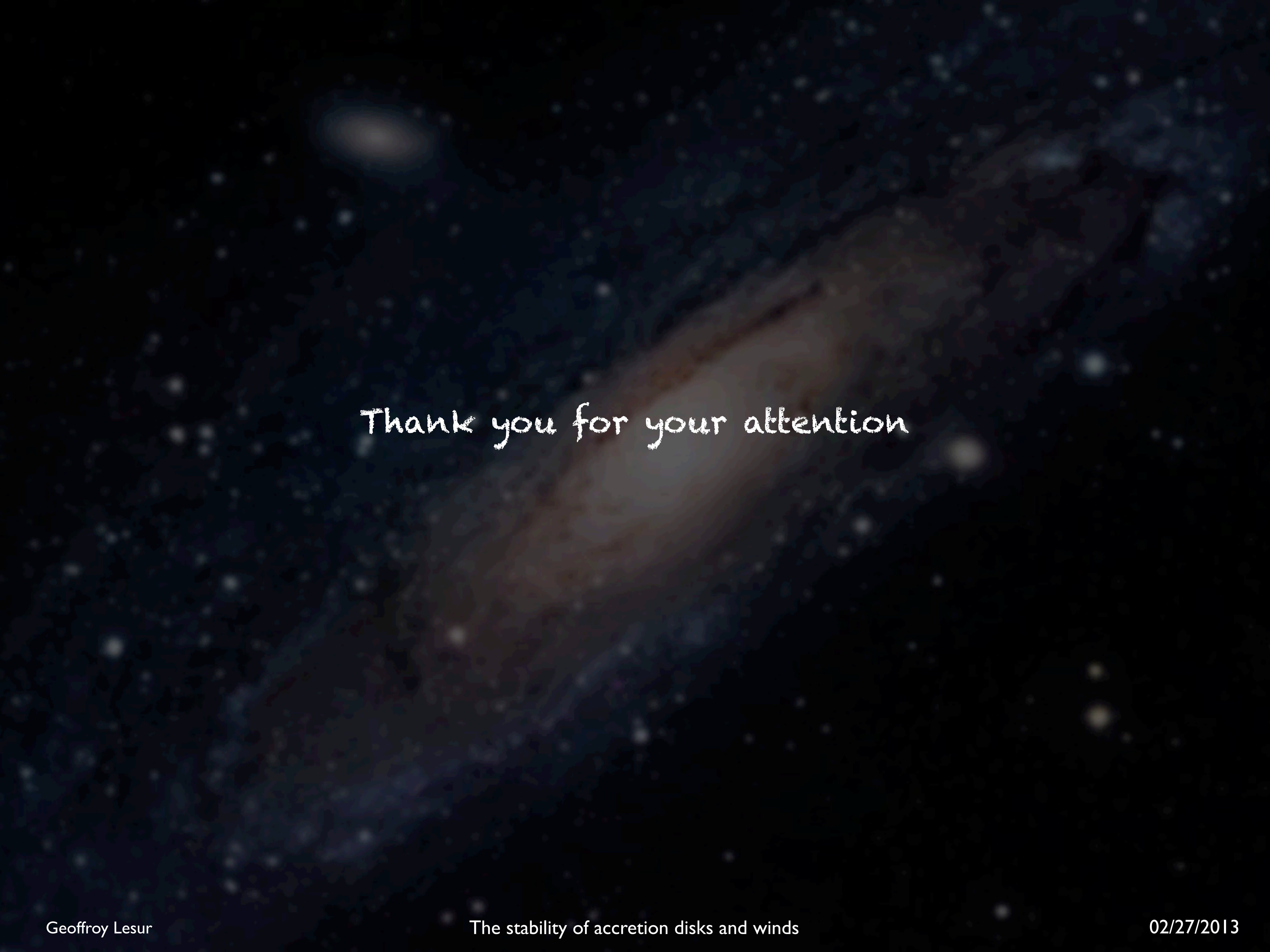


Fig. 15. y component of the vorticity in the stationary outflow solution.

refs: Moll et al. 2012, Lesur et al. 2013



Thank you for your attention